2010 Space Telescope Science Institute Calibration Workshop
Cover page and artwork designed by Pam Jeffries. The multi-color image of Omega Centauri was obtained during the Service Mission Observatory Verification (SMOV) after SM4. It is a composite of data obtained with the WFC3 UVIS channel, using the F814W, F606W, F438W, F336W, F275W and F225W filters. The COS spectrum (upper) is of 30 Dor016, a massive O2-type star. This star was observed as part of the SMOV, using the G130M and G160M gratings in the COS Far UV channel. A STIS spectrum of HDE 2699810, a near twin, is also shown (lower). These COS data were published by Evans et al. 2010, ApJ 715, L74.
The 2010 STScI Calibration Workshop

Hubble after SM4
Preparing JWST

Proceedings of a Workshop held at the
Space Telescope Science Institute
Baltimore, Maryland

21-23 July 2010

Edited by Susana Deustua and Cristina Oliveira

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3700 San Martin Drive, Baltimore, MD 21218, USA
For Rodger E. Doxsey

The proceedings of this workshop are dedicated to the memory of Rodger E. Doxsey, champion of the Hubble Space Telescope and former Head of the Hubble Mission Office at the Space Telescope Science Institute. Rodger’s tireless efforts to ensure that Hubble remain at the forefront of astronomical research by constantly improving upon the quality of science data calibrations challenged and inspired us all. Though we miss him dearly, we carry on with this tradition of Hubble data improvements and share that information with you in this volume, just as he would have wanted.
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Preface

We thank all the participants and organizers for making this workshop a success. The workshop was sponsored by the Instruments Division at STScI, and we thank all the members of the Organizing Committee, as well as Robin Auer, Roz Baxter, Steve Dignan, Pam Jeffries, Thomas Marufu, Jeff Nesbitt and his crew, Greg Pabst, Susan Rose, Dixie Shipley, Calvin Tullos, Anna Maria Valenzuela, and Mike Wiggs, for all their logistical and administrative support, without which, the workshop would not have been so smoothly run. Thanks also to Irene Stein and the Café Azafran folks for providing the tasty food, coffee and tea during the workshop.

We also thank NASA Headquarters, the HST Project at the Goddard Space Flight Center, and the Space Shuttle Program support staff at the Johnson Space Flight Center and Kennedy Space Center for their outstanding support of all HST servicing activities.

We dedicate these proceedings to the memory of Rodger Doxsey.

The Editors
Susana Deustua and Cristina Oliveira
January 2011
The Organizing Committee

Susana Deustua co-chair
Cristina Oliveira co-chair
Rachel Anderson
Azalee Bostroem
Ralph Bohlin
Marco Chiaberge
Colin Cox
Rosa Diaz
Parviz Ghavamian
Karl Gordon
Tony Keyes
Knox Long
Aparna Maybhate
Sami Niemi
Massimo Robberto
Elena Sabbi
Michael A. Wolfe
Participant List

Thomas Ake          CSC & Space Telescope Science Institute
Alessandra Aloisi   Space Telescope Science Institute
Jay Anderson        Space Telescope Science Institute
Rachel Anderson     Space Telescope Science Institute
Barbara Anthony-Twarog University of Kansas
Amber Armstrong     Space Telescope Science Institute
Thomas Ayres        University of Colorado
Sylvia Baggett      Space Telescope Science Institute
Brian Baptista      Indiana University
Elizabeth Barker    Space Telescope Science Institute
Tracy Beck          Space Telescope Science Institute
Stéphane Béland     University of Colorado at Boulder
William Blair       The Johns Hopkins University & STSCI
Ralph Bohlin        Space Telescope Science Institute
Tiffany Borders      Space Telescope Science Institute
Benoit Borguet      Virginia Tech.
Azalee Bostroem     Space Telescope Science Institute
Art Bradley         Spacecraft System Engineering Services
Marc Buie           Southwest Research Institute
Howard Bushouse     Space Telescope Science Institute
Ivo Busko           Space Telescope Science Institute
Sean Carey          Spitzer Science Center, Caltech
Pierre Chayer       Space Telescope Science Institute
Nathan Cole         The Johns Hopkins University
Alberto Conti       Space Telescope Science Institute
Colin Cox           Space Telescope Science Institute
Misty Cracraft      Space Telescope Science Institute
Gustavo Adolfo Cruz Díaz Instituto Nacional de Técnica Aeroespacial, CAB
Tomas Dahlen        Space Telescope Science Institute
Emily Dare          Tufts University
Tony Darnell        Space Telescope Science Institute
Guido De Marchi     European Space Agency
Nadia Dencheva      Space Telescope Science Institute
Pieter Deroo        Jet Propulsion Lab
Susana Deustua      Space Telescope Science Institute
Rosa Diaz           Space Telescope Science Institute
Van Dixon           The Johns Hopkins University
Timothy Dolch       Space Telescope Science Institute
Linda Dressel       Space Telescope Science Institute
Mike Droettboom     Space Telescope Science Institute
Michael Dulude      Space Telescope Science Institute
Lorenzo Faccioli    Lawrence Berkeley National Laboratory
Paul Feldman        The Johns Hopkins University
Henry Ferguson       Space Telescope Science Institute
Kathy Flanagan      Space Telescope Science Institute
Scott Friedman      Space Telescope Science Institute
Andrew Fruchter     Space Telescope Science Institute
Patrick Fry         Space Science & Engineering Center
Suvi Gezari         The Johns Hopkins University
Parviz Ghavamian    Space Telescope Science Institute
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<td>Kyan Schahmaneche</td>
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<td>Vithal Tilvi</td>
<td>School of Earth &amp; Space Exploration, ASU</td>
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<td>Wei Zheng</td>
<td>The Johns Hopkins University</td>
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Part 1. Workshop Schedule
Workshop Schedule

1. Wednesday, July 21

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<tr>
<td>9:00 am</td>
<td>Matt Mountain</td>
<td>Opening Remarks</td>
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<tr>
<td>9:15 am</td>
<td>J. C. Howk</td>
<td>The Legacy of the Hubble Space Telescope Spectrographs</td>
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<td>10:00 am</td>
<td>J. Green</td>
<td>Science Capabilities of the Cosmic Origins Spectrograph</td>
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<td>10:20 am</td>
<td>A. Aloisi</td>
<td>The On-Orbit Performance of the Cosmic Origins Spectrograph</td>
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<td>Coffee Break Posters</td>
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<td>11:10 am</td>
<td>T. Ake</td>
<td>COS FUV Flat Fields and Signal-to-Noise Characteristics</td>
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<td>11:25 am</td>
<td>R. Osten</td>
<td>COS Sensitivity Trends in Cy 17</td>
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<td>11:40 am</td>
<td>S. Osterman</td>
<td>Observing With HST Below 1150 Å: Extending COS/G130M and G140L Coverage to 905 Å</td>
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<td>11:55 am</td>
<td>C. Proffitt</td>
<td>The Space Telescope Imaging Spectrograph After SM4 Repair</td>
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<td>12:10 pm</td>
<td>T. Ayres</td>
<td>Ironing Out the Wrinkles in STIS</td>
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<td>12:25 pm</td>
<td>J. Walsh</td>
<td>Slitless Spectroscopy with HST Instruments</td>
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<td>E. Pancino</td>
<td>An Insight on the Gaia BP/RP and G-band Flux Calibration</td>
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<td>2:45 pm</td>
<td>K. Schahmaneche</td>
<td>SINDICE, a Calibrated Multi-Wavelength Light Source for Optical Telescope Calibration With a Stability and a Precision of $10^{-4}$</td>
</tr>
<tr>
<td>3:00 pm</td>
<td>J. Biretta</td>
<td>WFPC2 Closeout Calibrations</td>
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<tr>
<td>3:20 pm</td>
<td>P. L. Lim</td>
<td>WFPC2 Filters After 16 Years on Orbit</td>
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<tr>
<td>3:35 pm</td>
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<td>Coffee Break Posters</td>
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<tr>
<th>Time</th>
<th>Speaker</th>
<th>Talk</th>
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<tbody>
<tr>
<td>4:05 pm</td>
<td>A. Koekemoer</td>
<td>The NICMOS Legacy Archival Calibration Project</td>
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<tr>
<td>4:40 pm</td>
<td>G. Schneider</td>
<td>NICMOS Coronography: Recalibration and the NICMOS Legacy Archive PSF Library</td>
</tr>
<tr>
<td>4:55 pm</td>
<td>P. Ripoche</td>
<td>Calibration of the HST NICMOS F110W Using High Redshift Red-Sequence Galaxies</td>
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<tr>
<td>5:10 pm</td>
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<td>Posters &amp; Opening Reception</td>
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2. **Thursday, July 22**

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<tr>
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<tr>
<td><strong>Session V – Chair: Brad Whitmore</strong></td>
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<tr>
<td>9:00 am</td>
<td>B. Rauscher</td>
<td>An Overview of Detectors</td>
</tr>
<tr>
<td>9:45 am</td>
<td>R. O’Connell</td>
<td>New Scientific Capabilities of the HST WFC3</td>
</tr>
<tr>
<td>10:05 am</td>
<td>J. MacKenty</td>
<td>Performance and Calibration of the Wide Field Camera 3</td>
</tr>
<tr>
<td>10:25 am</td>
<td>Coffee Break</td>
<td>Posters</td>
</tr>
<tr>
<td>10:55 am</td>
<td>J. Kalirai</td>
<td>The Photometric Performance of the WFC3 UVIS and IR Cameras</td>
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<tr>
<td>11:10 am</td>
<td>E. Sabbi</td>
<td>WFC3 UVIS and IR Channel Flat Fields</td>
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<tr>
<td>11:25 am</td>
<td>M. H. Wong</td>
<td>Fringing in the WFC3/UVIS Detector</td>
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<tr>
<td>11:40 am</td>
<td>N. Pirzkal</td>
<td>WFC3 IR ”Blobs”, IR Sky Flats and the Measured IR Background Levels</td>
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<tr>
<td>11:55 am</td>
<td>H. Kuntschner</td>
<td>HST/WFC3 In-Orbit Grism Performance</td>
</tr>
<tr>
<td>12:10 pm</td>
<td>B. Hilbert</td>
<td>WFC3 IR Detector Behavior</td>
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<tr>
<td>12:25 pm</td>
<td>Lunch</td>
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<td><strong>Session VII – Chair: Max Mutchler</strong></td>
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<tr>
<td>1:25 pm</td>
<td>K. Long</td>
<td>Persistence in the WFC3 IR Detector</td>
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<tr>
<td>1:40 pm</td>
<td>H. Bushouse</td>
<td>WFC3 Image Calibration and Reduction Software</td>
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<tr>
<td>1:55 pm</td>
<td>A. Riess</td>
<td>First On-Orbit Measurements of the WFC3-IR Count Rate Non-Linearity</td>
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<tr>
<td>2:10 pm</td>
<td>D. Golimowski</td>
<td>ACS After SM4: New Life for an Old Workhorse</td>
</tr>
<tr>
<td>2:30 pm</td>
<td>J. Anderson</td>
<td>A Pixel-Based Correction for Imperfect CTE in ACS’s WFC</td>
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<tr>
<td>2:45 pm</td>
<td>P. Feldman</td>
<td>Red Leak Effects in Observations of Solar System Objects with ACS/SBC</td>
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<tr>
<td>3:00 pm</td>
<td>M. Loose</td>
<td>Application of the SIDECAR ASIC as the Detector Controller for ACS and the JWST Near-IR Instruments</td>
</tr>
<tr>
<td>3:15 pm</td>
<td>Coffee Break</td>
<td>Posters</td>
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<tr>
<td><strong>Session VIII – Chair: Jay Anderson</strong></td>
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<tr>
<td>3:35 pm</td>
<td>W. V. Altena</td>
<td>The Astrometric Context of HST in 2010</td>
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<tr>
<td>4:20 pm</td>
<td>V. Kozhurina-Platais</td>
<td>Multi-Wavelength Geometric Distortion Solution for WFC3/UVIS and IR</td>
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<tr>
<td>4:35 pm</td>
<td>E. Nelan</td>
<td>Calibration of HST’s Fine Guidance Sensor</td>
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<tr>
<td>4:55 pm</td>
<td>S. T. Sohn</td>
<td>A New Technique for Measuring Absolute Proper Motions with HST: Using Background Galaxies as Positional References</td>
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<tr>
<td>5:10 pm</td>
<td>N. Cole</td>
<td>Cross-Matching the Hubble Legacy Archive</td>
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<tr>
<td>5:25 pm</td>
<td>Cash Bar</td>
<td>Posters</td>
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<tr>
<td>6:25 pm</td>
<td>Conference Dinner</td>
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### Workshop Schedule

#### Friday, July 23

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<tr>
<td><strong>Session IX – Chair: Guido de Marchi</strong></td>
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<tr>
<td>9:00 am</td>
<td>S. Carey</td>
<td>Cross Calibration of Space Infrared Telescopes</td>
</tr>
<tr>
<td>9:45 am</td>
<td>J. Mather</td>
<td>JWST Overview</td>
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<tr>
<td>10:05 am</td>
<td>M Greenhouse</td>
<td>The JWST Integrated Science Instrument Module: Status and Test Plans</td>
</tr>
<tr>
<td>10:25 am</td>
<td>R. Jedrzejewski</td>
<td>JWST Pipeline: Discussion of Data Reduction</td>
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<td>10:40 am</td>
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<td>Coffee Break Posters</td>
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<tr>
<td><strong>Session X – Chair: Kathy Flanagan</strong></td>
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<tr>
<td>11:10 am</td>
<td>K. Gordon &amp; R. Bohlin</td>
<td>JWST Absolute Flux Calibration Planning</td>
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<tr>
<td>11:30 am</td>
<td>T. Nakos</td>
<td>Calibrators for the In-Orbit Spectrophotometric Calibration of the Medium Resolution Spectrograph of MIRI Onboard the JWST</td>
</tr>
<tr>
<td>11:45 am</td>
<td>T. Beck</td>
<td>JWST Science Operations: The Phase I &amp; Phase II Process</td>
</tr>
<tr>
<td>12:00 pm</td>
<td>A. Fruchter</td>
<td>A Redesign of Multidrizzle</td>
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<td>12:30 pm</td>
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<td>Lunch Book Signing/Meet the Author - Ed Weiler</td>
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<tr>
<td><strong>Session XI – Chair: John Grunsfeld</strong></td>
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<tr>
<td>2:00 pm</td>
<td>E. Weiler</td>
<td>Hubble - A Journey Through Space and Time</td>
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<tr>
<td>3:00 pm</td>
<td>Splinter Session I</td>
<td>Slitless Spectroscopy</td>
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<tr>
<td>3:00 pm</td>
<td>Splinter Session II</td>
<td>Count Rate Non-Linearity in the IR</td>
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Part 2. Invited and Contributed Talks
The Legacy of the Hubble Space Telescope Spectrographs

J. Christopher Howk

Univ. of Notre Dame

Abstract.

The spectrographs of the Hubble Space Telescope have each offered order of magnitude improvements over their predecessors, including those space-based instruments before HST. They have provided a wealth of information on the local and distant universe, opening up new fields of study that were not possible prior to the launch of HST. They have, of course, come with their own set of design and use challenges. I will discuss the legacy of HST spectroscopy and look forward to the types of spectroscopy we might wish to see in future space missions.
Science Capabilities for the Cosmic Origins Spectrograph

James Green

University of Colorado

Abstract. I will highlight the observational capabilities of COS, and the increase in sensitivity it provides in various observing situations. The scientific return of the instrument in its first year of operation will be quantified and presented, along with selected observations from the first year of COS data.
The On-Orbit Performance of the Cosmic Origins Spectrograph


Space Telescope Science Institute, Baltimore, MD 21218

J. Green, C. Froning, S. Béland, E. Burgh, K. France, S. Osterman, S. Penton

Center for Astrophysics and Space Astronomy, University of Colorado, Boulder, CO 80309

J. McPhate

Space Sciences Laboratory, University of California, Berkeley, Berkeley, CA 94720

T. Delker

Ball Aerospace Technologies Corporation, Boulder, CO 80309

Abstract. The Cosmic Origins Spectrograph (COS) was installed on the Hubble Space Telescope (HST) in May 2009 as part of the most recent Servicing Mission 4. COS is a fourth-generation instrument that has significantly extended HST UV spectroscopic capabilities. This paper highlights the current instrument performance, with particular emphasis on the initial on-orbit characterization during the Servicing Mission Observatory Verification (SMOV) and the subsequent calibration during Cycle 17. Plans for future improvements in the COS on-orbit calibration are also included.

1. Introduction

The Cosmic Origins Spectrograph (COS; Green 2000) is Hubble’s fourth-generation ultraviolet spectrograph. It was installed on Hubble on May 16, 2009, during the third extra-vehicular activity of Servicing Mission 4. COS is designed to perform high-sensitivity, medium- (R \( \sim \) 20,000) to low-resolution (R \( \sim \) 2,000-3,000) FUV and NUV spectroscopy of astronomical point sources in the 1150-3200 Å wavelength range. COS significantly expands the spectroscopic capabilities of Hubble at ultraviolet wavelengths, providing observers with unparalleled opportunities for observing faint point sources at ultraviolet light. The primary science objectives of the mission are the study of the origins of large-scale structure in the Universe, the formation and evolution of galaxies, the origin of stellar and planetary systems, and the cold interstellar medium.

\textsuperscript{1}Science Programs, Computer Science Corporation
\textsuperscript{2}Department of Physics and Astronomy, Johns Hopkins University
\textsuperscript{3}European Space Agency
The COS optical design achieves high performance, particularly in the FUV, by minimizing the number of reflections in the optical path and through the use of large format detectors, thus maximizing the wavelength coverage per exposure. Each channel has its own detector and a selection of gratings. An Optics Select Mechanism (OSM) is commanded to select either the FUV or NUV channel for an observation (Fig. 1). In addition, the user specifies the aperture, grating, central wavelength, grating offset position (FP-POS), and data collection mode (TIME-TAG or ACCUM) for each observation. In TIME-TAG mode individual photon events from the detector are tabulated into a list, while in ACCUM mode the photon events are integrated on-board into an image. COS is optimized to perform in TIME-TAG mode, although ACCUM mode is fully supported in the pipeline processing, and both TIME-TAG and ACCUM data are calibrated similarly.

The FUV channel spans the wavelength region extending from 1150 to 1775 Å. This is a single reflection system where a high-efficiency, first-order, aspheric holographic grating (G130M, G160M, or G140L) completely corrects the beam in the dispersion direction but has low spatial resolution perpendicular to dispersion. This channel uses a large format, solar blind cross delay line detector (XDL). This is a two segment photon counting detector with microchannel plates feeding a XDL anode in a 16384 × 1024 detector element format for each segment. The NUV wavelength range extends from 1750 to 3200 Å. To retain efficiency without the use of a large spectroscopic format NUV detector, multiple cameras simultaneously image multiple, fully aberration-corrected, spectra (in G185M, G225M, G285M, or G230L) on the Space Telescope Imaging Spectrograph (STIS) flight spare CsTe (1024 × 1024) MAMA detector. The NUV channel has also some limited broad-band imaging capability that is mostly used for target acquisition.

The initial on-orbit characterization of COS was performed during the servicing mission observatory verification (SMOV) activities, which were completed at the end of September 2009. This milestone marked the commissioning of COS for observations by the astronomical community. At that time, COS operational responsibilities were handed off to the Institute from the COS instrument development team (IDT), which is led by the principal investigator, Dr. James Green of the University of Colorado. Since commissioning, the In-
The On-Orbit Performance of the Cosmic Origins Spectrograph

2. COS Sensitivity

2.1. Initial On-Orbit Sensitivity

Initial on-orbit measurements of the sensitivity of the COS detectors were performed during SMOV (Fig. 2). For most of the NUV modes, the sensitivities resulted to be within 10-15% of the values predicted from ground testing (Massa et al. 2010a). The two NUV bare-Al gratings (G225M and G285M) that displayed time-dependent sensitivity loss on the ground, had initial on-orbit measurements that were consistent with this pre-launch degradation. In the NUV, COS sensitivity is 2 to 3 times that of STIS. The NUV detector suffers from some vignetting when observing external targets, which can decrease the sensitivity by as much as 20% at one edge of the detector. Vignetting is uniform for all three stripes, and the effect is now included in the NUV flat-field reference file (Ake et al. 2010a).

The initial on-orbit sensitivities of all COS FUV modes also resulted to be within 10-20% of the values expected from ground testing (Massa et al. 2010b). In this wavelength region, COS is more sensitive than STIS by factors of 10 to 30.
2.2. Sensitivity below 1150 Å

Initial on-orbit measurements with Segment B (short wavelength segment) of the FUV detector with the G140L grating have also confirmed some sensitivity at $R \sim 2,000$ at wavelengths shorter than 1150 Å and all the way down to the Lyman limit ($\sim 900$ Å). The sensitivity in this unusual wavelength regime for HST is due to the fact that the reflectivity of the MgF$_2$ coatings is still non-zero (McCandliss et al. 2010). This new extreme FUV mode, G140L@1280 Å Segment B, has been calibrated as part of the Cycle 17 calibration program and is fully supported starting from Cycle 18.

Exploratory observations with two new FUV settings that cover the 900-1150 Å spectral range, G130M@1055 Å and G130M@1096 Å, were also taken as part of the Cycle 17 calibration program. These new modes provide greater sensitivity below 1050 Å at a similar resolution ($\sim 2,000$) compared to G140L@1280 Å Segment B. They also have an effective area that is pretty constant over one segment (Segment A), contrary to G140L@1280 Å Segment B where the large increase in effective area at wavelengths longer than 1050 Å may easily cause a local rate violation depending on the spectral energy distribution of the source. By switching off Segment B in these two new G130M settings, much fainter FUV sources can be observed in the 900-1050 Å region (see Fig. 3 and Osterman et al. 2010 for more details). The two new G130M settings will be offered to the HST observers starting from Cycle 19 and will be fully calibrated as part of the COS Cycle 19 calibration program.

2.3. Time Evolution of the Sensitivity

On-orbit monitoring of the COS NUV and FUV sensitivity is an important part of the COS Cycle 17 and Cycle 18 calibration programs. This activity has produced some interesting results over the first year of COS on-orbit operations.
Before launch, the bare-Al gratings G225M and G285M in the NUV channel showed a wavelength-independent sensitivity degradation of about 1.6% and 4.5% per year, respectively. This degradation had been attributed to the growth of a thin oxide layer on the grating surface during the nitrogen purge that kept humidity out of the instrument before launch. Expectations were that this growth would have stopped once on orbit and that the sensitivity decline of the G225M and G285M gratings would have ceased. However, this is not the case since the throughput of the bare-Al gratings continues to decline at a rate similar or slightly higher than pre-launch rates (up to ~10% per year). While the continued on-orbit degradation of the NUV bare-Al gratings G225M and G285M was not anticipated before launch, this effect only marginally impacts a very limited amount of COS science. The gratings are used very little during Cycle 17 and Cycle 18 (less than 2-3% of the total COS time). They still provide a throughput that is slightly higher or comparable to the throughput of the STIS medium-resolution gratings covering the same wavelength range (E230M and G230M). The NUV MgF$_2$-coated gratings G185M and G230L, instead, showed no evidence of significant throughput changes over time before launch, and continue to show no significant changes in orbit.

The FUV gratings are also periodically monitored as part of the COS Cycle 17 and Cycle 18 calibration programs. The throughput of these gratings is declining faster than we expected based on UV contamination only. As of July 2010, the degradation in throughput is wavelength-dependent, with the longest wavelengths manifesting the largest decrease. Also, comparable throughput losses are observed in the same wavelength region covered by
different gratings, independently of the pixel location in the dispersion direction. Each of the two segments of the XDL detector show slightly different rates, with Segment B (shorter wavelengths) declining faster than Segment A. These characteristics of the FUV throughput degradation suggest a detector effect, particularly aging of the CsI photocathode. This aging is not localized to the most heavily used areas of the FUV detector, i.e., it is not related to the fluence or the total number of photons that hit the detector, but is a global effect. The rate of the FUV throughput loss is between 4-6% and 12% per year, depending on the wavelength. This rate is higher than ever observed in any other CsI photocathode flown on Hubble or on any other space-astronomy mission. This difference could be due to some gas contamination, because the COS FUV detector is windowless, instead of being a sealed tube. The Institute’s COS team, in collaboration with the COS IDT and the Hubble project at Goddard Space Flight Center, is currently investigating the physical causes of the FUV throughput loss and possible mitigating actions. Despite this larger than expected FUV loss, COS is still the most sensitive spectrograph ever flown on Hubble.

The decline in throughput of all the FUV gratings and the two bare-Al NUV gratings (Fig. 4) have been included into the COS exposure time calculator (ETC). ETC 18.2 was released to the public in April 2010 to support preparation of Phase II of the COS Cycle 18 approved programs. This version uses the measured trends in throughput extrapolated to the mid-point of Cycle 18 (March 2011). The time dependence of the COS NUV and FUV sensitivity has also been included into CALCOS, the COS calibration pipeline, with the delivery of new TDS reference files in mid July 2010. For more details on the time dependence of the COS sensitivity see Osten et al. (2010a,b).

3. Evolution of the Gain of the FUV XDL Detector

In TIME-TAG mode, in addition to time and location of each photon that falls on the detector, the FUV detector records the pulse-height amplitude (PHA). This gives a measure of the gain of the detector, i.e., how many photo-electrons are produced as the result of one photon striking the detector at that position.

By analyzing the FUV images with all the counts accumulated within a certain period of time since the beginning of on-orbit operations and by evaluating the corresponding maps of the pulse-height distributions as a function of time, it has been noticed that the peak of the pulse-height distribution at a certain x (dispersion) position on the detector has been steadily shifting to smaller gain values (see Fig. 5 and Sahnow et al. 2010).

This expected gradual gain degradation is due to charge extraction and is most apparent in regions of the detector subject to the largest photon accumulations. On Segment B of the FUV detector, these are the regions where geocoronal Lyα emission from the most used central wavelengths and FP-POS positions of G130M are located. Since the data are filtered by pulse height in the standard processing with CALCOS, events falling below the current minimum PHA threshold of 4 are removed from the events that contribute to the extracted spectra. As a consequence, several artificial “absorption features” due to this gain sag are now apparent in the science data. The localized flux losses associated with these features are up to ~ 5-10% as of July 2010.

The gain degradation of the XDL detector in the first year of on-orbit operations has been faster than pre-launch expectations, but is already slowing down significantly in the most heavily exposed areas. This behavior seems to be consistent with some kind of tiny gas contamination at or around launch time and a subsequent on-going on-orbit “re-scrubbing” of the detector by UV photons from the observed external targets.

Short-term strategies being implemented to reduce the localized effects of the gain sag on the science data, include lowering the minimum PHA threshold and marking the low gain regions with a data quality flag. Long term options are also investigated, including
Figure 5: Evolution of the gain versus x (dispersion) position on the detector at a variety of times since COS was installed on HST. The “modal gain” is the peak of the pulse-height distribution for events in the spectral extraction region. The large jump between the top green curve and the lower blue curve is the result of a lowering of the operational high voltage of the detector on August 15, 2009. Since then the gain has been steadily decreasing, and pulse height minima are appearing which correspond to regions of enhanced geocoronal Lyα emission.

Figure 6: Calculated LSFs for the COS FUV medium-resolution gratings that include the effects of the MFWFEs are shown as solid lines. The dotted line shows a Gaussian LSF with a 6.5 pixel FWHM. This Gaussian profile is consistent with the typical FUV LSF observed during ground testing in 2006.

raising the operational high voltage back to the original on-orbit value or moving to a new y (cross-dispersion) position on the detector.

4. On-Orbit Spectroscopic Line Spread Function

Analysis of SMOV data indicates that the COS on-orbit spectroscopic line-spread function (LSF) differs from that observed during pre-launch ground testing (Ghavamian et al. 2009, 2010). While the COS optics successfully correct for the spherical aberration of the HST primary mirror, mid-frequency wavefront errors (MFWFEs) due to zonal (polishing) errors in the HST primary and secondary mirrors result in an LSF with extended wings and a core that is slightly broader and shallower than expected. These MFWFEs could not be simulated during ground testing.

The MFWFE effects are particularly noticeable in the FUV. While the pre-launch FUV LSF is well described by a Gaussian, the on-orbit FUV LSF has up to ~ 40% of its total power distributed in non-Gaussian wings. The power in these wings is largest at the shortest wavelengths covered by the COS medium-resolution gratings (~ 1150 Å). The effect diminishes with increasing wavelength but remains non negligible even at the longest wavelengths (Fig. 6).
The effects of the MFWFEs are also present in the COS NUV LSF, particularly for the shorter wavelength gratings (G185M and G225M), although at a lower level than in the FUV. At wavelengths longer than $\sim 2500$ Å, the broad wings of the NUV LSF are instead dominated by the response of the MAMA detector (similar to what is seen for the STIS MAMA LSFs).

When a substantial fraction of the power in an LSF is transferred from the core to its extended wings, traditional measures of resolution, such as the FWHM of the line core, can be misleading. Nevertheless, the FWHM is a convenient tool to describe the resolution $R$, that in the case of the observed on-orbit COS LSF is defined as the empirical width at half the peak. With this caveat in mind, we can state that the MFWFEs reduces the COS resolving power by up to $\sim 20\%$ in the most extreme cases (shortest wavelengths of G130M). This in turn reduces the detectability of faint, narrow spectral features, leading to an increase between 20\% and 40\% of the minimum detectable equivalent width (see Ghavamian et al. 2009 for more details).

5. On-Orbit Dark Rates and their Evolution with Time

5.1. FUV XDL Detector

The average dark rate measured during SMOV for the FUV XDL detector away from the South Atlantic Anomaly (SAA) meets the pre-launch expectations of $\sim 1.5-2 \times 10^{-6}$ counts s$^{-1}$ pixel$^{-1}$ when a PHA filtering between 4 and 30 is applied. Weekly monitoring shows that the average dark rate is quite stable with time.

Deep FUV detector dark images revealed that the dark rate varies as a function of position on the detector. Segment A appears relatively featureless, while segment B exhibits several regions with higher count rates (Fig. 7). In most cases, these features will have a negligible effect on the extracted spectra, since they are quite faint. In TIME-TAG mode, these features are nearly eliminated by the default pulse-height filtering (4-30) used by CALCOS when the data are processed. They cannot be removed from ACCUM mode exposures, however, because pulse-height information is not available. Since ACCUM mode is used only for bright targets, these features should constitute a negligible fraction of the total counts in this case (see Sahnow et al. 2010 for more details).

5.2. NUV MAMA Detector

The NUV MAMA detector dark images are pretty featureless, with an average dark rate measured during SMOV away from the SAA that has a value of $\sim 6 \times 10^{-5}$ counts s$^{-1}$ pixel$^{-1}$. This value is significantly below the pre-launch predictions of $\sim 20 \times 10^{-5}$ counts s$^{-1}$ pixel$^{-1}$.

Since the early SMOV measurements, however, the dark current has continuously increased as indicated by the weekly monitoring. By late June 2010, it has risen to a level of $\sim 35 \times 10^{-5}$ counts s$^{-1}$ pixel$^{-1}$ which is slightly higher than pre-launch expectations. The NUV MAMA dark rate appears to be rising linearly with time at a rate of $\sim 2.8 \times 10^{-4}$ counts s$^{-1}$ pixel$^{-1}$ per year. In addition to this trend, there also appears to be a correlation between the dark rate and the detector temperature (Fig. 8). It is possible that the increasing dark current is due to window phosphorescence, similar to that observed in the STIS NUV MAMA detector. However, the COS NUV detector dark rate is still several times lower than both the pre-SM4 STIS NUV dark rate ($\sim 1.3 \times 10^{-3}$ counts s$^{-1}$ pixel$^{-1}$) and the current enhanced STIS NUV dark rate ($\sim 4 \times 10^{-3}$ counts s$^{-1}$ pixel$^{-1}$). The most up-to-date value of the COS NUV dark current has been included into the COS ETC version 18.2 (see Zheng et al. 2010 for more details).
Figure 7: Sum of a large number of on-orbit dark exposures. Each segment is displayed as an image and as a projection onto the x axis (units are counts s\(^{-1}\) pixel\(^{-1}\) \(\times\) 10\(^6\)). While segment A appears relatively featureless, segment B shows four pseudo-emission lines, known as arcs and knots (and labeled A and K, respectively), as well as a pair of divot/clod features, which result when their large pulse heights cause photons to be mis-registered. No pulse-height screening is employed here; when the default TIME-TAG screening is used, the Segment B features are significantly reduced.

Figure 8: Dark current (top panel) of the COS NUV MAMA as a function of time from SM4 until the end of June 2010. The dark current is still increasing linearly with time, at a rate of \(\sim 2.8 \times 10^{-4}\) counts s\(^{-1}\) pixel\(^{-1}\) per year (solid line in the top panel). Notice that the dark current also shows some correlations with the temperature of the MAMA tube (bottom panel).
6. On-Orbit Flat Fields

6.1. NUV MAMA Detector

The flat field of the COS NUV MAMA detector obtained on orbit during SMOV with the internal deuterium lamp showed consistency with the much higher S/N ground flat field that was obtained during thermal vacuum testing as a combination of internal and external deuterium lamp exposures (alignment better than $\sim 1$ pixel ; Ake et al. 2010a).

The COS calibration pipeline, CALCOS, is currently performing flat field calibration for the NUV on-orbit data as the default. The file currently used for this calibration was built from the combination of the ground and on-orbit flat-field exposures in order to improve in the final S/N compared to the pre-launch flat field file. This file was also updated to include a vignetting correction (depression up to $\sim 15-20\%$ in the first 150-200 pixels) needed for external observations. Separate vignetting corrections are currently applied for the M (G185M, G225M, and G285M) and L (G230L) gratings. SMOV observations with high S/N demonstrated that these new on-orbit COS MAMA flat fields can allow to routinely achieve S/N of 100 or more per resolution element in combination with the 4 recommended FP-POS positions (see Ake et al. 2010a for more details).

6.2. FUV XDL Detector

Two-dimensional (2D) flat fields of the COS FUV detector were obtained during ground testing. A deep flat field of the whole detector with a S/N $\sim 100$ per resolution element was obtained in 2001 at Berkeley as part of the final laboratory calibration before integration of the XDL device into the spectrograph. The purpose of the Berkeley 2D flat field was not to calibrate the relative pixel-to-pixel response of the detector, but rather to serve as a baseline for estimating on-orbit variations (Vallerga et al. 2001).

Additional pre-launch 2D FUV flat fields were obtained with the spectrograph assembled during the thermal vacuum campaign of 2003. The two internal flat-field calibration D$_2$ lamps were used extensively for this purpose. Since the light of these lamps goes through the flat-field calibration aperture (FCA) and illuminates the detector through the optics in a way similar to an external target, only the part of the detector where the source spectra land were illuminated. A deep 2D flat field with a S/N $\sim 130$ per resolution element was produced. However, this flat field was not deep enough to improve the ground-testing data beyond what could be achieved by simply using FP-POS.

Additional FUV flat field data were acquired during SMOV. Since the D$_2$ lamps are not bright enough to properly map out the flat field at FUV wavelengths, properly designed on-orbit observations of external targets, in particular bright white dwarfs, were instead used for this purpose. However, once again these data do not have adequate S/N to provide an on-orbit 2D flat field that could validate the ground flats and/or be used for flat-field calibration.

As of late July 2010, COS FUV exposures are currently not corrected for flat field by the COS calibration pipeline. However, data retrieved from the archive after March 2010 are calibrated with CALCOS in a way that the bad detector regions and grid wires are properly masked when combining data taken at different FP-POS positions into an $x1dsun$ spectrum. While this process cannot remove the $\sim 5\%$ pixel-to-pixel variations of an FUV flat field, it does a reasonable job removing the fixed pattern noise due to the grid wires.

Work is also in progress to release in the near future a post-calibration tool that combines data taken at different FP-POS positions into one summed spectrum. This tool would simultaneously fit the data with an initial guess of the flat field and of the science spectrum as inferred from the data, and will continue to iterate this process until the best fit is found for both. This one-dimensional (1D) iterative technique combined with the 4 recommended FP-POS positions looks promising and allows to achieve S/N of 100 or more per resolution element, as demonstrated on high S/N data acquired during SMOV.
We also expect to fully integrate 1D flat-field corrections for FUV observations into CALCOS by Spring 2011. This awaits the full analysis with the 1D iterative technique of the on-orbit data taken for all FUV gratings as part of Cycle 17 calibration program 12086 (see Ake et al. 2010b for more details).

7. Wavelength Calibration

Thermal vacuum testing in 2003 revealed that the dispersion solutions of the Primary Science Aperture (PSA) and Wavelength Calibration Aperture (WCA) differ from each other. During this testing dispersion solutions for all NUV and FUV gratings and settings were obtained for both PSA and WCA. Additional thermal vacuum testing in 2006 with a different set-up showed that the difference in the dispersion solution between WCA and PSA depends on the optical alignment. New data were obtained during SMOV to infer on-orbit lamp templates and measure on-orbit zero points compared to pre-launch calibrations.

NUV and FUV dispersion reference files and lamp templates to be used with CALCOS were updated for on-orbit data in January 2010. This update substantially improves the zero-point of the wavelength calibration for most COS spectroscopic modes (G185M is still to be completed), and allows an observer to achieve an accuracy of the on-orbit absolute wavelength calibration that meets the requirements (15 km/s for NUV and FUV medium-resolution gratings, and 150 km/s and 175 km/s for G140L and G230L, respectively). Monitoring during Cycle 17 indicates that the PSA-to-WCA on-orbit offsets are stable with time. We are still investigating the nature and frequency of some residual localized distortions up to 10 pixels, particularly in the FUV wavelength scale, which are not corrected by the current dispersion solution (see Oliveira et al. 2010a, 2010b, 2010c for more details).

8. COS-to-FGS Alignment

The COS-to-FGS alignment was optimized during SMOV and early Cycle 17. The tabulated position of the COS aperture relative to the FGSs was last updated in the science instrument aperture file (SIAF), based on on-orbit measurements of an astrometric target, at the beginning of March 2010. The revised aperture position should be accurate to $\sim 0.1^\prime$. The accuracy of COS initial pointing should now be limited by the accuracy of the individual stellar positions from the GSC2 ($\sim 0.2^\prime$) and of the user-supplied coordinates (that need to be better than $\sim 0.4^\prime$), not by the accuracy of the COS-to-FGS offset. Following this update, monitoring of target acquisitions has confirmed that the targets are usually well centered into the COS aperture. The Institute has revised the policy requiring that each visit starts with an ACQ/SEARCH target acquisition (TA) mode. In most circumstances, observers with approved Cycle 18 Phase I programs will be allowed to drop ACQ/SEARCH and use the time previously allocated to this TA mode for longer integrations of the science exposures (see Keyes & Penton 2010 and Penton et al. 2010 for more details).

9. Summary

Overall, COS is performing within expectations and has enabled a vigorous UV spectroscopic scientific program that has already produces several press releases. While COS on-orbit characterization continues, timely updates and documentation will be made available through our web pages (http://www.stsci.edu/cos).
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COS FUV Flat Fields and Signal-to-Noise Characteristics

Thomas B. Ake\textsuperscript{1}, D. Massa

\textit{Space Telescope Science Institute}

S. Béland, K. France, S. V. Penton

\textit{University of Colorado, Boulder}

D. Sahnow

\textit{Johns Hopkins University}

J. McPhate

\textit{University of California, Berkeley}

\textbf{Abstract.} The COS FUV channel employs a detector comprised of two microchannel plate (MCP) segments with cross delay line anodes. The detector shows several types of non-uniformities due to the hexagonal and moire patterns in the MCPs, dead spots, gain variations, and shadows from the wire grid installed in front of the MCPs to increase quantum efficiency. These features induce fixed-pattern noise in FUV spectra. The effects of these artifacts can be reduced by dividing the data by a flat field and combining exposures taken at different grating settings. A spectral iterative technique, similar to that used for GHRS and FOS, shows that $S/N > 100$ can be achieved in extracted spectra. Although flat field observations were obtained during SMOV using white dwarfs, a two dimensional flat field of sufficient quality for standard CALCOS processing was not achieved. Other methodologies are being explored for flat field correction and are expected to be installed in CALCOS to improve the $S/N$ of data incrementally. As an initial step, CALCOS currently ignores grid wire regions when creating a summed spectrum from exposures taken at different FP-POS positions. Average one-dimensional flats generated through spectral iteration have been investigated to correct individual exposures and show promise as an alternate flat fielding methodology. These may require separate flat fields for different cross-dispersion locations. An important result is that the flat fields and flux calibrations used by CALCOS are dependent on each other and should be derived together.

\textbf{1. Introduction}

The COS instrument was designed to use two main techniques to improve the signal-to-noise ratio of observations, flat fielding and fixed pattern offsets. Flat fielding corrects for pixel-to-pixel variations in the two-dimensional image of a single exposure prior to spectral extraction, while fixed pattern offsets smooth the variations when individual extracted spectra, taken at different locations on the detector, are merged into a combined spectrum.

\textsuperscript{1}Computer Sciences Corporation
For the latter technique, an observer controls the placement of the spectrum when preparing an observing proposal. A spectral region of interest is positioned through the use of FP-POS steps with a particular central wavelength (CENWAVE), or if there is sufficient wavelength overlap, through adjacent CENWAVE settings. Either of these causes a small rotation of the grating wheel, shifting the spectrum on the detector. The COS calibration pipeline, CALCOS, automatically merges spectra taken at different FP-POS settings for a CENWAVE (the XIDSUM files), but observers need to merge spectra taken at different CENWAVEs themselves. Flat fielding, on the other hand, is not performed through an observational sequence designed by the observer. The technique relies on the availability of a high quality flat field reference file that can be aligned with detector images at the pixel level. Although flat field observations were obtained during ground system tests for both COS detectors, they did not improve the S/N of other exposures for the FUV. Further work was deferred until after COS was installed in HST and checked out on-orbit.

We discuss here our investigations into on-orbit flat fielding techniques for the FUV detector. In Section 2 we summarize the COS Servicing Mission Observatory Verification (SMOV) program, which used a white dwarf star rather than the internal flat field lamps to create two-dimensional flat field images. Section 3 discusses an iterative technique used on extracted spectra that simultaneously cleans detector features from the merged spectrum and obtains a one-dimensional flat field. In Section 4, we use the 1-D flats from this technique to create average flats for the G130M and G160M gratings using SMOV and Cycle 17 calibration programs. Section 5 presents an analysis of the S/N that is achievable using the merged 1-D flat fields and FP-POS stepping. In Section 6 we discuss future directions in the investigations.

2. SMOV Flat Field Program for the FUV Detector

COS SMOV program 11491 observed WD0320–539 at five cross dispersion positions with G130M and two each with G160M and G140L to map the full science region of the FUV detector. Different CENWAVEs and FP-POS settings were used to separate the spectral and detector features. Each grating and FUV segment was analyzed independently. The time-tagged data were converted into images, exposures were added together in pixel space to accumulate the maximum number of photons, then the composite spectral and detector slope were removed by fitting each row with a sixth-order polynomial, avoiding the edges of the detector. The slope-corrected image was normalized to unity in a featureless 500-pixel region near the middle of the detector. Figure 1 shows the result for segment A with the G130M grating.

To evaluate how well these flats improved the S/N, we used observations from SMOV program 11494, which obtained high S/N observations of WD0947+857 (G130M, G140L) and WD1057+719 (G160M) capable of supporting S/N $> 60$ per resel. Different correction methods were tested to study S/N improvements. First, individual exposures were processed by dividing the SMOV flat field images into the 2-D exposure data, then spectra were extracted from the corrected images, as in the CALCOS design. This was found to remove the grid wire shadows, but induced some structure due to the low S/N per pixel of the flat and, perhaps to some degree, misalignment of the flat to the data. To improve the S/N of the flat field correction, a second type of processing was examined. An extraction slit was run across the flat field image similar to the extraction of the WD spectrum, then this 1-D flat was divided into the 1-D unflattened spectrum. This was found to be somewhat better than the 2-D processing for individual exposures. When different FP-POS spectra were merged into a final spectrum, the 2-D method often did not work substantially better than FP-POS summing with no flat fielding, but the 1-D processing always showed some improvement. A third method, using a spectral iterative technique (Section 3), yielded the highest S/N. Figure 2 illustrates the results for G130M segment A.
Figure 1: G130M flat field image for segment A from program 11491. Vertical bars are wire shadows from the QE grid. Hexagonal patterns arise from multi-fiber packing and pore deformation during microchannel plate (MCP) manufacturing. Moire stripes arise from offsets of pore locations in the stacked MCPs. Dead spots appears as white holes.

Figure 2: Examples of four FP-POS exposures for WD0947+857 corrected and merged by various techniques. Top spectrum is a straight sum of the data in wavelength space (the X1DSUM spectrum). The middle two spectra are merges after flat fielding is performed on each exposure, by the SMOV 2-D reference file or 1-D extracted flat. The bottom spectrum is the result of self-correction through the iterative technique, without using the flat field image at all. Spectra are offset by an arbitrary amount for clarity.
Since the on-orbit flats from program 11491 did not improve the S/N, no reference file updates were made to CALCOS and flat fielding for the FUV detector continues to be disabled. A separate change was made to pipeline processing instead. To eliminate the effects of the grid wire shadows on FP-POS summed spectra, the default SDQFLAGS keyword was changed so that these regions are excluded when the data are merged (Ake et al. 2010). The shadows are 50 pixels wide and ~20% deep in an individual exposure, occurring every ~840 pixels. Originally when CALCOS coadded FP-POS exposures, the wire shadows were reduced in depth but appeared in more places. For the maximum four FP-POS steps, the shadows became only 5% deep, comparable to the fixed pattern noise, but then corrupted 30% of the spectrum. In the current CALCOS products, ignoring the grid wires causes those regions to have lower S/N than surrounding areas since fewer exposures contribute in the sum, but no pixels are corrupted and the S/N of the spectrum on average improves. This change is a temporary measure until an acceptable flat fielding strategy is found for the FUV detector.

3. Spectral Iteration

Since a 1-D flat field correction appeared to do better than a 2-D one from the SMOV analyses, we investigated whether average 1-D flats generated from spectral iteration could improve the flat fielding instead. The iterative technique was developed for high S/N observations obtained with the GHRS (Cardelli & Ebbets 1993; Lambert et al. 1994) and has been applied to FOS and STIS spectra as well (e.g., Gilliland 1998). The algorithm takes advantage of the property that, when spectra are taken at different locations on the detector, features belonging to the star are constant when comparing data in wavelength space, while features arising from the detector are constant in pixel space. The first step of the iteration merges all exposures in wavelength space to obtain an initial best estimate of the stellar spectrum. This is then divided into the individual exposures, with the resulting residuals being then shifted and merged in pixel space to obtain an estimate of the flat field. This average flat is divided back into the spectra in pixel space, which results in an improved estimate of the stellar spectrum when they are merged in wavelength space. By switching between wavelength and pixel space in merging and ratioing the data, the stellar spectrum and underlying fixed pattern noise are solved simultaneously. For COS, the SHIFT1A and SHIFT1B FITS header keywords provide the conversion between spaces.

Several SMOV and Cycle 17 programs made observations at different grating settings, either with various FP-POS locations for one CENWAVE or with multiple CENWAVEs. Table 1 lists those for the G130M and G160M used for this study. Each grating still has to be analyzed separately since their spectra nominally lie at different cross-dispersion (Y) locations. Locations for different CENWAVE and FP-POS exposures for a grating vary

<table>
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<tr>
<th>ProgID</th>
<th>Program</th>
<th>Target</th>
<th>Grating</th>
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<th>FP-POS</th>
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<td>WD0320–539</td>
<td>G130M</td>
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<td>1577,1589,1600, 1611,1623</td>
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Figure 3: Stellar spectrum (left) and flat fields (right) resulting from a spectral iteration for the G130M observations in program 11494. See text for further explanation. Details of the region near Si IV $\lambda 1393$ in segment A can be seen in the bottom spectrum in Fig. 2.

with Y, but only by 2–3 pixels. Target acquisition errors can also cause spectra to fall at different Y positions.

Figure 3 illustrates the results of an iteration of eight FP-POS G130M exposures for the SMOV high S/N program, 11494. The upper left panel is the final merged spectrum of WD0947+857 for both FUV segments, which are iterated separately. The bottom left panel shows the total counts that went into making the spectrum, illustrating how the near-edge regions have fewer exposures contributing to the final spectrum. The right panels show the resultant 1-D flat field for each segment and an estimate of the S/N in the flats assuming that half the information in the data goes to making the spectra and half to the flat. Except for the right hand side of segment A and the region around Ly$\alpha$ in segment B, which have the fewest contributing counts from the star, the S/N achieved indicates most of the variations in the flats are due to fixed pattern noise. The prominent regular dips in the flats are the grid wire shadows.

The situation when iterating on multiple CENWAVEs is a little more complex. As with other HST detectors, the flat field for COS can be considered to be the product of two components: a P-flat, which characterizes the pixel-to-pixel sensitivity variations of the detector, and an L-flat, which accounts for larger scale, low-frequency variations. In CALCOS, the flat field reference file is a P-flat. The L-flat correction is essentially folded into the sensitivity curves, which are specified for each grating and central wavelength. The iterative algorithm, which attempts to equalize the fluxes in the wavelength overlap regions, will create a combined P- and L-flat along with a correction for the relative difference in sensitivity between CENWAVEs.

Figure 4 illustrates the effect for program 11491, where WD0320–539 was observed at two FP-POS each with CENWAVEs $\lambda 1291$ and $\lambda 1309$. The upper two panels show an iterated solution for the segment A exposures processed separately by CENWAVE. Spectra from the two settings do not match exactly, likely because the large grating rotation between them causes a small, but apparent, change in the blaze function at the detector. When all
the exposures are iterated together, the fluxes now match, but the flat field becomes curved (lower panels). These results illustrate that ultimately the flat field and flux calibrations are dependent on each other and should be derived together.

4. Generation of 1-D Flats

The spectral iterative technique was executed on each of the programs in Table 1. For the Cycle 17 sensitivity monitoring program, 11897, flats were created for each epoch of the observations individually since the shape of the grating response curves have been found to change over time (Osten et al 2010). The flats were combined for each grating and segment, weighting by the \((S/N)^2\) estimates as in Fig. 3. A second-order polynomial, ignoring the grid wire shadows, sufficiently removed the L-flat curvature for multiple CENWAVE iterations.

As an example, Figure 5 shows the final 1-D flats for the G130M grating. The S/N is estimated to be between 50–75 per pixel, or an error of roughly 1.5–2%. This is smaller than the variation seen in the flats, indicating that we are resolving the fixed pattern noise in the detector, albeit as an average within the spectral extraction slit height. The data sets are not of high enough S/N to allow us to determine any differences of the flats with CENWAVE, which could occur due to the spectra falling at different cross-dispersion locations.

To evaluate the stability of the flats, we can perform a consistency check by dividing the final flat into each of the contributing programs and calculate statistics on the residuals. In Figure 6 and 7 we show the results for the G130M flats. First we can see that the grid wire shadows are nicely corrected in each program flat, leaving no visible artifacts above the overall noise level. Detector dead spots do leave residuals since the individual spectra have contributions from different cross-dispersion positions. These regions, such as the one
at X ∼ 9400 in segment A, were never expected to be correctable by a flat field and are marked by CALCOS with a bad data quality flag. The long wavelength end of segment A, beginning at X ∼ 11000, shows variations that increase towards the end of the segment. This may be evidence of misalignment of the flats or may be due to low S/N effects since the long wavelength ends of the WD exposures have the fewest counts, caused by the downward slopes of both the grating sensitivity and stellar energy distribution. Further observations are necessary to investigate the discrepancy.

Table 2 lists computed statistics of the G130M residuals. The RMS variation in the final flat for regions between the grid wires, $\sigma_{\text{FLAT}}$, is 3.5%, and including the shadows, 5%. The program flats range from 5–7% and 3.5–6.4%, with and without the grid wire regions. The residual RMSs, after dividing the individual flats by the final one ($\sigma_{\text{RESID}}$), are cut roughly in half, demonstrating that the fixed pattern structure occurring in the final flat is stable and well-aligned from star to star. We also note that $\sigma_{\text{RESID}}$ values are the same with and without the grid wires, indicating that the grid wire shadows have been well-corrected by the 1-D flat.

5. Signal-to-Noise Evaluation

The ultimate criterion of the usefulness of flat fields is their ability to improve the S/N for any observation. We have determined the S/N ratios achievable in spectra corrected by the 1-D flats, contrasting them to unflattened exposures taken at one FP-POS position and merged spectra from multiple FP-POS and CENWAVEs with no flat fielding. First we reprocessed the data sets listed in Table 1 using the final 1-D flat fields. Each exposure was divided by the flat appropriate to its grating and segment mode, then were combined by weighting by the exposure time. Spectra were made for both single FP-POS settings, and thus only corrected by flat fielding, and as fully merged spectra, with both flat fielding and FP-POS smoothing. Then to simulate the output products from the current version of
Figure 6: The G130M segment A final 1-D flat field (black line) and results of dividing it into the individual contributing program flats (colored lines, offset by arbitrary amounts for clarity).

Figure 7: The G130M segment B final 1-D flat field and residuals, as in Fig. 6.
Table 2: G130M Flat Field Comparison Statistics

<table>
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<th>With Grid Wires</th>
<th>Without Grid Wires</th>
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<td>Segment B</td>
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</table>

CALCOS, merged spectra were created similarly using a unity flat and ignoring the grid wire shadows. A difference here is that we combined exposures at different CENWAVES, which CALCOS does not do, to maximize the S/N that can be obtained with the data available. Finally, we combined, without flat fielding, observations that had multiple exposures at one FP-POS position, to determine the maximum S/N obtainable with exposures taken at only one grating setting. All spectra were binned to the size of an FUV resolution element, six pixels, for the S/N measurements.

The S/N ratios were determined in the same way for all spectra. Because the WD spectra in Table 1 are relatively featureless, except in areas around sharp interstellar lines, we fit the continua with polynomials and calculated the RMS in 10 Å sections along the spectrum. This allowed us to evaluate the S/N compared to the Poisson limit at different exposure levels, such as in the wings of Lyα on segment B. We take the S/N to be the inverse of the RMS value. For segment A, we ignored the longest wavelength regions since the flat fields are problematic there.

Figure 8 shows the measured S/N for the different processings compared to the Poisson limit determined by the total counts in each 10 Å section. We have averaged both gratings (G130M and G160M) and both segments together. We find that the maximum S/N achievable with a single grating setting is $\sim 20$, consistent with prelaunch measurements and the $\sigma_{\text{FLAT}}$ values in Table 2. The CALCOS merging of exposures can reach S/N~50 for observations with four FP-POS positions totaling at least 6000 counts per resel. With flat fielding, S/N~40 can be reached for a single grating setting, and with four FP-POS positions, S/N~100 is possible.

One caveat with this analysis is that we are using the same data to assess the flats as what went into making them. In this case, the maximum expected S/N would no longer be the Poisson limit, but should be somewhat higher. Although different targets, CENWAVES, and FP-POS steps were involved with the investigation, we have not sufficiently evaluated the usefulness of applying the 1-D flats to other targets. The programs in Table 1 were the highest S/N data available to us and we deemed it better to utilize all the data to make flats better than 2%, rather than dividing the data sets into two populations, one to create the flats and one to evaluate them.
Figure 8: Average S/N ratios measured for G130M and G160M spectra created by different kinds of data processing. Open circles are for exposures taken at one grating setting, filled circles for merges of multiple FP-POS and/or CENWAVE positions. Unflattened, single position exposures are in red. Violet symbols are merges of multiple setting, unflattened exposures ignoring grid wire shadows, as in current CALCOS processing. Blue symbols are for spectra corrected by the 1-D flat fields.

6. Summary

Although 2-D flat field images made from the FUV SMOV 11491 program were unable to improve the S/N ratio of FUV data significantly, we find that 1-D flats show promise. The spectral iterative technique is a robust method to generate them for high S/N data, simply requiring that features in the source spectrum are accurately aligned in wavelength between exposures and are non-varying in shape (e.g., geocoronal Lyα cannot be corrected by this technique). With the limited data available to us, the flats we have created are not expected to be universally applicable to all COS observations. They have been derived from well-centered point sources for a limited set of CENWAVEs. We have discovered that a self-consistent flux and flat calibration is needed to perform an L-flat correction, either through a separate reference file or through the sensitivity curves. We have not analyzed G140L spectra yet, which may require changes to the iterative routine since spectra only cover part of the detector segments and second-order light appears at the long wavelength end of segment A.

More high S/N observations are planned for Cycle 18, particularly to characterize the flat fields at other CENWAVEs, and hence, different cross-dispersion locations. More investigation is also needed into why the G130M 1-D flats at the long wavelength end of segment A appear to be inconsistent. Lastly, checks should be made on the stability of the flats over time.
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Space Telescope Science Institute, 3700 San Martin Drive, Baltimore, MD 21218

S. V. Penton, S. Osterman
CASA, University of Colorado, Boulder, CO 80309

D. Sahnow
The Johns Hopkins University, Baltimore, MD 21218

Abstract. After the initial on-orbit determination of the absolute flux calibration of the Cosmic Origins Spectrograph was performed, we have been monitoring the instrument’s spectroscopic sensitivity regularly. The bare Aluminum gratings G225M and G285M exhibited sensitivity degradation on the ground relative to the Al+MgF$_2$-coated gratings, measured during semi-annual grating efficiency tests. The rates of degradation were about 1.8% per year for the G225M and 5.4% per year for the G285M grating relative to the G230L grating. Observations of spectroscopic white dwarf standard stars have shown a decline in sensitivity in the G225M and G285M gratings on the NUV channel, while the gratings with Al+MgF$_2$ coatings (G230L and G185M) appear stable. The trends seen in external targets confirm those seen using internal exposures, and appear to be wavelength-independent but grating dependent. All gratings on the FUV channel are experiencing wavelength-dependent sensitivity degradation, which is worse at longer wavelengths. At the shortest FUV wavelengths the sensitivity decline is around 5% per year, increasing to about 12%/year at 1800 Å.

The sensitivity changes on the NUV and FUV channels appear to be uncoupled. The sensitivity decline of the NUV bare Al gratings appears to be caused by continued growth of an oxide layer, either from additional deposition of atomic oxygen in space or migration of existing atomic oxygen in the system to the outside. Detector QE loss due to localized exposure of the FUV cross-delay line detector to UV irradiation can be ruled out by several tests which examine the sensitivity decline versus total counts in specific regions of the detector. The wavelength dependence of the FUV sensitivity degradation is the opposite sense for contamination to be an issue. The source of the FUV sensitivity decline does not appear to be due to shifts in the pulse-height distribution from gain sag. The loss of quantum efficiency shows the classic signature of photocathode degradation.

1. Introduction

Aloisi et al. (2010) provide a summary of the on-orbit performance of the Cosmic Origins Spectrograph (COS). This paper is a summary of results from the COS spectroscopic sensitivity monitoring calibration programs for the first 9 months of Cycle 17, from September 2009 through early May 2010. This topic is discussed in more detail in Osten et al.
2010. Early results indicate changes in sensitivity for some modes, only some of which were anticipated before installation on HST.

2. Internal Monitoring of Spectroscopic Sensitivity

The internal monitoring of the near-UV spectroscopic sensitivity is performed via grating efficiency tests. These use controlled wavelength calibration lamp exposures and compute ratios of emission lines in common to a pair of gratings. They were designed in 2003 after thermal vacuum testing suggested that the bare Aluminum gratings G225M and G285M on the NUV channel were declining in efficiency. These tests were done on the ground starting in 2003 and performed semi-annually until just before launch in 2009. They were repeated in orbit twice during the time period under consideration here. The grating efficiency test is a relative measure of spectroscopic sensitivity since it uses ratios of lines from different gratings. A summary of GET results from ground data is shown in Figure 1. Each data point plotted is the slope of the fit to the efficiency ratio versus time for an emission line at a particular wavelength. Only ratios relative to the G230L grating are plotted. Both the G185M and G230L gratings are Al coated with MgF$_2$. The G185M/G230L grating ratio appears stable in time, having a slope near 0. The bare Al gratings show a decline relative to G230L, and this decline is a different value for the two bare Al gratings: for G225M relative to G230L it is $-1.8 \pm 0.9\%$/year, and for G285M relative to G230L it is $-5.4 \pm 0.5\%$/year. The results for the G225M and G285M gratings reveal that the trend is independent of wavelength.

Bare Al can react with oxygen and develop a thin oxide coating. The NUV detector was in a nitrogen purge environment, enabling the possible continued growth of an oxide layer. It was initially thought that this growth would cease once in orbit. The two on-orbit GET measurements appear to be consistent with a continuation of the trends observed on the ground.

3. External Monitoring of Spectroscopic Sensitivity

The external monitoring of spectroscopic sensitivity uses spectrophotometric white dwarf standard stars. Table 1 lists the white dwarf standards and which gratings they are used to monitor. After sensitivity declines in the FUV gratings were detected in early 2010 the monitoring frequency of the medium-resolution FUV gratings was increased to monthly from quarterly. The time-dependent sensitivity (TDS) trends are computed independently for each grating, central wavelength, and detector segment or stripe. Spectral ratios are taken relative to the first spectrum for a given grating/central wavelength/segment or stripe, and then averaged over suitable wavelength ranges to avoid the edges of the spectra or other features like Lyman $\alpha$. These are plotted as a function of time, and linear fits are computed. These slopes constitute the time-dependent sensitivity. A summary is shown in the top panel of Figure 2 for all COS FUV and NUV gratings. Each data point refers to a stripe or segment of a particular configuration. The bottom panel of Figure 2 displays a close-up view of the FUV gratings. For the FUV channel segments A (B) are indicated with open (closed) symbols, respectively. The target acquisition for the G140L+G130M monitoring visits and G160M monitoring visits was changed in February and March 2010, respectively, to use an NUV imaging target acquisition scenario. Prior to those dates dispersed light target acquisitions had been used. The change results in a small displacement of the spectra within the aperture but the effect on throughput is negligible.

The NUV sensitivity decline seen on the ground in the bare Al gratings continues in orbit. The weighted mean change of the G230L sensitivity from external target is $-1.1 \pm 0.4\%$/year, and the G185M weighted mean change is $-0.8 \pm 0.4\%$/year. The G225M grating shows a sensitivity decline of $-3.3 \pm 0.3\%$/year, and the G285M grating shows a decline of
Figure 1: Summary of results from ground-based grating efficiency test. The bare-Aluminum gratings exhibit an efficiency decline relative to the G230L grating which is independent of wavelength, while the G185M grating shows no efficiency decline relative to the G230L grating.
−10.8 ± 0.2%/year. An unanticipated result was the appearance of sensitivity decline in the FUV channel. The G130M grating has a weighted mean sensitivity change of −5.6 ± 0.2%/year, and the G160M grating it is −11.8 ± 0.3%/year. For G140L the decline ranges from a weighted mean of −7.2 ± 1.5%/year from 1300 < λ(Å) < 1400 to −10.9 ± 3.5%/year from 1700 < λ(Å) < 1800. There is a difference of about 3% between FUV gratings at the same wavelength and segment, as well as different gratings and segment at the same wavelength.

<table>
<thead>
<tr>
<th>Detector</th>
<th>Grating</th>
<th>WD Monitor</th>
<th>Monitoring Frequency in Cycle 17</th>
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<td>COS/FUV</td>
<td>G140L</td>
<td>WD0947+857</td>
<td>monthly</td>
</tr>
<tr>
<td></td>
<td>G130M</td>
<td>WD0947+857</td>
<td>quarterly initially, monthly since March 2010</td>
</tr>
<tr>
<td></td>
<td>G160M</td>
<td>WD1057+719</td>
<td>quarterly initially, monthly since March 2010</td>
</tr>
<tr>
<td>COS/NUV</td>
<td>G230L</td>
<td>WD1057+719</td>
<td>monthly</td>
</tr>
<tr>
<td></td>
<td>G185M</td>
<td>G191B2B</td>
<td>quarterly</td>
</tr>
<tr>
<td></td>
<td>G225M</td>
<td>G191B2B</td>
<td>monthly</td>
</tr>
<tr>
<td></td>
<td>G285M</td>
<td>G191B2B</td>
<td>monthly</td>
</tr>
</tbody>
</table>

4. Tests of Global vs. Local Sensitivity Declines

The rapid rate of FUV sensitivity decline was unexpected. This rate is larger than what other space FUV detectors have experienced (Davidsen et al. 1992, Bohlin 1999, Dixon et al. 2007). The wavelength dependence of the FUV sensitivity decline suggests that it is an issue with the photocathode. Contamination would be expected to have the opposite trend with wavelength, i.e. worse at shorter wavelengths. It is known that cross-delay line detectors such as the one used on COS can experience detector quantum efficiency loss due to localized exposure to UV irradiation (Tremsin & Siegmund 2001). In order to test whether the sensitivity decline was a global or local phenomenon, a test (CAL/COS program 12096) was performed which obtained spectra of the sensitivity monitor at positions offset from the default spectral location. Figure 3 shows the TDS trends for the monitoring data obtained through early May, along with spectra obtained at 1.2 and 3.0″ from the nominal location. The offset spectra show count rates within 1% of the nominal position, indicating that the sensitivity decline is a global phenomenon.

The regions of the detector where the largest number of photons have accumulated are the regions where geocoronal Lyman α emission fall. There was no evidence of localized sensitivity declines in these regions. However, in the course of examining these regions Sahnow et al. (2010) discovered that the peak of the pulse height distribution is steadily shifting to smaller values in these regions. The gain sag is not related to the time-dependent decline in spectroscopic sensitivity; although the size of the charge cloud of photoelectrons is reduced, this does not affect the probability of detecting the photon.

5. Conclusions

The NUV bare Al sensitivity decline was not unexpected; the general trends in ground-based GETs are being mirrored in on-orbit performance. The cause of the FUV sensitivity loss is still under investigation. Laboratory tests are being conducted to examine the effect of atomic oxygen on the CsI photocathode as one possible cause.

On-orbit measurements of the time-dependent sensitivity have been implemented in TDS reference files, which were delivered on July 14, 2010 for both the NUV and FUV gratings. These files correct for the NUV grating using a wavelength-independent value.
Figure 2: (top) Summary of time-dependent sensitivity trends in COS data obtained through early May 2010. The Y axis gives the values of the linear fit to temporal relative count rate data for each configuration noted. Low resolution grating data are shown in black, medium-resolution NUV data in red, medium-resolution FUV data in blue. For the FUV data, segment A (B) is shown with open (filled) symbols. (bottom) Close-in view of COS FUV TDS results. Symbols are as in the top panel.
Figure 3: Test of global vs. local sensitivity decline. The black stars are sensitivity monitoring data in the G140L/1230 configuration of the white dwarf standard WD0947+857. The red and blue stars correspond to spectra obtained at 1.2 and 3.0" from the nominal location, and do not exhibit any different change in sensitivity.
The correction for FUV data is segment- and grating-dependent, as well as wavelength dependent. The reference file should correct fluxes for time-dependent effects to ±2%. Version 18.2 of the ETC has been updated with sensitivities projected midway through Cycle 18. Further updates to the reference files will be made based on inspection of later data.

6. Recent Developments

Since the HST calibration workshop, there has been a change in the FUV TDS trends. While the sensitivity continues to decline, the wavelength dependence has disappeared and the segment- and grating-dependent differences have disappeared. Data taken since 2010.2 are now consistent with a decline of 4–6 %/year across all wavelengths. These new trends will be incorporated into an updated reference file; check the COS website (http://www.stsci.edu/hst/cos) and STScI Analysis Newsletter for updates.

References

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Observing with HST below 1150Å: Extending the Cosmic Origins Spectrograph Coverage to 900Å

Steve Osterman, Steven V. Penton, Kevin France, Stéphane Béland

Center for Astrophysics and Space Astronomy, Astrophysics Research Lab,
University of Colorado, Boulder, CO, 80309

Stephan McCandliss,
Department of Physics and Astronomy, Johns Hopkins University, Baltimore, MD 21218

Jason McPhate
University of California, Berkeley Space Sciences Lab, Berkeley, CA 94720

Derck Massa
Space Telescope Science Institute, Baltimore, MD 21218

Abstract. The far-ultraviolet (FUV) channel of the Cosmic Origins Spectrograph (COS) is designed to operate between 1130Å and 1850Å, limited at shorter wavelengths by the reflectivity of the MgF2 protected aluminum reflective surfaces on the Optical Telescope Assembly and on the COS FUV diffraction gratings. However, because the detector for the FUV channel is windowless, it was recognized early in the design phase that there was the possibility that COS would retain some sensitivity at shorter wavelengths due to the first surface reflection from the MgF2 coated optics. Preflight testing of the flight spare G140L grating revealed ~ 5% efficiency at 1066Å, and early on-orbit observations verified that the COS G140L/1230 mode was sensitive down to at least the Lyman limit with 10-20 cm² effective area between 912Å and 1070Å, and rising rapidly to over 1000 cm² beyond 1150Å. Following this initial work we explored the possibility of using the G130M grating out of band to provide coverage down to 900Å. We present calibration results and ray trace simulations for these observing modes and explore additional configurations that have the potential to increase spectroscopic resolution, signal to noise, and observational efficiency below 1130Å.

1. Introduction

The Cosmic Origins Spectrograph (COS), installed in the Hubble Space Telescope in May, 2009, was intended to provide high sensitivity, moderate to low resolution spectroscopy between 1130Å and 3200Å (Green 2003). In addition to meeting this goal, COS has demonstrated sensitivity down to wavelengths approaching 900Å (McCandliss 2010), providing coverage in the Far Ultraviolet Spectroscopic Explorer (FUSE) band at sensitivities comparable to individual FUSE channels. The nominal G140L/1230 mode provides coverage from the detector cutoff at 1850Å down to <910Å, and two new modes, G130M/1096 and /1055, provide higher sensitivities and potentially much higher signal to noise at these wavelengths.

The COS light path is shown in figure 1. Light from the Optical Telescope Assembly (OTA) enters COS through an oversized aperture admitting the entire aberrated wavefront
Figure 1: COS Light Path. The FUV channels are distinguished by having only one reflection and a windowless, two segment detector. The holographically-ruled aspheric FUV diffraction gratings provide dispersion, reimaging, and astigmatism and aberration correction in a single reflection. The two segment detector allows for one segment to be disabled to prevent overlight if asymmetric illumination is anticipated. (Froning 2009)

from a point-like source. The aperture is windowless so that any short wavelength (> 1130Å) light remaining after the two reflections in the OTA will travel unobstructed to the FUV grating. The gratings perform diffraction, aberration correction and focus in a single reflection in order to minimize reflections, maximizing short wavelength sensitivity. Light then travels to the windowless FUV detector. Laboratory testing indicated that the detector retains relatively high (>30%) quantum efficiency down to at least 800Å. While the COS diffraction gratings and the OTA mirrors are coated with MgF\textsubscript{2} protected aluminum (typically used for wavelengths longer than 1150Å), these optics were expected to retain some first-surface reflectivity below the MgF\textsubscript{2} transmission cutoff at ∼1150Å. This was verified for the G140L-C (flight spare) grating in laboratory testing (fig.2) (Osterman 2002), and for the OTA during COS on-orbit testing.

2. G140L Modes

2.1. G140L/1230

The FUV detector on COS is composed of two independently commandable 10 × 85mm segments, referred to as segments A and B (B is the shorter wavelength segment in all modes). Although the B segment is typically maintained a reduced voltage (HV-Low) for G140L/1230 setting, effectively disabling the short wavelength half of the detector, segment B was intentionally left at the operating voltage (HV-Nom) during the first calibration tests. It was immediately obvious that the instrument and the OTA retained significant sensitivity down to approximately 900Å. Measured sensitivity is shown in fig. 3, dotted line + triangles). Spectroscopic resolution (λ/Δλ) is expected to drop from ∼ 2300 at 1250Å to no better than 2000 at 900Å (McCandliss 2010).

2.2. G140L/800

The success of the G140L/1230-B observations suggested the possibility of shifting the wavelength scale to the blue so that the central wavelength (falling on the detector gap) is ∼ 800Å. This would place the entire G140L pass band on a single detector segment and eliminate the need for multiple observations to obtain full coverage. This requires a relatively large focus mechanism move, placing the focus mechanism (discussed below)
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Figure 2: G140L-C (flight spare) laboratory test results showing ~5% total efficiency (groove efficiency × reflectivity) at 1066Å.

Figure 3: Effective area of G140L and G130M modes at wavelengths below 1200Å. The G140L/1230-B segment (dotted line + triangles) provides coverage from >920Å-1170Å with sensitivity increasing rapidly above 1075Å. Instrument configurations using the G130M for sub-1150Å observations (solid line + circles, dashed line + stars) provide narrower coverage per detector segment. By placing the long wavelength edge of the G130M B segment below 1075Å we avoid the large variation in sensitivity across a single detector segment evidenced in the G140L/1230 mode. This permits observations of brighter objects, enabling higher signal to noise observations at shorter wavelengths. FUSE effective area ranged from ∼ 7–25 cm² per channel (Sahnow 2000).
outside of the nominal focus range for the G130M and G160M gratings, but it could be argued that this does not represent an unacceptable risk given the increased observational efficiency that this mode could provide.

2.3. Red Leak

One difficulty encountered in the G140L/1230-B mode (and anticipated for the G140L/800 mode) is the greater than two order of magnitude increase in the effective area with wavelength across the B segment of the detector, rising from $\approx 10 \text{ cm}^2$ at the shortest wavelengths to over $1300 \text{ cm}^2$ at the long wavelength edge. As a result of this sensitivity variation only relatively dim targets can be observed without triggering the bright object protection despite the low sensitivity at the blue end of the band pass, significantly reducing the obtainable signal to noise. While the G140L grating could be repositioned so as to place the higher effective area portions of the spectrum on the A (disabled) detector segment to obtained more uniform sensitivity, this would bring zero-order light onto the B segment (as is the case with the G140L/1105 setting), undoing any attempts to flatten the countrate on the B segment.

3. G130M Modes

The instrument development team had not originally considered any short wavelength configurations for the G130M grating because we anticipated significantly degraded resolution as we moved farther and farther out of band and because we had no model or test data for the grating efficiency at wavelengths far from blaze. However, in light of the better than expected performance in the G140L/1230-B segment, we began exploring the possibility driving the short wavelength cutoff of the G130M spectrum down to 900 Å. This had the potential to mitigate the unwieldy variation in sensitivity exhibited in the G140L/1230 mode and expected with the G140L/800 mode.

3.1. G130M/1055 and /1096

The COS FUV gratings are located $\sim 178 \text{ mm}$ from the center of rotation of the grating select mechanism (fig. 1). In addition to providing grating selection, this mechanism permits small adjustments to the band for each grating. This flexibility ensures that the full spectrum can be covered despite the gap introduced by the physical separation of the A and B detector segments. The grating select mechanism is in turn mounted on a linear translation mechanism to take out motion of the grating along the optical axis and to approximately accommodate changes to the location of the focal surface introduced by grating rotation. Larger than nominal moves are possible, but require substantial granting mechanism translations to correct for the focus offset. Large wavelength offsets may require a greater focus adjustment than the mechanism can provide, and even then would not fully recover the nominal instrument performance since the focal plane will no longer be coincident with the detector face. Also, outside of the nominal band the cross dispersion height is compromised.

Nevertheless, if reduced resolution is acceptable the gratings can be configured to support observations significantly outside of the design wavelength range. We proposed observing with the G130M grating rotated so that 900Å light would fall on the short wavelength edge of B detector segment and with a second position offset from the first by +41Å to ensure full coverage across the detector gap and overlap with the G130M/1291 mode. This configuration requires rotating the grating mechanism approximately 2.8°; the corresponding focus mechanism position that returns the best possible resolution is roughly a factor of two beyond the mechanism hard stops. Given the shallow slope of line width versus focus mechanism position, we chose to limit the focus offset to within the range used by existing modes. This represents a small reduction in resolution, but ensures that in the
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Figure 4: G130M/1055 and /1096 observations of GD561 obtained by COS.

Figure 5: Ray trace results for G130M/1055 and /1096 show (a) resolution decreasing at longer wavelengths and (b) cross dispersion width peaking near the central wavelength. Pixels are approximately 6 μm wide (dispersion direction) by 25 μm tall. OTA mid frequency wavefront error is not included in the geometrical ray trace model so that actual resolution is expected to be somewhat lower than model predictions.

event of a focus mechanism failure during an observation in one of these new configurations performance would not be as severely degraded in the in the nominal G130M modes as it would with the mechanism at the extreme range of motion.

The two new modes, G130M/1055 and /1096, provide continuous spectral coverage from 900 Å to the short wavelength edge of the existing G130M modes. Observations of GD561 were carried out over the third quarter of 2010 (fig. 4) and the effective area was determined to be higher than for the G140L/1230-B mode (fig. 3). Spectral resolution in these modes is under evaluation and appears to be consistent with ray trace predictions, dropping from approximately 2500 at the shortest wavelengths to 1400 at longer wavelengths (fig. 5 and table 1). While the large sensitivity variation is still apparent, with the G130M/1055 and /1096 modes the A segment of the detector can be disabled permitting observation of brighter objects to obtain higher S/N observations.

3.2. G130M-1215

In addition to the G130M/1055 and /1096 modes, we have performed ray trace modeling of a third new G130M mode, extending short wavelength coverage down to 1065 Å. This mode does not require the large grating rotation and corresponding large refocusing move demanded by the G130M/1055 and /1096 configurations as the short wavelength edge is now only 55 Å beyond the blue edge of G130M-1291. This off-nominal geometry compromises the best possible resolution and cross dispersion height, but to a much lesser degree than for the more extreme cases already explored, maintaining better than 16,000 resolution across
the band. This mode has the added advantage of placing geocoronal Ly-α on the detector gap.

4. Conclusion

The Hubble Space Telescope has provided spectacular imaging and spectroscopy longwards of 1150Å for over 20 years. The Cosmic Origins Spectrograph, while originally intended to provide spectroscopic coverage from 1130Å to 3200Å, has now extended the usable wavelength range of HST to the Lyman limit, providing spectroscopic access to wavelengths unobservable since the end of the FUSE mission in 2007. The capabilities of these new and proposed modes are summarized in Table 1.

<table>
<thead>
<tr>
<th>COS Mode</th>
<th>Wavelength Range</th>
<th>Effective Area</th>
<th>Modeled Resolution</th>
<th>Background (cts/resl/ksec)</th>
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<tbody>
<tr>
<td>G140L/1230</td>
<td>&lt;920-1160Å</td>
<td>~8-10 cm² (at 1000Å)</td>
<td>~2100 (1000Å)</td>
<td>0.3</td>
</tr>
<tr>
<td></td>
<td>1230-1850Å</td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>G140L/800</td>
<td>&lt;920-1850Å (a segment)</td>
<td>~8-10 cm² (at 1000Å)</td>
<td>~2100 (1000Å)</td>
<td>0.2</td>
</tr>
<tr>
<td>G130M/1096</td>
<td>940-1081Å</td>
<td>~15-25 cm² (b segment)</td>
<td>~2300 (1000Å)</td>
<td>1.3</td>
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<tr>
<td></td>
<td>1096-1238Å</td>
<td></td>
<td></td>
<td></td>
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<tr>
<td>G130M/1055</td>
<td>900-1041Å</td>
<td>~15-25 cm² (b segment)</td>
<td>~1800 (1000Å)</td>
<td>1.8</td>
</tr>
<tr>
<td></td>
<td>1055-1196Å</td>
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<td></td>
<td></td>
</tr>
<tr>
<td>G130M/1215</td>
<td>1065-1205Å</td>
<td>~30-2000 cm² (b segment)</td>
<td>~16,000 (1100Å)</td>
<td>0.1</td>
</tr>
<tr>
<td></td>
<td>1220-1360Å</td>
<td></td>
<td></td>
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</table>

Underlined modes are available or will be made available to observers in cycle 19. Modes in italic have not been tested and performance projections are based on modeling and on similar modes.

By expanding coverage to these shorter wavelengths, we make possible a range of investigations not previously accessible to HST, including studies of the Lyman continuum escape fraction from low redshift galaxies, of the Gunn Peterson effect at redshifts between 2 and 2.8 along multiple lines of sight, and observations of the O VI λλ 1032, 1038 doublet.

The higher resolution of the G130M/1215 mode could support observation of atomic and molecular diagnostics that could be used to study winds and atmospheres of massive stars, as well as the bulk of the mass in the translucent ISM out of which those stars form.

References

Performance of the Space Telescope Imaging Spectrograph after SM4

Charles R. Proffitt¹, A. Aloisi, R. C. Bohlin, K. A. Bostroem, C. R. Cox, R. I. Diaz, W. V. Dixon², P. Goudfrooij, P. Hodge, M. E. Kaiser², M. D. Lallo, D. Lennon³, S. Niemi, R. V. Osten, I. Pascucci³, E. Smith, M. A. Wolfe, B. York, W. Zheng²,

Space Telescope Science Institute, 3700 San Martin Drive, Baltimore MD 21212

T. R. Gull⁴, D. J. Lindler⁵, and B. E. Woodgate⁴

Code 667, Goddard Space Flight Center, Greenbelt Maryland 20771

Abstract.

On May 17, 2009, during the fourth EVA of SM4, astronauts Michael Good and Mike Massimino replaced the failed LVPS-2 circuit board on the Space Telescope Imaging Spectrograph (STIS), restoring this HST instrument to operation after a nearly 6 year hiatus. STIS after this 2009 repair operates in much the same way as it did during the 2001-2004 period of operations with the Side-2 electronics. Internal and external alignments of the instrument are similar to what they had been in 2004, and most changes in performance are modest. The STIS CCD detector continued to experience radiation damage during the hiatus in operations, leading to decreased charge transfer efficiency (CTE) and an increased number of hot pixels. The sensitivities for most modes are surprisingly close to what was expected from simple extrapolation of the 2003-2004 trends, although the echelle modes show somewhat more complex behavior. The biggest surprise was that the dark count rate for the NUV MAMA detector after SM4 has been much larger than had been expected; it is currently about 2.5 times bigger than it was in 2004 and is only slowly decreasing. We discuss how these changes will affect science with STIS now and in the future.

1. Introduction

STIS was designed as a replacement for the two original HST spectrographs, the Goddard High-Resolution Spectrograph and the Faint Object Spectrograph. These previous generation spectrographs had been limited to small 1-D sensors (512 pixel Digicons), while STIS was specifically designed to take advantage of availability of large, sensitive 2-D detector arrays.

STIS is a complex instrument with a large number of modes and options. The instrument includes three independent detectors (see Table 1). While only one of these detectors can be used at any given time, switching between them takes a minimal amount of time,

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¹Science Programs, Computer Sciences Corporation
²Department of Physics and Astronomy, Johns Hopkins University
³European Space Agency
⁴National Aeronautics and Space Administration
⁵Sigma Space Corporation
Figure 1: A detail from the STIS Early Release Observation OAC6010D0 of the nebula around the superluminous blue variable star η Car taken in June of 2009, using the 7283 CENWAVE setting of the G750M grating with the 52X0.1 aperture, illustrates the utility of spatially resolved long slit observations. The prominent acute-shaped emission lines near 7450Å are [Fe II] and [Ni II] emission lines from the Little Homunculus, ejecta from the 1890s first discovered by HST/STIS (Ishibashi et al 2003). In this figure, wavelength increases along the x-axis, and the y direction shows spatial structure along the length of the long slit.

so it is straightforward to combine observations using the different channels. Two of these detectors are Multi-Anode Microchannel Arrays (MAMAs), while the third is a CCD.

<table>
<thead>
<tr>
<th>Detector</th>
<th>type</th>
<th>dimensions (pixels)</th>
<th>≈ wavelength range (Å)</th>
<th>FOV (&quot;)</th>
<th>plate scale (&quot;)</th>
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<tbody>
<tr>
<td>FUV MAMA</td>
<td>CsI</td>
<td>1024 × 1024</td>
<td>1150 - 1700</td>
<td>25 × 24</td>
<td>0.025</td>
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<tr>
<td>NUV MAMA</td>
<td>Cs₂Te</td>
<td>1024 × 1024</td>
<td>1600 - 3200</td>
<td>25 × 24</td>
<td>0.025</td>
</tr>
<tr>
<td>CCD</td>
<td>SITe</td>
<td>1024 × 1024</td>
<td>1650 - 11,200</td>
<td>25 × 24</td>
<td>0.0507</td>
</tr>
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</table>

A number of gratings can be used with each of these detectors, and for each grating a large number of apertures are available. The first order gratings provide resolutions that vary between ≈ 600 and ≈ 1400 (depending on mode and wavelength), while the echelle modes provide resolution between ≈ 25,000 and 114,000, although with special observing techniques, resolutions of up to 200,000 can be obtained with the STIS E140H and E230H echelle modes (e.g., Jenkins, & Tripp 2001).

Most commonly, the first order grating modes are used with the long slits, allowing spatially resolved spectroscopy (Fig. 1), while the echelle modes are used with small apertures to minimize the overlap between the multiple spectral orders projected onto the detector (Fig. 2).

1.1. STIS History

STIS was originally installed into HST on February 14, 1997 during HST Servicing Mission 2 (SM2). The initial on-orbit performance of STIS is detailed in Kimble et al. (1998). The instrument operated using its primary Side-1 electronics until May 16, 2001 (42,000 hours), when a malfunction disabled the Side-1 electronics. This malfunction is believed to have been a hard short in a tantalum capacitor, and this failure rendered all of the Side-1 electronics unusable.

STIS resumed observations on July 7, 2001, using the redundant set of Side-2 electronics. There were two notable differences in the operation of STIS after the switch to Side-2.

One of these changes was the presence in the CCD channel of additional electronic read-noise in the form of a herringbone pattern with approximately 1 e⁻ RMS. The behavior
Figure 2: The flat fielded image for the STIS E140M echelle observation OB8709020 of the standard star BD+28° 4211 is shown using a logarithmic stretch. This figure illustrates the multiple spectral orders that can be simultaneously observed using the STIS echelle modes.

of this additional read-noise and techniques for removing it are discussed by Jansen et al. 2003.

The STIS Side-2 electronics also lack the ability to measure the temperature of the STIS CCD detector. On Side-1, the thermo-electric cooler (TEC) was operated in a closed loop to maintain the detector at a constant set point of −83 C, but the lack of the appropriate sensor makes this impossible on Side-2. Instead, the TEC is operated at constant current, and the temperature of the CCD sensor fluctuates as the thermal environment of the aft shroud influences the STIS instrument. These temperature fluctuations lead to significant variations in the detector dark current, including the brightness of hot pixels, on orbital time scales. However, there is a working sensor on the CCD housing, and it has proved possible to use this housing temperature to scale the dark current for these orbital variations (see Brown 2001 and Brown & Davies 2003). This scaling is applied in the STIS data reduction pipeline when constructing the STIS CCD dark reference files and when subtracting these reference darks from individual STIS CCD exposures.

Operations on Side-2 continued until August 3, 2004 (an additional 27,000 hours of operation), when a 5V power converter failed. The detectors were undamaged, but because this converter is necessary for the operation of essentially all the STIS instrument’s moving parts, there was no way to adjust the gratings and apertures to allow light to reach the detectors.

Following the 2004 failure, a plan was conceived to replace the Side-2 circuit board containing the failed converter during HST Servicing Mission 4 (SM4). Technical details of the planned repair are described in Rinehart et al. (2008). On May 17, 2009, this repair was successfully carried out by astronauts Michael Good and Mike Massimino, and STIS was returned to operation. Since only the Side-2 electronics were repaired, the performance changes described above that are unique to Side-2 continue during the new post-SM4 period of operations.
2. Changes between 2004 and 2009

Most internal and external alignments, the overall focus, and the mechanical and electronic functioning of STIS are similar to those seen during previous Side-2 operations. Some small zero-point offsets are compensated for by normal acquisition and wavecal procedures.

Significant fading of the STIS PtCr/Ne wavelength calibration lamps has been noted at the very shortest wavelengths (Pascucci et al. 2010ab). For a small number of STIS CENWAVE settings, the lamp choice and/or exposure times are being adjusted to compensate for these declines.

The STIS instrument does exhibit some degradation from its time on-orbit. Radiation damage is continuing to accumulate on the STIS CCD detector, resulting in increased dark current, more hot pixels, and increased effects from the decline in charge transfer efficiency. The throughput for most STIS channels showed modest declines after the hiatus, but not more than would have been expected from a simple extrapolation of previous trends.

The biggest surprise after STIS repair was the very high, but slowly declining, dark rate seen in the NUV MAMA detector.

Below we go into greater detail on some of these changes.

2.1. STIS CCD Performance

A detailed update on STIS CCD performance can be found elsewhere in these proceedings (Wolfe et al. 2010), so here we will concentrate on discussing the effects that these changes have on observers.

The herringbone pattern that appeared with the switch to the Side-2 electronics is still present with essentially the same characteristics and amplitude. However, the overall read-noise of the STIS CCD does appear to have increased by \( \approx 0.3 \) e\(^-\). This change appears to affect all gain setting and amplifiers, and unlike the herringbone noise does not appear to have a coherent pattern that would allow its removal via Fourier techniques.

The CCD dark rate has continued to increase, not only because of the accumulation of additional radiation damage, but also because temperatures inside STIS are higher than they were in the past due to the gradual degradation of HST’s thermal insulation. In 2010, the average CCD housing temperature has been about 4 degrees C higher than it was in 2001. Given the observed 7%/degree dependence of the dark rate on the housing temperature, this effect alone results in a 31% increase in the dark rate.

Even after correction for the temperature variations, the CCD dark current after SM4 shows considerably more scatter than previously (Fig. 3). The causes of this additional scatter are currently under investigation.

For observations targeting faint objects, the most serious effect of radiation damage may be the “tails” that appear on hot pixels and cosmic rays. These are caused by the reappearance of charge from short time-scale traps. Effectively, some of the charge lags a bit behind during the CCD readout, i.e., the charge transfer efficiency (CTE) is < 1 and decreasing over time. The tail size and length increases with the number of charge transfers performed during the readout, so these effects increase dramatically with increasing distance from the readout register. These tails are obvious when looking at a typical STIS CCD dark reference file image (see the vertical streaks in Fig. 4; the readout register is along the top in this figure). Note that because CTE losses in the serial readout register are small, there are no tails visible along the rows (horizontal direction in Fig. 4).

Because of CTE effects, there is a significant decrease in the effective background and noise if the spectrum can be placed closer to the serial readout, especially for faint objects. To this end, special “E1” aperture positions along the long slit were defined to put the spectrum near row 900 of the detector; this reduces the number of parallel transfers during the readout by about a factor of 4. There are a few disadvantages to the E1 positions. Since they lie between the long slit fiducial bars and the top of the detector, there is a smaller
Figure 3: Median CCD dark rate after rejection of hot pixels as determined from monitoring observations throughout the on-orbit lifetime of STIS.
unobstructed distance available along the slit. Also, there is some vignetting near the top of
the detector, that varies between gratings, making the absolute flux calibration less certain.
Finally, the quality of the G750L and G750M IR fringe flats for point sources at the E1
positions will not be as good. Near row 512, the 0.1x0.09 aperture can be used to provide
fringe flats with a PSF very similar to that of point source observed in the long slit. If the
slit wheel were to be simply rotated to put this aperture near row 900, the rotation of the
slit wheel will have shifted it too far along the dispersion direction to allow the resulting
fringe flat to be well aligned with the external spectrum.

2.2. Instrument Throughputs
The dependence of STIS throughputs as a function of time and wavelength have been
tracked since 1997, using the same set of standard stars. The observed changes are believed
to result from a combination of optical contamination (primarily organic molecules darkened
by exposure to UV light) and possibly some degradation of the detectors or small changes
in the grating tilts. These time-dependent sensitivity (TDS) changes are monitored by
periodic observations of selected standard stars. Results from before the 2004 STIS failure
were summarized in Stys et al. (2004). While the detailed dependance at a given wavelength
Performance of STIS after SM4

Figure 5: Time dependent changes in the throughput of STIS low dispersion modes.

varies with channel, for a given detector, applying the degradation derived from the low-dispersion modes to the higher resolution gratings used with that detector has been an adequate approximation to past behavior.

Figure 5 shows the history of the time dependent sensitivity changes for a typical wavelength bin of each of the STIS low dispersion modes. The upper two panels show the results for the MAMA G140L and G230L gratings, while the lower three panels show results for the STIS CCD G230LB, G430L, and G750L gratings. Starting around 2002, the rate of decline of the STIS UV throughputs appears to have flattened, and degradation between 2004 and the post-SM4 period appears to be consistent with a simple projection of those trends. Further details are given in Osten et al. 2010.

The behavior of the STIS E140H grating may be a possible exception to the trends shown in Figure 5. This grating initially seemed to show substantially lower throughput (by 15 - 20 %) than a simple extrapolation of the previous trends would have suggested. However, the grating throughput appears to have recovered from this anomaly within a few months after SM4. Additional details of this behavior, along with a discussion of changes in the echelle blaze function alignment, can be found in Bostroem et al. (2010) elsewhere in these proceedings.

2.3. NUV MAMA Dark Rate

The dark rate seen by the STIS NUV MAMA detector is believed to be dominated by a phosphorescent window glow. A simple model for this glow was initially presented by Kimble (1997) and also by Jenkins (1997), and was further discussed by Ferguson & Baum (1999). Metastable states resulting from impurities in the window become populated by charged particle impacts; these states do not decay, but instead are depopulated by thermal transitions to unstable states, which produce the UV photons that cause the background. This process is illustrated in Figure 6. The need to thermally excite the meta-stable state to depopulate causes the dark rate over short time scales to be an exponential function of the window temperature. However, over longer time scales, keeping the detector window cold causes a large population of meta-stable states to build up, and when the detector is
subsequently warmed back up, a large, but temporary, increase in the window glow results until a new equilibrium is reached. Based on pre-SM4 modeling, it had been expected that the dark rate would be initially enhanced by a factor of 2 to 3 above the pre-SM4 long term equilibrium mean rate of about 0.0013 counts/pixel/s, but would come back to near that equilibrium rate within two to three weeks. Instead the dark rate was found to be as high as 0.015 counts/pixel/s, and the subsequent decline has been much slower than expected. As of fall 2010, the mean dark rate was about 0.0034 counts/s/pixel, about 2.5 times the mean rate of seen in 2004, and the time scale for further decreases appears to be very long. Additional details of the history of the STIS NUV MAMA dark rate, as well as a comparison with the COS NUV MAMA, can be found in Zheng et al. (2010) elsewhere in these proceedings.

3. Future Calibration Plans

Following the completion of the initial post-SM4 verification observations, the Cycle 17 STIS calibration program was initiated to perform monitoring and ongoing calibration of instrument performance and to fill in any gaps left by the initial calibrations. Full details of each Cycle’s STIS calibration plan can be found at \url{http://www.stsci.edu/st/stis/calibration/}.

In addition to the routine monitoring observations, there are a number of ongoing activities of particular note:

- Flux recalibration of all STIS echelle settings to allow shifts in the echelle blaze function to be determined – see Bostroem et al. (2010) and HST CAL/STIS program 11866 for further details.

- Improvements to CCD dark current scaling and subtraction are underway (see also Jansen et al. 2003 for a discussion of the STIS herringbone pattern noise). In the near future, we also hope to consider whether pixel-based corrections for charge transfer efficiency losses, similar to those being developed for ACS by Anderson & Bedin (2010), might also be usefully applied to STIS CCD detector data.
• New observations to determine the stability of the MAMA pixel-to-pixel flats and to derive new STIS CCD pixel-to-pixel flats (Niemi et al. 2010, and HST CAL/STIS proposals 11852, 11861, and 11862).

• Improvements to our understanding of STIS wavelength calibration lamp performance and dispersion relations (Pascucci et al. 2010ab, and Ayres 2010).

4. Summary

Overall, the performance of STIS after its repair during SM4 is, in most ways, close to that seen during the previous period of operations between 2001 and 2004 using the Side-2 electronics. Most of the changes seen are indeed close to extrapolations of previous trends. It is hoped that STIS will remain an important part of HST’s instrument complement for many years to come.

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Ironing Out the Wrinkles in STIS

Thomas R. Ayres

Center for Astrophysics and Space Astronomy
University of Colorado, Boulder, CO 80309

Abstract.

The echelle wavelength scales of Space Telescope Imaging Spectrograph were evaluated based on a novel high-density line list for the on-board Pt/Cr–Ne emission lamps. The new reference wavelengths were obtained by a bootstrapping technique that exploited the space-borne instrument as its own “laboratory spectrometer.” A number of strategies were explored to mitigate subtle wavelength scale deviations identified in the process (known from earlier work), either by modifying the pipeline dispersion relations directly, or by a post-facto distortion correction. The main conclusion is that the STIS echelle wavelengths can be significantly improved with only modest changes to the current dispersion model.

1. Overview

Since its installation in Hubble more than a decade ago, Space Telescope Imaging Spectrograph (STIS: Woodgate et al. 1998) has been the premier high-resolution ultraviolet spectrometer in space. As such, it has conducted a wide diversity of observing programs on planets, stars, gaseous nebulae, interstellar matter, and even extragalactic sightlines. STIS is especially capable of measuring accurate radial velocities of narrow emission and/or absorption lines, and recording often subtle profile distortions in the spectra of astrophysical objects shaped by a variety of kinematic processes.

Because of its high resolving power, however, STIS is crucially reliant on a robust wavelength calibration, so that the high-precision velocity measurements attainable in principle, can be achieved in practice. For that purpose, the instrument carries a set of wavelength calibration lamps: Pt/Cr–Ne hollow-cathode discharge sources that emit a rich spectrum of very sharp lines in the 1150–3200 Å range where the STIS echelles operate.

The “Wrinkles Project” (aka, “Ironing out the Wrinkles” – Cycle 17 AR–11743) was designed to test the performance of the STIS wavelength scales and identify improvements if warranted. Wrinkles is a follow-on from a Cycle 13 archival program “The Deep Lamp Project,” which examined STIS echelle calibration exposures (“wavecals”) to assess internal precision and absolute accuracy of the pipeline-assigned wavelength scales (Ayres 2008). The method in both projects was to process long-duration exposures of the on-board lamps as if they were ordinary science data; measure positions of high-S/N lines in the calibrated flux density tracings; and compare the recorded values with laboratory wavelengths. The earlier study identified systematic distortions — sometimes subtle, sometimes not — in many of the supported grating ‘tilts.’ An approach had been developed (for the Cycle 14 Legacy Archival project “StarCAT”\(^1\); Ayres 2010) to correct the distortions post facto, but a central goal of Wrinkles was to evaluate whether an upgraded dispersion model could be implemented in the pipeline itself.

\(^1\)See: http://archive.stsci.edu/prepds/starcat/
An important advance proposed for Wrinkles was to exploit new laboratory line lists derived specially for the STIS lamps by a partnership between the U.S. National Institutes of Standards and Technology (NIST) and the Space Telescope European Coordinating Facility’s STIS Calibration Enhancement group (see Kerber et al. 2006). This was viewed as critically important for the NUV band, where a rich spectrum of Cr I and Cr II lines appears, in addition to the first and second spectra of Pt and Ne in traditional “platinum lamps,” but could not be utilized directly owing to lack of supporting laboratory wavelengths. Unfortunately, the ST-ECF effort appears to be in stasis as of this writing.

2. Laboratory Self-Calibration

Because of lack of progress on the ST-ECF side, the Wrinkles project had to compensate for the pivotal missing Cr I and Cr II calibration material by deriving its own “laboratory” spectrum. This was accomplished as an integral part of the overall analysis, as follows.

First, all the available STIS wavecal exposures longer than 60 s in duration were de-archived (including those taken for re-commissioning STIS following its successful repair during Servicing Mission 4 in May 2009). Examples of representative raw frames for E140H (λ1234) and E230H (λ1763) are illustrated in Figure 1. These tilts are at the extreme short ends of the FUV and NUV bands, respectively. Each echellegram covers ~300 Å in ~40 partially overlapping “orders.” The lamp lines (dark spots) were superimposed on a WD spectrum (light stripes) to show the orders more clearly. The shortwavelength ends of each frame (bottom for E140H, top for E230H) can be recognized by increased crowding of the orders. In both cases, the exposures are not deep enough to capture many lamp features in these high orders, so the dispersion solutions would be less well constrained there.

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2STIS has two general working wavelength ranges for echelle spectroscopy: FUV (1150–1720 Å) and NUV (1650–3200 Å), corresponding to the two MAMA cameras in the instrument.

3Chromium was introduced into the STIS lamps to increase the calibration line density at longer wavelengths.
There were 320 datasets in all, counting subexposures: 61 for the single E140M tilt, 69 distributed among the eleven E140H settings, 68 for the six E230Ms, and 122 for the twenty-six E230Hs. About a quarter of these proved to be unsuitable for the wavelength analysis because they had been taken through inappropriate apertures.

Figure 2 depicts the available exposure depth for each setting, counting only wavecals taken through the narrow “spectroscopic” apertures. The total cumulative exposure is 37 hours. Each block corresponds to an individual mode/tilt: the width indicates the wavelength coverage, and the height is the relative exposure depth. Dark shaded, color-coded blocks in each of the four sections refer to the prime settings. Thinner, yellow-shaded tilts are ones judged to have poor depth (and candidates for new calibration exposures).

Next, a slightly modified version of the original Deep Lamp pipeline was used to process the datasets, including the cases of multiple subexposures (avoided for simplicity in original Deep Lamp). This involved rewriting the FITS headers of the o*.raw wavecal files to reset specific processing keywords so that calstis treated the lamp exposure as if it were a normal science image, but taking care to deactivate certain options, like Doppler compensation for the spacecraft motion, which normally are invoked for observations of external targets. Fittingly, the “WAVELINE” (o*.wav file, a short [5–10 s] exposure of the lamp preceding a science exposure to set the wavelength zero-point) is a copy of the wavecal itself (first subexposure in case of multiple subexposures). The general procedure has been described by P. Hodge and colleagues (1998).

In the calstis processing, a new cross-correlation template initially was implemented, based on the preliminary full-coverage H-mode lamp “atlas” constructed as a byproduct of the Deep Lamp project. In tests, described below, only minor differences were found compared with the historical template used in calstis, although the default version does exhibit conspicuous defects in a number of places, as illustrated later (Fig. 4).

The post-processing scheme developed for StarCAT then was applied, in various guises specialized to subsequent analysis steps. The simplest level retained the full 2D order-resolved format of the x1d file, either with or without the StarCAT wavelength distortion correction (as described by Ayres 2010). In both cases, the x1d spectra were re-sampled to twice the spectral density (to now four points per resolution element [resol]), and the photometric error curve, $\sigma_\lambda$, was modified slightly to agree better with Poisson expectations at low counts (as described by Ayres 2010). The uncorrected 2D spectral datasets were the raw material for a new distortion correction, while the corrected echellegrams were used to judge the success of the distortion compensation over the full 2D format.

The second level of post-processing involved merging the orders of the distortion-corrected x1d pipeline file into a coherent 1D spectrum for the particular setting, along the lines described in the StarCAT study. The order-overlap zones benefited directly in higher S/N, but more subtly in averaging over presumably uncorrelated residual wavelength distortions, and fixed pattern noise, at opposite sides of the echellegram.

Subsequent steps in the StarCAT protocol involved, first, co-adding subexposures, registered by cross-correlation; then, combining like exposures taken in different epochs. This scheme was applied to the 2D and 1D spectral collections alike. A final processing step, solely for the 1D traces, spliced together all of the individual H-mode wavelength segments (or the M-modes) to yield a coherent “atlas” of the calibration lamp over the full 1150–3150 Å range.

4The default slits are 0.20×0.09 (in arcseconds) for H settings and 0.20×0.06 for M tilts, although the nonstandard 0.10×0.09 aperture often was used for E230H to mitigate order crowding. In several cases, particularly the 3 prime M settings, the 0.20×0.20 “photometric” slot was utilized. These exposures tend to be brighter, owing to $\sim 3\times$ higher throughput compared with the default “spectroscopic” slit, but the line profiles also are broader and more “U-shaped” than their sharp Gaussian cousins from the narrow apertures.

5Collection of traced spectra for each echelle order of a setting, the main output product of calstis.
Figure 2: Wavecal exposure depth in each supported STIS echelle mode/setting.
When combining like spectra from different epochs, the wavelength registration step was skipped, because there was no reason to prefer one epoch over the others. In the initial subexposure co-addition, the registration was performed, because the reference “WAVE-LINE” for each subexposure defaulted to the leading one of the set, which then would not account for any thermal drifts of the subsequent ones. The post facto cross-correlation procedure takes care of this. (However, any exposure that had an anomalously large shift compared with the majority of a group was excluded from consideration.) Further, in the splicing procedure that combined, for example, the 37 individual high-resolution echelle settings (11 FUV E140H, 26 NUV E230H), there also was no cross-segment wavelength registration (unlike in StarCAT), to avoid biasing the final spectrum by tagging it to positions of a few specific lines across the full range. At most wavelengths the H-spectrum is a combination of several independent echelle settings, recorded on different sets of pixels at different positions on the camera, and each individually corrected for wavelength distortions. Because of this, the final spliced spectrum in principle should have higher wavelength precision than any of the individual segments.

Another important departure from the StarCAT protocols involved the treatment of spectral flux densities at the different layers of co-addition. In StarCAT, the calibrated $f_{\lambda}$ were weighted in a average by $\sigma_{\lambda}^{-2}$ which can be shown to be equivalent to weighting by net counts (see Ayres 2010). This assumes constant intrinsic flux densities (i.e., no variability), regardless of exposure time. For the lamp spectra, however, the intrinsic fluxes are not independent of exposure time, epoch, or aperture; the former two because the lamp output is not constant (changing as the discharges ‘warm up’), and the last because the lamp is a diffuse source. To counteract any bias, the flux densities, $f_{\lambda}$, and photometric uncertainties, $\sigma_{\lambda}$, of a group of lamp exposures selected for co-addition were multiplied by a scale factor to bring into agreement a specific bright line in common to the exposures. The scale factor was set relative to the observation having the highest S/N at the reference feature. A consequence of the procedure is that the independent spectral segments might be offset in flux density from one another, to varying degrees. These offsets were cleared, however, in the final splicing step according to a bootstrap fluxing procedure developed for StarCAT. The flux bootstrapping worked very well, especially for the highly overlapping H settings, because there always were sufficient bright lines in the overlap zones to determine an accurate flux density ratio between adjacent segments.

Figure 3 illustrates a page from an atlas of the resulting H-mode spectrum. The tracings were slightly Gaussian smoothed for display purposes. Note the highly logarithmic ordinate scale. This is required to simultaneously show the strongest lamp lines, together with those at the horizon of detectability, in the face of the remarkable dynamic range achieved in these co-added spectra. Red vertical ticks mark the measured features. Annotations in black are from a hybrid line list derived from NIST laboratory measurements; red are Cr I lines from Wallace & Hinkle (2009); and blue are preliminary identifications based on Atomic Line List v2.04 (more detailed descriptions of the atomic data sources are provided later). The green curve is a heavily smoothed $5\sigma$ photometric noise level. Only features with peak S/N at or exceeding that threshold were measured in the H tilts.

Figures 4 compares the original calstis cross-correlation template (red curve) with one derived from the new H-spectrum (dots), illustrating a number of conspicuous flaws in the former (note lines missing in the old template, e.g., $\lambda 1426$). To be sure, this is the worst example of defects in the pipeline template; for the most part, the agreement between new and old is good. When constructing the new template, the H-spectrum was interpolated onto the wavelength scale of the original pipeline version, and the flux densities were scaled by the ratio of the total integrated intensity of the original spectrum to that of the new one. Thus, the new template has the same intensity units as the original, but a somewhat different (likely more accurate) distribution over wavelength.

Following spectral processing, every sharp lamp line with peak S/N $\geq 5$ was measured in the two sets of 2D records (i.e., the 44 individual frames distortion-corrected or not),
Figure 3: Page from an atlas of highest-resolution version of STIS lamp spectrum.
Figure 4: Comparison of new and old cross-correlation templates.
and the single 1D spliced H-spectrum, utilizing an autonomous Gaussian fitting “robot.” The automaton was taught to ignore overly broad features, and weak peaks on the flanks of stronger lines. In general, it was very successful in capturing the truly sharp, isolated emissions prized for calibration, without direct operator intervention (except to ‘tweak’ the governing parameters).

The robo-fitting of the H-spectrum yielded an extensive, refined tabulation of the STIS lamp features. A hybrid laboratory list based on the Sansonetti et al. (2004) measurements of STIS-type lamps (including Cr I and Cr II lines) below 1800 Å, and the earlier NIST study by Reader et al. (1990) of GHRS-style lamps (lacking appreciable chromium) for the longer wavelengths, contains 4335 entries, but some of the features are below the S/N$\geq 5$ threshold in the co-added Wrinkles lamp atlas, and thus are not useful for defining distortion corrections. The new Wrinkles line positions then were matched to those in common to the hybrid Sansonetti/Reader list (lacking Cr measurements above 1800 Å) and a more recent study of specifically Cr I by Wallace & Hinkle (2009) utilizing the high-precision Kitt Peak Fourier transform spectrometer (FTS), although their wavelengths were restricted to the mid- and long-NUV portion of the STIS band, above 2360 Å.

In comparison to the laboratory measurements, the Wrinkles atlas lines displayed long range systematic behavior, albeit at a low level (few hundred m s$^{-1}$). The systematic behavior was assumed to be intrinsic to the Wrinkles material, rather than to the laboratory wavelengths, because there was good internal agreement between the deviations displayed by the NIST-classified lines (Pt and Ne) on the one hand, and the Wallace & Hinkle lines (Cr I) on the other, over the wavelengths in common ($\lambda > 2360$ Å, as mentioned above). A high-order, low-amplitude polynomial was fitted to the deviations and applied to all the Wrinkles wavelengths to put them on the NIST/W–H scale. The corrected Wrinkles list coincides with the corresponding NIST and W–H features (i.e., the average velocity difference is zero) with a standard deviation of just $\sim 200$ m s$^{-1}$ ($1\sigma$), about twice as good as the quoted uncertainties in the laboratory work (suggesting that the latter are conservative).

The procedure was repeated to obtain reference wavelengths for M-mode features, although instead of using the M-spectrum directly (one E140M plus six E230Ms), the high-precision H-spectrum was utilized, after Gaussian smoothing it to the equivalent M-mode resolution. This ensures that reference wavelengths applied to derive distortion corrections for the M modes have the heritage of the heavily averaged high-resolution spectrum, which itself can be tied directly to the original laboratory measurements recorded at comparably high resolving power. Accordingly, the final M line list was corrected for the subtle long range systematic trend identified in the corresponding H-spectrum list.

The smoothing procedure required separate resolution matching for the FUV and NUV regions, owing to the $\sim 50\%$ higher M resolution of the former. The break point was taken at 1650 Å. In addition, a bright spike at 1187.5 Å was edited out of the H-spectrum prior to smoothing. This feature is an unflagged intermittent detector hot spot on the FUV MAMA, described in the StarCAT study, which survives averaging because there are only two independent spectra that contribute to the sum at these short wavelengths. It probably is present at other wavelengths in the merged spectra, but less obvious because of more widely spaced orders (bright spot perhaps falls in the inter-order background) and the larger number of independent spectra contributing at each final wavelength.

Beyond the numerous matches with the two fundamental reference laboratory sources, the Wrinkles list contains many entries that are not in the NIST and W–H tabulations, mainly the Cr I and Cr II lines that the ST-ECF effort was intended to measure. Most of

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6 The S/N$\geq 5$ criterion translates to a maximum “photometric” measurement error of about 500 m s$^{-1}$ in the line positions in equivalent velocity units. The quoted uncertainties of the lab wavelengths are at the level of 400 m s$^{-1}$, although there is evidence that the true 1$\sigma$ uncertainties are smaller.
the “new” features can be assigned preliminary identifications by reference to Cr line lists from atomic databases such as maintained by the University of Kentucky (Atomic Line List v2.04\textsuperscript{7}). These new features, regardless of their exact identities, can be utilized — treating the observed wavelengths as “laboratory” values — together with the ones corresponding to the “original” NIST features (but now using the new H-spectrum wavelengths) to refine the distortion maps.

In essence, STIS became its own laboratory spectrometer; bootstrapping from the existing line lists to obtain a better distortion correction, and from that an improved spliced H-spectrum, and from that new, more accurate measurements of the lamp lines. The process was iterated a few times to ensure optimum results. It worked because: (1) the STIS echelle dispersion relations are very linear (\S 4 below); and (2) the apparent distortions that survive the pipeline processing are large scale and systematic (as shown later).

Figure 5 illustrates schematically the improvement obtained for the H settings (2D spectra). The top panel represents uncorrected \texttt{calstis}, as one would routinely retrieve from the MAST archive. The image depicts the density of points (wavelength deviations, \(\lambda_{\text{obs}} - \lambda_{\text{lab}}\) expressed in equivalent Doppler shift), as binned in wavelength and velocity shift. Solid symbols represent averages over 50 consecutive measurements, to illustrate any systematic behavior. The middle panel is for uncorrected \texttt{calstis}, but with the new cross-correlation template described earlier. The significant deviation at 1750 Å in the top panel is muted somewhat. The bottom panel is for distortion-corrected \texttt{calstis} with the new template. The deviations now are smaller and more uniform, and likely reflect mainly unavoidable measurement errors on both the STIS and laboratory sides. Note that the H mode resol (2.6 km s\(^{-1}\)) is comparable to the full extent of the y axis in each panel.

### 3. Enhanced Distortion Correction

Based on experience developed during this analysis, it was decided to modify the way in which the distortion maps were constructed and applied. Previously, the maps were based on a linear grating parameter “\(k\)” (\(k \equiv m\lambda\)) on the \(x\) axis (the grating parameter varies, in principle, exclusively in the dispersion \([x]\) direction) and a linear order “\(m\)” scale on the \(y\) axis. However, the orders in the actual echellegram have a nonlinear \(y\) spacing, which should be maintained in the correction maps to properly capture any coherent spatial distortions related to the detector geometry. A new mapping was undertaken, replacing \(m\) with a spatial coordinate \(\tilde{y}\) that reflects the true nonlinear order spacing. The \(\tilde{y}(m)\) function was determined by measuring mean order positions (the orders are slightly tilted) in calibration spectra of the DA white dwarf G191-B2B (which had been recorded in all 44 supported echelle settings, and has a bright, nearly line-free continuum spectrum). Further, both independent variables were treated differentially: \(\Delta k = k - k_0\) (with a range of approximately \(\pm 1000\)) and \(\Delta \tilde{y} = \tilde{y} - \tilde{y}_0\) (with a range 0–1023 [low-res pixels]), where \(k_0\) is a constant for each mode and \(\tilde{y}_0\) is a constant specific to each setting. These new values were further normalized by a factor of 500 to put the difference variables into ranges of order unity, to promote more stable numerical solutions. (The dependent variable describing the local distortions was equivalent velocity shift, \(v\) [km \(^{-1}\)].)

The distortion maps were modeled empirically by fully (bi-)cubic functions (16 terms) for the H settings, and by 5th-order polynomials (36 terms) for the denser (in orders per tilt) M settings. In addition, the 2D polynomial corrections were “damped out” beyond certain limits in \(\Delta k\) and \(\Delta \tilde{y}\), such that at least ten measurements contributed to the solutions beyond the cutoffs in each axis to avoid edge effects (i.e., wild variations of the polynomials beyond where they might be constrained by measurements, but still within the valid wavelengths

\textsuperscript{7}See: http://www.pa.uky.edu/~peter/atomic/
Figure 5: Effects of different processing on 2D order-resolved wavelength deviations.
of the order where the corrections might be applied). Figure 6 illustrates the approach for two representative settings: one M, the other H.

In these figures, the top frame depicts the distribution of measured lamp lines in $\Delta k$ (x-axis) and $\Delta \tilde{y}$ (y-axis). The size of the symbol is proportional to the absolute velocity shift, and the color codes whether the shift was negative (blue), positive (red), or below a threshold (small open symbols). The velocity cutoffs were 250 m s$^{-1}$ for the E140H and E230H tilts, 500 m s$^{-1}$ for E140M-1425, and 750 m s$^{-1}$ for the E230M settings. These are approximately the limiting velocities that could be measured in principle for a feature of $S/N \sim 10$. The maximum velocity shifts typically are only about a third of a resol, $\sim 1$ km s$^{-1}$ for H settings, and a few km s$^{-1}$ for M tilts (see, e.g., Fig. 5 for H). The middle frame visualizes the polynomial distortion model derived from the spatial distribution of velocity offsets in the upper panel, where again blue refers to negative velocities and red to positive. The color scales saturate at $\pm 4\times$ the velocity cutoffs mentioned above. The bottom panel illustrates the residual displacements after application of the distortion correction, with the same scaling as the top panel. The orange outlined area indicates boundaries of the region beyond which the models were forced to take on the values along the periphery. Figure 7 summarizes analogous distortion maps for the 44 supported STIS echelle settings. Orientations of the individual frames are the same as in Figs. 1 and 6.

4. Improvements to the Standard Dispersion Model

The next step was to explore whether improvements might be achieved simply by introducing new terms into the existing calstis dispersion model, or whether a completely different approach altogether was warranted. As a point of reference, below is the polynomial model incorporated in calstis,

$$S = a_0 + a_1 (m \lambda) + a_2 (m \lambda)^2 + a_3 m + a_4 \lambda + a_5 m^2 \lambda + a_6 m \lambda^2 + a_7 (m \lambda)^3$$

where $S$ is the “sample” coordinate (echellegram “x” axis) in low-resolution pixels, $m$ is the order number, and $\lambda$ is the wavelength (in ˚A). Only coefficients up to $a_6$ are populated in the STIS calibration library for echelle settings. Note, also, that the model is not fully bi-quadratic because $m^2$ and $\lambda^2$ are missing. Smith (1990) has discussed the venerable IUE dispersion relations, upon which the STIS (and GHRS before it) polynomial models were based, including motivation for specific terms from the echelle grating equation.

To carry out dispersion modeling experiments, one needs the distributions of lamp line pixel positions, $S$, as a function of order. Conveniently, calstis traces the spectrum in raw pixel space, then assigns wavelengths to the pixels by inverting the dispersion relation to infer $\lambda(m) = f(S,m)$. Thus, it is simply a matter of re-fitting each lamp spectrum in the native pixel coordinates, rather than assigned wavelengths. The derived line pixel positions should then be in the same coordinate system as the calstis dispersion relations, modulo perhaps a constant shift.

As a first experiment, motivated by the Deep Lamp study, the calstis model was reformulated using the ostensibly orthogonal independent variables $k \equiv m \lambda$ and $m$ (see Smith 1990). The grating parameter $k$ in principle varies exclusively along the x-axis, and the order number varies (trivially) exclusively along the echellegram y-axis, leading to the desired orthogonality. However, a fully bi-quadratic solution (with $\Delta k = k - k_0$, where $k_0$ was taken as constant for each mode) did not significantly improve $\chi^2$ with respect to a fully bi-quadratic model with terms in $\lambda$ and $m$ (which is the calstis formulation dropping $[m\lambda^3$, but adding $m^2$ and $\lambda^2$). The reason for the somewhat disappointing performance will became apparent later.

Figure 8a illustrates the success, or lack of same, of progressively more complex polynomial dispersion models (in $k,m$) applied to the STIS echelle settings. Each shaded box depicts $\chi^2$ (actually, the “reduced” $\chi^2$ [per degree of freedom]) of the fit for the particular
Figure 6: Left: 5th-order distortion map, E140M-1425. Right: 3rd-order for E230H-2762.
Figure 7: Distortion maps for the 44 supported STIS echelle settings.
setting. Prime tilts are indicated in red at left. The lighter the $\chi^2$ shade, the better the fit. The best examples approach the optimum $\chi^2 = 1$. The first column ("orig calstis") was obtained by taking the normal pipeline dispersion coefficients and calculating the $S$ values according to $A$ and $m$ from the 2D measurements of each setting, then comparing the predicted pixel values with the measured ones (as described above). Before evaluating $\chi^2$, the average deviation was subtracted, to account for the fact that the constant coefficient in the pipeline dispersion relations is essentially arbitrary because, in practice, that coefficient is adjusted during the processing to match the zero point indicated by the associated WAVELINE exposure (by the template cross-correlation procedure alluded to earlier).

In general, the $\chi^2$ for the prime settings appear to be the best (with the possible, although possibly glaring, exception of E140H-1416), but for the secondary tilts generally are worse (and in some cases, very bad). Apparently, coefficients for several of the secondary settings with poor ground calibrations were simply "extrapolated" from neighboring, better characterized prime tilts. In principle, that is a reasonable strategy, but in practice it does not appear to have worked very well. The reasons for this will be described shortly.

The second column ("new calstis") was the result of re-deriving the dispersion coefficients for the pipeline polynomial, but using the new, more extensive sets of on-orbit measurements now available for the different settings. This provided a fairer baseline for later comparisons than the original calstis coefficients, and illustrates the extent to which that model is capable of reproducing the dispersion properties of each setting, given the spe-
pecific (new) wavecal material available. The subsequent columns illustrate the improvement, if any, achieved with higher order polynomials, up to 5th-order. Note that for nearly all the settings, there was not much improvement going from the calstis model to fully bi-quadratic (2 additional terms). The 3rd-order polynomial made a noticeable difference for a few of the tilts, mainly the E230Ms. But, aside from E140M, going to 4th- or 5th-order did not improve the $\chi^2$ appreciably (note that it can even worsen slightly at high $n$ owing to the decrease in the degrees of freedom tied to the rapidly increasing number of coefficients in the polynomial). The conclusion is that the relatively simple pipeline dispersion model actually performs pretty well, at least when the coefficients are derived explicitly for each setting.

Figure 8b illustrates various permutations of the calstis polynomial, itself, to test whether modest changes might significantly improve performance. The first column is “new calstis” (same as second column in Fig. 8a, but now with a more compact color table). The second column of Fig. 8b turns on the currently unused (at least for the echelle settings) eighth term in the calstis model ($[m \lambda]^3$). Except for a few of the E230M tilts, this term apparently is not very potent. The third and fourth columns illustrate what happens when either $m^2$ or $m^3$ is added to the pipeline model. Both of these terms appear to be equally improving for many of the settings, especially the numerous E230Hs, but have not much influence on the tilts where $(m \lambda)^2$ was more effective. The fifth column activates $(m \lambda)^3$ and $m^2$ together (equivalent to adding a single new term to calstis, if the coefficients for the existing $[m \lambda]^3$ are populated). This does as well, or in some cases better, than a fully bi-quadratic model (last column). (The bi-quadratic model would add $m^2$ and $\lambda^2$ to calstis, but take away the $[m \lambda]^3$ term.) The conclusion here is that a simple modification to the existing calstis dispersion relation could significantly improve performance for all the tilts.

5. Orthogonal Variables and Global Models

There still, however, was the lingering issue of why the supposedly orthogonal independent variable set $(k, m)$ did not produce the level of improvement over the decidedly non-orthogonal variables $(\lambda, m)$ anticipated from general principles. The reason became obvious with a simple consideration of the behavior of the dispersion properties of the individual settings. Figure 9 illustrates the representative case of E140H. The left panel displays, order by order, parabolic fits to line centroids ($S$: low-resolution pixels) as a function of $k - k_0$,

$$
\left[ \frac{S}{500} \right] = a_0 + a_1 \left[ \frac{(k - k_0)}{500} \right] + a_2 \left[ \frac{(k - k_0)}{500} \right]^2,
$$

(2)

for all orders with ten or more measurements (to ensure an accurate polynomial solution) in all eleven tilts of that mode. If the $k$ values truly were orthogonal to $m$, the curves would fall exactly on top of one another, which clearly they do not.

The lower right panel depicts the coefficients of the quadratic fits. Note that the $S(k - k_0)$ relation is dominantly linear, with a small but non-negligible parabolic term. Both the linear and parabolic terms appear to be nearly independent of $m$ in a given setting, and among the settings as well. The constant term, in contrast, displays a clear dependence on $m$ for each tilt (slanted curves at top of lower right panel).

Another way of looking at this is illustrated in the upper right panel, which depicts the variation of the inferred central $k_0$ with $m$ for each setting. (The parameter $k_0[m]$ is the set of values for a tilt that would force all the curves for that setting in the left panel of Fig. 9 to coincide with one another.) Not only does the “central blaze peak” depend significantly on order number in each tilt, but the whole format apparently shifts systematically (in $k$ direction) from setting to setting. Here, it appears that $k_0$ has at least a parabolic dependence on $m$, so a bi-quadratic dispersion model in $k - k_0$ and $m$, and with $k_0 = \text{const}$, would effectively be missing terms in $m^3$ and $m^4$. At the same time, the apparent dependence of $k_0$ on $m$ can be measured, and compensated to restore the proper
orthogonality of $k$ and $m$. In principle, one could exploit this to solve for a global fit over all the orders of all the tilts of a mode simultaneously.

A robust global model would be valuable for the settings currently lacking sufficient exposure depth to achieve a well constrained 2D dispersion solution (yellow shaded in Fig. 2). Without the global model, the next best option simply would be to obtain additional observations in those settings to bring up the exposure depth to a more useful level.

An initial foray into global modeling applied a fully bi-cubic polynomial (16 terms) to all the wavecal lines in all the orders of all the tilts of a given mode (14,000 separate measurements for E230H, for example), but compensating for the apparent $k_0$ displacements according to relations like those illustrated in the upper right panel of Fig. 9 for E140H. In principle, the $k_0(m)$ correction should put all the replicas of a given wavecal spot in a given order $m$ exactly on top of each other in $S$, regardless of setting, thereby effectively boosting $S/N$ in a collective fit. It would be as if one could supply a tall MAMA detector that would simultaneously image all the orders of a mode (as E140M does), instead of just a subset. Here, the $m$ variable was replaced by a spatial variable $\tilde{y}$ that accounted for the non-linear $y$ spacing of the orders, and $\Delta k$, as in eq. 2 above, again was used, but now taking explicitly $k_0 = k_0(m)$ as described earlier. This effectively is a “dispersion-only” model, because there is no memory concerning the original location of each measurement in detector coordinates.

Despite the high degree of the model, and promising $k_0$ correction, the resulting global fits were disappointing. The $\chi^2$ was 44 for the E230H case, for example. There were large, mixed polarity residuals in diagrams analogous to Fig. 6 (but with $y$ values extending up to 5700 for E230H, for example, to accommodate all the orders), especially on the extreme left and right edges. The mediocre fits indicated something was missing from the model, probably a systematic behavior tied specifically to the detector reference frame.

A second global experiment attempted to compensate for these putative detector frame effects by introducing six terms to a bi-quadratic form of the dispersion-only model (now,
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15 terms total, one short of the fully bi-cubic model used in the initial global experiment. The new terms utilized an independent variable $\Delta \tilde{y} \equiv \tilde{y} - \tilde{y}_0$, where $\tilde{y}_0$ is the $y$-position of the first order of a setting at the bottom of the detector. Thus, $\Delta \tilde{y}$ spans the detector $y$-axis (0–1023, in low-res pixels) for each setting, and keeps track of where an order falls physically on the detector for each specific tilt. Again, $\Delta \tilde{k} = k - k_0(m)$ served as the $x$-axis independent variable, but for simplicity $m$ itself was used for the order-like variable, rather than the $\tilde{y}$ in the initial global experiment. The final polynomial was:

$$S = a_0 + a_1 m + a_2 m^2 + a_3 \Delta \tilde{k} + a_4 \Delta \tilde{k}^2 + a_5 m \Delta \tilde{k} + a_6 m \Delta \tilde{k}^2 + a_7 m^2 \Delta \tilde{k} + a_8 m^2 \Delta \tilde{k}^2 + a_9 \Delta \tilde{y} + a_{10} \Delta \tilde{y}^2 + a_{11} \Delta \tilde{y} \Delta \tilde{k} + a_{12} \Delta \tilde{y} \Delta \tilde{k}^2 + a_{13} \Delta \tilde{y}^2 \Delta \tilde{k} + a_{14} \Delta \tilde{y}^2 \Delta \tilde{k}^2.$$  

Astonishingly, this simple approach actually worked. The $\chi^2$ (now $\sim 3$ for previous worst case E230H) was in the range achieved by a bi-quadratic model applied to individual settings having good line densities. The fact that the modeling achieved a dramatic improvement over the initial global fitting efforts (which did not include detector-specific terms) must mean that the large residuals seen earlier do have a very systematic component in the detector frame.

This possibility was investigated by collecting the parabolic fit coefficients for $S = f(\Delta \tilde{k})$ (the “$a$” coefficients from eq. 2, but now taking $k_0 = k_0[m]$), averaging them over all the settings of a mode, applying the average transformation to each of the settings individually, and then plotting the residuals in detector coordinates (the $\Delta \tilde{y}$ mentioned earlier). The results are illustrated in Figure 10.

The apparent quadrupolar distortion pattern is consistent between the M and H modes of each camera, and very similar between FUV and NUV. It is quite distinct from the common types of optical distortions (pincushion and barrel), which would produce a bilaterally symmetric map (redshifts on one side, blueshifts on the other). This probably eliminates the optical path as the culprit and points to the MAMAs themselves. In fact, the inferred distortion maps qualitatively are very similar to that reported in Walsh et al. (2001: their Fig. 11) for the FUV camera, although their companion figure for NUV superficially is quite different. (Note that the Walsh et al. study measured the full vector displacements, whereas here only those parallel to the dispersion [detector $x$ axis] can be resolved. Also, the Walsh et al. results were for direct imaging with the MAMAs, without the full pass through the complex echelle optics.)

From a practical point of view, introducing an appropriate spatial distortion correction step into calstis would allow the use of a simplified — but potentially more accurate — dispersion model, along the lines sketched earlier, and would permit more reliable “extrapolation” of the model to those settings that have poorer line densities.

6. Summary

The central objective of the Wrinkles project was to explore ways to improve precision and accuracy of the STIS echelle mode wavelength scales. Extensive measurements of deep Pt/Cr–Ne lamp spectra in each of the 44 supported modes were utilized to derive 2D distortion maps to correct for large scale systematics traced to limitations of the polynomial dispersion relations embedded in the calstis pipeline. The corrected x1d high-resolution (H-mode) echelle spectra then were merged, averaged, and spliced to produce an improved global atlas of the calibration lamp. This pseudo-laboratory spectrum was measured by an automated fitting ‘robot’ to yield a refined list of the sharp lamp lines, especially the

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8 An important component of accuracy, for normal science programs, is the quality of the target centering.
Figure 10: Collective spatial distortion maps for the four STIS echelle modes.

many unidentified features longward of 1800 Å attributed to Cr I or Cr II. The new line list was corrected for low-amplitude, long-range systematics by reference to the subset of features for which accurate laboratory measurements were available. In parallel, the high-quality H-spectrum was substituted for the cross-correlation “template” in the pipeline; the original reference file has a number of conspicuous defects. The high-density line list then was applied to derive second generation distortion maps, and the overall cycle was repeated. Several iterations of the process resulted in a factor of $\sim 2$ improvement in the (already very good) precision of the STIS echelle wavelength scales, especially for the many secondary settings that are less well calibrated than the primary ones.

A companion analysis considered possible enhancements of the existing pipeline polynomial dispersion relations. The largest improvement was found by simply re-deriving the coefficients for each individual echelle setting, exploiting the now higher density of reliable lamp wavelengths. Further improvement was obtained by adding the term $m^2$ to render the polynomials almost fully bi-quadratic (only the term $\lambda^2$ still would be missing), while also activating the vestigial term $(m\lambda)^3$ (whose coefficients currently are not populated in the calibration library for the echelle settings).

Although it was argued previously that replacing the pipeline independent variables $\lambda$ and $m$ with properly orthogonal variables $k(=m\lambda)$ and $m$ should improve the solutions, experiments showed, counter-intuitively, that this was not the case. The reason turned out to be that the $k$ variable, as defined, was not completely orthogonal to the orders
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$m$, because the central value for each order, $k_0$, curved slightly along the detector $y$ axis. Deriving the curvature terms empirically allowed a reformulation in terms of $k - k_0(m)$, which not only solved the orthogonality issue, but also allowed the modeling of all the settings simultaneously for a particular mode.

Global bi-parabolic solutions then were obtained, with the simple addition of a few extra terms to account for spatial distortions associated with the detector frame. The quality of the global model was comparable to similar-degree fits to individual settings with the highest line densities. Furthermore, applying average values of the $k - k_0(m)$ parabolic fits to the individual settings of each mode allowed the spatial (detector frame) distortions to be isolated. The apparent quadrupolar warping is identical in the two cameras, and is similar to independently derived spatial distortions reported for at least the FUV MAMA. If a suitable spatial correction could be introduced into calstis, a simplified set of dispersion relations in principle could be utilized.

Third-, fourth-, and fifth-order dispersion models achieved further improvement over the modified calstis model. Except for E140M, however, the gains were modest. Indeed, there is some merit to applying the modified calstis formulation as an initial wavelength transformation, then following with a subsequent third-, or higher, order distortion correction to account for any low-amplitude residuals. (The latter step might not be necessary if the detector frame spatial correction, mentioned above, could be implemented.)

The main recommendations to the STIS project, in order of increasing importance, if not also increasing difficulty, are:

1. Re-derive the calstis dispersion coefficients using the updated lamp line lists and the existing on-orbit long duration wavecal exposures. (As of this writing, new coefficients have been derived by the author and are undergoing tests.)

2. Improve the wavecal exposure depth for the many modes with relatively poor coverage at present. This includes especially the short FUV end of E140H, and the short NUV end of E230H, where the existing line densities are comparatively low, and the dispersion solutions are less well constrained. (This work is planned for the author’s approved Cycle 18 GO calibration program “Deep Lamp Too.”)

3. Consider extending the pipeline polynomials to a ninth term ($m^2$), and populating the coefficients for the existing eighth term ($[m\lambda]^3$). (The STIS calibration team has agreed to explore implementation of these changes, and suitable coefficients have been supplied by the author for testing purposes.)

4. Re-cast the dispersion modeling into two pieces: a de-warping correction followed by a simplified polynomial in orthogonal variables like the $k - k_0(m)$ approach. With this strategy, the dispersion coefficients for each setting of a mode should be very nearly identical, thereby allowing more legitimate extrapolations to settings with poor line densities. A post facto, high-order distortion correction specific to each tilt could clean up small residuals, if any, from the two-step dispersion correction. (The two-step approach, while the best justified from a “physical instrument” perspective, probably is overkill, given the level of improvement anticipated with step [3], above.)

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Slitless Spectroscopy with HST Instruments

J. R. Walsh, M. Kümmel, H. Kuntschner

Abstract. Slitless spectroscopy is generally regarded as a niche, perhaps ‘difficult’, astronomical observation technique. However the low background from space and the high spatial resolution offered by HST instruments has enabled it to become a powerful survey tool, with some applications to single object work. The ST-ECF has been involved with slitless spectroscopy from NICMOS, through ACS to WFC3. Techniques and software have also been developed for the bulk extraction of NICMOS and ACS slitless spectra, which are available through the Hubble Legacy Archive.

A short overview of the slitless spectroscopy modes on HST is provided and the essential elements that distinguish slitless from slit or multi-slit spectroscopy practice are detailed. From these elements the strategy for reducing slitless spectra is developed and its implementation in the aXe software package is summarised with examples from ACS and WFC3 spectra. Extension of the basic technique to deal with dithered spectra and contamination are outlined. An application of the decomposition of spectra to the case of a complex slitless spectroscopy scene is described.

1. HST slitless spectroscopy modes

HST has had slitless spectroscopy capabilities since launch and of the Cycle 1 instruments WF/PC-1 had grisms and FOC had prisms, both designed for use without a slit. However on account of the spherical aberration these elements were not used for scientific observations. After correction by COSTAR, the FOC prisms were used for occasional slitless spectroscopy (e.g. Jacobsen et al. 1993). With Servicing Mission 2 (SM2) in 1997, two instruments were flown with slitless spectroscopic capability. NICMOS was named for its multi-object spectroscopic capability through slitless grisms and STIS could also be used without a slit in many of its modes. Table 1 lists the slitless spectroscopy modes on HST over its orbital life. NICMOS produced many results from late type stars (e.g. Greissl et al. 2007) to distant galaxies (e.g. Freudling et al. 2003) and an archive of G141 grism spectra was released through the Hubble Legacy Archive (HLA); see Freudling et al. (2008).

With SM3B in 2003, the Advanced Camera for Surveys (ACS) was installed: all modes of this camera had slitless facilities with both prisms and a grism (see Table 1). Pasquali et al. (2006) provide an overview of ACS slitless grism spectroscopy. Deep slitless surveys were completed with the red G800L grism and ACS Wide Field Channel (GRAPES: Pirzkal et al. 2004; PEARS: e.g. Straughn, et al. 2009). In the latest servicing mission (SM4), the Wide Field Camera 3 (WFC3) was installed, with both the UV-optical and near-infrared channel having slitless grism capabilities. The overview in Table 1 should be supplemented by the respective instrument handbooks for more detailed descriptions of the slitless capabilities.
Table 1: Overview of HST slitless spectroscopy modes

<table>
<thead>
<tr>
<th>Instrument</th>
<th>Orbital life</th>
<th>Disperser</th>
<th>$R = \lambda / \Delta \lambda$ @ $\lambda$</th>
<th>$\lambda$ range</th>
</tr>
</thead>
<tbody>
<tr>
<td>WFPC1</td>
<td>1990-1993</td>
<td>G200 grism</td>
<td>75 @ 1800Å, 40 @ 4500Å, 35 @ 8000Å</td>
<td>1300-2000Å, 3000-6000Å, 6000-10000Å</td>
</tr>
<tr>
<td>FOC</td>
<td>1990-2002</td>
<td>FUVOP prism</td>
<td>50 @ 1200Å, 100 @ 2500Å</td>
<td>1250-3000Å, 1600-4000Å</td>
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<tr>
<td>STIS</td>
<td>1997-present</td>
<td>NUV prism</td>
<td>900 @ 1200Å</td>
<td>1150-3000Å</td>
</tr>
<tr>
<td></td>
<td></td>
<td>All 1st order gratings, e.g. G800L</td>
<td>800 @ 8000Å</td>
<td>5420-10270Å</td>
</tr>
<tr>
<td>NICMOS</td>
<td>1997-present</td>
<td>G096 grism</td>
<td>100 @ 1.0μm</td>
<td>0.8-1.2μm</td>
</tr>
<tr>
<td></td>
<td></td>
<td>G141 grism</td>
<td>100 @ 1.7μm</td>
<td>1.1-1.9μm</td>
</tr>
<tr>
<td></td>
<td></td>
<td>G200 grism</td>
<td>100 @ 2.0μm</td>
<td>1.4-2.5μm</td>
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<tr>
<td>ACS</td>
<td>2002-present</td>
<td>PR110L prism</td>
<td>80 @ 1500Å</td>
<td>1150-1800Å</td>
</tr>
<tr>
<td></td>
<td></td>
<td>PR130L prism</td>
<td>100 @ 1500Å</td>
<td>1250-1800Å</td>
</tr>
<tr>
<td></td>
<td></td>
<td>PR200L prism</td>
<td>60 @ 2500Å</td>
<td>1700-3900Å</td>
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<td></td>
<td></td>
<td>G800L grism</td>
<td>100 @ 8000Å</td>
<td>5500-10500Å</td>
</tr>
<tr>
<td>WFC3</td>
<td>2009-present</td>
<td>G280 grism</td>
<td>70 @ 3000Å</td>
<td>1900-4000Å</td>
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<td>G102 grism</td>
<td>160 @ 1.04μm</td>
<td>0.8-1.15μm</td>
</tr>
<tr>
<td></td>
<td></td>
<td>G141 grism</td>
<td>120 @ 1.30μm</td>
<td>1.1-1.7μm</td>
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</table>

2. The elements of slitless spectroscopy

The essence of slitless spectroscopy is that the sky positions of the targets, and their sizes and shapes, define the ‘slits’ in analogy to slit or multi-slit spectroscopy. One can consider that an object, whose light is dispersed in a slitless image, defines its own virtual slit that modulates the height of the 2D spectrum and the resulting spectral resolution through the width of the object in the dispersion direction. For non-circular objects, a preferred axis, such as the major axis of a galaxy, can define a tilted virtual slit. Since the positions of the virtual slits are set by the pattern of objects on the sky, then supposition of the spectra is bound to occur commonly as the spectra are many times longer than the target in the dispersion direction. This occurrence is called contamination and in contrast to slit spectroscopy occurs predominantly in the dispersion direction. If a grism is used as the slitless dispersing element, then the zeroth and higher orders (that can be both positive and negative) present an additional source of contamination in contrast to prisms. Depending on the design of the grism, the zeroth and higher orders have lower transmission than the strongest order (generally designated +1). Figure 1 shows a very typical slitless grism image, here taken with the ACS WFC G800L grism, of a moderately crowded field in the Chandra Deep Field South (CDFS) exemplifying all these effects of virtual slits, contamination and multiple spectral orders. Note also that objects outside the field also give rise to partial spectra (at the left edge of the detector in the case of WFC3 and ACS) since the position of the dispersing object is displaced from the low wavelength end of the +1st order spectrum.

A further characteristic of slitless spectra is the sky background level that is distinct from direct imaging and slit spectroscopy. Since the effective slit width is that of the detector, then the background level is many times larger than for a slit spectrometer. Since the spectral width of the grism is that of its passband then it acts like a very broad filter, with a passband of about 3300Å in the case of the ACS WFC G800L grism (see Table 1, column 5). The background count level is thus high and has structure. The structure arises
mainly from the cut-off of the sky spectrum from the different orders at the edges of the field, which typically are slightly beyond the detector area. Figure 2 shows, as an example, the WFC3 G141 master background image; the knee to the right edge of the detector, for example, is caused by the zeroth order moving out of the unvignetted field.

3. Reducing slitless spectroscopy data

The elements of slitless spectra described in the previous section contribute to the commonly held notion that slitless spectra rate as ‘difficult’ to reduce. However providing accurate information on the positions, sizes and spectral shape of the objects whose spectra are dispersed, together with the requisite calibration data, allows a straightforward pipeline approach for slitless spectra. Indeed given the large number of detectable spectra collected on deep ACS or WFC3 slitless images, from hundreds to over a thousand, at least a semi-automatic approach has to be adopted to extracting this many spectra from the slitless images.

Operationally each slitless image, or set of slitless images at an HST single guide star pointing, is accompanied by a direct image, usually, but not necessarily, in a filter included in the passband of the slitless spectra. A morphological and photometric catalogue of this image, such as made with the SExtractor task (Bertin & Arnouts, 1996), then provides all the virtual slit information required to localise the spectra. With the positional mapping from the direct image to the slitless image, the reference position of an object on the slitless image (equivalent to the slit position) and the extent of the spectra for all its orders can be determined. With the photometric information, the level of contamination of overlapping spectra can be quantitatively assessed, given simple assumptions about the spectral shape. With multiple filters a simple SED can be constructed, so that the model
of the contaminating spectra can be improved. Zeroth orders present a particular case of contamination, because, although slightly dispersed (by the prism on which the grooves are ruled), they can resemble an emission line for small and point source objects.

With the information on the reference position and dimensions of the dispersing object, the extent of all the order(s) can be traced from the known pre-calibrated properties of the prism or grism; then the wavelength assignment to pixels along the spectral traces can be made. Once the wavelength extent of each pixel is known, then a flat field correction can be applied. In slitless spectroscopy, potentially any wavelength in the passband of the dispersing element can be recorded by every pixel so that the flat field information must be described by a cube, having two spatial dimensions for the detector and the wavelength dependence. Operationally the flat field cube is parametrised by a polynomial fit of the variation of flat field with wavelength for each pixel (see Kümmel et al. 2009b).

The array of pixels belonging to a given object and order can be combined (‘extracted’) in two ways - as a rectified 2D long slit with perpendicular dispersion and cross-dispersion axes or the pixels summed in the cross-dispersion direction over the virtual slit (possibly with cross-dispersion profile weighting) to form a 1D spectrum. Absolute sensitivity can be applied to the 1D spectrum using a calibration spectrum established on spectrophotometric standard stars for conversion to absolute flux. For flux calibration of extended sources, a correction to the sensitivity curve established for a point source is required based on the shape and size of the extended object in the dispersion direction.

4. The aXe software package and its calibration files

All these steps have been combined in the software package aXe (the name is not an acronym). The individual steps have been coded in C for speed and then the set of tasks scripted in python, so that for example with the "axecore" task one can proceed from an input slitless image and a catalogue of objects on a companion direct image to a set of extracted 2D, wavelength calibrated, spectra and fully wavelength and flux calibrated 1D
spectra. The aXe software is described in Kümmerl et al. (2009b) and a full user manual is available (Kümmerl et al. 2010a). Full instrument independence of this software is provided through input of a configuration file which stores all of the instrument-specific information, such as on the trace and wavelength solutions (including their 2D dependence on position of the reference position on the detector) and the required calibration files such as flat field cube and sensitivity file. The configuration file and associated files are the set of reference files for a given HST slitless mode (though not included in CDBS).

Slitless spectroscopy precludes the use of wavelength calibration lamps, so that the dispersion solution as a function of the reference position must be calibrated in advance, either on-ground or in space, or both. Compact targets with a good grid of emission lines (such as planetary nebulae, emission line stars - Be, Wolf-Rayet or chromospherically active - and AGN) are typically chosen (see Pasquali et al. 2006). For grisms many orders may be detected that can be potentially be extracted and scientifically exploited. In practice this has not been done since, even though the dispersion is improved, the efficiency in the higher orders is generally low. In addition, the higher orders are increasingly out of focus on account of the optical path through the grism, so the theoretically expected increase in spectral resolution is not fully realised.

The variation of the trace over the field can be mapped in slitless exposures from spectra of the many stars in low Galactic latitude fields. The flat field cube contains both the large scale and pixel-to-pixel flat field properties (combined L-flat and P-flat) and can be constructed from ground or in-orbit flat data. In the case of ACS the flat field cube was formed by combining the wavelength dependence of the in-orbit direct imaging filter flats (Walsh & Pirzkal, 2005); for WFC3 good ground flats were available (Kuntschner et al. 2009). The fringing of the detector can also be considered in the reduction process (see Walsh et al. 2003).

Spectrophotometric standards provide flux calibration as a function of position so that spatial dependence of throughput can be corrected. Kuntschner et al. (2010, this volume) provide an overview of calibration activities for WFC3 slitless modes and Kümmerl et al. (2010c, this volume) step through an example reduction of WFC3 G102 and G141 spectra taken in the CDFS as part of the Early Release Science (Straughn et al, 2010) with the aXe software.

Slitless spectroscopy is by virtue a survey mode and detectable spectra of hundreds to greater than a thousand targets are typically available for scientific exploitation on a deep ACS or WFC3 slitless grism image. aXe also provides facilities for generating web pages for rapid and comfortable scanning of many slitless spectral extractions enabling quality assessment, search for objects of interest (e.g. with emission lines), etc. Walsh & Kümmerl (2004) provide a description of this add-on facility to the aXe software.

The aXe software was originally developed for ACS slitless spectra and was then applied to NICMOS and most recently to WFC3. A fully automatic pipeline (PHLAG) was developed in the framework of the Hubble Legacy Archive (HLA http://www.stecf.org/archive/hla and http://hla.stsci.edu) and 1923 NICMOS G141 spectra were extracted and released (Freudling et al. 2008). PHLAG was further developed for ACS G800L spectra (Kümmerl et al. 2009c) and 47919 spectra were released into the HLA (Kümmerl et al. 2010b, in prep.). Figure 3 shows an example preview of one of these ACS spectra that includes its associated direct image stamp.

5. Further development of aXe

The aXe software has continued to be developed and now uses the drizzle software to combine multiple dithered exposures. Since the spectra are differentially distorted across the field by the instrument optics, simple combination of slitless images is not possible, but the extracted 2D images with the wavelength and cross-dispersion corrected can be combined.
Figure 3: An example spectrum taken from the HLA ACS slitless spectroscopy archive showing a preview of an $i=22.8$ mag. galaxy in the HUDF.

aXedrizzle now performs this and an example of its application can be found in Kümmel et al. (2010c, this volume). Cosmic ray rejection, comparable to the implementation in multidrizzle in STSDAS, has been built into aXedrizzle (in aXe version 2.1).

Since the zero point of the wavelength of any given spectrum depends on an independent image, then issues of repeatability of filter and grism wheel placement and spacecraft jitter enter into the zero point determination. One approach that has been successfully implemented in the PHLAG pipeline is to catalogue stars on an image and cross-correlate their spectra against a library of templates. Selecting the template that best matches allows the computation of the zero point shift for each spectrum; given enough stars on an image then a robust selection of the shift for the whole image can be made. Application of this shift improves the wavelength zero point for all extracted spectra on that image (see Kümmel et al. 2010b).

Great care has to be taken in treating slitless spectra of extended objects taken at multiple spacecraft rolls. The orientation of the virtual slit defined by the object itself and the direction of the dispersion can result in very different spectra of the same source, such that they should not be combined. To combine spectra with different rolls a restoration approach can be applied (e.g. Freudling 1997) or individual features in extended objects can be extracted as point sources (e.g. HII regions, see Kümmel et al. 2010b). Care should also be applied in comparing slit and slitless extraction for extended objects since it is the object shape which controls the light distribution. The Appendix of Freudling et al. (2008) considers the correction to the slit direction based on the goal of choosing the virtual slit that minimizes the variations of wavelength; this correction was applied to the spectra in the ACS HLA release (http://www.stecf.org/archive/hla/acs_g800l_release.php).

Contamination is an intrinsic limitation of all slitless spectroscopy. Using the simple SED constructed from the direct image photometry allows a continuum model of the spectra.
be formed, enabling estimation of the contamination contribution. This is very useful, for example, for assessing if an emission line is actually caused by a zeroth order of a nearby object. However a simplified SED cannot, for example, flag when an emission line from one spectrum is contaminating another. To go further an iterative scheme would need to be contemplated whereby the initial estimate of the contaminating spectrum is replaced (perhaps only partially) by the extracted spectrum, and some stopping criterion is implemented. This procedure would be computer intensive as, in principle, the complete extraction step, involving many hundreds of spectra, has to be repeated for each contamination iteration; it has not been implemented in aXe to date.

The process of modelling the contamination from direct image photometry is analogous to simulating a slitless spectroscopy image from a direct image catalogue. Simulation can play an important role in planning slitless spectroscopic observations, such as for deriving the optimal roll angle for observations of specific objects; one example could be slitless spectroscopy of strong lenses in galaxy clusters. The ability to simulate slitless spectra of course plays a role in signal-to-noise estimation, in complement to the more direct HST Exposure Time Calculator for slitless modes. The slitless simulator built on aXe provides close compatibility with aXe itself, so that for example the same configuration file can be used as for extracting the spectra from real data (Kümml et al. 2009a). Although aXe output products provide fully propagated statistical errors based on the instrument pipeline error arrays, the ability to simulate spectra and add them to real images, in an analogous way to artificial star tests in photometric reduction, can be useful for extracting spectra from complex fields.

6. Disentangling slitless spectra

Even aside from contamination, slitless spectra of composite objects, such as a spiral galaxy with a bulge and HII regions, can present a complex spectral scene. However the independent data presented by the direct imaging provide additional information that can be used to extract multiple objects. This is most amenable to decomposition techniques for spectra such as implemented for long slit spectra by Lucy & Walsh (2003). The direct image can provide a prior for the positions, shapes and fluxes (spectra) of the components of the image that can then be iteratively restored to match the slitless spectrum. An implementation of the technique described in Lucy & Walsh (2003) for the case of spatially complex background (implemented in stecf.specres.specinholucy) has been coded and demonstrated on the slitless spectrum of a strong lens system in a field elliptical galaxy, discovered by Blakeslee et al. (2004) in ACS data. Figure 4 shows the direct image with the knots indicated and the fairly confused slitless spectrum. The ten lensed knots were represented as slightly extended points sources, all with the same PSF, and the spectrum of each was extracted. The figure shows the combined lensed nuclear and arc spectrum. Such an analysis procedure is highly interactive and would not be contemplated on a large scale but does demonstrate that useful area spectroscopy is achievable with the slitless technique.

7. Conclusions

One of the aims of this contribution was to try to demonstrate that slitless spectra are not more ‘difficult’ to exploit than more conventional slit spectra. A well-defined approach to the extraction of slitless spectra has been outlined and is implemented in the aXe software. This software is robust and flexible in its instrument applicability and has been proved on ACS, NICMOS and WFC3 data from HST and on ground-based data. It has been included in a fully automatic pipeline which has enabled many tens of thousands of spectra to be released through the HLA.
Figure 4: Example of a Lucy iterative decomposition of ten strong lensed knots in a field elliptical galaxy. Upper left is shown the colour image from ACS WFC F435W, F606W, F775W and F850LP filter images and lower left is shown the designation of the knots on the image with the galaxy removed (Blakeslee et al. 2004). The ACS G800L slitless image is shown to the upper right and the individual extracted spectra of all the knots to the lower right. The sum of the spectra of the seven knots in the arc and the three in the nuclear lensed knots are indicated as is the sum of all the spectra, at lower right.
8. Acknowledgements

Many members of the ST-ECF have contributed to the support of HST slitless spectroscopy over the years 1997-2010, including Wolfram Freudling and Richard Hook. Former ST-ECF members who played a large role were Soeren Larsen, now at the University of Utrecht, Anna Pasquali now at Max-Planck Institute for Astronomy, Heidelberg and Nor Pirzkal, currently at STScI.

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An insight into the flux calibration of Gaia G-band images and BP/RP spectrophotometry

E. Pancino

INAF – Bologna Observatory, Via Ranzani 1, I-40127 Bologna, Italy

Abstract.

The Gaia mission is described, focussing on those technical aspects that are necessary to understand the details of its external (absolute) flux calibration. On board Gaia there will be two spectrophotometers, the blue one (BP) and the red one (RP) covering the range 330-1050 nm, and the white light (G-band) imager dedicated to astrometry. Given the fact that Gaia’s focal plane will constitute 105 CCDs and the sources will cross the the focal plane at constant speed, at different positions in each of the foreseen passages (on average 70–80, but up to 350) in the mission lifetime, the “simple” problem of calibrating the integrated BP/RP and G-band magnitudes and the low resolution BP/RP spectra flux turns into a very delicate and complicated issue, including CTI effects, LSF variations across the focal plane and with time, CCD gating to avoid saturation and the like. The calibration model requires a carefully selected set of $\approx 200$ Spectrophotometric Standard Stars (SPSS) with a nominal precision of a few $\%$, with respect to Vega.

1. The Gaia mission

Gaia is a cornerstone mission of the ESA Space Program, presently scheduled for launch in 2012. The Gaia satellite will perform an all-sky survey to obtain parallaxes and proper motions to $\mu$as precision for about $10^9$ point-like sources and determine astrophysical parameters ($T_{\text{eff}}, \log g, E(B-V)$, metallicity etc.) for stars down to a limiting magnitude of $V \approx 20$, plus 2-30 km/s accuracy (depending on spectral type), radial velocities for several millions of stars down to $V < 17$.

Such an observational effort has been compared to the mapping of the human genome for the amount of collected data and for the impact that it will have on all branches of astronomy and astrophysics. The expected end-of-mission astrometric accuracies are almost 100 times better than the HIPPARCOS dataset (see Perryman et al. 1997). This exquisite precision will allow a full and detailed reconstruction of the 3D spatial structure and 3D velocity field of the Milky Way galaxy within $\approx 10$ kpc from the Sun. This will provide answers to long-standing questions about the origin and evolution of our Galaxy, from a quantitative census of its stellar populations, to a detailed characterization of its substructures (as, for instance, tidal streams in the Halo, see Ibata & Gibson, 2007, Sci. Am., 296, 40), to the distribution of dark matter.

The accurate 3D motion of more distant Galactic satellites (as globular clusters and the Magellanic Clouds) will be also obtained by averaging the proper motions of many thousands of member stars: this will provide an unprecedented leverage to constrain the mass distribution of the Galaxy and/or non-standard theories of gravitation. Gaia will determine direct geometric distances to essentially any kind of standard candle currently used for distance determination, setting the whole cosmological distance scale on an extremely firm basis.
Table 1: Expected numbers of specific objects observed by Gaia.

<table>
<thead>
<tr>
<th>Type</th>
<th>Numbers</th>
<th>Type</th>
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</tr>
</thead>
<tbody>
<tr>
<td>Extragalactic supernovae</td>
<td>20,000</td>
<td>Extra-solar planets</td>
<td>15,000</td>
</tr>
<tr>
<td>Resolved galaxies</td>
<td>$10^6$–$10^7$</td>
<td>Disk white dwarfs</td>
<td>200,000</td>
</tr>
<tr>
<td>Quasars</td>
<td>500,000</td>
<td>Astrometric microlensing events</td>
<td>100</td>
</tr>
<tr>
<td>Solar system objects</td>
<td>250,000</td>
<td>Photometric microlensing events</td>
<td>1000</td>
</tr>
<tr>
<td>Brown dwarfs</td>
<td>$\geq$50,000</td>
<td>Resolved binaries (within 250 pc)</td>
<td>$10^7$</td>
</tr>
</tbody>
</table>

As challenging as it is, the processing and analysis of the huge data-flow from Gaia is the subject of thorough study and preparatory work by the Data Processing and Analysis Consortium (DPAC), in charge of all aspects of the Gaia data reduction. The consortium comprises more than 400 scientists from 25 European institutes. Gaia is usually described as a self-calibrating mission, but it also needs external data to fix the zero-point of the magnitude system and radial velocities, and to calibrate the classification and parametrization algorithms. All these additional data are termed auxiliary data and have to be available, at least in part, three months before launch. While part of the auxiliary data already exist and must only be compiled from archives, this is not true for several components. To this aim a coordinated program of ground-based observations is being organized by a dedicated inter CU (Coordination Unit) committee (Ground Based Observation Group), that promotes synergies and avoids duplications of effort.

1.1. Science goals and capabilities

Gaia will measure the positions, distances, space motions, and many physical characteristics of some billion stars in our Galaxy and beyond. For many years, the state of the art in celestial cartography has been the Schmidt surveys of Palomar and ESO, and their digitized counterparts. The measurement precision, reaching a few millionths of a second of arc, will be unprecedented. Some millions of stars will be measured with a distance accuracy of better than 1 per cent; some 100 million or more to better than 10 per cent. Gaia’s resulting scientific harvest is of almost inconceivable extent and implication.

Gaia will provide detailed information on stellar evolution and star formation in our Galaxy. It will clarify the origin and formation history of our Galaxy. The results will precisely identify relics of tidally-disrupted accretion debris, probe the distribution of dark matter, establish the luminosity function for pre-main sequence stars, detect and categorize rapid evolutionary stellar phases, place unprecedented constraints on the age, internal structure and evolution of all stellar types, establish a rigorous distance scale framework throughout the Galaxy and beyond, and classify star formation and kinematical and dynamical behavior within the Local Group of galaxies.

Gaia will pinpoint exotic objects in colossal and almost unimaginable numbers: many thousands of extra-solar planets will be discovered (from both their astrometric wobble and from photometric transits) and their detailed orbits and masses determined; tens of thousands of brown dwarfs and white dwarfs will be identified; tens of thousands of extragalactic supernovae will be discovered; Solar System studies will receive a massive impetus through the observation of hundreds of thousands of minor planets; near-Earth objects, inner Trojans and even new trans-Neptunian objects, including Plutinos, may be discovered.

Gaia will follow the bending of star light by the Sun and major planets over the entire celestial sphere, and therefore directly observe the structure of space-time – the accuracy of its measurement of General Relativistic light bending may reveal the long-sought scalar correction to its tensor form. The PPN parameters $\gamma$ and $\beta$, and the solar quadrupole
moment \( J_2 \), will be determined with unprecedented precision. All this, and more, will be obtained through the accurate measurement of star positions.

We summarize some of the most interesting object classes that will be observed by Gaia, with estimates of the expected total number of objects, in Table 1. For more information on the Gaia mission: http://www.rssd.esa.int/Gaia. More information for the public on Gaia and its science capabilities are contained in the Gaia information sheets\(^1\). An excellent review of the science possibilities opened by Gaia can be found in Perryman et al. (1997).

### 1.2. Launch, timeline and data releases

The first idea for Gaia began circulating in the early 1990, culminating in a proposal for a cornerstone mission within ESA’s science program submitted in 1993, and a workshop in Cambridge in June 1995. By the time the final catalogue will be released approximately in 2020, almost two decades of work will have elapsed between the original concept and mission completion.

Gaia will be launched by a Soyuz carrier (rather than the originally planned Ariane 5) in 2012 from French Guyana and will start operating once it reaches its Lissajous orbit around \( L_2 \) (the unstable Lagrange point of the Sun and Earth-Moon system), a month later. Two ground stations will receive the compressed Gaia data during the 5 years\(^2\) of operation: Cebreros (Spain) and Perth (Australia). The data will then be transmitted to the main data centers throughout Europe to allow for data processing. We are presently in technical development phase C/D, and the hardware is being built, tested and assembled. Software development started in 2006 and is presently producing and testing pipelines with

\(^1\)http://www.rssd.esa.int/index.php?project=GAIA\&page=Infosheets\_overview.

\(^2\)If – after careful evaluation – the scientific output of the mission will benefit from an extension of the operation period, the satellite should be able to gather data for one more year, remaining within the Earth eclipse.
the aim of delivering to the astrophysical community a full catalogue and dataset ready for scientific investigation.

Apart from the end-of-mission data release, foreseen around 2020, some intermediate data releases are foreseen. In particular, there should be one first intermediate release covering either the first 6 months or the first year of operation, followed by a second and possibly a third intermediate release, that are presently being discussed. The data analysis will proceed in parallel with observations, the major pipelines re-processing all the data every 6 months, with secondary cycle pipelines – dedicated to specific tasks – operating on different timescales. In particular, verified science alerts, based on unexpected variability in flux and/or radial velocity, are expected to be released within 24 hours from detection, after an initial period of testing and fine-tuning of the detection algorithms.

1.3. Mission concepts

During its 5-year operational lifetime, the satellite will continuously spin around its axis, with a constant speed of 60 arcsec/sec. As a result, over a period of 6 hours, the two astrometric fields of view will scan across all objects located along the great circle perpendicular to the spin axis (Figure 2, left panel). As a result of the basic angle of 106.5° separating the astrometric fields of view on the sky (Figure 1), objects transit the second field of view with a delay of 106.5 minutes compared to the first field. Gaia’s spin axis does not point to a fixed direction in space, but is carefully controlled so as to precess slowly on the sky. As a result, the great circle that is mapped by the two fields of view every 6 hours changes slowly with time, allowing repeated full sky coverage over the mission lifetime. The best strategy, dictated by thermal stability and power requirements, is to let the spin axis precess (with a period of 63 days) around the solar direction with a fixed angle of 45°. The above scanning strategy, referred to as “revolving scanning”, was successfully adopted during the Hipparcos mission.

Every sky region will be scanned on average 70-80 times, with regions lying at ±45° from the Ecliptic Poles being scanned on average more often than other locations. Each of the Gaia targets will be therefore scanned (within differently inclined great circles) from a minimum of approximately 10 times to a maximum of 250 times (Figure 2, right panel). Only point-like sources will be observed, and in some regions of the sky, like the Baade’s window, ω Centauri or other globular clusters, the star density of the two combined fields of view will be of the order of 750 000 or more per square degree, exceeding the storage capability of the onboard processors, so Gaia will not study in detail these dense areas.
1.4. Focal plane

Figure 3 shows the focal plane of Gaia, with its 105 CCDs, which are read in TDI (Time Delay Integration) mode: objects enter the focal plane from the left and cross one CCD in 4 seconds. Apart from some technical CCDs that are of little interest in this context, the first two CCD columns, the Sky Mappers (SM), perform the on-board detection of point-like sources, each of the two columns being able to see only one of the two lines of sight. After the objects are identified and selected, small windows are assigned, which follow them in the astrometric field (AF) CCDs where white light (or G-band) images are obtained (Section 1.5.). Immediately following the AF, two additional columns of CCDs gather light from two slitless prism spectrographs, the blue spectrophotometer (BP) and the red one (RP), which produce dispersed images (Section 1.6.). Finally, objects transit on the Radial Velocity Spectrometer (RVS) CCDs to produce higher resolution spectra around the Calcium Triplet (CaT) region (Section 1.7.).

1.5. Astrometry

The AF CCDs will provide G-band images, i.e., white light images where the passband is defined by the telescope optics transmission and the CCDs’ sensitivity, with a very broad combined passband ranging from 330 to 1050 nm and peaking around 500–600 nm (Figure 4). The objective of Gaia’s astrometric data reduction system is the construction of core mission products: the five standard astrometric parameters, position ($\alpha$, $\delta$), parallax ($\pi$), and proper motion ($\mu_{\alpha}$, $\mu_{\delta}$) for all observed stellar objects. The expected end-of-mission precision in the proper motions is expected to be better than 10 $\mu$as for G<10 mag stars, 25 $\mu$as for G=15 mag, and 300 $\mu$as for G=20 mag. For parallaxes, considering a G=12 mag star, we can expect to have distances at better than 0.1% within 250 pc, 1% within 2700 pc, and 10% within 10 kpc.

To reach these end-of-mission precisions, the average 70–80 observations per target gathered during the 5-year mission duration will have to be combined into a single, global, and self-consistent manner. 40 Gb of telemetry data will first pass through the Initial Data Treatment (IDT) which determines the image parameters and centroids, and then perform object cross-matching. The output forms the so-called One Day Astrometric Solution (ODAS), together with the satellite attitude and calibration, to sub-milliarcsecond accuracy. The data are then written to the Main Database.
The next step is the Astrometric Global Iterative Solution (AGIS) processing. AGIS processes together the attitude and calibration parameters with the source parameters, refining them in an iterative procedure that stops when the adjustments become sufficiently small. As soon as new data come in, on the basis of 6 months cycles, all the data in hand are reprocessed together from scratch. This is the only scheme that allows for the quoted precisions, and it is also the philosophy that justifies Gaia as a self-calibrating mission. The primary AGIS cycle will treat only stars that are flagged as single and non-variable (expected to be around 500 millions), while other kinds of objects will be computed in secondary AGIS cycles that utilize the main AGIS solution. Dedicated pipelines for specific kinds of objects (asteroids, slightly extended objects, variable objects and so on) are being put in place to extract the best possible precision. Owing to the large data volume (100 Tb) that Gaia will produce, and to the iterative nature of the processing, the computing challenges are formidable: AGIS processing alone requires some $10^{21}$ FLOPs which translates to runtimes of months on the ESAC computers in Madrid.

1.6. Spectrophotometry

The primary aim of the photometric instrument is mission critical in two respects: (i) to correct the measured centroids position in the AF for systematic chromatic effects, and (ii) to classify and determine astrophysical characteristics of all objects, such as temperature, gravity, mass, age and chemical composition (in the case of stars).

The BP and RP spectrophotometers are based on a dispersive-prism approach such that the incoming light is not focussed in a PSF-like spot, but dispersed along the scan direction in a low-resolution spectrum. The BP operates between 330–680 nm while the RP between 640–1000 nm (Figure 4). Both prisms have appropriate broad-band filters to block unwanted light. The two dedicated CCD stripes cover the full height of the AF and, therefore, all objects that are imaged in the AF are also imaged in the BP and RP.

The resolution is a function of wavelength, ranging from 4 to 32 nm/pix for BP and 7 to 15 nm/pix for RP. The spectral resolution, $R=\lambda/\delta\lambda$ ranges from 20 to 100 approximately. The dispersers have been designed in such a way that BP and RP spectra are of similar sizes (45 pixels). Window extensions meant to measure the sky background are implemented. To compress the amount of data transmitted to the ground, all the BP and RP spectra – except for the brightest stars – are binned on chip in the across-scan direction, and are transmitted to the ground as one-dimensional spectra. Figure 4 shows a simulated RP spectrum, unbinned, before windowing, compression, and telemetry.
The final data products will be the end-of-mission (or intermediate releases) of global, combined BP and RP spectra and integrated magnitudes $M_{BP}$ and $M_{RP}$. Epoch spectra will be released only for specific classes of objects, such as variable stars and quasars, for example. The internal flux calibration of integrated magnitudes, including the $M_G$ magnitudes as well, is expected at a precision of 0.003 mag for $G=13$ stars, and for $G=20$ stars goes down to 0.07 mag in $M_G$, 0.3 mag in $M_{BP}$ and $M_{RP}$. The external calibration should be performed with a precision of the order of a few percent (with respect to Vega).

1.7. High-resolution spectroscopy

The primary objective of the RVS is the acquisition of radial velocities, which combined with positions, proper motions, and parallaxes will provide the means to decipher the kinematical state and dynamical history of our Galaxy.

The RVS will provide the radial velocities of about 100–150 million stars up to 17th magnitude with precisions ranging from 15 km s$^{-1}$ at the faint end, to 1 km s$^{-1}$ or better at the bright end. The spectral resolution, $R=\lambda/\delta\lambda$ will be 11 500. Radial velocities will be obtained by cross-correlating observed spectra with either a template or a mask. An initial estimate of the source atmospheric parameters will be used to select the most appropriate template or mask. On average, 40 transits will be collected for each object during the 5-year lifetime of Gaia, since the RVS does not cover the whole width of the Gaia AF (Figure 3). In total, we expect to obtain some 5 billion spectra (single transit) for the brightest stars. The analysis of this huge dataset will be complicated, not only because of the sheer data volume, but also because the spectroscopic data analysis relies on the multi-epoch astrometric and photometric data.

The covered wavelength range (847-874 nm) (Figure 5) is a rich domain, centered on the infrared calcium triplet: it will not only provide radial velocities, but also many stellar and interstellar diagnostics. It has been selected to coincide with the energy distribution peaks of G and K type stars, which are the most abundant targets. In early type stars, RVS spectra may contain also weak Helium lines and N, although they will be dominated by the Paschen lines. The RVS data will effectively complement the astrometric and photometric observations, improving object classification. For stellar objects, it will provide atmospheric parameters such as effective temperature, surface gravity, and individual abundances of key
elements such as Fe, Ca, Mg, Si for millions of stars down to G ≃ 12. Also, Diffuse Interstellar Bands (DIB) around 862 nm will enable the derivation of a 3D map of interstellar reddening.

1.8. The DPAC

ESA will take care of the satellite design, build and testing phases, of launch and operation, and of the data telemetry to the ground, managing the ESAC data center in Madrid, Spain. The data treatment and analysis is the responsibility of the European scientific community. In 2006, the announcement of opportunity opened by ESA was successfully answered by the Data Processing and Analysis Consortium (DPAC), a consortium that presently consists of more than 400 scientists in Europe (and outside) and more than 25 scientific institutions.

The DPAC executive committee (DPACE) oversees the DPAC activities; work has been organized among Coordination Units (CU) in charge of different aspects of data treatment:

- **CU1. System Architecture** (manager: O’Mullane), dealing with all aspects of hardware and software, and coordinating the framework for software development and data management.

- **CU2. Data Simulations** (manager: Luri), in charge of the simulators of various stages of data products, necessary for software development and testing.

- **CU3. Core processing** (manager: Bastian), developing the main pipelines such as IDT, AGIS and astrometry processing in general.

- **CU4. Object Processing** (managers: Pourbaix & Tanga), for the processing of objects that require special treatment such as minor bodies of the Solar system, for example.

- **CU5. Photometric processing** (manager: van Leeuwen), dedicated to the BP, RP, and $M_G$ processing and calibration, including image reconstruction, background treatment, and crowding treatment, among others.

- **CU6. Spectroscopic Processing** (managers: Katz & Cropper), dedicated to RVS processing and radial velocity determination.

- **CU7. Variability Processing** (managers: Eyer & Evans & Dubath), dedicated to processing, classification and parametrization of variable objects.

- **CU8. Astrophysical Parameters** (managers: Bailer-Jones & Thevenin), developing object classification software and, for each object class, software for the determination of astrophysical parameters.

- **CU9. Catalogue Production and Access** (to be activated in the near future), responsible for the production of astrophysical catalogues and for the publication of Gaia data to the scientific community.

These are flanked by a few working groups (WG) that deal with aspects that are common among the various CUs, such as the GBOG (Ground Based Observations Group), which coordinates the ground based observations for the external calibration of Gaia) or of general interest (such as the Radiation task force, serving as the interface between DPAC and industry in all matters related to CCD radiation tests).
2. The flux calibration of Gaia data

Calibrating (spectro)photometry obtained from the usual type of ground based observations (broadband imaging, spectroscopy) is not a trivial task, but the procedures are well known (see e.g., Bessell, 1999) and several scientists have developed sets of standard stars appropriated for the more than 200 photometric systems known, and for spectroscopic observations. Generally, magnitudes are calibrated to a standard system with equations in the form

\[ M = m + ZP + \alpha(\text{color}) + \beta(\text{airmass}) \]

where \( M \) is the calibrated magnitude in a chosen photometric band, \( m \), the instrumental magnitude in the same (or very similar) band, \( \alpha \) is the color term and \( \beta \), the extinction coefficient due to the Earth’s atmospheric extinction. For the spectra, the instrumental effect on the observed spectral energy distribution (SED) is parametrized as

\[ S_{\text{obs}}(\lambda) = R(\lambda) S(\lambda) \]

where the observed SED, \( S_{\text{obs}}(\lambda) \), is the result of the convolution of the “true” SED, \( S(\lambda) \), with all the instrumental (transmissivity, quantum efficiencies) effects, which are empirically determined in the form of a response curve, \( R(\lambda) \) through the use of spectrophotometric standard stars (SPSS). In the case of Gaia, several instrumental effects – much more complex than those usually encountered – redistribute light along the SED of the observed objects. In particular these are: the TDI integration mode, the large focal plane, radiation damage and resulting CTI (charge transfer inefficiencies), and that the whole instrumental model is well known only before launch.

2.1. Challenges

The most difficult Gaia data to calibrate are the BP and RP spectra, requiring a new approach to the derivation of the calibration model (Section 2.3.) and to the SPSS needed to perform the actual calibration (Section 3.). The large focal plane with its large number of CCDs makes it so that different observations of the same star will be generally on different CCDs, with different quantum efficiencies. Also, each CCD is in a different position, with different optical distortions, optics transmissivity and so on. Therefore, each wavelength and each position across the focal plane has its (sometimes very different) PSF (point spread function). The TDI and continuous reading mode, combined with the need of compressing the data before on-ground transmission, make it necessary to translate the full PSF into a linear (compressed into 1D) LSF (line spread function), which of course add complications. In-flight instrument monitoring is foreseen, but never comparable to the full characterization that will be performed before launch, so the real instrument – at a certain observation time – will be different from the theoretical one assumed initially, and this difference will change with time.

Radiation damage deserved special mention as it is one of the most important factors in the time variation of the instrument model. It has particular impact on the BP and RP dispersed images since the objects travel along the BP and RP CCD strips in a direction that is parallel to the spectral dispersion (wavelength coordinate). Radiation damage causes traps that subtract photons from each passing object at a position corresponding to a certain wavelength. Slow traps release the trapped charges once the object is already passed, while fast traps can release the charges within the same object, but at a different wavelength. Given the low resolution, one pixel can cover as much as 15–20 nm (depending on the wavelength) and therefore the net effect of radiation damage can be to alter significantly the SED of some spectra. Possible solutions under testing are the equivalent of CCD pre-flashing, the statistical modeling of the traps behavior and the fact that different transits for the same object will be affected differently by CTI effects, allowing for a certain degree of correction through average or median spectra. Finally the PSF/LSF itself is generally
larger than one Gaia pixel in the BP/RP spectra, introducing a large LSF smearing effect, i.e., the spread of photons with one particular wavelength into a large range of wavelengths.

In this paper, we will adopt the current Gaia calibration philosophy, where most of these instrumental effects are taken into account during the so-called internal flux calibration. A large number of well behaved stars (internal standards) observed by Gaia will be used to report all observations to a reference instrument, on the same instrumental flux and wavelength scales. All transits for each object observed by Gaia will be then averaged to produce one single BP and RP spectrum for each object, with its integrated instrumental magnitudes: \( M_G \), \( M_{BP} \), and \( M_{RP} \). Only for specific classes of objects will epoch spectra and magnitudes will be released, with variable stars as an obvious example. The mean and epoch spectra will be mostly free from many of the problems examined just above, but they will still contain residuals due to the imperfect knowledge of the real instrument at each precise moment of time, and the most significant effects are expected to be the LSF smearing and the CTI effects.

In this paradigm, the internal and external flux and wavelength calibrations are treated as two entirely separate and consecutive pieces of the CU5 photometric pipeline, with different calibration models. We always start from internally calibrated BP/RP spectra and \( M_G \), \( M_{BP} \), and \( M_{RP} \) magnitudes, without giving importance to the exact way they are produced. Presently, two alternative approaches are being considered to maximize the precision of the global calibration procedure: the first one is a hybrid model that partially combines internal and external models (Montegriffo et al. 2010), while the second is the so-called full forwarding model (Carrasco et al. 2010, in preparation), using the same calibration model for both the internal and external calibration.

2.2. The external calibration teams

Two development units (DU) within CU5 (Photometric processing), are dedicated to the external calibration of Gaia photometry. They are DU13: Instrument absolute response characterization: ground-based preparation, coordinated by E. Pancino and DU14, Instrument absolute response characterization: definition and application’, coordinated by C. Cacciari. They are both based in Bologna, Italy, in collaboration with the Bologna, Barcelona, and Groningen Universities. The actual team members at the time of writing are: G. Altavilla, M. Bellazzini, A. Bragaglia, C. Cacciari, J. M. Carrasco, G. Cocozza, L. Federici, F. Figueras, F. Fusi Pecci, C. Jordi, S. Marinoni, P. Montegriffo, E. Pancino, S. Ragaini, E. Rossetti, S. Trager.

2.3. Dispersion Matrix basic definition

If we concentrate now on the mean, internally calibrated BP/RP spectra calibration, we can write:

\[
S_{\text{obs}}(\kappa_I) = \sum_{\lambda_i=0}^{N} T(\lambda_i) L_{\lambda_i}(\kappa_I - \kappa_P(\lambda_i)) S_{\text{true}}(\lambda_i)
\]

where \( S_{\text{obs}} \) and \( S_{\text{true}} \) are the observed and “true” SEDs respectively, expressed the first in Gaia pixels \( \kappa_I \) and the second in wavelength intervals \( \lambda_i \) corresponding to the actual sampling of the SPSS used in the flux calibration process. \( T(\lambda_i) \) is a combination of all the instrument and telescope transmissivity functions and aperture, while \( L \) is the LSF at a certain \( \lambda_i \), centered at the appropriate \( \kappa_I \) pixel, but of course calculated over the whole wavelength interval from \( \lambda_i=0 \) to \( N \) (the total number of samples in the tabulated SPSS spectrum).

Such an equation can be written in its much simpler matrix form:

\[
S_{\text{obs}} = D \times S_{\text{true}}
\]
where $D$ is called a “Dispersion Matrix”, an object that can be determined if $S_{\text{obs}}$ and $S_{\text{true}}$ are known, i.e., using a well defined set of SPSS observed by Gaia, that also have well known SED (see below). Once $D$ is properly determined, it can be inverted to convert each mean, internally calibrated BP/RP spectrum\(^3\) ($S_{\text{obs}}$) into a flux calibrated spectrum $S_{\text{true}}$

$$S_{\text{true}} = D^{-1} \times S_{\text{obs}}$$

The main advantage of this approach is that $D$ contains (and therefore corrects empirically for) the actual effects of LSF smearing – even if the real shape of the LSF is not perfectly known a priori. More than that, the effective LSF – as determined with the chosen SPSS set – at each wavelength can be extracted from each column of the matrix. The matrix rows represent instead the effective passbands corresponding to each Gaia pixel, including the full effect of LSF smearing. This peculiar property of the dispersion matrix makes it the best (and possibly only) solution to the external calibration of Gaia BP/RP spectra. By definition, the dispersion matrix $D$ contains also the actual dispersion function, which can be seen in Figure 6 as the curved structure close to the diagonal of the matrix.

Finally, an important by-product of the described calibration model is the absolute wavelength calibration of the BP/RP spectra to a precision of at least a few tenths of a Gaia pixel\(^4\) (Montegriffo & Bellazzini 2009b), which is automatically performed together with the absolute flux calibration.

There are a few problems in the use of the dispersion matrix as proposed. We will discuss in the three following Sections the two most important ones and their proposed solutions: (1) the matrix is rectangular and its inversion is not so straightforward; (2) the matrix needs a set of independent vectors to be determined in a non-degenerate way (which also implies that the set of SPSS must be carefully chosen).

### 2.4. Smoothing the input SPSS spectra

Clearly, the dispersion matrix is a rectangular matrix: the Gaia observations have a smaller number of samples (pixels) than the SPSS spectra used to build their calibration model (wavelength sampling). Inverting a rectangular matrix is a non-trivial task so if we want

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\(^3\)Incidentally, for some object classes that need it, such as variable stars, single transits – the so called epoch spectra – will be published. The described calibration model can be applied also to single epoch spectra once they have been internally calibrated, i.e., reported to a common, instrumental scale of flux and wavelength.

\(^4\)This first estimate of the wavelength zero point and scale precision is based on a slightly outdated calibration model formulation, an therefore has to be considered just as an upper limit to a more realistic uncertainty (Montegriffo, 2010, private communication).
Figure 7: On both panels, grey dots are simulated spectra of different metallicity and reddening (this explains the parallel sequences). Green dots are white dwarfs and hot subdwarfs, while red symbols are different types of red stars with significant absorption features, i.e., molecular bands. Abscissae represent the BP–RP color, ordinates are the difference between the “known” magnitude of the used SEDs and the “calibrated” ones obtained with a dispersion matrix. Left: all points are calibrated with a matrix built only with hot spectral types (green dots): the reddest stars are calibrated with an error of 0.15 mag and more. Right: all points are calibrated with a matrix determined using also 10 red SPSS with absorption features (red symbols): all stars with SEDs similar to the 10 red stars (2 mag<BP–RP<4 mag) are calibrated with an error of less than 1%. ©ESA

to use the inverted matrix to calibrate our data we must find a method to reduce the dimensionality of the input SPSS flux tables. A few different methods have been considered such as the B splines representation (Montegriffo & Bellazzini, 2009a), Gaussian smoothing with variable width Gaussians (Montegriffo, 2010, private communication), and smoothing through rectangular functions corresponding to the Gaia pixels (Montegriffo et al. 2010). The idea is that the $S_{true}$ of each SPSS is observed from the ground with a resolution that exceeds the Gaia BP/RP one by at least a factor of 4-5, and then compressed in a way that minimizes information losses. Once the SPSS spectra are smoothed, the dispersion matrix becomes a square effective dispersion matrix, $D_e$

$$S_{obs} = D_e \times S_{smooth}$$

However, $D_e$ can still be non-diagonal or degenerate and before proceeding we must find the criteria to build the best possible $D_e$ with the data in hand.

2.5. The ideal SPSS set

One reason why the dispersion matrix is non-diagonal, is that the SPSS adopted set can never be an orthogonal set of independent calibrators: stars are all similar to each other, they have all a black-body like continuum with some features (absorption and emission lines or bands). As a result, if a dispersion matrix is built with a particular set of SPSS, such as white dwarfs and hot subdwarfs (the ideal calibrators in the classic spectroscopic observations), it will be able to calibrate properly only objects with similar spectra, i.e., relatively smooth, with some absorption lines in the blue part of the spectrum.

An example of the above case is shown in Figure 7, where two different dispersion matrices are used to calibrate the same set of simulated Gaia observations. In the first
case (left panel of Figure 7), a dispersion matrix is built using only white dwarfs and hot subdwarfs, with a minority of solar type stars, and it can clearly be seen that red stars with deep absorption bands are calibrated with an error of 0.15 mag at least, failing the specified requirements. In the second case (right panel of Figure 7), a small number (10) of red stars with deep absorption bands are included in the SPSS set used to build a second dispersion matrix. The second dispersion matrix is able to calibrate all red stars with absorption features to better than 1%, exceeding the requirements.

This example shows the importance of spectral features in the SPSS set used in construction of the dispersion matrix. Hot stars have prominent absorption lines in the blue, but no features in the red. The addition of a few red stars with absorption bands “trains” the matrix in the calibration of stars with features in the red (effectively reducing degeneracy). Similarly, problems are encountered in the calibration of emission line objects (peculiar hot stars and quasars, for example). But it is quite difficult to include these objects into the SPSS set since they are often variable.

Even if several types of objects are included when determining the dispersion matrix, other effects can have a large impact on the degeneracy, such as edge effects. For these reasons, the accurate choice of the SPSS set is crucial, but does not solve the problem of degeneracy once and for all.

2.6. Nominal dispersion matrix

To further reduce degeneracy of the effective dispersion matrix $D_{e}$, we can use other constraints such as the fact that we know most aspects of the instrument from pre-launch characterization. These include the quantum efficiency of the CCDs, the optical layout and transmissivity, the nominal LSF at various positions along the focal plane and at different wavelengths, the nominal dispersion function and its variation along the focal plane. The slow change of these with time can also be monitored to a certain extent, and included in the models.

We can therefore separate the dispersion matrix in a part that is theoretically modeled based on pre-launch instrument description and on its (partially reconstructed) variation with time, which we call $D_{n}$, or nominal dispersion matrix, and in a part that is completely
unknown, which can be considered as a correction matrix $K$, made of the residual corrections after the nominal model is taken into account (Montegriffo et al. 2010)

$$D_e = K \times D_n$$

The nominal matrix will be clean: diagonal and non-degenerate (see Figure 8). The correction matrix will be partially degenerate, but all signal that lies far away from the diagonal can be safely considered spurious (the system varies in a continuous way, the corrections must be “small” compared to the nominal system), and the part of the correction matrix close to its diagonal can be easily modeled.

To summarize all the previously defined steps, once an appropriate SPSS set is chosen, the calibration model becomes

$$S_{obs} = D_e \times S_{smooth} = K \times D_n \times S_{smooth}$$

and the matrices involved can be easily inverted to calibrate all Gaia observations since they are all square and (almost completely) diagonal.

### 2.7. Integrated magnitudes

A classical approach can be adopted for the absolute flux calibration of integrated $M_G$, $M_{BP}$, and $M_{RP}$ magnitudes (Ragaini et al., 2009a,b) in the form

$$M = m + ZP$$

where $M$ is the calibrated magnitude, $m$ the internally calibrated one observed by Gaia, and $ZP$ is the required zero-point. No significant color term appears necessary.

However, if we consider that an integrated magnitude $M$ is the convolution of the spectral distribution $S_{true}$ and the effective passband $B$, we can calibrate integrated magnitudes with the same approach adopted for BP/RP spectra, with a much more homogeneous procedure from the point of view of pipeline code writing

$$M = S_{true} \times B$$

Since generally the passband $B$ is sampled differently than the SPSS flux table $S_{true}$, we must smooth to one or both $S_{true}$ and $B$. Similarly to the case of BP/RP spectra, we can split $B$ into two components

$$B = K' \times B_n$$

where $K'$ is a correction vector, made of the actual residuals to a theoretically known – or nominal – effective passband $B_n$, known before launch and slowly varying with time due to several causes, the most important one being the decrease in CCD quantum efficiency due to radiation damage. With this kind of treatment, the problem becomes a simple least square fitting problem to derive the unknown $K'$ vector (Ragaini et al., 2010, in preparation).

### 2.8. RVS calibration

The possibility of flux calibrating the RVS spectra has been so far considered a secondary problem, since both radial velocities and astrophysical parameters can be derived without the need of an absolute flux scale attached to the spectra. A preliminary set of considerations (Trager, 2010) shows that in principle the SPSS grid for the calibration of Gaia G-band and BP/RP data, that is presently under construction, should be sufficiently sampled to ensure a flux calibration of RVS spectra as well. We expect the topic to be further explored by CU6 (Spectroscopic processing) in the near future, but we will not consider RVS spectra flux calibration in this paper.
3. The Gaia grid of spectrophotometric standard stars

From the above discussion, it is clear that the Gaia SPSS grid has to be chosen with great care. The Gaia SPSS, or better their reference flux tables (corresponding to $S_{\text{true}}$ in the previous Sections) should conform to the following general requirements (van Leaven et al. 2010):

- Resolution $R = \lambda/\delta\lambda \simeq 1000$, i.e., they should over sample the Gaia BP/RP resolution by a factor of 4–5 at least;
- Wavelength coverage: 330–1050 nm;
- Typical uncertainty in the absolute flux scale, with respect to the assumed calibration of Vega, of a few percent, excluding small troubled areas in the spectral range (telluric bands residuals, extreme red and blue edges), where it can be somewhat worse.

The total number of SPSS in the Gaia grid should be of the order of 200–300 stars, including a variety of spectral types. Clearly, no such large and homogeneous dataset exists in the literature yet\(^5\). It is therefore necessary to build the Gaia SPSS grid with new, dedicated observations. We describe the characteristics of the Gaia SPSS and of the dedicated observing campaigns in the following sections.

3.1. SPSS Candidates

We have followed a two step approach (Bellazzini et al. 2007) that first creates a set of Primary SPSS, i.e., well known SPSS that are calibrated on the three Pillars of the CALSPEC\(^6\) set, described in Bohlin et al. (1995,2007), and tied to the Vega flux calibration

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\(^5\)The CALSPEC database (Bohlin, 2007) is not large enough for our purpose, especially considering the strict criteria described below. Its extension to more than 100 SPSS is eagerly awaited, but still not available to the public.

\(^6\)http://www.stsci.edu/hst/observatory/cdbs/calspec.html
Figure 10: The preliminary reductions (no telluric correction, only TNG (Telescopio Nazionale Galileo) observations, library extinction curve, and so on) of star GD 71 (top panel) are compared with the CALSPEC spectrum (bottom panel). Prominent telluric features are marked in both panels. Except for the spectral edges – which will need to be reconstructed with the use of models – the main body of the spectrum is always close to the CALSPEC one within 1% or better. ©ESA

by Bohlin & Gilliland (2004) and Bohlin et al. (2007). An example of the kind of spectra obtained for Pillar GD 71 (with DoLoRes at the TNG is shown on Figure 10. The Primary SPSS will constitute the ground-based calibrators of the actual Gaia grid, and need to conform to the following requirements (van Leeuwen 2010):

- Primary SPSS have spectra as featureless as possible;
- Primary SPSS shall be validated against variability;
- Primary SPSS have already well known SEDs;
- The magnitude of each Primary SPS results in S/N≈100 per pixel over most of the wavelength range when observed from the ground with 2m class telescopes;
- The location of Primary SPSS is in non crowded areas of the sky;
- Primary SPSS cover a range of RA and Dec to ensure all-year-round ground based observations from both hemispheres.

The Primary SPSS candidates set is described in more detail in Altavilla et al. (2008), and some of the most important sources for Primary candidates are the CALSPEC grid, Oke (1990), Hamuy et al. (1992,1994), Stritzinger et al. (2005) and others. The actual Gaia SPSS grid, or Secondary SPSS, conforms to a different set of requirements (van Leeuwen 2010):

- Secondary SPSS have spectra as featureless as possible (but see below for exceptions);
- Secondary SPSS shall be validated against variability;
- The magnitude and sky location (i.e., number of useful, clean transits, see Carrasco et al. 2006, 2007) of Secondary SPSS grants a resulting $S/N \approx 100$ per sample over most of the wavelength range when observed by Gaia (end of mission);

- Secondary SPSS cover a range of spectral types and spectral shapes, as needed to ensure the best possible calibration of all kinds of objects observed by Gaia.

As already mentioned, Secondary SPSS will be mostly hot and featureless stars but will include a small number of selected spectral types, to ensure that the calibration model can work on all object types. More details on Secondary SPSS can be found in Altavilla et al. (2010), including a long list of literature catalogs and online databases from which the candidates are extracted. Clearly, all the Primary SPSS that, at the end of the data reductions, will satisfy also the criteria for Secondary SPSS, will be included in the Gaia SPSS grid.

Additionally, special members of the Secondary SPSS candidates are: (1) a few selected SPSS around the Ecliptic Poles, two regions of the sky that will be repeatedly observed by Gaia, in the first two weeks after reaching its orbit in L2, for calibration purposes; (2) a few M stars with deep absorption features in the red; (3) a few SDSS stars that have been observed in the SEGUE sample (Yanny et al., 2009), since the SEGUE sample has the potential of being extremely useful in the Gaia flux calibration (Bellazzini et al. 2010), both internal (relative) and external (absolute); (4) a few well known SPSS that are among the targets of the ACCESS mission (Kaiser et al., 2010), dedicated to the absolute flux measurement of a few stars besides Vega.

3.2. Observation strategy and campaigns

A basic consideration when starting the observations of such a large campaign, is that the traditional spectrophotometry techniques require too much observing time: each SPSS should be spectroscopically observed in perfect photometric conditions, ideally more than once. No TAC (Time Allocation Committee) would grant such a large amount of observing time to a proposal that does not contain any cutting edge science in it.

We therefore chose (Bellazzini et al. 2007) to split the problem into two parts: (1) spectra are taken in good sky conditions, but not necessarily perfectly photometric; they
are calibrated with the help of a Pillar or Primary SPSS thus recovering the correct spectral shape; (2) absolute photometry in the B, V, and R (sometimes I) Johnson-Cousins bands is taken in photometric sky conditions and used to fix the spectral zero-point by means of synthetic photometry. This is motivated by the fact that absolute photometric night points are faster to obtain than spectra. A subset of SPSS candidates will be spectroscopically observed in photometric sky conditions, to check the whole procedure.

Besides the Main Campaign just described, it is necessary to monitor candidate SPSS for constancy (Auxiliary Campaign), since very few of them have systematically been monitored in the literature, and there are illustrious examples of stars that showed unexpected variability (Landolt & Uomoto, 2007). An example of a different kind of problem, that could greatly benefit from good quality dedicated imaging, is star HZ 43. It was initially chosen by Bohlin et al. (1995) to be one of the Pillars, and later rejected because of an optical companion lying 3' away, only visible in the V band, that made it useless as an SPSS from the ground.

Most of our SPSS candidates are WDs close to the instability strip, and, which, sometimes have poorly known magnitudes, so it is necessary to monitor them for short-term variability on scales of 1–2 hours (Figure 11). Binary systems are frequent and can be found at all spectral types, so we also monitor all our candidates for long-term variability (3 years) collecting approximately 4 night points per year. These two monitoring campaigns rely on relative photometry (using stars in the field of view) to derive variability curves. An SPSS candidate is considered constant if it does not vary with an amplitude larger than a few milli-mags.

The facilities that are being used for the two observing campaigns are (Federici et al., 2007, Altavilla et al. 2010):

- EFOSC@NTT (New Technology Telescope) La Silla, Chile (primarily Main campaign);
- ROSS@REM (Rapid Eye Munt Telescope), La Silla, Chile (primarily Auxiliary campaign);
- LARUCA@1.5m, San Pedro Martir, Mexico (primarily Auxiliary campaign and absolute photometry);
- BFOSC@Cassini, Loiano, Italy (primarily Auxiliary campaign);
- CAFOS@2.2m, Calar Alto, Spain (primarily Main campaign);
- DoLoRes@TNG (Telescopio Nazionale Galileo), La Palma, Spain (primarily Main campaign);

Observations started in the second half of 2006, comprising more than 35 accepted proposals. We have been awarded a total of 230 observing nights approximately, at the rate of 33 per semester. More than 50% of this time resulted in at least partially useful data. Given the large number of facilities involved, and of different observers, it has become necessary to establish strict observing protocols (Pancino et al. 2008,2009). The campaigns are now more than 50% complete, with the spectroscopy observations 75% complete, and we expect to complete all our observing campaigns around 2013.

3.3. Data reduction and analysis

The large amount of data collected needs to be reduced and analyzed with the maximum possible precision and homogeneity. An initial set of data is collected for each CCD/instrument/telescope combination and an Instrument Familiarization Plan (IFP) is
conducted, to derive shutter times, linearity, calibration frames and lamps stability, photometric distortions, 2\textsuperscript{nd} order contamination of spectra (Figure 12), and so on. This plan is now almost complete, and the protocols are presently being finalized and written.

The data reduction is regulated by strict data reduction protocols, that are presently being finalized. While the data reduction methods are fairly standard, care must be taken in considering the characteristics of each instrument, as determined during the IFP, to extract the highest possible quality from each instrument. Semi-automatic quality check (QC) criteria are defined for each kind of observation (minimum & maximum S/N, seeing and roundness requirements of images, presence of bad columns, companions, and so on). Only frames that pass the QC are reduced. For imaging, we term “data reduction” the removal of the instrument characteristics (dark, bias, flat-field, fringing), QC, and the measurement of aperture photometry with SExtractor (Bertin & Arnouts, 1996). The data products are 2D reduced images and aperture magnitude catalogues. For spectroscopy, we term “data reduction” the removal of instrumental features (dark, bias, flat-field, illumination correction, wavelength calibration, 2\textsuperscript{nd} order contamination correction, relative flux calibration, telluric features removal), followed by QC and spectra extraction. The data products are 2D reduced frames, 1D extracted and wavelength calibrated spectra, 1D flux calibrated spectra, 1D telluric absorption corrected spectra.

The data reduction procedures are well advanced for photometry (almost half of the data reduced) and have just started for spectra (10\% of the data reduced) at the moment of writing.

The data analysis is presently in the design and testing phases. The study of short-term variability curves is proceeding (10\% of the data analyzed, see Figure 11). Absolute photometry and relative spectroscopy procedures are being refined: for example, preliminary end-to-end reduction of photometric imaging nights have been performed for TNG and NTT observations, to allow us to identify those nights that were actually photometric and did not need to be repeated. Preliminary extinction curves have been determined for TNG and CAHA (Centro Astronómico Hispano Alemán) (spectroscopic observations, allowing us to see
that extinction varies in a grey manner (within a few percent) even in the case of some Calima (desert dust) in the sky over La Palma.

The final data products for the Auxiliary campaign will be relative magnitudes and lightcurves for all the monitored candidate SPSS; for the Main campaign, absolute magnitudes and errors will be released together with their uncertainties and flux tables (Figure 10) in the form \((\lambda\text{nm})F_{\lambda}\text{(photons s}^{-1}\text{m}^{-2}\text{nm}^{-1})\). Possibly, also other intermediate data products will be released (see above).

3.4. Data availability

All the data SPSS ground-based observations, along with the collected literature information and measurements, are stored in the CU5-DU13 local Wiki pages in Bologna (Wiki-Bo)\(^7\). Wiki-Bo also contains all our technical documentation, internal reports, observation status and data products, along with literature references and sources, observing proposals and the like. The raw and reduced data products are stored in a local archive\(^8\) for internal purposes (Figure 13).

In the future, when CU9 begins (Catalogue production and access), it is foreseen that all the ground-based data used for the calibration of Gaia data (radial velocity standards, SPSS, spectral libraries, Ecliptic pole observations, observations of Gaia itself from the ground, and so on) will be published as well, although no decision on the format and type of data products has been made yet.

4. Conclusions

The Gaia mission and its data reduction is a challenging enterprise, carried out by ESA and the European scientific community. As an example of the DPAC (Data Processing and Analysis Consortium) tasks, I have briefly summarized the problem of the external (absolute) flux calibration of spectrophotometric Gaia data, and more specifically of the BP/RP low resolution spectra and the integrated G-band and BP/RP magnitudes. An innovative calibration model is presently under study and testing, and a large (\(\simeq 200 – 300\))

\(^7\)http://yoda.bo.astro.it/wiki, guest username and password can be obtained from E. Pancino.

\(^8\)http://spss.bo.astro.it, guest username and password can be obtained from E. Pancino.
grid of SPSS with 1–3% flux calibration with respect to Vega is being built from multi-site ground-based observations.

But once the Gaia data will become available, a greater challenge will have to be faced: the impact in almost all fields of astrophysics require that the scientific community (and not only the European one) be adequately prepared to extract the most scientific output from the data. The training of a new generation of scientists, and the collection of complementary data, necessary to answer key questions when combined with Gaia data, should start now. The challenge requires that large groups of scientists get efficiently organized and ready to collaborate on large and comprehensive datasets.

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ACCESS: Mission Overview, Design and Status


Johns Hopkins University, Baltimore, MD, USA 21218

Bernard J. Rauscher, Jeffrey W. Kruk, Randy A. Kimble, Dominic J. Benford, Roger Foltz, Jonathan P. Gardner, Robert J. Hill, Duncan M. Kahle, Eliot Malumuth, D. Brent Mott, Augustyn Waczynski, Yiting Wen, Bruce E. Woodgate

NASA Goddard Space Flight Center, Greenbelt, MD USA 20771

Ralph C. Bohlin, Susana Deustua

Space Telescope Science Institute, Baltimore, MD, USA 21218

Robert Kurucz

Harvard Smithsonian Center for Astrophysics, Cambridge, MD, USA 02139

Michael Lampton

Space Sciences Laboratory, Berkeley, CA 94720

Saul Perlmutter

University of California, Berkeley, Berkeley, CA 94720

Edward L. Wright

University of California, Los Angeles, Los Angeles, CA 90024

Abstract.

ACCESS, Absolute Color Calibration Experiment for Standard Stars, is a series of rocket-borne sub-orbital missions and ground-based experiments designed to enable improvements in the precision of the astrophysical flux scale through the transfer of absolute laboratory detector standards from the National Institute of Standards and Technology (NIST) to a network of stellar standards with a calibration accuracy of 1% and a spectral resolving power of 500 across the 0.35 – 1.7 µm bandpass.

Establishing improved spectrophotometric standards is important for a broad range of missions and is relevant to many astrophysical problems. Systematic errors associated with problems such as dark energy now compete with the statistical errors and thus limit our ability to answer fundamental questions in astrophysics.

The ACCESS payload and ground calibration components currently span a range of readiness levels extending from the design phase, through procurement, fabrication, and component test phases. The strategy for achieving a <1% spec-

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1Joint appointment with Space Telescope Science Institute
trophotometric calibration accuracy, a description and status of the instrument and the ground calibration system, and the NIST traceability components are discussed.

1. Motivation

A variety of fundamental problems in astrophysics need a precise (better than 1%) network of astrophysical flux standards spanning a wide dynamic range across the visible and near-infrared (NIR) bandpass to address critical questions with the required uncertainty. Understanding the expansion history of the universe, and its implications, is one such key problem.

The discovery of the accelerated expansion of the universe (Riess et al., 1998, Perlmutter et al., 1999) compared the standardized brightness of high redshift \( (0.18 < z < 1.6) \) Type Ia supernovae (SNe Ia) to low-redshift SNe Ia, showing that at a given redshift the peak brightness of SNe Ia is fainter than predicted (Riess et al. 2004a, Riess et al. 2004b, Knop et al. 2003, Tonry et al. 2003, Hamuy et al. 1996). This implies they are further away than expected, indicating a period of accelerated expansion of the universe and the presence of a new, unknown, negative-pressure energy component - dark energy.

Accurate testing of dark energy models through observation of SNe Ia depends on the precise determination of the relative brightness of the SNe Ia standard candles. For each supernova, its redshift, \( z \), is plotted against its rest-frame B-band flux to determine the SNe Ia Hubble brightness-redshift relationship. Cosmological and dark energy parameters are determined from the shape, not the absolute normalization, of this relationship. Since the rest-frame B-band is seen in different bands at different redshifts, the relative zero-points of all bands from 0.35 \( \mu m \) to 1.7 \( \mu m \) must be cross-calibrated to trace the supernovae from \( z = 0 \) to \( z \sim 1.5 \). The term “absolute color calibration” is defined as the slope of the absolute flux distribution versus wavelength.

Controlling the systematic errors to this level of accuracy and precision is required, not only for the absolute color calibration, but also for other sources of systematic errors which themselves depend on the color calibration (e.g. extinction corrections due to the Milky Way, the SN host galaxy, and the intergalactic medium, in addition to the K-corrections which provide the transformation between fluxes in the observed and rest-frame pass-bands).

Using SNe Ia to distinguish dark energy models from one another levies a requirement for 1% precision in the cross-color calibration of the SNe Ia flux across a bandpass extending from 0.35 – 1.7 \( \mu m \).

However the systematic errors in the flux calibration network spanning the visible through the NIR currently exceed 1%. The astrophysical flux scale is ultimately pinned to a single star, Vega (Hayes & Latham 1975, Hayes 1985) through ground based measurements of a fundamental standard, a melting point black body furnace (Oke & Schild 1970). Since that fundamental calibration in 1970, there have been significant technological advances in detectors, telescopes and instruments, and the precision of the fundamental laboratory standards used to calibrate these instruments. These technologies have not yet been transferred to the fundamental astrophysical flux scale across the visible to NIR bandpass. In addition, Vega, the ultimate fundamental stellar standard for the astrophysical flux scale, has proven to be a poor choice for the following reasons. Vega is a pole-on rotator (Hill et al. 2010, Yoon et al. 2008, Aufdenberg et al.2006). As a result, it presents a range of effective temperatures and is difficult to model uniquely. It exhibits an infrared excess that arises from its circumstellar debris disk. And, it is too bright to be observed with today’s premier optical telescopes.

Recently, programs have been initiated to approach facets of this problem from various perspectives, primarily to account for the impact of the Earth’s atmosphere on observations through dedicated simultaneous monitoring of the atmosphere with ground-based compan-
ion telescopes (Burke et al. 2010, Stubbs et al. 2007) or LIDAR observations (Zimmer et al. 2010, McGraw et al. 2010). A program is in progress to compare ground-based visible measurements of stars with on-site measurements of fundamental standards to establish a set of stellar flux standards (Smith et al. 2009). This paper will describe the ACCESS program’s strategy to obtain a set of precise (<1%) absolute flux measurements spanning the visible into the NIR for a set of stellar flux standards through observations from a rocket-borne telescope with a calibration tied to National Institute of Standards and Technology (NIST) fundamental irradiance (detector) standards.

2. ACCESS Overview

ACCESS - “Absolute Color Calibration Experiment for Standard Stars”, is a series of rocket-borne sub-orbital missions and ground-based experiments that will enable the absolute flux for a limited set of primary standard stars to be established using calibrated detectors as the fundamental metrology reference (Kaiser et al. 2010, Kaiser et al. 2008). These experiments are designed to obtain an absolute spectrophotometric calibration accuracy of <1% in the 0.35 – 1.7 µm bandpass at a spectral resolution greater than 500 by directly tracing the observed stellar fluxes to NIST irradiance standards. Transfer of the NIST detector standards to our target stars will produce an absolute calibration of these standards in physical units, including the historic absolute standard Vega, the Sloan Digital Sky Survey (SDSS) standard BD+17°4708 (Fukugita et al. 1996), and the Spitzer IRAC standard HD 37725 (Reach et al. 2005).

The ACCESS payload and ground calibration components currently span a range of readiness levels extending from the design phase, through procurement, fabrication, and component test phases (Kaiser et al. 2010). This paper will present the strategy for achieving a <1% spectrophotometric calibration accuracy, a description and status of the instrument and the ground calibration system, and the NIST traceability components.

3. Observing Strategy

ACCESS has adopted a multi-faceted strategy for reducing systematic errors and establishing the first members in a network of standard stars calibrated to 1% precision.

First in this multi-pronged approach is the elimination of the Earth’s atmosphere as a spectral contaminant. One of the prime impediments to obtaining a precise flux calibration in the NIR is the plethora of hydroxyl (OH) lines that are formed high (∼70 km - ∼89 km) in the Earth’s atmosphere (Moreels et al. 1977). These spectral lines are strong, numerous - extending across 0.85 µm < λ < 2.25 µm, and highly variable on short timescales. Correcting ground based observations to 1% precision for absorption by the Earth’s atmosphere in the NIR is prohibitive. Spectral contamination by OH emission remains daunting even at balloon altitudes.

As a result, the ACCESS observations will execute from a sounding rocket platform completely above the Earth’s atmosphere. The typical launch trajectory for a Black Brant IX sounding rocket carrying a 1000 lb payload yields an apogee of 300 km with the time above 100 km in excess of 400 seconds. Observing at an altitude exceeding 100 km eliminates the challenging problem of measuring the residual atmospheric absorption and strong atmospheric emission seen by ground-based observations and even by observations conducted at balloon altitudes. The number, strength, and variability of the OH lines was an important factor in the selection of a sounding rocket platform for the ACCESS observations.

Next is the judicious selection of standard stars. Stars were selected as potential ACCESS flux standards based upon their ability to serve as precise calibration standards in both the visible and the NIR spectral region. This required that they (1) have existing photometric heritage, (2) present a stellar atmosphere capable of being precisely modelled,
and are (3) bright enough to obtain a spectrum with S/N of 200 in a single rocket flight. Since our program includes observations of Sirius and Vega, which are 12 hr apart on the sky, this also imposed the requirement that (4) additional targets be within 45° of the zenith on either the Sirius or Vega flight. To optimally utilize the flight time and to eliminate a step in the calibration transfer to larger telescopes, we required that (5) all targets other than Vega and Sirius be fainter than 6th magnitude. After satisfying these constraints, (6) candidate targets were prioritized according to their ability to further our understanding of stellar physics through model atmosphere calculations. The current set of ACCESS targets is presented in Table 1.

Third in the reduction of systematic errors is the design of a spectrograph with a bandpass extending from the visible through the NIR to eliminate the cross-calibration between these historically distinct bandpasses. Key to achieving this broad bandpass is the use of WFC3 heritage HgCdTe detector whose sensitivity extends from below 4000 Å to ~1.7 µm.

Another component of the calibration strategy is to establish an a priori error budget and track it.

Also essential, is the use of multiple methods to determine the payload sensitivity through NIST traceable sub-system and end-to-end payload calibrations. A key component of the calibration strategy is the use of NIST calibrated photodiode detector (irradiance) standards to establish NIST traceability. This is a fundamental component of the calibration. The NIST traceable photodiode irradiance standards will be used to establish the artificial star used to calibrate the integrated telescope and spectrograph. In addition to both an absolute and relative calibration of the payload, and hence the stars, the irradiance standards will establish the calibration system in fundamental physical units. Standard detectors have stability heritage exceeding 15 years in the visible for silicon photodiodes and 10 years for the NIR (e.g. Ge, InGaAs) photodiodes.

End-to-end payload calibrations will be performed with both NIST continuum and emission line radiance standards as cross-checks for systematic errors.

Knowing the performance of the instrument immediately prior to and after observations of the target stars is also a critical component of the calibration. Consequently, ACCESS includes an On-board Calibration Monitor (OCM) to track the payload performance in the field prior to launch and while parachuting to the ground post-observation. The OCM will be calibrated at the same time as the payload and re-calibrated and monitored between flights.

Current target list for the ACCESS program.

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<td>HD 37725</td>
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<td>+18 5 33.2</td>
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Table 1: Current list of the ACCESS program targets.
4. ACCESS Telescope and Spectrograph

The prime disadvantage of the sounding rocket platform is the short observing time. Since the amount of time above the Earth’s atmosphere is a function of payload mass, the selection of a sounding rocket platform for the ACCESS observations requires a compact and lightweight telescope and spectrograph design to maximize the time the payload is above the atmosphere.

The ACCESS telescope is a Dall-Kirkham cassegrain with Zerodur mirrors. The primary mirror has a 15.5-inch diameter. The telescope mirrors will be coated with aluminum with a protective MgF$_2$ overcoat. This coating was chosen in order to retain sensitivity in the 3500−4000 Å range where the higher order Balmer lines provide leverage for stellar atmosphere modeling. The telescope feeds a low-order echelle spectrograph with a cooled, substrate-removed, HgCdTe detector.

The spectrograph is a Rowland circle design configured as an echelle (Fig. 1) used in 1st (9000−19000 Å), 2nd (4500−9500 Å) and 3rd orders (3000−6333 Å), with some sensitivity in the 4th order. It consists of just two figured optical elements: a concave diffraction grating with a low ruling density, and a Fery prism by Ferre1911 with spherically figured surfaces placed in the converging beam. The prism provides astigmatism correction and cross-disperses the spectrum.

The flight grating (Fig. 2) is a mechanically ruled (blazed), quad-partite, concave grating with a radius of curvature of 400.7 mm, a groove density of 45 lines per mm, and a groove depth of ~660 nm. It is on a BK7 substrate coated with aluminum and a SiO$_x$ protective overcoat. The need for a compact spectrograph design to fit within a 17.26-inch outer diameter of the rocket skin motivated the Rowland mount for the grating. The spectrograph design is optimized to use the grating through a range of small diffraction angles. This restricts the astigmatism generated by the Rowland circle mount configuration. Component level efficiency measurements will be made under a nitrogen gas purge when the ground calibration fixture has been completed.

The prism needs to provide high, relatively uniform, transmission across the bandpass extending from 3500 Å to 1.8 μm. It must also provide sufficient dispersion to separate orders and isolate the wings of the line spread function of adjacent orders on the detector (Fig. 3 Right). Based on optical models for the prism using NBK10 glass, the minimum separation expected between adjacent orders on the detector is 24 pixels (437 μm).
Figure 2: The flight grating. The grating is ruled along its full length in the spatial (figure horizontal) direction. The dispersion direction is in the vertical direction. The grating is ruled in four partites to maintain efficiency.

The resolution of the spectrograph depends on the telescope point spread function (PSF) and the size of the detector pixels. For the telescope PSF of 1.17″ (as achieved on recent flights with a similar design) the 18 µm pixels of the detector provide critical sampling and produce constant wavelength resolution elements in each order, giving a resolving power ranging between 500 and 1000 (Fig. 3).

The spectrograph housing (McCandliss et al. 1994) will be evacuated and mounted to the back of the primary mirror base plate. An angled mirrored plate with a 1 mm aperture in the center located at the telescope focus serves as the slit jaw, allowing light to enter the spectrograph while reflecting the region surrounding the target into an image-intensified video camera for real-time viewing and control by the operator on the ground.

The optical elements are sealed in a stainless steel vacuum housing to provide for contamination control, thermal stability, and calibration. A fused silica entrance window sits behind the slit jaw and provides the seal for the spectrograph vacuum housing. The grating and cross disperser are mounted inside along with a set of baffles. The detector is mounted on a bellow that is used as part of the focus adjustment mechanism. The spectrograph vacuum is maintained by a non-evaporable getter and is monitored by an ion gauge. The typical vacuum is \(<10^{-6}\) Torr.

The focal plane detector array is a HgCdTe device with a 1024×1024 pixel format and 18 µm × 18 µm pixels. The detector array is indium bump-bonded to a Hawaii 1-R multiplexer resulting in 1014×1014 active imaging pixels and 5 rows and columns of reference pixels at each edge. The reference pixels are connected to capacitive loads rather than active imaging pixels; they track the effects of thermal drift and low frequency noise.

The detector composition is tailored to produce a long-wavelength cutoff at \(\sim 1.7\mu m\). The CdZnTe growth substrate is removed to provide high near-IR quantum efficiency (QE) and response through the visible to 3500Å. The flight detector QE (Fig. 4) is \(\sim 20\%\) at 5000Å. This device was developed for the HST/Wide Field Camera 3 (WFC3) program by Teledyne Imaging Sensors (TIS). The band gap corresponding to the 1.7 µm cutoff yields low dark current at operating temperatures near 140K and results in a detector that is relatively insensitive to thermal background radiation (Fig. 4). A cold baffle at \(\sim 200K\) is mounted in front of the detector. The inclusion of additional thermal baffles is under evaluation.

The flight detector was selected based upon initial characterization tests. The QE (Fig. 4) and dark tests have been repeated to ensure that detector performance has not
Figure 3: Spot diagrams for the optical layout at a selection of wavelengths spanning the ACCESS bandpass. Each box is labelled in microns and each spot is referenced to the chief ray. In general, the ray trace yields images that are within 2 pixels in the dispersion direction, \(\sim 4\) pixels in the cross-dispersion direction, with a minimum separation between orders of 24 pixels (18x18 \(\mu\)m pixels).

Figure 4: Left: Flight detector quantum efficiency (with the interpixel capacitance correction applied). Right: Dark current at an operating temperature of 145K.
changed and to extend QE measurements to the shorter wavelengths in the ACCESS band-pass. Reciprocity failure testing has been conducted to measure the count rate dependent non-linearity (Hill et al. 2010).

Characterizing the throughput of the ACCESS instrument using the a mix of predicted and measured component efficiencies for the grating design, transmissive optics, reflective coatings, and the flight detector QE (Fig. 5: Left) indicates that a S/N of 200 per spectral resolution element over the full spectral bandpass will be achieved easily for the 0th V magnitude stars Sirius and Vega (Fig. 5: Right). Subsecond integration times will be required to avoid saturation of the detector for these bright stars. Proven algorithms for sub-array readouts of the detector will be used to accomplish this.

For the fainter targets (e.g., BD+17°4708) a 400 second observation yields a S/N of 200 per spectral resolution element down to the Balmer edge. Additional binning can further increase the background subtracted signal-to-noise ratio of the acquired spectrum.

5. ACCESS Calibration

5.1. Calibration Overview

The ACCESS calibration program consists of five principal components.

1. Establish a standard candle that can be traced to a NIST detector-based irradiance standard.

2. Transfer the NIST calibrated standard(s) to the ACCESS payload - calibrate the ACCESS payload with NIST certified detector-based laboratory irradiance standards.

3. Transfer the NIST calibrated standard(s) to the stars - observe the standard stars with the calibrated ACCESS payload.

4. Monitor the ACCESS sensitivity - track system performance in the field prior to launch, while parachuting to the ground, and in the laboratory, to monitor for changes in instrument sensitivity.

Figure 5: Left: Component sensitivities for the grating design, transmissive optics, reflective coatings, and the flight detector QE across the full spectral bandpass. The efficiency for each order is also plotted. Right: The flux limit for a signal-to-noise of 200 in a 400 second rocket flight is shown for each of the four ACCESS orders imaged on the detector. The flux for three selected targets is over plotted for comparison. Except at the very shortest wavelengths for BD+17°4708, a signal-to-noise ratio of 200 is achievable at a resolving power of 500 in a single rocket flight.
Figure 6: Auto-collimator subsystem used to calibrate the reflectivity of the collimator primary and secondary mirrors. From the right, a light source (not shown) feeds a vacuum monochromator, which is fiber fed into the auto-collimator housing to an order blocking filter and input to an integrating sphere. The output of the integrating sphere is baffled to match the collimator f-ratio. The light then passes through a pinhole at the focus of the collimator and is incident on the NIST standard photodiode detector which is positioned to measure the beam entering or exiting the auto-collimator. The auto-collimating flat mirror (yellow) is on the left. This assembly calibrates the collimator that will be feeding the artificial star-at-infinity to the telescope.

5. Fit stellar atmosphere models to the flux calibrated observations - confirm performance, validate and extend standard star models.

The crux of the ACCESS instrument sensitivity is knowing the ratio of the total number of photons entering the telescope aperture to the total number of photons detected by the spectrograph detector as a function of wavelength.

Determination of the number of photons entering the telescope requires a source with a known number of photons in a beam matched to the entrance aperture of the telescope. For ACCESS, this source (Fig. 7 Bottom) consists of a stellar simulator and a collimator to provide the "star-at-infinity" required by the telescope. The stellar simulator (Fig. 7) is comprised of an illuminated pinhole placed at the collimator focus where the light illuminating the pinhole originates with a quartz tungsten halogen (QTH) continuum lamp fed to an integrating sphere to erase the spatial signature of the tungsten filament. The light output from this integrating sphere is fiber fed to the vacuum monochromator. The output from the monochromater is fiber fed, through an evacuated tube, into the 22-inch cylindrical vacuum housing. There it illuminates an order blocking filter, followed by an integrating sphere to generate a spatially uniform, unpolarized beam that is input to a baffle box to match the focal ratio of the stellar simulator beam to the F/12 collimator.

The total number of photons in the "star-at-infinity" output beam of the collimator will be provided by two measurements. The first is a measurement of the reflectivity of the auto-collimating flat mirror using a NIST calibrated photodiode detector standard in a nitrogen gas purge environment. The second is a measurement of the intensity of the light from the stellar simulator that is input to and reflected by the auto-collimator. Figure 6 illustrates the auto-collimator calibration configuration used to determine the reflectivity of the primary and secondary mirrors. The absolute calibration transfer is obtained through the use of a NIST calibrated photodiode detector transfer standard to measure the input and output images at the auto-collimator focal plane.

The end-to-end calibration of the telescope with spectrograph (Fig. 7: Bottom) is then performed as a function of wavelength by measuring the intensity of the simulated star, measuring the count rate at the spectrograph detector, correcting for the collimator attenuation of the simulated star by the reflectivity of the collimator primary and secondary
mirrors, and dividing the calibrated radiant flux by the illuminated area of the primary mirror.

Although simple in principle, systematic effects, such as the uniformity of reflective coatings, matching of the collimator and telescope apertures, the spatial uniformity of the photodiode detectors, the transmission of the slit, the scattered light determination, the determination of the area of the primary and secondary telescope mirrors, the stability of the light source, etc., must be closely tracked if this process is to yield the required precision and accuracy.

In addition to the calibration of the ACCESS payload to a fundamental irradiance standard, ACCESS will be calibrated to both continuum and emission line fundamental radiance standards. Using emission from tunable lasers, the NIST SIRCUS facility (Brown et al. 2006) will provide an end-to-end radiance calibration transfer to ACCESS while the use of a spectral light engine will calibrate ACCESS using a continuum spectral energy distribution similar to the spectral energy distribution of the stellar targets (Brown et al. 2006).

5.2. Calibration Monitoring

An On-board Calibration Monitor (OCM) will track sensitivity changes in the telescope and spectrograph as a function of time after the absolute calibration data have been collected (Kaiser et al. 2010, Kaiser et al. 2008, Kruk et al. 2008). The current baseline design (Fig. 8) of the OCM uses 8 pairs of feedback stabilized LEDs, with central wavelengths spanning the ACCESS bandpass, to illuminate the telescope by scattering off a diffuser mounted on the interior of the telescope cover (Kruk et al. 2008). The LEDs are located in a multi-layer annular assembly mounted as a collar around the nose of the star tracker positioned behind the secondary mirror of the telescope. This assembly does not increase the central obscuration of the cassegrain telescope. Feedback stabilization is achieved through brightness monitoring of each LED by an adjacent dedicated photodiode. The OCM will monitor instrument performance during the end-to-end transfer of the NIST calibration of the diode standards to ACCESS in the laboratory. This will provide the necessary transfer in sensitivity to the spectrograph detector to compare against subsequent monitoring observations of the OCM during the various I&T phases. This light source will provide the capability to switch on-off during an observation to check the detector dependence on count rate. The use of the OCM will provide real time and up-to-date knowledge of the ACCESS sensitivity.

6. Status

The ACCESS instrument and calibration hardware are in an active fabrication phase (Kaiser et al. 2010). The telescope and spectrograph flight optics are in-hand with the exception of the prism. The auto-collimating flat mirror has been fabricated, polished, and delivered. The telescope optics and the auto-collimating flat mirror are ready for coating. The housing for the collimator and auto-collimating flat mirror systems are in fabrication. The collimator optics are ready for integration once the collimator assembly housing is delivered. The flight detector has been selected and a subset of performance tests have been presented here. Electronics and calibration equipment are being defined, procured and fabricated. The detector subsystem assembly is being finalized by GSFC and delivery to JHU is expected within the next few months.
Figure 7: Left: The ACCESS telescope and spectrograph (not all components are shown). Right: Assembled ground calibration configuration of ACCESS with the collimated artificial star. From the right, a light source (not shown) feeds a dual-monochromator, which is fiber fed to an order blocking filter and input to an integrating sphere. The output of the integrating sphere is baffled to match the collimator f-ratio. The light then passes through a pinhole at the focus of the collimator. This resulting collimated beam is the calibrated light source for the ACCESS instrument.
Figure 8: Onboard Calibration Monitor used to track sensitivity changes in the field and while parachuting to the ground. The OCM LEDs are mounted in an annular ring about the star tracker behind the secondary mirror and they illuminate a diffuser on the inside of the telescope cover which in turn illuminates the telescope primary with an angular distribution of rays.

7. Summary

ACCESS is an active sub-orbital program with a payload development component followed by four launches over the subsequent two years. This experiment will enable a fundamental calibration of the spectral energy distribution of established bright primary standard stars, as well as standard stars 10 magnitudes fainter, in physical units through a direct comparison with NIST traceable irradiance (detector) standards. Each star will be observed on two separate rocket flights to verify repeatability to <1%, an essential element in establishing standards with 1% precision. The first launch is scheduled for late 2011.

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K. Schahmaneche, for the SNDICE Collaboration
(http://supernovae.in2p3.fr/sndice)

LPNHE, CNRS-IN2P3 and Universités Paris 6 & 7, 4 place Jussieu, F-75252 Paris Cedex 05, France

Abstract.
Cosmological topics such as the study of Dark Energy, require a better photometric precision than what has been achieved until now. An accuracy better than 1% is at least necessary. To go in that direction, we have designed, constructed and installed at the Canada-France-Hawai‘i Telescope (CFHT), a calibrated light source. The goal is to monitor and calibrate MegaCam, a wide field imager used by projects such as the CFHT Legacy Survey and the Supernova Legacy Survey (SNLS). The aim is to obtain an accuracy better than 0.1%.

1. Introduction

The need for a precise photometric calibration has grown over the last few years. Several projects, such as the measurement of the Dark Energy Equation of State using type Ia supernovae, have now reached a point where the precision of the photometric calibration must be better than one percent [Astier06]. Usual astrophysical calibration procedures relying on standard star observations [Regnault09] are limited by the knowledge of the emission spectra of the reference standards (Vega or the HST white dwarf calibrators). More precisely, since the cosmological parameters are measured by comparing fluxes of nearby supernovae (measured in the blue bands of the imagers, around 400 nm) to fluxes of high redshift supernovae (measured in the red, around 800 nm), it is necessary to control the inter-calibration between the imager passbands. Instrumental devices have been developed to study, monitor and model the transmission curves of the astronomical detectors (telescope + camera) with a precision of the order of one per-thousand (Stubbs07, Doi10, Stubbs10). In this paper, we present a similar device called SNDICE (SuperNovae Direct Illumination Calibration Experiment) based on the concept of the direct illumination of the instrument by LEDs, whose emitted light can be controlled with a precision of about 10^{-4}. The opportunity offered by the progress of LED technologies for supporting CCD photometry was underlined in a first paper which described SNDICE-type systems [Juramy08]. The light beams emitted by these LEDs are monitored by photodiodes located along the light path. The device is calibrated on a precision test bench with respect to a NIST photodiode. The SNDICE project was initiated in the Spring 2007 and was installed at CFHT in February 2008 in order to test the concept, and eventually cross-calibrate the pass-bands of the MegaCam wide field imager (Boulade03). We present here first results obtained with this prototype.

The calibrated source is described elsewhere [Juramy08]. Its main features (Figure 1) are listed below:

Light is produced by 24 LEDs whose central wavelength sample the full bandwidth of MegaCam. The bandwidth of each LED is between 7% and 9% of the central value. A source of current with a precision of $10^{-5}$ allows to achieve comfortably a stability of the light beam better than $10^{-4}$.

Due to the “flat-top” design of the selected LEDs, the emitted LED beam is an isotropic beam. Using a series of two consecutive masks, the resulting conical beam, after reflection on CFHT primary mirror, covers the entire field of view of MegaCam.

Redundant control measurements to monitor the stability of the calibrated light source are also performed. The current flowing through the LEDs is measured (sampled at a frequency of 30 kHz). For each LED, the light emitted is controlled at the output of the LEDs source (i.e. 25cm after the LED) by off-axis control photodiodes. Another device, located close to the focal plane, a few centimeters from the MegaCam CCDs, permits to monitor the LED light flux. Since this detector must be as sensitive as the CCDs which equip MegaCam, we designed a Cooled Large Area Photodiode (CLAP), which consists in a two-level micro-cryostat encapsulating a 1 cm$^2$ photodiode and continued by an ultra-low current amplifier ASIC. The whole device is small enough to be placed inside MegaCam.

2. Bench calibration of SNDICE

The LED source was calibrated on a precision bench at LPNHE (Paris). The calibration was transferred from an NIST photodiode to the LED light beams. The calibration was then transferred to the CLAP. This calibration procedure was performed in two steps. A “spectrometric” calibration was obtained using a monochromator to measure the LED spectra at fixed currents (the ones to be used at CFHT for MegaCam calibration). A “photometric” calibration was then performed by mapping each LED light field at different distances from the LED, using a small area calibrated photodiode delivered by Gigaherz-Optik (Figure 2).

During this calibration procedure, we studied the stability of the LED light beams. The main variations were found to be due to thermal effects. A first study of this effect shows that the variation of the LED spectra can be modeled as a function of the temperature with a simple linear law at an accuracy of $2.5 \times 10^{-4}$. A similar first order linear modeling of
3. SNDICE beam quality at CFHT

The SNDICE device (LED light source and the CLAP module) was installed in Hawaii in February 2008. The CLAP module, first located close to the primary mirror, was installed in August 2008 just in front of MegaCam filters.

MegaCam images taken under LED illumination display a high reproducibility. At the pixel scale, for an illumination corresponding to approximately 10000 ADU (for 5 seconds exposure time and an electronic gain around 2), the fluctuations observed between two images taken back to back is slightly larger than the photon noise, and is around 1%. The mean flux measured on a larger scale, namely on 128x128 superpixels, shows a reproducibility down to $10^{-4}$.

For longer time intervals between exposures, around one hour, one can measure a decrease of the flux measured by MegaCam. This effect of a few per thousand is due to thermal variations: MegaCam calibration runs are done during day-time, in the morning and the ambient temperature generally increases during the run. This implies an expected variation of the LED light flux. Introducing a simple linear parametrization of this effect reduces the dispersion of the global LED light flux measured by MegaCam at $1.8 \times 10^{-4}$ (Figure 4).

A first longer time scale analysis, over 4 month in 2010 (4 SNDICE calibration runs from March to July) shows a stability of the response of MegaCam to SNDICE illumination better than 0.5% (Figure 5).

We finish this stability study with a linearity test. By increasing MegaCam exposure times from 1s to 5s, we can check the linearity of the response of the imager in the illumination range [from 3000 ADU / pixel to 13000 ADU / pixel], using the high stability of the LED light source (Figure 6). The residuals to the linear fit are all below 0.01%.
Figure 3: Variation of LED light flux measurements as a function of temperature (on the left, in arbitrary units). A linear parametrization of that temperature dependency allows us to obtain an accuracy of $5 \times 10^{-5}$ (on the right).

Figure 4: a) Mean values in ADU over the MegaCam mosaic (72 amplifiers) for several exposures taken during one hour, the response displays a small drift. b) A linear fit of this drift yields a residual distribution with $1.8 \times 10^{-4}$ RMS attributed to residual temperature variations.
Figure 5: Mean values in ADU for one MegaCam amplifier as a function of MJD for several calibration runs at CFHT (circles). The drift is due to seasonal thermal variations. By taking into account the thermal dependency we are able to reduce the dispersion below $5 \times 10^{-3}$ (black stars).
Figure 6: The linearity of MegaCam response is checked by using the high stability of the LED beam and by increasing the exposure time from 1s to 5s. On the top, mean ADU in one half-CCD as a function of the exposure time in seconds. On the bottom, residual to the linear fit. The residuals are below 1 ADU, i.e.: $1 \times 10^{-4}$.

4. MegaCam monitoring using SNDICE beam

A first study of MegaCam can be performed. The LEDs sample in wavelength the open transmission of the instrument (Figure 7) or any filter (Figure 8).

The knowledge of the LED field map gives access also to the spatial variations of the response functions. The filter transmission as a function of the position on the focal plane is obtained by simply dividing two MegaCam frames: one taken with a filter, another taken without any filter. The result displays the spatial variations of the filter transmission at a given wavelength (Figure 9).

One can also sample the instrument response at different spatial scales to study a given CCD or a given amplifier (Figure 10).

The limitation of this measurement is due to other aspects discussed in this paper: the thermal variation of LED spectra and the internal reflections (see following section).

4.1. Internal reflections

SNDICE is a unique type of light source: a point source at a finite distance. The light beam falling into a given pixel goes through a thin spatial path. This is why a MegaCam frame under SNDICE illumination displays diffraction patterns due to dust or defects located along the light path. On science or twilight exposures, all the diffraction patterns generated by dust or defects are averaged out.

This specificity of the SNDICE beams makes it a unique tool to investigate the light path and, for example, the reflections that are usually averaged. These internal reflections have to be taken into account for precise photometry. Indeed, in usual “flat-field” frames, their effect is to create fake inefficiencies areas of the mosaic. Due to these internal reflec-
Figure 7: MegaCam open transmission (no filter) sampled with all LEDs. This is a relative measurement: a global coefficient was used to adjust the measurement for the "central" LED (\( \lambda = 600 \) nm).

Figure 8: MegaCam \( u^*, g', r', i', z' \) sampling with LEDs. The filters’ transmission curves are the ones delivered by the constructor (REOSC-SAGEM).
Figure 9: Comparison of the spatial variations of $g'$ filter transmission for several LED wavelength: SNDICE measurement (top), simulation based on the constructor (REOSC-SAGEM) data (bottom). The central inefficiency is a well known feature of the MegaCam $g'$ filter. For the highest wavelength it is compensated by a central higher efficiency.

In the process, "flat-field" frames are polluted and in turn affect the uniformity of the photometry of astrophysical sources.

For example, one can easily see a pincushion pattern on some MegaCam/SNDICE frames showing the reflection of the beam on the CCD plane or on the filter, if there is one, then again on the last surface of the Wide Field Corrector located in front of MegaCam and composed of 4 lenses (Figure 11).

To study more precisely these reflections, a central LED channel, initially designed for SNDICE-MegaCam geometrical alignment (called the “planet” channel because, contrary to the other LEDs, its illumination gives only a spot a few thousand pixels wide), is an extremely efficient tool for reflections investigation. The MegaCam/SNDICE frames taken with this specific channel provide unique data to fine tune an optical simulation of the instrument (Figure 12).

4.2. Surfaces studies

The diffraction pattern caused by dust or defects along the light path can be used to study the different surfaces encountered by the SNDICE beam.

The CCDs can be illuminated by the same SNDICE light field but with different relative positions of SNDICE with respect to the telescope (i.e.: different “impact parameters” of the optical axis of the LED light beam w.r.t. the optical axis of the telescope). Using this possibility, one can “track” the same diffraction pattern in different illumination positions, and deduce from this the optical surface where the defects or dust grains are located (Figures 13). This has been done by showing the accumulation of dust/defects on the different optical surfaces of the Wide Field Corrector or on the primary mirror (Figure 14).
Figure 10: Mean effective gain (electronics gain times CCD QE at $\lambda_{LED}$) for 72 amplifiers measured using SNDICE LEDs (arbitrary units).

Figure 11: Evidence for Fresnel reflection on a MegaCam image (whole mosaic) obtained under LED beam (on the left) and its simulation using ray tracing (on the right).
Figure 12: Spot generated by the specific alignment "planet" channel (on the left with a linear grey-scale). Multiple reflections of this spot are visible on a log colored-scale (on the right). This special channel is unique to ray trace testing.

Figure 13: Surface defects on the mirror or any lens of the Wide Field Corrector, generate diffraction patterns on the focal plane. The location of these patterns varies with the geometrical relative position of SNDICE w.r.t. telescope axis. By changing this position one can locate the defect in the optical path.
Figure 14: The distribution of the "displacement" (measured in superpixels = 16x16 pixels) of the diffraction pattern on the frame shows peaks corresponding to the different surfaces where the defects are located. The main peak corresponds to defect/dust accumulated on the primary mirror. The two other peaks correspond to shorter "displacement" and so shorter distance of the diffraction source w.r.t. the focal plane, i.e., correspond to both of the extreme optical surfaces of the Wide Field Corrector located just in front of MegaCam.

5. Actual limitations of this first prototype

The thermal variations measured on the calibration bench at LPNHE, span a temperature range of a few degrees (from 16°C to 24°C). These measurements showed that it not possible to extrapolate our calibration measurements (spectra and light mapping) to the CFHT dome temperatures (close to 0°C) with a sufficient precision. Instead of extrapolating, it is necessary to interpolate and to do so, to calibrate the LEDs light in the temperature range of their utilization in situ. This is under development: we are modifying the calibration bench to be able to cool it down to 0°C.

Another limitation is due to the internal reflections showed previously. Using SNDICE specific alignment channel (the “planet”) we will be able to check a fine tuned model of the telescope optics by taking into account these reflections and modeling the SNDICE beam.

At the moment, SNDICE is not yet used as an absolute calibrated light source. Once the two developments mentioned above will be achieved, we will have an absolutely calibrated light source in hand.

6. Conclusion

We have shown that SNDICE can be calibrated with a precision better than $10^{-4}$ in the lab. In the MegaCam environment, we could measure a short term reproducibility of the MegaCam response at $10^{-4}$. This gives access to a precise relative calibration and monitoring of MegaCam at CFHT.

Improvements of our calibration bench (to parametrize precisely flux and spectral variations due to thermal variation) and improvements in the SNDICE beam simulation (to
take into account reflections on the different MegaCam surfaces) will lead to an absolute calibration of MegaCam on CFHT.

In addition to the calibration of MegaCam at CFHT, we are currently designing a second generation device for the wide field imager SkyMapper (Siding Spring Observatory, NSW, Australia).

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WFPC2 Filters after 16 Years on Orbit

P. L. Lim¹, M. Quijada², S. M. Baggett¹, J. Biretta¹, J. MacKenty¹, R. Boucarut², S. Rice², and J. del Hoyo²,³

Abstract.
Wide Field Planetary Camera 2 (WFPC2) was installed on Hubble Space Telescope (HST) in December 1993 during Servicing Mission 1 by the crew of Shuttle Mission STS-61. WFPC2 replaced Wide Field Planetary Camera 1 (WFPC1), providing improved UV performance, more advanced detectors, better contamination control, and its own corrective optics. After 16 years of exceptional service, WFPC2 was retired in May 2009 during Servicing Mission 4 (SM4), when it was removed from HST in order to allow for the installation of Wide Field Camera 3 (WFC3).

WFPC2 was carried back to Earth in the shuttle bay by the crew of Shuttle Mission STS-125. In a joint investigation by Goddard Space Flight Center (GSFC) and Space Telescope Science Institute (STScI), the Selectable Optical Filter Assembly (SOF A) of WFPC2 was extracted and the filter wheels removed and examined for any on-orbit changes. The filters were inspected, photographed and scanned with a spectrophotometer at GSFC. The data have been analyzed at STScI with a view towards understanding how prolonged exposure to the HST space environment affected the filters and what the resultant impacts are to WFPC2 calibrations.

In this paper, we present our early results from these post-SM4 laboratory studies, including comparisons of pre- to post-mission filter transmission measurements for F343N, F160W, F160BW, F450W, and F170W.

1. Introduction

In the 16 years of WFPC2 mission, over 180,000 science exposures were taken, all of which are now public and available from STScI Archive¹. WFPC2 was brought back safely to the ground during SM4. We take this opportunity to study the filters that have been exposed to the space environment for nearly two decades and use the results to improve science calibrations.

The main goal is to look for changes in filter transmissions or bandpass wavelengths, if any. Due to resource limitations, we are unable to scan all 48 filters. However, we have selected the filters of interest by prioritizing them according to usage, wavelength coverage, etc. The identified elements are tabulated in Table 1. The filters chosen for post-mission scanning are those with known on-orbit changes, high and regular utilizations, UV filters for which we wish to check the red leak, and others that might be interesting to examine for various reasons, such as unique technology (Wood’s filters) or anomaly seen during inspection. If there is a need from the scientific community to scan an unlisted filter, a

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¹Space Telescope Science Institute, Baltimore, MD
²Goddard Space Flight Center, Greenbelt, MD
³University of Arizona, Tucson, AZ
¹http://archive.stsci.edu/hst/search.php

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General Observer could contact Lim et al. We will give the request full consideration in the presence of strong scientific justifications and resources.

<table>
<thead>
<tr>
<th>Category</th>
<th>Filters</th>
</tr>
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<tbody>
<tr>
<td>Known on-orbit changes(^a)</td>
<td>F122M, F160BW, F343N</td>
</tr>
<tr>
<td>UV filters(^b)</td>
<td>F170W, F185W, F218W</td>
</tr>
</tbody>
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\(^a\)See Section 1..
\(^b\)Red leak only.
\(^c\)Included based on lab inspection, etc.

Through on-orbit monitoring programs, there are several filters with known changes. For instance, Biretta & Gonzaga (2008, private communications) reported a $\sim$20-25% drop in transmission for F122M. F160BW was found to have exponential pinhole growth (Biretta & Verner, 2009). Last but not least, the transmission of F343N declined linearly throughout the mission, as much as 50% lower in 2009 compared to 1997 (Gonzaga & Biretta, 2009). Despite these cases, most of the filters are expected to have changed less than a few percent.

We discuss our analysis methods in Section 2.. Early results for F343N, F160AW, F160BW, F450W, and F170W are presented in Sections 3.-6.. Lastly, we summarize our work in Section 7..

2. Methods

All 12 wheels were removed from the SOFA for inspection and scanning at GSFC. Follow-up analysis is primarily done in STScI. Cosmetic inspections to look for surface features (e.g., stains, spots, and pinholes) were completed in July 2010 by J. MacKenty (main inspector), S. Baggett, J. Biretta, P. L. Lim, and M. Quijada. Such features might or might not affect the filter transmissions.

Scanning of filter transmissions by M. Quijada and J. del Hoyo are on-going. We use a spectrophotometer with a double-beam monochromator. This differential beam measurement enables us to measure on-orbit filter performance in a non-vacuum lab environment. Scanning is performed at room temperature ($\sim 25^\circ$C), which is slightly warmer than on-orbit ($\sim 12-14^\circ$C), but we do not expect this to have significant effect on our results.

Due to UV absorption by atmospheric O$_2$, we are unable to obtain good signal-to-noise for wavelengths less than $\sim$2000Å. The scanner wavelength resolution ranges from $\sim$0.5-50Å. We find it sufficient to use higher resolution for in-band transmissions and lower for out-of-band. The lower limit of measurable transmission is $\sim 10^{-6}$-10$^{-8}$.

To study the spatial homogeneity of the transmissions, we scan a filter at five positions, as illustrated in Figure 1. For square filters (most of them), the positions are top left, top right, center, bottom left, and bottom right. For round filters (F160AW and F160BW), they are top, left, center, right, and bottom. These correspond to the locations on a filter when the wheel (Wheel 1 is also shown in the figure) is rotated on its rim with the smooth side (opposite of the clamp side that is shown) facing the incoming scanner beam and the filter is at the bottom. We note that for Wheels 1-11, it was the clamp side that faced the on-orbit incoming beam in HST, but this does not affect our scanned results.
WFPC2 Filters after 16 Years on Orbit

Figure 1: (a) A photograph of Wheel 1 clamp side in the lab. Clockwise from CLEAR (at 1 o’clock position): F953N, F160AW, F160BW, and F122M. (b) Five-point scan locations for a square filter. Top left (blue diamond), top right (green square), center (black circle), bottom left (yellow square), and bottom right (red triangle) correspond to the orientation where the wheel is standing on its rim with the smooth side (opposite of clamp side) facing the incoming scanner beam and the filter is at the bottom of the wheel. (c) Five point scan locations for a round filter (e.g., F160BW), with top (green square), left (yellow square), center (black circle), right (red triangle), and bottom (blue diamond). Orientation is the same as (b).

Once we have the scanned transmission curves (hereafter, post-flight), we compare them with pre- and in-flight data. Pre-flight is defined as the first version of the transmission curve taken in the lab prior to the WFPC2 mission. In-flight is the latest version used during the WFPC2 mission, mostly adjusted by Baggett et al. (1997) based on photometric calibrations. Both pre- and in-flight transmission scans are publicly available from STScI2. For a homogeneous filter, we will only present the averaged post-flight transmission from all the five points in this paper.

3. F343N

Figure 2 shows our results for F343N, a narrowband filter for [NeV]. The pre-flight transmission scan was rescaled by Baggett et al. (1997) to the in-flight version in order to bring the on-orbit spectrophotometric standard star observations into agreement with predictions and provide observers with the proper photometric calibration (i.e., conversion from counts to flux) for each image. Post-flight scanning confirms a further transmission decline, as well as a bandpass shape change, a peak wavelength redshift by $\sim$2A, and full-width at half-max (FWHM) broadening by $\sim$3A. These changes are effectively identical at all five points scanned.

The overall transmission decline is in agreement with a later analysis of standard star observations with WFPC2 by Gonzaga & Biretta (2009). Figure 3 compares observed PC1 and WF3 fluxes to in- and post-flight filter transmissions. Note that the observed fluxes are, by necessity, compared to predictions based upon a convolution of the throughput information for all elements along the light path such as the HST mirror, WFPC2 optics, filter

transmission, and detector quantum efficiency. As expected, post-flight filter transmission data agrees well with the late WFPC2 standard star observations. Hence, in order to accurately predict observed fluxes and obtain the proper photometric calibration throughout WFPC2’s lifetime, an interpolation between in- and post-flight filter transmission curves will be needed.

4. **F160AW & F160BW**

During the actual talk given in this workshop, we assumed that the slot positions of F160AW and F160BW (Wood’s filters) on Wheel 1 were correctly tabulated in the WFPC2 Instrument Handbook (McMaster et al., 2008, Table 3.1). However, we later discovered that the positions are reversed based on actual telescope commanding data (Welty 2010, private communication), i.e., F160AW resides in Slot 2 and F160BW in 3. This has been corrected here and therefore, might not reflect the information given during the talk.

F160AW was never used in-flight due to known pinholes. Figure 4 shows pre- and post-flight transmissions. It is apparent that the pinholes contribute to significant red leak that varies from $10^{-3}$-$10^{-5}$ across the filter. Unfortunately, due to the lack of on-orbit observations in this filter, we are unable to study its pinhole characteristics over time. However, this does not affect any science calibration.

Figure 5 shows in- (i.e., pre-) and post-flight filter transmissions for F160BW. The red leak shown ($10^{-6}$) represents an upper limit due to spectrophotometer limitations (see Section 2.). Unlike F160AW, no inhomogeneity is found. It appears that the pinhole growth reported in Biretta & Verner (2009) does not significantly increase the red leak. This is consistent with the Lim et al. (2009, Section 5.6.2) analysis that found negligible red leak in F160BW cross-filter observations as recently as May 2009.
Figure 3: Similar to Gonzaga & Biretta (2009, Figure 1) for F343N but with post-flight filter transmission comparison also added. The figure shows the ratio of observed to predicted fluxes over time. Standard star observations using PC1 and WF3 are shown as filled circles and crosses, respectively. Blue and red data points compare the observations to the in- and post-flight filter transmission curves, respectively. Ratio of 1 indicates that observation agrees with prediction.

Figure 4: Filter transmissions across wavelength for F160AW in log-linear scale. Blue line is the pre-flight lab filter scan. As F160AW was never used on-orbit, there is no in-flight version. Green, red, cyan, magenta, and yellow lines are post-flight data for the five points, as given in the legend. Post-flight red leak is inhomogeneous and ranges from $\sim 10^{-3}$ (“Top”) to $\sim 10^{-5}$ (“Right”).
Figure 5: Filter transmissions across wavelength for F160BW in log-linear scale. Blue and green lines are in- (same as pre-) and post-flight measurements, respectively. The $\sim 10^{-6}$ scanned red leak is at the instrument limit and thus, only reflects the upper limit of actual values.

5. F450W

Figure 6 shows in-band and red leak transmissions in linear and log-linear scales, respectively, for F450W, a Wide B filter. The post-flight in-band measurement agrees with pre- and in-flight to within a few percent.

As for the red leak, pre-flight data had transmission set to $10^{-5}$ (mixed in with some negative values) for wavelengths longer than $\sim 7500\,\text{Å}$. This region was re-adjusted in the in-flight version. Post-flight filter scan data have considerable scatter and negative values due to instrument settings, hence, they only reflect the transmission upper limit for wavelengths greater than 7500Å. Otherwise, they are also consistent with pre- and in-flight data. Remeasurement in this region with better sensitivity will improve the results.

6. F170W

Figure 7 shows pre-, in-, and post-flight red leak transmissions for F170W, a broadband UV filter. After the talk for this workshop, we rescanned this filter with improved sensitivity. The pre-, in-, and post-flight data are all in excellent agreement. The post-flight data shown in this paper supersedes the version shown during the talk.

7. Summary

Given the 16 years of significant and unprecedented WFPC2 contributions to the scientific community, it is imperative that we study its filters flown back from orbit during SM4 in order to improve science calibrations.

We inspected the filter surfaces and scanned them at five points to assess inhomogeneity. F343N shows $\sim 50\%$ transmission decline, which is consistent with on-orbit standard star observations, as well as an overall bandpass shape change. Significant inhomogeneity and
Figure 6: Filter transmissions across wavelength for F450W in-band (top) and red leak (bottom, log-linear scale). Blue, green, and red lines are pre-, in-, and post-flight measurements, respectively. Post-flight red leak only reflects the upper limit beyond $\sim$7500Å. Gaps in pre- and post-flight red leaks are negative values.
red leak increase are seen in F160AW, but this does not impact WFPC2 science because the filter was never used. We are only able to obtain the upper limit of F160BW red leak, which is consistent with non-detection in on-orbit calibrations despite reported pinhole growth. F450W and F170W agree very well with existing data, although the post-flight red leak of the former could be improved using better instrument sensitivity.

We will scan the rest of the filters from Table 1 and post the post-flight transmissions on the WFPC2 website tentatively in October 2010. In addition, our results will be published in a WFPC2 Instrument Science Report.

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The Legacy of the Hubble Space Telescope Spectrographs

Anton Koekemoer

Space Telescope Science Institute

Abstract.
The legacy recalibration and reprocessing of the entire archive of NICMOS data, consisting of more than 100,000 observations obtained since its installation on HST in 1997, has led to the production of a final archive that includes all the calibration and processing improvements that have been obtained as a result of our improved knowledge of the instrument performance during this time. This archive is unique in being the only set of HST observations in the 1 - 2 micron regime for over a decade until the installation of WFC3 in 2009, and provides the foundation for a wide range of follow-up science. The infrared detectors are susceptible to a wide range of calibration issues that are specific to infrared detectors, including persistence, temperature and bias-dependent effects, and an extensive program of calibration and software development has been carried out to correct these. These have been applied in a full recalibration and reprocessing of the entire archive of NICMOS data, including also MultiDrizzle combination and release in the HLA, resulting in a dataset with dramatically improved scientific value that will serve as a unique and extensive archival legacy in its own right, in addition to benefiting future infrared instruments and science programs.
Update on NICMOS Reference Files

Tomas Dahlen and the NICMOS Team

Space Telescope Science Institute, Baltimore, MD 21218

Abstract.

The response of the three NICMOS detectors when observing a standard star showed variations of 2-4% during both the era before the installation of the NICMOS Cooling System, NCS (1997-98) and during the era of NCS operation 2002-08. If uncorrected, these variations could affect the quality of reduced NICMOS images, introducing errors in the photometry. Here we discuss steps taken to mitigate these effects, including the creation of dynamic reference files that takes into account the variations in detector response. This includes the photometric calibration tables and the flat-field and dark reference files. Furthermore, updates have been made to the calnica calibration software to accommodate these new reference files. Using these new reference files will in some cases improve the photometric accuracy of NICMOS images by ~2-3%.

1. Introduction

Two sets of reference files are used when calibrating NICMOS data. One set for data taken during the operation of the NICMOS cryocooler 1997-98, when the operating temperature was ~61K. The other set is used for the era after the installation of the NICMOS Cooling System (NCS) 2002-2008 when the operating temperature was ~77K. The different detector temperatures during these eras is the reason for the need for two sets of reference files. However, within both there eras, the original reference files were static. By studying the photometry of standard stars during both these periods, slight but significant changes in the response in e-/s for a constant source were detected. For the 1997-98 period, the response increased by a few percent, while for the 2002-08 period, the response showed a decrease by a few percent. If uncorrected, these changes may introduce errors in the measured photometry of NICMOS images by ~2-3%.

In this write-up, we quantify the changes in the detector response and the consequences these have on NICMOS calibrations. We present a new set of NICMOS reference files, i.e., photometric reference tables, flat-field images and dark current images, that are corrected for the varying detector response when NICMOS images are calibrated using the calnica software.

2. Response of NICMOS Detectors

In Figure 1, we show the relative response of the three NICMOS cameras as a function of time for the pre-NCS 1997-98 era (left panels) and for the post-NCS era after 2002 (right panels). Different colored lines show the response for different filters. Black arrows symbolize the overall trend. There are significant deviations from unity in both eras.

The reason for the increase in response in the 1997-98 era is the increase in detector temperature as the solid nitrogen evaporated. The cause behind the decrease in detector response the post-NCS era is most likely a gradual decrease in the bias voltage with time.
3. Creating Dynamic Reference Files

With a response that changes with time for the NICMOS detectors, accurate calibrations of NICMOS images will be difficult. In particular, photometric calibration will be affected, but also the flat-fields, dark current and possibly also the number of hot/bad pixels in the NICMOS data quality masks will be affected by a non-constant response of the detectors. In an ideal case, we would like to apply calibration files that are valid at the particular detector response when the images were taken. Fortunately, there is a way to estimate the response of individual NICMOS images. The bias level of the zeroth read of NICMOS MULTIACCUM images is correlated with the response of the detectors. For each image, the NICMOS pipeline derives a measurement of the bias level which is quantified as an “effective temperature” called biastemp (which is assigned to the TFBTEMP keyword in the image header of NICMOS images). See Pirzkal et al. 2009 for more information on the CalTempFromBias task that is used in the pipeline to calculate the biastemp. Biastemp is not a physical temperature but depends both on detector temperature and the detector bias voltage. In Figure 2, we show the evolution of biastemp with time (green symbols) together with the actual detector temperature as measured by the mounting cup sensor (red symbols). For the pre-NCS era, there is a qualitative agreement between biastemp and the detector mounting cup temperature, implying that the change in biastemp is due to a change in the true detector temperature. For the post-NCS era, the detector mounting cup temperature is stable, while the biastemp shows a significant decline with time. This suggests that the change in biastemp is not due to a change in the physical detector temperature, but is instead caused by a decrease in the detector bias voltage.

Comparing Figures 1 and 2 indicates that the change in detector response is correlated with the change in biastemp. It is therefore possible to get an estimate of the response of individual NICMOS images at the time of exposure by measuring the biastemp. In order to include biastemp dependent calibrations of NICMOS images, we have created a set of calibration files that depends on biastemp. We have also modified the NICMOS calibration
software, calnica, to allow calibrations that depends on biastemp (see Dahlen et al. 2009). Calnica is used in the OTFR pipeline at STScI when images are requested from the archive, but is also available in the NICMOS package of STSDAS for users who want to make their own customized calibrations of NICMOS data.

3.1. Photometric calibration

A main concern with a varying detector response is the accuracy of the measured photometry in NICMOS exposures. Figure 1 shows a change in response of 2-4% in the 1997-98 era, and 2-3% in the post-NCS era after 2002. If uncorrected, this translates to an uncertainty in the derived NICMOS photometry by \( \pm 2\% \).

The PHOTCALC step in calnica reads the photometric keywords from the PHOT-TAB reference file and populates the image header of the \(*_{\text{raw}.fits}\) images with e.g., photometric zero-points. There are basically two ways to correct for the change in the response with biastemp when running calnica. Either the photometric zero-point can be made biastemp dependent, or the image could be scaled to account for the biastemp dependence while leaving the zero-point unchanged. We have used the latter approach, because using zero-points that vary with temperature will result in different zero-points for each individual NICMOS exposure. Such a scenario would make it difficult to combine separate images into a common image or mosaic. Instead, a scaling factor is calculated that corrects the image counts (or count rate) for the temperature dependence of the response and puts all images on the same zero-point (for each camera and filter combination).

To derive the scaling factor, we have used available photometric standard star observations to derive the relative response as a function of biastemp for individual NICMOS images during both the pre-NCS and post-NCS eras. The relative response is thereafter fitted to a straight line to derive the scaling factor, ZSCALE:

\[
ZSCALE = PHOTF.C1 \times (TFBTEMP - PHOTREFT) + PHOTF.C0
\]

here PHOTF.C1 is the slope of the relation and PHOTF.C0 is the offset, i.e., the two parameters that are fitted. TFBTEMP is the measured “effective temperature”, while
PHOTREF T is the reference temperature at which the tabulated zero-point in the reference file is valid (i.e., for which the scaling is unity).

After the scaling factor is calculated, the science extension of the NICMOS image is multiplied by this factor in the pipeline calibrations. The scaling factor is also assigned to the ZPSCALE keyword in the primary header. A scaling factor is only calculated when the TFBTEMP keyword is populated in the primary header of the *.raw.fits file. If this is not the case, the ZPSCALE keyword is set to unity. Furthermore, if biastemp is outside a valid range then the scaling is also set to unity. For the three polarizers in NIC1 and NIC2, as well as for the three grisms in NIC3, the scaling is always set to unity. Besides populating the ZPSCALE keyword, PHOTCALC step in calnica also sets the keyword PHOTFERR. This gives an estimate of the error in the zero-point due to the uncertainty in the temperature dependence of the photometry. The size of this uncertainty is typically 0.015 mag.

The new PHOTTAB reference tables have been included in the OTFR since June 4, 2009, for 1997-98 data, and since June 2, 2009, for data taken 2002-08.

3.2. Flat-field Corrections

In the FLATCORR step in calnica, the *.raw.fits images are corrected for pixel-to-pixel variations in the detector sensitivity by multiplying by a flat-field image. The flat-field image is, by construction, normalized to unity and inverted (therefore, the multiplication and not the usual division). With the change in the response due to the variations in biastemp, the flat-fields also change. These changes are, however, of second order, i.e. any uniform change over whole the detector will not affect the flat-field, since such a change will disappear when the flat-fields are normalized to unity. A uniform change over the whole detector due to the change in biastemp is instead corrected for by the scaling performed in the PHOTCALC step in calnica described in the previous subsection. There are, however, second order effects of the changes in the response that do affect the flat-fields. I.e., because different pixels have a relative response that changes with temperature or bias voltage, the structure of the flat-field changes with biastemp, even though the overall normalization is unity. To account for this change in structure, a set of five different flat-fields has been created for each filter in both the pre-NCS and post-NCS eras. Each of these five flat-fields is constructed using data from a particular biastemp range, and these five ranges cover the observed biastemp span for each era. In the FLATCORR calibration step, calnica reads the biastemp from the TFBTEMP keyword and then selects the flat-field that closest matches the biastemp when it performs the flat-field correction.

Figure 3 illustrates how the flat-field structure depends on biastemp. The figure shows the flat-field structure at five different biastemp values normalized to a flat-field constructed using the median biastemp. Since the latter is close to the value for the third of the five flat-fields, this flat is close to unity over the whole field. If the change in biastemp caused a uniform change in the response in all pixels over the whole detectors, then all five normalized flat-fields shown in the figure would be unity. If uncorrected, the change in the shape of the flat-fields due to the change in biastemp may introduce photometric errors of up to ~1%.

The multiple extension biastemp dependent flat-fields images are used on NICMOS images retrieved from the HST Archive after November 19, 2008 for post-NCS data and after January 23, 2009 for the pre-NCS data. Details on how these flat-fields were constructed and how they are implemented in calnica are given in the contribution by Thatte & Dahlen in these proceedings and Thatte & Dahlen 2009.

3.3. Dark Corrections

Dark current corrections are performed in the DARKCORR step when the calnica task is run. The header for the *.raw.fits NICMOS images contains keyword information on two different types of dark reference files that can be used in the DARKCORR step. The
DARKFILE keyword contains a static dark reference file with extension *drk.fits. Each NICMOS sampling-sequence has its own DARKFILE and when NICMOS data is retrieved from the archive, this keyword is populated with the reference file that is appropriate for the combination of sampling sequence (given in keyword SAMP_SEQ) and the number of reads used in that particular sampling sequence (NSAMP keyword). Most *drk.fits darks were made in 1997 and are based on synthetic models. The TEMPFILE keyword points to a dynamic dark reference file (extension *tdd.fits) that is constructed so that it can be made dependent on both biastemp and the temperature measured by the mounting cup sensor. There is one pre-NCS and one post-NCS TEMPFILE for each camera. By default, calnica uses the file given by the TEMPFILE keyword for the dark current subtraction. Only in the case where the TEMPFILE keyword is missing or has a “N/A” value, will the DARKFILE be used. All files retrieved from the archive after April 9, 2002 have the TEMPFILE keyword populated. However, an actual temperature dependence was initially only implemented for data taken in the pre-NCS era. The biastemp/temperature dependence of the post-NCS *tdd.fits files was derived and implemented in 2009. For old data lacking the TEMPFILE keyword, we recommend retrieving the data again from the OTFR to get the most up-to-date reference files and calibrated files.

The TEMPFILE contains information on the three separate components of the NICMOS dark current: the linear dark, the amplifier glow, and the shading component. Below, we describe how calnica uses these separate components in the biastemp dependent calibrations.

**Linear Dark**
The linear dark current component is found in the first extension (EXT=1) of the *tdd.fits
file, i.e. the file given by the TEMPFILE keyword. Figure 4 shows the linear dark extension for NIC3 in the post-NCS era. There are both pixel-to-pixel variations, creating the “salty” pattern, as well as large-scale gradients. The linear dark extension also includes the coefficients $C_{0,\text{LIN}}$ and $C_{1,\text{LIN}}$, that can be used to calculate a scaling factor, $\text{LINSCALE}$, that corrects the linear dark current for its dependence on biastemp. The logarithm of the scaling factor is given by:

$$\log(\text{LINSCALE}) = C_{1,\text{LIN}}/\text{TFBTEMP} + C_{0,\text{LIN}}.$$  \hspace{1cm} (2)

We use this parametric form, because, at least in principle, the dark current should be proportional to the voltage over each pixel (which can be seen as a diode) for which $V \propto e^{(1/T)}$. In the DARKCORR step, calnica multiplies the linear dark component with $\text{LINSCALE}$ before subtracting that component.

Even though the dark current should be dependent on both detector temperature and bias voltage, investigations of the biastemp dependence of the linear dark current of the three NICMOS cameras have shown that the any detectable trend in the linear dark with temperature is smaller than the overall scatter in the dark current. Figure 5 shows the measured dark current as a function of biastemp for the three NICMOS cameras in the post-NCS era. Fitting the data points to a straight line shows no significant dependence of the dark current on the temperature (the same hold if fitted to the parametric form in Equation 2). Error-bars in the figure represent the rms scatter for each data point.

Therefore, at this point, no temperature dependence of the linear dark component is implemented. The coefficients are set to $C_{0,\text{LIN}}=0.0$ and $C_{1,\text{LIN}}=0.0$. To date no temperature dependence of the linear dark is implemented; but calnica is now prepared to include this should future investigations deem it necessary.

**Amplifier glow**

The amplifier glow is a signal that is added to NICMOS images every time the cameras are read out. The signal is most likely due to radiation from the amplifiers situated in each of
the four corners of the three NICMOS cameras. The amplifier glow component (normalized to one read-out) is found in the second extension (EXT=2) of the *tdd.fits file given by the TEMPFILE keyword. Figure 6 shows the NIC3 amplifier glow image.

In the DARKCORR step one amplifier glow image is subtracted for each read-out of the science image. Investigation of how the amplifier glow varies with biastemp has found a ∼2% change in the signal in the post-NCS era. This change maybe significant enough to affect the quality of the calibrated files, leaving unwanted residuals. Figure 7 shows the counts of the amplifier glow signal in the corners of the NIC1 camera as a function of biastemp for about 100 individual images. The counts are normalized to the median of all images.

Due to the observed change in amplifier glow signal, a biastemp dependent scaling of the amplifier glow component has been implemented in calnica. The scaling factor is calculated using a straight line fit to the count vs. biastemp relation, as indicated by the black line in Figure 7 for NIC1. The fit determines the coefficients \(C_0\_AMP\) and \(C_1\_AMP\) that gives the scaling factor according to:

\[
AMPSCALE = C_1\_AMP \times (TFBTEMP - REFTEMP) + C_0\_AMP.
\]  

where TFBTEMP is the biastemp and REFTEMP is the reference temperature for which the median ampglow image is created (indicated by blue dot in the Figure). The coefficients are read by calnica from the header of the ampglow extension (EXT=2) of the dark reference file. Before the amplifier glow signal is subtracted during the DARKCORR step in calnica, the amplifier glow image is multiplied by AMPSCALE to take into account the dependence on biastemp.

There only exists sufficient calibration data to derive the \(C_0\_AMP\) and \(C_1\_AMP\) coefficients for the data taken after the installation of the NCS, 2002-08. For the pre-NCS era,
Figure 6: Amplifier glow image for NIC3 in the post-NCS era. The elevated signal in the corners from the heat of the amplifiers is clearly visible. Individual bright spots indicate hot pixels.

Figure 7: Relative signal of individual ampglow images as a function of biastemp. Black line shows a straight line fit to data, while the blue point shows the median of all data.
the coefficients are set so that the scaling is always unity, AMPSCALE=1.

**Shading profile**
The shading is a noiseless signal that shows up as a gradient across each of the four detector quadrants in the three NICMOS cameras. The shading is caused by a gradual change of the bias level in the pixels with time as the quadrant is read out. Figure 8 shows an example of the shading for a NIC3 image.

The shading varies between read-outs and depends on the time since the last read-out, which is called the DELTATIME. In total, there are twelve different DELTATIMEs used by the MULTIACCUM sampling mode. The twelve different shading profiles associated with each of the DELTATIMEs are described in EXT=3 to EXT=26 in the *tdd.fits file given by the TEMPFILE keyword. Each DELTATIME is associated with two of the extensions. One is an image of a static shading profile valid for that DELTATIME, and the other is a binary table including coefficients needed to calculate the temperature dependence of the shading correction. For more information on these files, see Jedrzejewski (2002).

In the DARKCORR step of calnica, the temperature of the detector is read and then the shading profiles for each of the DELTATIMEs used in the exposure is scaled using the coefficients in the *tdd.fits file before the shading component of the dark is subtracted.

Investigations of the pre-NCS data showed a strong correlation between the shading signal and the mounting cup temperature (Monroe & Bergeron 1999). Figure 9 shows an example of the NIC2 shading profile for a 64s DELTATIME readout at two different temperatures. In this case, the difference in shading signal is up to 10 counts per Kelvin.

The temperature dependence of the post-NCS data has been investigated using both biastemp and the mounting cup temperature. No significant trend in the shading profile with biastemp has been detected. If the shading was dependent on the biastemp in a similar way to the mounting cup results from the pre-NCS era, such dependence would have been easily detected because the change in biastemp is equivalent to a change in temperature by more than 2K in the post-NCS era. The shading profile may not depend on biastemp in a way that is similar to the amplifier glow and photometry.

Therefore, the mounting cup temperature is used to derive the coefficients for the temperature dependence of both the pre-NCS and post-NCS shading signal. However, the
mounting cup temperature was very stable in the pre-NCS era (see right panel of Figure 2) and it has not been possible to derive trend in shading profile vs temperature for this era. Therefore, the coefficients in the dark reference files for the post-NCS era are set so that no temperature dependent scaling of the shading profile is performed when calibrating post-NCS data.

3.4. Data Quality Masks

The first data quality masks for the post-NCS era were produced in 2002 using the data from the first NICMOS calibration programs. Since then, calibration programs run annually during the whole period 2002-08. With the change in detector response, as well as aging of the detectors, it is likely that there have been changes to the number of pixels that should be flagged in the DQ masks. Therefore, we have used available calibration data to re-derive the NICMOS data quality masks for the post-NCS era. In particular, we have looked at the number of pixels flagged as bad, i.e., hot, cold, or dead pixels and the number of pixels affected by grot. The latter are pixels with lower sensitivity most likely due to small paint flecks on the detector that obscures the light.

In summary, the new masks have a slightly higher number of bad pixels compared to the 2002 masks, while the number of grot affected pixels is somewhat smaller. There is only a weak temporal variation in the number of flagged pixels, therefore only a single new DQ map has been created for each camera for the post-NCS era. These new masks have been used in the OTFR pipeline processing starting April 7, 2009. For more details on the creation of these data quality masks, see the contribution by Barker & Dahlen in these proceedings and in Barker & Dahlen 2009.

4. Conclusions

New NICMOS reference files take into account the variation in the detector response observed in both the pre-NCS era (1997-98) and the post-NCS era (2002-08). These variations

Figure 9: Shading profiles at two different mounting cup temperatures for a 64s DELTA-TIME readout in NIC2. Figure taken from Monroe & Bergeron 1999.
are caused by a change in the detector temperature (pre-NCS) and by a drift in the bias voltage (post-NCS). The quantity “biastemp” is introduced as a measurement of bias level of the zeroth read of NICMOS images. The biastemp is correlated with the detector response and can, therefore, be used to determine the response for individual images.

The new reference files that takes into account the variations in response include:

- photometric calibration tables that corrects the photometry for the variation in detector response
- flat-field images for both eras that consists of multiple flat-fields extensions, each valid for a different range in biastemp
- dark reference files that depend on biastemp

Also, to accommodate these new reference files, a number of updates in the calnica calibration software have been made. Altogether, these new reference files improve the photometry of calibrated NICMOS images by up to $\sim$2-3%. We also present new data quality masks for the post-NCS era.

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NICMOS Coronagraphy: Recalibration and the NICMOS Legacy Archive PSF Library

Glenn Schneider  
*Steward Observatory, The University of Arizona, Tucson, AZ 85721*

Murray D. Silverstone  
*Eureka Scientific, Inc., Oakland, CA 94602*

Elizabeth Stobie  
*National Optical Astronomy Observatories, Tucson, AZ 85719*

Joseph H. Rhee  
*University of California, Los Angeles, CA 90095*

Dean C. Hines  
*Space Telescope Science Institute, Baltimore, MD 21218*

**Abstract.**  
NICMOS coronagraphy, with well-matched template Point Spread Function (PSF) subtraction, probes the closest environments of occulted targets with the highest imaging sensitivity in intrinsically high contrast fields at the smallest radial distances uniquely afforded by HST. NICMOS PSF-subtracted coronagraphy has been invoked in a wide variety of HST programs with science themes as divergent as detecting and characterizing disks of circumstellar material in neo-natal stellar environments, to studying faint nebulosity associated with luminous active galaxies, to searching for planetary-mass companions in extrasolar planetary systems recently born and in the “stellar graveyard.” The investment in HST time in the execution of these and other NICMOS coronagraphic programs has met with mixed returns. Stunning scientific advances contrast with more common failed analyses due to the lack of suitable template PSFs that are required to produce high-fidelity, photometrically robust, high contrast coronagraphic images. To remedy this situation, we are undertaking a rigorous, homogeneous, and complete recalibration and analysis of the full archival set of raw NICMOS coronagraphic images (through HST Cycle 15) to create a Legacy library of template PSFs enabling the recovery of the large body of science otherwise lost. This PSF library and enhanced data recalibration processes, along with our generically applicable analysis software, will: (1) critically augment the needs of observational programs reliant on high fidelity PSF subtractions, (2) increase survey yields and improve photometric efficacy, (3) reduce HST orbit allocations that are otherwise required for near-contemporaneous reference PSF observations, and (4) greatly enrich the yet-unrealized potential of the many NICMOS coronagraphic observations already acquired from the broad spectrum of science programs previously executed.
1. Introduction

In comparison to the other HST coronagraphic instruments, NICMOS coronagraphy with “well matched” Point Spread Function (PSF) subtraction simultaneously provides: (a) the highest imaging contrast, (b) the smallest Inner Working Angle (IWA), (c) full circumferential (azimuthal) spatial coverage, and (d) very high dynamic range (via multiaccum) sampling. Residual energy in the non-central region of the point source PSF, diffracted beyond the 0′′.3 radius image plane mask of the NICMOS camera 2 coronagraph, may be further reduced by coronagraphic PSF template subtraction (CPSFTS) to reveal much fainter target circumferential light of astronomical origin (e.g., from light-scattering orbiting dusty debris or near-IR emissive giant planets in close angular proximity to much brighter host stars; e.g., Fig 1 left). Coronagraphic starlight suppression in combination with flux-density renormalized (scaled) and astrometrically registered CPSFTS can provide per resolution element (resel) to star image contrasts of app. $10^{-4}$ at 0′′.3, few $10^{-6}$ at 1″, and $5 \times 10^{-7}$ at 2″ from the occulted target (e.g., Fig 1 right). However, these performance levels are often difficult to achieve because of a combination of: (a) chromaticity from differences in target vs. PSF template stellar spectral energy distributions (SED) under the passband of the filter used, (b) “wiggling” (metrologic instability) of the cold pupil (Lyot stop) in the dewar (e.g., Krist et al. 1998), (c) “breathing” of the HST optical telescope assembly (differential wavefront errors and PSF instability on sub-orbit to orbit timescales; e.g., Schneider et al. 2001), and (d) target (and PSF template) acquisition (decentration) precision. Contemporaneously obtained “reference” PSFs often do not sufficiently replicate these (a-d) variability phase spaces to enable efficacious PSF subtraction. Additionally, the instrumental calibration of the target and PSF template imaging data must be carried out with extreme rigor (beyond what the OPUS pipeline typically delivers) to remove low-level instrumental signatures that can otherwise be conflated with (or mistaken for) true features in PSF subtracted imaging data. Such calibrations are also needed for “roll subtractions” (a strategy for faint companion detection; e.g., Lowrance et al. 2005) where the target is used as its own PSF by observing at two or more celestial orientation (roll) angles.

To reach the above levels of CPSFTS detection sensitivity, a large ensemble of high fidelity candidate PSF template images are required to tile the dominant parameter spaces of PSF instability and variability. Because HST orbits are a precious commodity, HST science programs have not, individually, observed a sufficiently diverse set of PSF template stars (and many have observed only one – or none). By drawing on the large and diverse data set of raw NICMOS coronagraphic images in the MAST, with reprocessing, we have created well-calibrated and characterized library of both target and candidate PSF images across all NICMOS coronagraphic programs through Cycle 15 (and partially for Cycle 16) to fill this need. With such a library of high quality and well characterized coronagraphic PSFs, as provided by our NICMOS Legacy Archive PSF Library (LAPL: produced under HST AR program 11279), science attempted, but previously unfulfilled or with “negative” results, can be recovered and re-examined. The LAPLACE (LAPL And Circumstellar Environments) re-calibration project provides this ensemble with:

1) the BEST calibrated basic data available for NICMOS coronagraphic imaging.

2) the BEST PSF-template “matching” library for coronagraphic imaging data possible from NICMOS raw archival data reprocessed from the Multi-mission Archive at Space Telescope (MAST).

3) these data and augmenting analysis software (S/W), as made available through STScI and the MAST.
2. Impediments to Efficacious PSF Subtraction Addressed by the LAPL

Template PSFs subtracted from target images that are imperfectly matched in their structural details produce artifacts in difference images that impede unambiguous interpretation and hinder accurate photometry. The principle causes simultaneously contributing to diversities in PSF structures and CPSFTS artifacts are generally not observationally controllable, but with the underlying causalities understood “well matched” PSF templates may be found from ensemble LAPL reprocessed images to attack the problem.

2.1. Chromaticity

Image artifacts on many spatial scales are produced in target minus template PSF subtractions where the point-source filter-band SEDs are not identical (e.g., Fig. 2A), even when the in-band flux densities have been perfectly scaled and the images astrometrically co-aligned with appropriate rigor. This chromatic problem is obviated with two- or multi-roll PSF subtractions for point-source detections; but, generally, this observation strategy is inapplicable for imaging spatially extended structures such as circumstellar disks. However even in this case, other causal mechanisms can be the “tall pole” resulting in PSF mis-matches even when using the target (in roll subtractions) as its own template. For this reason, even point-source (e.g., planet-finding) detection programs conducted in this manner will benefit from the LAPL. Many observers had programmed PSF star observations of “similar” colors to their target stars for circumstellar or AGN studies, but many were not sufficiently close in color (particularly under the broad NICMOS F110W and F160W filters) due to other observational or operational selection constraints. The LAPL data products provide J, H, and Ks flux data (and thus J-H and H-Ks color indices) on all targets, mostly derived from 2MASS and transformed through spectral templates into the NICMOS filter band used for each observation. With that, as a proxy to full in-band spectra for science targets and candidate templates, the LAPL can be searched for candidate PSF templates “well matched” in color and of sufficient brightness.

2.2. Absolute (and Differential) Coronagraphic Centering Precision

The NICMOS flight software in concert with the HST Pointing Control System is responsible for locating, acquiring, and centering coronagraphic targets at the “sweet spot” of the
coronographic optics. Generally, this is accomplished (per a level 3 instrument requirement) to a precision of appx 1/8 pixel (∼ 10 mas) RSX for most target acquisitions. The LAPL data products carry with them the location of the target obscured in the coronagraph to a typical precision of appx 1/100th pixel RSX, predicated upon target acquisition image centroids and spacecraft (target recentering) slew data telemetered with the embedded downlinked engineering telemetry. This information can be used for target:template image alignment before resampling and PSF subtraction, automatically with the LAPL enhanced IDP3 (Stobie & Ferro 2006) S/W, which mitigates the first-order effects of differential mis-centerings with post-facto image processing. However, acquisition decentration at these levels alters the detailed intrinsic structure of the coronagraphic PSF in both the residual diffractive and instrumental scattering components in the halo of the PSF, leaving artifacts due to differential mis-centerings after subtraction (e.g., Fig. 2B). Target:template images obtained with small differential miscenterings minimize these residual artifacts, as evidenced through exploration using down-selected PSF candidate observations from the LAPL.

2.3. **OTA “Breathing”**

Metrological instabilities in the HST Optical Telescope Assembly (OTA) induce orbit-phase and pointing dependent differential wavefront errors (WFE; primarily in focus) arising from desspacing (“breathing”) of the secondary mirror (Suchkov & Hershey 1998) that changes the structure of PSF (Makidon et al. 2006). These changes in PSF structure act on both intra- and inter- orbit timescales and induce residual artifacts in CPSFTS (e.g., Fig 2C). The WFE-correlated “breathing” phase and amplitude of PSF variations are generally not controllable by observation design. The LAPL candidate PSF templates, as an ensemble, stochastically tiles the OTA breathing phase spaces and provides “breathing matched” coronagraphic template PSFs to subtract from similarly breathing-affected science images.

2.4. **Cold Pupil Mask (Lyot Stop) Motions**

The NIC2 Lyot stop is designed to mask the telescope central obscuration, spider vanes, primary mirror hold down pads, and primary mirror outer support ring. Unfortunately, due to the dewar deformation from solid nitrogen (SN₂) expansion and zero gravity release, the mask itself is both slightly decentered and subject to small, thermally induced, motions. As a result small parts of the pupil intended to be masked are differentially revealed and covered with changes in the instrument thermal state causing changes in PSF structure. The pupil stop motion is primarily in the V2/V3 plane causing the diffraction spikes in CPSFTS images to become very dark or very light (e.g., Fig 2D), as well other effects, with differential pupil mask offsets. This represent yet another PSF “instability space” that is tiled, by the LAPL.

3. **LAPL Image Reprocessing - Improvements over the OPUS Pipeline**

To enable high fidelity PSF subtraction using the LAPL, we have undertaken a rigorous reprocessing of ∼ 8,800 raw science data images obtained in 46 NICMOS coronagraphic programs¹ executed in Cycles 7, 7N, 11-15 (all inclusive) and selective targets from Cycle 16. The LAPL recalibration of these high dynamic range science target and PSF template observations (including also supplemental non-coronagraphic imaging obtained in some of these programs) improve the photometric efficacy and image quality over the OPUS pipeline and comprise the LAPL.

¹7038, 7052, 7157, 7179, 7220, 7221, 7226, 7227, 7233, 7248,7329, 7418, 7441, 7808, 7828, 7829, 7834, 7835, 7857, 7897, 7924, 8079, 8979, 8983, 8984, 9693, 9768, 9834, 9845, 10147, 10167, 10176, 10177, 10228, 10244, 10448, 10464, 10519, 10527, 10540, 10560, 10599, 10847, 10849, 10852, 11155
Figure 2: The four principle causes for target:template mis-matches in PSF structures giving rise to image artifacts in PSF subtracted images over a large range of spatial scales are separately illustrated above. The large pool of candidate PSF templates provided by the LAPL that, as an ensemble, tiles these parameter spaces and enables PSF subtraction that is otherwise compromised; e.g. see Fig 3.

3.1. Dark Frames

The two-dimensional dark structure in NICMOS science data readouts depends not only upon the integration time and details of how the detector is clocked (i.e., what “sample sequence” is used) with multiple non-destructive readouts, but also upon the detector bias voltage sourced by electronics far removed from the focal plane array (FPA) and upon the temperature at which the device is operating. Changes in temperature due to self-thermalization with the operation of the FPA readout amplifiers, instabilities from other sources, and secular drift over time during the SN2 and NICMOS Cooling System (NCS) eras, altered both the dark current and also other image artifact structures. The OPUS pipeline synthesizes model dark frames characterized from (often sparse) on-orbit calibration data to reduce these instrumental signatures, but sometimes fails to sufficiently mitigate zero-point (DC or bias) offsets, amplified flat-field imprints and other artifacts. Such artifacts are often apparent in OPUS calibrated images and, differentially (with non-repeatabilities), become very prominent in difference (PSF subtracted) images. To mitigate this problem, whenever possible, we cull from the totality of all available on-orbit dark frames (of the same sample sequences), matched in detector temperature by $\leq 100$ mK in Cy 7/7N and $\leq 50$ mK in Cy 11-16 to create a dark calibration reference frame for every science image to be calibrated. We provide these calibration reference dark frames in the LAPL along with the science image data products. In some cases where suitable on-orbit dark data are unavailable (notably for some sample sequences at some epochs in Cycles 11-16), we rely as well on model darks, but in other circumstances significant improvements in dark calibrations are realized (e.g. see Fig 4).
Figure 3: Due to chromatic, breathing, centering, and cold mask mis-alignment effects, CPSFTS images exhibit a large diversity in image quality and unmitigated optical artifacts. Here, with images (from upper right to lower left) ordered in decreasing J-H color, the HD 181327 (J-H = +0.22) debris ring is well revealed in only a few images (outlined in green), but obliterated in many more (e.g., a few highlighted in red).
Figure 4: Representative improvement in dark (re)-calibration with LAPL “custom reference dark” creation (left) compared to processing with the OPUS pipeline using synthesized model darks (right), both identically stretched in this figure. Pixel-to-pixel medians and variances within the three subarrays unaffected by flux from the target are used to assess the level of improvement, though zero-point offsets, flat field imprints, and the strong appearance of the “photometrically challenged” column (128) are visually apparent in the OPUS processed image. (Dataset: n41k23nkm)

3.2. Flat Fields at Small Inner Working Angles

One of the major strengths of the NICMOS coronagraph is its ability to push to inner working angles as small as 0\prime\prime.3 (the angular size of its image plane obscuration, aka the coronagraphic “hole”, projected onto the camera 2 detector) from an occulted target. Because of metrological instabilities in the NICMOS dewar system, the location of the obscuration (and targets acquired into it) projected on the detector focal plane can vary by several pixels\(^2\) on time scales of days to weeks. This is not an issue for coronagraphic centering and acquisition on orbit or sub-orbit timescales, but is fundamentally incompatible with STScI’s flat field calibration program that obtains data on monthly to (for many filters) as long as multi-year time scales. The high SNR camera 2 flat field reference files, available through CDBS and used by OPUS, have static imprints of the corongraphic hole that often do not match its location when science (and PSF) observations were made. While applicable elsewhere in the FOV, mis-registered hole imprints create image artifacts when these flats are used for calibration that are deliterious to imaging at small coronagraphic IWAs (e.g., see Fig 5). To mitigate this problem, we create contemporaneous flat field reference files for every coronagraphic observation, derived from “hole finding flats” that are autonomously acquired as part of the target acquisition process\(^3\). Hole finding flats are always taken in only the F160W filter, and thus we perform a pixel-by-pixel detector-temperature dependent band-transformation to the science data filter band based upon the instrumental calibration to create contemporaneous flats to match each science observation. These flats

\(^2\)While these motions are typically \(\leq\) 2 pixels, the photometric reach is greater with a vignetting signature extending \(~ 0\prime\prime.5\) from the obscuration edge as it is slightly afocal w.r.t. the HST OTA f/24 focal plane.

\(^3\)Because of the coronagraphic hole motion over time, it is necessary to locate its position on the detector focal plane through internal flat field observations analyzed by the on-board flight S/W. These acquisition support images are downlinked with the science data images and are used advantageously in LAPL reprocessing.
Figure 5: Left: Two PSF-subtracted images of RY Tau (using the same PSF template) obtained in the same spacecraft orbit but at field orientations differing by 30° as flat-fielded by the OPUS pipeline exhibit nearly identical rotationally invariant bright spots exterior to the r=0''3 coronagraphic hole, belying an instrumental origin. Middle: Multiplying the contemporaneous (unity normalized) “hole finding” with the CDBS reference flat used in the OPUS pipeline calibration reveals a misplaced imprint of the coronagraphic hole (detailed in the right panel). Red circle is the location of the coronagraphic hole at the epoch of the science observation, it’s (opposite parity) imprint in the CDBS flat is displaced by about half an arcsecond. Color scale in the linear stretch indicates photometric multiplicative error factors which arise using the OPUS calibrated image.

remove the “hole imprint” artifacts induced by the use of CDBS flats in the OPUS pipeline. LAPL calibrated science data images named *.calc.fits use these observation-matched contemporaneous flats in calibration that we also provide in the LAPL.

3.3. Flat Fields at Larger Angular Distances

The target acquisition “hole finding flats” are exposed to a per-pixel SNR ~ 50, as opposed to several hundred to ~ 1200 in the medium or broad band filter calibration reference flats created by STScI that reside in the CDBS and used by OPUS. Except in the regions (generally of greatest interest in coronagraphic programs) near the coronagraphic obscuration, the CDBS flats would be superior to use. However, secular changes in the flat-field response in different filters (separately within the SN2 and NCS eras of operations) have been noted that result in photometric errors after flat-fielding. For each filter we have re-examined all available on-orbit internal flats to assess field (detector location) dependent variations in sensitivity (photometric response) with both time and temperature. From this we create, by exposure binning and/or interpolation as is appropriate, an “epochal” flat field calibration reference file for every science data image that is also used to produce a LAPL calibrated image (denoted *.cale.fits) applicable everywhere in the FOV except close to the coronagraphic obscuration for reasons discussed above. Thus, the LAPL provides two calibrated data images for each exposure, separately, with hole-edge artifact rejection using lower SNR flats, and higher precision photometry at angular distances unaffected by the presence of the coronagraphic hole. The LAPL does not synthesize these into a single image, as the applicability depends upon the metrical needs (and goals) of any particular investigation.
3.4. Bad (uncalibratable) Pixels - Rejection and Correction

The OPUS pipeline uses static “bad” pixel maps to identify defective or inoperable pixels. Pixels that are deemed as bad, by adopted criteria, are left uncalibrated in the calnica pipeline process (and those pixel values set to zero). The criteria used in the STScI OPUS pipeline to identify “hot” pixels (with excessively high dark current) or under-responsive pixels (possibly by shadowing from contaminants on the detectors themselves) are very conservatively set – appropriate for very faint object imaging in sparse fields and where image dithering is a viable strategy for recovery. These criteria, however, are not so well suited for high dynamic range imaging where the target region of interest is filled with bright flux and dithering (with coronagraphy) is not a viable observation technique. Hence, in LAPL processing we developed and applied static bad pixel maps (separately for the SN₂ and NCS eras) with identification criteria specifically applicable to bright object coronagraphy. In particular, we define a pixel as hot if it contributes noise from dark current in 64 seconds that is more than half the read noise. With this criteria the static number of excessively hot pixels during the SN₂ era excluded from LAPL calibration was 36 and during the NCS era was 196. Excessively under-responsive pixels were evaluated from F222M internal flat field lamp (and background) exposures defined as $\leq -5\, \sigma$ less responsive than the median of good neighbors in a 9x9 pixel box about every pixel. With that criteria we found 143 and 214 pixels uncalibratable due to under-responsivity in the SN₂ and NCS eras, respectively.

We document the calibration status (good, dead, saturated, under-responsive, excessively hot, etc.) of every pixel in a data quality (DQ) extension on all LAPL image data products. As in the OPUS pipeline, pixels that are found to be defective, by the criteria summarized above, are not calibrated - but in LAPL processing are “repaired” post-facto. Bad data reparation is required with coronagaphic data that in application for PSF subtraction are (later) digitally re-binned and resampled with template-to-target astrometric alignment; bad pixels if “set to zero” (as in the OPUS pipeline) would adversely affect a ring of neighbors when resampled. Uncalibratable pixel values are estimated by two-dimensional Gaussian-weighted bi-cubic interpolation of good neighbors, using sinc-function apodization with a kernel width matched to the FWHM of the bandpass employed for the observation (to suppress interpolative “ringing”, particularly if undersampling the PSF which is the case in the F110W filter in camera 2). LAPL data products which carry the name *calf have had uncalibratable pixel values estimated in this manner (for an example, see Fig 6). While no bad (zero) pixel values appear in the LAPL *calf files, the DQ extensions reflect the state of the pre-repaired pixels, so investigators can know which pixel values have been estimated rather than directly measured.

4. Using the LAPL

To assist LAPL archival users in finding candidate NICMOS coronagraphic PSFs templates potentially suitable for their science targets of interest, we have attached a number of qualifying, descriptive, and metrical FITS file header keywords and corresponding values to characterize every image (with a recommendation to STScI that these be ingested into a meta-database in the MAST to facilitate data searches). We summarize these in Tables 1 and 2. We also have augmented, and make available through STScI, the IDL-based IDP3 image analysis S/W with added functionality to enable and enhance interactive exploration and quantitative data analysis of PSF subtractions using large image ensembles derived from the LAPL and other sources.
Figure 6: Example of LAPL “bad pixel” reparation (see main text; log10 image displays). Left: Pixels in black are defective and uncalibratable by the criteria adopted for the LAPL. Right: The same image after repairing uncalibratable pixels.

5. The Proof is in the Pudding

Many NICMOS coronagraphic science programs have reported negative results in source detections that are (now) attributable to the lack of suitable template PSFs as remedied by the LAPL. While yet to be systematically applied across the diversity of target classes, during the LAPL development a number of previously published “non detections” were revisited and turned into spatially resolved and photometrically reliable analysis quality data images with LAPL processing methods and PSF templates; for example the cases of the sub-arcsecond transition disks around the PMS stars MWC 480 and SAO 206462 shown in Figure 7. The same observing program succeeded in detecting the brighter, and larger (so less contrast challenging) circumstellar disk about the PMS star HD 100546. With LAPL reprocessing and PSF template subtraction, the improvement in image data quality will allow re-interpretation of these data with the image artifacts earlier affecting photometric isophotes now largely mitigated (Fig 8, panels A & B). Significant improvements in optically thin debris disk images are also demonstrated with original and LAPL CPSFTS imaging of the HD 181327 debris ring (Figure 8, panels C & D), for which the morphology of much more sharply bounded ring edges are now clearly seen.

6. Summary

The NICMOS LAPL will be made publically available through the MAST along with image analysis tools in the IDL/IDP3 environment. LAPL data keywords will enable archival science investigations and will leverage the large investment in orbits devoted to NICMOS coronagraphy with new results previously unobtainable. Good hunting!

Acknowledgement: Based on observations made with the NASA/ESA Hubble Space Telescope, obtained from the Data Archive at the Space Telescope Science Institute, which is operated by the Association of Universities for Research in Astronomy, Inc., under NASA contract NAS 5-26555. Support for program the LAPL program, AR 11279, was provided by NASA through a grant from the Space Telescope Science Institute.
Figure 7: During HST cycle 7, the close environments of four old pre-main sequence stars were probed with NICMOS coronagraphy in GO program 7857. Negative results for attempted scattered-light detections of circumstellar disks associated with two of those stars were reported by Augereau et al. (2001); but with a suitable PSF template ensemble and LAPL reprocessing, those disks were recently imaged and photometric surface brightness profiles measured.

Figure 8: Representative improvements in CPSFTS imaging of circumstellar disks. (A) HD 100546 (F160W) as reported Augereau et al. 2001, (B) with LAPL reprocessing and PSF templates. (C) HD 181327 (F110W) as reported by Schneider et al. 2005, (D) with LAPL reprocessing and PSF templates.
References

Makidon, R. B., et al. 2006, SPIE, 6270, 52
Stobie, E. & Ferro, A. 2006, ASPC, 251, 540
Suchkov, A. & Hershey, J. 1998, STScI ISR NICMOS 98-015

Table 1: Observation Description FITS Header Keywords.

<table>
<thead>
<tr>
<th>Name</th>
<th>Value</th>
<th>Description</th>
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<tr>
<td>TARGTYPE</td>
<td>POINT_SOURCE</td>
<td>Unresolved, Isolated, Point Source (Star)</td>
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<td></td>
<td>EXTENDED_SOURCE</td>
<td>Spatially Resolved Bright Source (e.g., circumstellar disk)</td>
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<td></td>
<td>DOUBLE_STAR</td>
<td>With Well Separated Components</td>
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<td></td>
<td>CORON_DOUBLE</td>
<td>≤ 0.3 overlapping PSF cores inside coronagraph</td>
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<td></td>
<td>FAINT_OBJ</td>
<td>Spatially Resolved Faint Source (e.g., QSO, AGN)</td>
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<td>TARGCNTR</td>
<td>CENTERED</td>
<td>Target Nominally Positioned in Coronagraph</td>
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<tr>
<td></td>
<td>DECENTRED</td>
<td>Target in Coronagraph But Decentered</td>
</tr>
<tr>
<td></td>
<td>PERIPHERY</td>
<td>Target at the Periphery of the Coronagraphian Mask</td>
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<tr>
<td></td>
<td>UNOCCULTED</td>
<td>Target in FOV but Not Coronagraphically Occulted</td>
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<tr>
<td></td>
<td>OUTFOV</td>
<td>Target Out of FOV but PSF Wings (Halo) Seen in FOV</td>
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<tr>
<td></td>
<td>NOTARGET</td>
<td>Valid Image Data but No Target Visible in FOV</td>
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<tr>
<td></td>
<td>UNSTABLE</td>
<td>Target Visible but Pointing Unstable (drifting, etc.)</td>
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<tr>
<td>PSFQUAL</td>
<td>OK</td>
<td>PSF Quality Nominal (useful as PSF Subtraction Template)</td>
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<tr>
<td></td>
<td>SATURATED</td>
<td>PSF Core Saturated (wings may be OK)</td>
</tr>
<tr>
<td></td>
<td>DEGRADED</td>
<td>PSF Optically Degraded (focus, image smear, other issues)</td>
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Table 2: Target Location and Photometry FITS Header Keywords.

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<td>2MASS Catalog J Magnitude</td>
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<td>H2MASS</td>
<td>6.380</td>
<td>2MASS H Magnitude</td>
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<td>HE2MASS</td>
<td>0.018</td>
<td>2MASS H Mag. Uncertainty</td>
</tr>
<tr>
<td>K2MASS</td>
<td>6.245</td>
<td>2MASS Ks Magnitude</td>
</tr>
<tr>
<td>KE2MASS</td>
<td>0.020</td>
<td>2MASS Ks Mag. Uncertainty</td>
</tr>
<tr>
<td>Fnnnn_EF</td>
<td>'Y'</td>
<td>Filter Band Transformation Error Flag</td>
</tr>
<tr>
<td>Fnnnn_JY</td>
<td>2.69286</td>
<td>NIC2 Filter Band Flux Density in JY</td>
</tr>
<tr>
<td>CPHOTSTR</td>
<td><code>rn(icat(k93models,5200.0,0.0,4.49), SYNPHOT Transformation Input band(h),6.380,vegamag)</code></td>
<td>SYNPHOT Transformation Input, NIC2 filter band</td>
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</table>

nnnn is NICMOS filter name, e.g. F180M

Error flag set to 'Y' (Yes) if transformation input parameters were not known

Stellar SED model and flux density used to transform to NIC2 filter band
Calibration of the HST NICMOS F110W Using High Redshift Red-Sequence Galaxies

Pascal Ripoche

Lawrence Berkeley National Laboratory, Berkeley, CA 94820

Abstract.

We present a new method to measure photometry zero-points using high redshift red-sequence galaxies. This method allows us to measure the HST NICMOS F110W zero-point close to the sky level and, thus, avoid calibration errors due to the poorly constrained HgCdTe non-linearity at these faint flux levels. This is the level at which almost all of the highest redshift supernovae observations are obtained with NICMOS camera. We combined HST optical and VLT near-infrared observations of distant red-sequence galaxies to constraint their spectral energy distribution and thus derived NICMOS calibration from the observed fluxes with NIC2 F110W. Using 23 red-sequence galaxies in three distant galaxy clusters, we determine the absolute NICMOS F110W zero point to 2.5% accuracy. This result shows that the non-linearity is over-corrected using the standard STScI pipeline leading to a significant zero-point offset at the sky level.

This work has been supported by the Office of Science, U.S. Department of Energy, through contract DE-AC02-05CH11231 and in part by NASA through grants associated with HST-GO-10496 and HST-GO-11799.
An Overview of Detectors

Bernard Rauscher

*NASA Goddard Space Flight Center, Greenbelt, MD*

**Abstract.**
Removing the instrument signature of the detectors is an important part of calibrating any imager or spectrograph. When doing this, it helps to understand how the detectors work, and the kinds of artifacts that can manifest themselves. This talk will provide an overview of the hybrid HgCdTe and Si:As detectors that will be used by the James Webb Space Telescope, with an emphasis on calibration.
New Scientific Capabilities of the HST WFC3

Robert O’Connell

*University of Virginia, Charlottesville, VA*

**Abstract.**

I will describe some of the new scientific capabilities enabled by the extended wavelength coverage (0.2-1.7um), large suite of specialized filters, and improved ”discovery efficiency” of HST’s Wide Field Camera 3. Some examples from Cycle 17 programs include the star formation histories of nearby disk and early-type galaxies (from both resolved sources and integrated light), discovery of multi-age pre-main-sequence populations in massive Local Group star clusters, the evolutionary state of galaxies at intermediate redshifts (z \( \approx \) 1-4), and identification of candidates for the highest redshift galaxies at z \( \approx \) 8-9 in the Hubble Ultra Deep Field.
Performance and Calibration of Wide Field Camera 3

John MacKenty

Space Telescope Science Institute, Baltimore, MD 21218

Abstract.

The Wide Field Camera 3 was installed in HST in May 2009 during SM4. Designed to provide a factor of 10 or better increase in the near-ultraviolet and near-infrared imaging capabilities of the observatory, WFC3 has surpassed these expectations by large margins and is operating well. This talk will provide an overview of the instrument’s design, its in-flight performance, and our strategy and priorities for its calibration.
The Photometric Performance and Calibration of WFC3

J. S. Kalirai and A. Rajan

Space Telescope Science Institute, 3700 San Martin Drive, Baltimore, MD, 21218

Abstract. We analyze multiple observations of five bright HST spectrophotometric standard stars obtained with WFC3/UVIS and WFC3/IR in order to establish the temporal stability of the instrument over its first year of science operations and its photometric sensitivity. Observations of the five stars, GD 153, G191B2B, GD 71, GRW+70d5824, and P330E were obtained at multiple dither positions at each epoch in all filters. The cadence of the observations was once per week shortly after the instrument was installed in Hubble in the summer of 2009 (i.e., during SMOV) to once every few months by the end of the first year. Analysis of the high signal-to-noise photometry in wide and medium-bands indicates that the instrument is photometrically stable, with measured variations that are <0.5%. As calibrated after SMOV, the absolute throughput of the instrument is substantially higher than predictions based on Thermal Vacuum 3 (TV3) tests. Preliminary photometric zero points have been calculated in UVIS and IR filters factoring in the measured on-orbit throughput, however this analysis still uses the WFC3 ground flat fields. A pending update will provide more accurate zero points based on the first on-orbit flat fields for WFC3 which have just been released.

1. Introduction and Background

Since its installation in Hubble in May 2009, the WFC3 instrument has quickly emerged as a powerful and unique tool for astrophysical investigations. The instrument offers two wide-field cameras that operate at high resolution and cover a broad wavelength range from the UV-visible-NIR (i.e., 200 – 1700 nm). A total of 77 narrow-, medium-, and wide-band imaging filters and 3 spectroscopic grisms on WFC3 ensure that a diverse set of scientific problems can be efficiently tackled. For example, during its first weeks of science operations, WFC3 captured images of an impact on Jupiter and also detected the highest redshift galaxies ever observed.

The success of scientific programs executed on WFC3 is intimately tied to the accuracy of the instrument’s calibration. During the past year, many WFC3 calibration programs have been executed to establish the instrument’s performance in orbit, several of which have the primary aim of characterizing the absolute sensitivity and stability of both the UVIS and IR cameras. The primary technique used to achieve the absolute calibration for HST instruments involves imaging bright spectrophotometric standard stars which have model spectral energy distributions that are well defined. For example, white dwarf stars contain pressure broadened Balmer lines that can be modeled to yield the temperature and gravity of the star in a non-degenerate way (e.g., Bergeron, Saffer, & Liebert 1992). The three primary spectrophotometric standards used by HST are the hot DA (hydrogen atmosphere) white dwarfs G191-B2B (V = 11.77), GD 153 (V = 13.35), and GD 71 (V = 13.03). For these three stars, the UV, visible, and NIR models agrees with the STIS spectra to better than 1%, indicating a superb absolute flux calibration. Further information on the calibration of these stars is provided in Finley, Koester, & Basri (1997), Bohlin (2000), and Bohlin, Dickinson, & Calzetti (2001).
In this paper, we describe the absolute calibration and photometric stability of WFC3 during its first year of science operations. Our goal is to provide a characterization of the end-to-end sensitivity of both instruments in all 77 imaging filters, taking into account the HST Optical Telescope Assembly (OTA) and instrument components such as the pickoff mirror, mirror reflectivity, inner/outer windows, detector quantum efficiency, and filter transmissions. Our final measurements are also dependent on the many other calibration reference files that are used to process raw HST data into flt and drz images, such as the geometric distortion solution and the flat fields. At the time of this calibration workshop, an alpha-release geometric distortion solution was available however on-orbit flat fields were not, and therefore the “final” on-orbit sensitivity measurements are likely to change (at most) by a few percent in the near future.

2. The First On-Orbit Sensitivity Measurements: SMOV4

The expected performance of WFC3 was estimated from ground tests during Thermal Vacuum 3 (TV3), using the “CASTLE” optical stimulus at NASA/GSFC (see Brown 2008). The results of these tests led to throughput tables which were populated in CDBS and SYNPHOT, and from which photometric zero points were measured. We performed the first on-orbit calibration of WFC3’s sensitivity shortly after the instrument was installed in Hubble. These first observations targeted GD 153 (one of the three primary spectrophotometric standards) on WFC3/UVIS1 in a subset of 37 widely used filters (based on Cycle 17 usage statistics). For WFC3/IR, we targeted both GD 153 and P330E (a solar analog secondary standard) in all 15 filters.

All of the exposure times were set to ensure a high signal-to-noise (e.g., ~100) detection of the star, which typically required only a few seconds. The measured count-rate of these stars was compared to the predicted count rate from convolving the TV3 total system throughput with the model spectrum of the stars. Surprisingly, all of our results indicated that the true sensitivity of both cameras is much higher than the ground calibration indicates (see Figures 1 and 2). This is possibly due to uncertainties in the “CASTLE” calibration or alternatively, could represent an error in the throughput table of the HST OTA. We note that the first HST/ACS on-orbit measurements reported by Sirianni et al. (2002) were also systematically higher than their ground tests.

Details of our photometric technique and the methods used to calibrate the WFC3/UVIS and IR instrument throughputs based on these SMOV results are presented in two extensive ISRs, Kalirai et al. (2009a) and Kalirai et al. (2009b). We specifically note here that our first calibration from these results used an iterative approach where we compare synthetic photometry, as calculated after folding in a smooth correction to SYNPHOT, to the measured counts. The advantage of this approach is that SYNPHOT will take the integral of the actual bandpass with the full throughput, including the on-orbit correction factor, and will yield output photometry that then determines a new correction factor. The total system throughputs from this analysis were used to produce photometric zero points for WFC3 in the VEGAMAG, ABMAG, and STMAG systems, and updated throughput tables were calculated and installed in the CDBS (e.g., for SYNPHOT and the ETC) and OPUS pipelines. These zero points are also available on the WFC3 web page along with enclosed energy curves.

3. The WFC3/UVIS and IR Cycle 17 Calibration

At the conclusion of SMOV4, a more detailed WFC3/UVIS and IR calibration program was implemented to measure the total system throughput using multiple standard stars on both CCDs of the UVIS camera and the IR detector. Observations were also obtained in an expanded set of configurations (e.g., subarrays and full frames) and in those filters which were not observed in SMOV4. With knowledge of the on-orbit throughput, the exposure
Figure 1: WFC3/UVIS measurements of a spectrophotometric standard star indicates that the total system throughput is 5 – 10% higher at blue and red wavelengths and 15 – 20% higher at 4000 – 7000 Å compared to ground tests.

Figure 2: WFC3/IR measurements of two spectrophotometric standards indicates that the total system throughput is 10 – 15% higher in all filters compared to ground tests.
times in the Cycle 17 calibration program were adjusted to fill the full well up to 90% of its capacity (thereby minimizing the random errors in the photometry). All of these data, as well as the earlier SMOV4 observations, have also been reprocessed with updated calibration reference files.

3.1. Photometric Temporal Stability
The temporal stability of WFC3 can be characterized from the multiple data sets obtained throughout the first year of operation. A detailed discussion of the photometric stability will be presented in Kalirai et al. (2010, in prep), which will include an analysis of all spectrophotometric standards at multiple locations on each of the detectors. The observations with the highest S/N on the UVIS detector are those targeting GD 153, whereas the observations with the best cadence are those of the contamination monitor programs which target the HST standard GRW+70d5824 (CAL11426 and CAL11907). For the IR detector, observations of GD153 and P330E have a combined cadence of roughly once per month. To minimize systematics from spatially varying the location of the stars (e.g., flat field variations), we restrict our analysis to separate subsets of data taken at the same location of the detector. As one example, we illustrate in Figure 3 that the GD 153 observations observed on the 512 pixel UVIS1-C512A-SUB subarray indicate temporal stability to >99.5% over a 320 day time frame. Our measurements taken at the other corners of the UVIS detector with multiple standards also demonstrate that temporal variations are less than 0.5%, often limited by the error in the measurement. We also measure a similar stability on WFC3/IR in all wide and medium-band filters.

3.2. Towards New Absolute Throughput Measurements
The stability of WFC3 over its first year implies that the SMOV4 absolute throughput calibration of the instrument’s UVIS and IR cameras is accurate. The calibration on the UVIS camera was based only on one star on one CCD, GD 153, whereas we have now observed G191B2B, P330E, and GD 71 on both CCDs. Similarly, for the IR camera, the calibration was based on both GD 153 and P330E and the other two stars have now been observed. Over the past year, several WFC3 reference files have changed (e.g., geometric distortion solution, bad pixel tables, dark, etc.), and so we have re-processed all of the SMOV4 and Cycle 17 data to re-measure the throughputs. Given the updated calibration to SYNPHOT based on the results shown in Figures 1 and 2, our expectation is that the measured counts of each of these stars will agree well with predictions from SYNPHOT.

We demonstrate the comparison between our aperture photometry and calibrated SYNPHOT predictions for WFC3/UVIS in Figure 4. Within each of the panels, the ratio of the measured to predicted counts is shown for a given standard, where all observations over multiple epochs have been averaged together. Observations in wide-band filters are shown with black open circles, narrow-band filters are in grey, and quad filters are in purple. All three of the stars show very similar trends with the ratios being near unity, as expected (the dotted lines mark +/-2% variations). Note specifically that observations in filters that deviate from unity, such as F336W or the red filters with $\lambda > 9000 \text{ Å}$, are consistent among the three stars. These deviations were hinted in the SMOV4 data as reported in Kalirai et al. (2009b), however, we intentionally did not calibrate out the effects.

In Figure 5 (top), we present a closer comparison of the three standards shown in Figure 4 in wide-band filters. This plot demonstrates that the overall throughput of WFC3/UVIS is about 1% higher than indicated by the SMOV4 calibration. The bottom panel shows the residuals from comparing GD 153 to G191B2B (red curve) and GD 153 to P330E (blue curve). The large tick marks on this plot denote 0.5% variations and therefore the two white dwarfs are in exquisite agreement with one another. P330E falls slightly below the two white dwarfs, which is also what is seen for ACS (see Bohlin 2007). This likely indicates a small error in the flux calibration of the STIS P330E spectrum.
Figure 3: The temporal stability of WFC3/UVIS as measured from higher signal-to-noise observations of the bright HST spectrophotometric standard GD 153 on UVIS1-C512A-SUB. At each of the epochs, four dither positions were obtained and averaged together, where the error bars reflect the 1σ error in the mean. All but a single wide-band filter (F218W, σ = 0.5%) shown here have standard deviations of <0.4% (note that the large tick marks represent 1% variations). Some of the observations in F606W and F814W shown in the bottom two panels were intentionally shifted to UVIS2 and are thus missing from the plot.

A comparison of Cycle 17 observations of standard stars on WFC3/IR also shows very similar results to the SMOV calibration. Similar to Figures 4 and 5, our calibration indicates consistency at the 1% level between three white dwarfs (including GD 71) and the solar analog P330E, in all 15 IR filters.

For new instruments, the absolute throughput is expected to change as new calibration reference files are produced. The largest correction is often anticipated from the flat field, and preliminary tests with WFC3 suggest that low frequency spatial variations affect the ground-based flats causing photometric errors of several percent (see Sabbi et al. 2009). These ground flats are currently the calibration files in the OPUS pipeline and also were used to make Figures 4 and 5. The first on-orbit flat fields for WFC3 have been released in August 2010 on the instrument web page, and they have been shown to improve spatial variations by up to 4.5% in some filters, leaving much smaller residuals.

Of particular concern for the absolute throughput measurements is the quality of the flat field at the location of the subarrays in which the standard stars are placed. For the IR camera, the subarrays are located at the center of the detector and the new on-orbit flat agrees with the ground flat to within 1%. However, for the UVIS camera, the subarrays are located in the corners of the mosaic, and there are presently few overlapping observations that can provide a reliable correction at these extremes. This will improve as the flat field calibration program continues to execute through Cycle 18, and we will make corresponding updates to the WFC3 zero points and SYNPHOT throughput tables. At the present time,
Figure 4: A comparison of the measured counts vs post-SMOV4 calibrated SYNPHOT predictions for three standard stars. Observations in wide-band filters are shown with black open circles, narrow-band filters are in grey, and quad filters are in purple. All three of the stars show very similar trends and agree with the predictions to within 2% in most filters. As discussed in the text, an updated calibration that factors in these results will be performed after on-orbit flat are used to reprocess the data.

it is important to note that the zero point of a filter such as F336W, which deviates from unity in Figures 4 and 5, is 5% higher than published.

4. Summary

The new WFC3 camera on Hubble is in full operation. The photometric throughput of the instrument’s UVIS and IR cameras is calibrated to $\sim 1 - 2\%$ in almost every filter, which will improve even more when we apply on-orbit flat field corrections to the standard star observations. The overall sensitivity of WFC3/UVIS and IR is substantially better than predicted from ground tests, which will enable new scientific investigations and enhance others.

WFC3/UVIS has a higher throughput than any HST instrument over the wavelength range extending from its blue cutoff (at 2000 Å) to $\sim 4000$ Å. Beyond this limit, the choice of which HST instrument is best suited for users depends on the details of the science requirements. Although the absolute throughput of ACS is higher at optical wavelengths, the highest WFC3 efficiency gains in our on-orbit observations also occur at 4000 – 7000 Å, and therefore the gap between the two instruments has been closed significantly. Specifically, some of the primary science observations with HST require several orbits where signal is coadded over multiple exposures. Relative to ACS/WFC, WFC3 has smaller pixels by 20%, a lower read noise by $\sim 50\%$, a smaller CTE correction, and a much lower dark current. Combined with our new efficiency measurements, the ability to better sample the PSF and naturally beat down the noise through multiple exposures can make WFC3 the
Figure 5: A closer look at the measured counts of three standards in the wide-band filters and their SYNPHOT predictions. The curves for all three standards agree well with one another, with residual variations between the two white dwarfs being <0.5%.

preferred instrument for even broadband F606W and F814W observations of faint sources. For example, the two instruments will both reach a limiting ABMAG of 27.9 (29.2) for a S/N = 10 point source detection in 1 hour (10 hours) in the F606W filter. The choice between the two instruments will require careful predictions from the respective ETCs, factoring in the detailed observational setup. Of course, WFC3 contains many more filters over its complete wavelength range than ACS/WFC, yet ACS offers a 50% larger field of view, both considerations being potentially important for users.

References

WFC3 UVIS and IR flat-fields

E. Sabbi\textsuperscript{1}, J. Mack\textsuperscript{1}, N. Pirzkal\textsuperscript{1}, A. Viana\textsuperscript{1}, J. S. Kalirai\textsuperscript{1}

Abstract. Flat fielding is a standard calibration step for astronomical data, which allows us to correct for variations in the local response of a detector and improve the accuracy of photometric analysis. As for other Hubble Space Telescope (HST) instruments, ground-based flat-fields are, and will be, the base to remove the pixel-to-pixel variations in the Wide Field Camera 3 (WFC3) data. Tests performed during the Servicing Mission Observatory Verification (SMOV4), that followed the installation of WFC3 on Hubble, indicate that ground-based flat-fields do not completely remove local variation in the response of the detector, but that low-frequency structures are still present in both the UVIS and IR data. As part of the WFC3 standard calibration process we have used observations of the rich globular cluster Omega Centauri to compare the flux of the same stars at different positions on the WFC3 detectors and derive a correction to remove the remaining low-frequency structures. Here we will present the characteristics of the ground-based high-frequency flat (better known as P-flat) and the low-frequency (or L-flat) flat fields, how these files are created and their impact on the processed astronomical data.

1. Introduction

WFC3 was installed on HST in May 2009, during the last Servicing Mission (SM4) to replace the Wide Field and Planetary Camera 2 (WFPC2). One of the most important features of WFC3 is its broad wavelength coverage from the near ultraviolet ($\sim 200$ nm) to the near infrared ($\sim 1700$ nm).

WFC3 consists of two independent channels, the optical/ultraviolet (UVIS) channel, with two $2051 \times 4096$ pixel CCDs with a spatial resolution of $0.04''$ pixel$^{-1}$ and operating from 200 to 1000 nm, and the near infrared (IR) channel with a HgCdTe, $1024 \times 1024$ pixel array operating between 800 and 1700 nm, with a spatial resolution of $0.13''$ pixel$^{-1}$.

Flat-fielding is one of the standard calibration steps of astronomical images. It is designed to normalize the system illumination path and remove pixel-to-pixel non-uniformities, so that count rates are independent of the position on the detector. In this proceeding we will discuss the characteristics of the WFC3 flat-fields for the UVIS and the IR channels, their accuracy and how they could be improved.

2. WFC3 flat-fields

As for several other HST instruments, WFC3 flat-fields were acquired from the ground under thermal vacuum conditions. In the case of WFC3 both the UVIS and the IR channels flat-fields where acquired at the Goddard Space Flight Center during the last WFC3 thermal vacuum campaign (TV3) at the beginning of 2008. The optical stimulus CASTLE (Figure 1) was used as a source of external illumination to reproduce the Optical Telescope Assembly (OTA).
During TV3, high signal-to-noise full frame flat-fields were acquired for each of the 77 filters (62 in the UVIS channel and 15 in the IR) of WFC3 by operating the camera at the same temperatures currently used in space (namely $-83^\circ$C for the UVIS and 145 K for the IR channel respectively, Sabbi et al. 2008; Bushouse 2008).

Figure 2 shows four examples of UVIS flat-fields from the four representative filters. It is evident that UVIS flat-fields are characterized by structures that are strongly wavelength dependent:

- At the shortest wavelengths the most striking feature is the cross-hatch structure most evident in chip 1. However these variations are less then 5% peak to peak.
- Moving toward the visible regime, the flat-fields are more uniform, allowing us to note two features that are common to all the UVIS filters: the dark spot to the bottom right side of chip 2 (highlighted by the red circle in Figure 2) corresponds to a region of lower coating thickness (Wong 2010), and the diamond like area that cross the UVIS field of view from the upper left corner of chip 1 to the center of right quadrant (quadrant D) in chip 2.
- The reddest UVIS flat-fields are characterized by strong fringing. Because of the particular patterns shown by the fringing crests in the dark spot, this region has been renamed “happy bunny” (for more details about the fringing refer to Wong 2010, and Wong’s contribution to this conference).
Figure 2: Examples of ground-based UVIS flat-fields in the filters F225W (top-left), F606W (top-right), F814W (bottom-left) and F953N (bottom-right). The red circles indicate the ‘happy bunny” region, other features are also highlighted in red.

Compared to the UVIS channel, flat-fields for the IR channel are more uniform (Figure 3), with the most noticeable features being the dark circular spot at the bottom of the lower left quadrant (known as the “death star”) and the dark arc that starts from the lower right corner of the detector, and has been called the “wagon wheel”. Both regions are characterized by very low sensitivity, and pixel variable response with time. These two features are poorly flat-fielded and therefore flagged in the bad pixel table. Appropriate dithering patterns minimize the impact of these structures on the photometry.

3. Flat-field components

Experience with previous HST instruments, such as ACS and STIS, show us that:

1. Detector response can spatially change with time;

2. Ground-based flat-fields can differ from the on-orbit data because of inaccuracies of the optical stimulus in reproducing the HST OTA.

As a consequence residual structures can still be present in the data after the flat-fielding. To reduce these inaccuracies, corrections to the ground-based flats can be applied. Figure 4 shows the portion of the header of a typical WFC3/UVIS image where the calibration files used by CALWF3 to process the raw images into calibrated .FLT images are identified. Three different flat-field files are listed in the header:

1. The pixel-to-pixel flat-field (PFLTFILE) corrects for the variation in pixel response, and is the flat-field acquired on the ground during TV3;

2. The delta-field (DFLFILE) is a correction that can be applied if the response of the detector(s) changes with time;

3. The low-order flat (LFLTFILE) is the correction that removes residual variations due to the difference between the telescope and the optical stimulus optics (i.e. low frequency structure).
Figure 3: Examples of ground-based IR flat-fields in the filters F110W (top-panel), and F160W (bottom panel).
Figure 4: Portion of a WFC3/UVIS image header that shows the calibration files used by CALWF3 to remove the artifacts that affects UVIS raw images. WFC/IR image headers have a similar structure.

For consistency with the header of WFC3 images, from now on we will refer to the three different kinds of flat-fields as Pflat, ∆flat and Lflat respectively. At the moment, WFC3 images retrieved from the Hubble archive are corrected only for the pixel-to-pixel variations. In the coming sections we will discuss the accuracy of the Pflats, and how they can be improved.

3.1. Pflat accuracy

In order to investigate the accuracy of the TV3 Pflats, during SMOV4 we observed two star clusters (ω Centauri and 47 Tucanae) with both the UVIS and the IR channels in several broad-band filters. For each filter we used a 9-point dither pattern, with a dither step of ∼1/4 of the field of view.

By comparing the magnitude of stars observed through the same filter, with the same exposure time, but on different positions over the detector, we identified residual structures of the order of 4–10% in the UVIS data (depending on the wavelength, Sabbi 2009) and of ∼2 % in the IR channel (Hilbert et al. 2009). Figure 5 shows the comparison between the average magnitude of stars detected in all the images and the magnitude of the same stars on one image for the F606W UVIS filters as function of the position.

As we said before, variations in the detector response with time can introduce new structures in the flat-fields. Both the WFC3 channels are equipped with internal lamps that allow us to illuminate the WFC3 detectors through all the available filters. These “internal flats” cannot be directly used to flat-field the WFC3 data, however they are periodically acquired when the telescope is in occultation (and therefore no scientific data can be acquired) to monitor the stability of the detector responses, and eventually provide
Figure 5: Difference between the average magnitude of the stars detected in the various WFC/UVIS F606W images of ω Centauri and the magnitude of the same stars in the central pointing as a function of their position on the central pointing.
3.2. Lflats

In order to remove the residual low order structures from the WFC3 flat-fields, during cycle 17 we observed the globular cluster ω Centauri using 10 broad-band UVIS filters (namely F225W, F275W, F336W, F390W, F438W, F555W, F606W, F775W, F814W, and F850LP) and 5 broad- or medium-band IR filters (the F098M, F110W, F125W, F139M, and F160W respectively).

For each filter we acquired 9 observations following a $3 \times 3$ pattern with a dither step corresponding to a quarter of the channel field of view. As already done for ACS (Van der Marel 2003) in each filter we compared the flux of stars on different positions of the detector and fitted the difference with a Legendre polynomial to determine the low-order correction. Figures 6 and 7 show examples of field corrections for both the UVIS and the IR channels. The Lflat corrections for the remaining filters were obtained through a linear interpolation of these filters. Figure 8 shows the comparison between the average magnitude of stars detected in all the images and the magnitude of the same stars on one image, after the Lflat correction was applied, both for the UVIS/F606W and the IR/F160W filters. α releases of both the UVIS and IR full frame Pflats corrected for the low-frequency structures (LPflats) can be downloaded from the WFC3 web-page at the urls:

http://www.stsci.edu/hst/wfc3/analysis/uvis_flats and
http://www.stsci.edu/hst/wfc3/analysis/ir_flats

Figure 6: Examples of UVIS channel Lflat corrections for the filters F275W, F336W, F438W, F606W, and F814W, derived from the observations of the globular cluster ω Centauri. Corrections are in magnitudes. Minimum and maximum values are given.

the necessary Δflat corrections. No significant variations have been identified so far (Rajan & Baggett, 2010; Hilbert, 2009).
Figure 7: Examples of IR channel Lflat corrections for the filters F110W, and F160W, derived from the observations of the globular cluster ω Centauri. Corrections are in magnitudes. Minimum and maximum values are given.

4. Summary and Conclusions

Flat fields for both the WFC3 channels were acquired from the ground during the last thermal vacuum campaign (TV3) at the beginning of 2008. Comparisons between ground-based internal flats and internal flats acquired from the sky indicate that the WFC3 detector responses have been stable over the first year of operations.

Observations of globular clusters with large dither steps show that even after the flat field is applied the flux of a source still depends on its position on the detector. These residual structures can be corrected to better than 1% by applying Lflat derived from the same observations. α releases of Lflats for both the WFC3 channels can be downloaded from the two following urls:

http://www.stsci.edu/hst/wfc3/analysis/uvis_flats and
http://www.stsci.edu/hst/wfc3/analysis/ir_flats

After one year of operations, several IR channel deep exposures characterized by a low level of star crowding to accumulate in the HST archive. The WFC3 team is stacking together these images to make external flat-fields from the sky (see Pirzkal contribution for more details). Observations of the Earth and of the Moon will also be acquired between the summer and fall of 2010 to make other external flats for both the channels.

References

Figure 8: Difference between the average magnitude of the stars detected in the nine images of ω Centauri and the magnitude of the same stars in the central pointing as a function of their position on the central pointing for the F606W filter after the Lflat correction was applied.
Figure 9: Same as Figure 8, but for the filter F160W of the IR channel.
Sabbi, E. 2009, WFC3-ISR 2009-19, “WFC3 SMOV Program 11452: UVIS Flat Field Uniformity”
van der Marel, R. 2003, ACS-ISR 2003-10: “Determination of Low-Frequency Flat-field Structure from Photometry of Stellar Fields”
Fringing in the WFC3/UVIS detector

Michael H. Wong

Astronomy Department, University of California, Berkeley CA 94720-3411

Abstract. In late 2010, a star cluster will be observed with narrowband red filters to determine the impact of fringing (position- and wavelength-dependent patterns of brightness variation) on flight data. Ground flat fields show peak-to-trough amplitude variations of 0.5% to 16% (among 12 affected filters). Different “fringe flat fields” created by models will be applied to the flight data to compare corrections. Two separate sets of ground test data provide different thickness maps (and thus create different fringe flats). The disagreement between the thickness maps is consistent with an error in the monochromatic illumination wavelengths in one of the ground test data sets, but an unexplained corner-to-corner slope across the detector remains even after correcting for the wavelength error. Flight data are needed to determine which thickness map produces the best correction for fringing, and improvements to fringe data analysis may be needed to improve the correction.

1. Introduction

Fringing is caused when multiple internal reflections within a detector lead to constructive and destructive interference. “Wood grain” patterns appear because the phase of the interference is strongly sensitive to detector thickness and wavelength of illumination. In WFC3’s UVIS channel, fringing is detectable only at long wavelengths, where silicon grows transparent enough that some incident photons can reflect off detector layer interfaces, before these photons are absorbed. Fringing is a common issue with astronomical CCD detectors such as WFC3/UVIS, in which the silicon substrates are thinned so that photoelectrons are generated close to the gate structures. Thinning improves detection efficiency at short wavelengths, but leads to more severe fringing effects at longer wavelengths.

In Section 2 I use long-wavelength narrowband UVIS flat fields to estimate the impact of fringing. Section 3 discusses two separate ground test data sets, obtained by illuminating the UVIS detectors with monochromatic light sources. These two data sets are being used to characterize the detector. These data are crucial inputs to models (Section 4) of the UVIS fringing effect. These models produce maps of the detector’s thickness across the entire CCD (Section 5). Reconciling the two ground test data sets has been hampered by an apparent error in the wavelength calibration of one of the data sets (Section 6). On orbit data are now being collected to compare fringe model solutions, and additional tricks may be required to achieve the best correction (Section 7).

2. Estimated Impact of Fringing

Figure 1 shows the effect of fringing by comparing flat fields for two filters in a single quadrant of the UVIS detector: F673N (left), in which fringing is not apparent, and FQ937N (right), which shows strong fringing effects due to the filter’s slightly narrower bandpass and much longer central wavelength.
Figure 1: Quadrant B of two ground flat fields: affected by fringing (right) and not affected (left). Black region of the FQ937N flat is masked to avoid areas affected by the quad filter edges; the same region of F673N is masked for consistency. Image scale runs from 0.85 (black) to 1.10 (white), with a square root stretch to emphasize low-level detail.

Figure 2: Histograms of the two flat field samples shown in Figure 1. Symbols correspond to fringe amplitude metrics listed in Table 1: rms deviation (triangles with error bars), full width at 20% maximum (circles with error bars), and bimodal histogram peaks (squares).
Analysis of these flat fields provides a generic estimate of the effect of fringing on observations using the narrowband red filters. This approach relies on characterizing histograms of pixel values. Four different metrics are listed in Table 1 and graphically presented in Figure 2. The root-mean-square deviation from the mean (triangles in Figure 2) and the full width at 20% maximum (circles) are the most general metrics. The rms deviation may be most useful in estimating photometric uncertainties due to fringing in on-orbit data, while the full width at 20% maximum may be more useful for exposure time calculations. In filters where fringing is very strong, two additional metrics are presented. The distance between bimodal histogram peaks (squares) may also be useful for exposure time calculations, while the manually measured variation in adjacent fringes on the detector (last column of Table 1) gives a sense of the maximum error that may be introduced by dithers of the same scale as the fringe patterns themselves. Values for a single type of fringing metric can be compared to assess the relative impact of fringing on the filters and quadrants listed. An entry for F606W is also listed as a reference case that is not affected by fringing.

Table 1: Estimates of fringing effects on WFC3/UVIS data, based on analysis of ground flat field data (corrects an error in filter names found in Wong, 2010b).

<table>
<thead>
<tr>
<th>Filter</th>
<th>Quadrant</th>
<th>rms deviation</th>
<th>Full width at 20% maximum</th>
<th>Distance between histogram peaks</th>
<th>Manual peak-to-trough</th>
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<td></td>
<td></td>
<td>(percent)</td>
<td>(percent)</td>
<td>(percent)</td>
<td>(percent)</td>
</tr>
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<td>A</td>
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<td>–</td>
<td>–</td>
</tr>
<tr>
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<td>3.0</td>
<td>–</td>
<td>–</td>
</tr>
<tr>
<td>F606W</td>
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<td>3.3</td>
<td>–</td>
<td>–</td>
</tr>
<tr>
<td>F606W</td>
<td>D</td>
<td>1.2</td>
<td>3.3</td>
<td>–</td>
<td>–</td>
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<tr>
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<td>A</td>
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<td>4.9</td>
<td>–</td>
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<td>4.5</td>
<td>–</td>
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<td>5.1</td>
<td>–</td>
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</table>
In Wong (2010b) the analysis of flat field data to characterize fringing is described in greater detail, along with images and histograms for every narrowband long-wavelength filter. The flat fields themselves are described in Sabbi et al. (2009).

Science target pixel distributions will differ from those in the flat fields if the spectral energy distribution (SED) of the target differs significantly from that of the calibration light source. Sources with narrow SEDs may be affected by fringing even if they are observed using broad-band filters. Sources with SEDs very similar to the calibration source will have most of the effects of fringing removed by flatfield division (as part of the standard calibration pipeline), but sources with significantly different SEDs may actually have additional error introduced by the pipeline. Section 7 discusses “fringe flats” which can be used to correct for fringing, if the SED of the target is known. Fringe flats can be created based on modeling of extensive sets of ground test data taken for this purpose.

3. The Ground Test Data

Two sets of ground test data were taken to measure fringe patterns as a function of wavelength in the UVIS detectors. The first set of tests were conducted at NASA Goddard Space Flight Center’s Detector Characterization Laboratory (DCL) in 2001. As part of the WFC3 Thermal Vacuum Test 3 (TV3) in 2008, a second set of data were acquired. Differences in experimental conditions between the DCL and TV3 tests are described in greater detail in Wong and Malumuth (2011), and the data are archived as described in Wong (2010a).

Three key differences distinguish the two data sets. The DCL data set contains about twice the number of frames as the TV3 set (~150 rather than 77), spanning a wider range of wavelengths. The detectors were illuminated at normal incidence in the DCL configuration, and at a flight-like incidence angle of 21° in the TV3 tests.

Figure 3 compares the ground test data wavelength range with the bandpasses of UVIS’s narrowband long-wavelength filters. Although the TV3 data span the wavelength range that is most strongly affected by fringing (longward of 880 nm; see Table 1), the DCL data have better coverage at the shortest wavelengths. Originally, it was thought that effects of fringing would be limited to wavelengths longer than about 750 nm (Bond and Kim Quijano, 2007). However, ground flats revealed that the cutoff fell closer to 600 nm (Sabbi 2008), with F656N being the shortest-wavelength filter showing visible signs of
4. Modeling Fringing

Malumuth et al. (2003a) developed a thin-film analytical model of the STIS CCD to predict and correct fringing effects. The code solves the Fresnel equations, modeling the detector as a series of layers, each characterized by a thickness and a complex index of refraction. The roughness of the interface between each pair of layers is also parameterized. The basic approach is to take a set of calibration data, using the model to find parameter sets that best simulate the data. These parameter sets can then be used to simulate any arbitrary data frame and to correct for fringing. Malumuth et al. (2003b) describes the adaptation of the STIS CCD fringing model to the WFC3/UVIS detector, and the technical details of the parameter fitting approach used in this work is described in Wong (2010a) and Wong and Malumuth (2011). Results of the UVIS modeling effort to date are described in Wong and Malumuth (2011).

Although the CCD assembly is modeled as a system of seven layers (eight including the semi-infinite vacuum above the detector stack) with seven interfaces, the thickness of the silicon detection layer has the greatest effect on fringing. Fits were tested to the TV3 and DCL calibration data for 8000 values of the detection layer thickness, in steps of 0.5 nm. Six sets of secondary parameter values (roughnesses and non-detection layer thicknesses) were also tested. Once the calibration data are used to find best fit model parameters, the parameters can be used to simulate fringing at intermediate wavelengths, or for an arbitrary SED.

Figure 4 compares the two data sets and four model calculations. The model calculations used best-fit parameters for either the DCL or TV3 data sets. The plot shows fringe
Figure 5: CCD detection layer thickness map based on TV3 data alone. For Chip 1, the median and standard deviation of thickness is 16.04 µm ± 0.23 µm, with 1% of pixels thinner than 15.51 µm and 1% thicker than 16.46 µm. For Chip 2, the thickness is 15.42 µm ± 0.58 µm, with 1% thinner than 13.87 µm and 1% thicker than 16.19 µm.

Figure 6: Same as Fig. 5, but for wavelength-corrected DCL data. For Chip 1, the median and standard deviation of thickness is 16.33 µm ± 0.23 µm, with 1% of pixels thinner than 15.81 µm and 1% thicker than 16.75 µm. For Chip 2, the thickness is 15.70 µm ± 0.58 µm, with 1% thinner than 14.15 µm and 1% thicker than 16.64 µm.
amplitude as a function of wavelength, for a single representative pixel. Similar results are seen for other pixels. The solid curves indicate that the TV3 (black) and DCL (red) data are well-fit by the models. The DCL data have already been corrected for a wavelength error described in Section 6.

Unfortunately, consistency between the DCL and TV3 data sets has not yet been achieved. The dashed lines in Figure 4 present models based on one set of retrieved parameters (e.g., the dashed red curve is based on TV3 data) but with fringe amplitudes appropriate for the incidence angle of the other data set (e.g., the dashed red curve is model output for 0° incidence as appropriate for comparison with DCL data). The disagreement between the dashed curves and the data points of the same color shows that even with the DCL wavelength correction, the two data sets do not agree. This disagreement is even more severe without the wavelength correction.

5. Thickness Solutions

Detection layer thickness maps are shown in Figures 5 and 6, as retrieved from the TV3 and DCL data sets respectively. Retrievals based on the two data sets result in slightly different thickness maps. Because fringing is so sensitive to detection layer thickness, models using the two thickness maps will predict very different fringing behavior. On-orbit cluster observations are being conducted in filters sensitive to fringing, to determine the best input parameters (thickness map, and secondary parameters map) for the model.

Section 6 describes a correction to the DCL wavelength calibration. With this correction (already applied to the DCL data used to retrieve the Figure 6 thicknesses), the model retrieves similar thickness maps using the TV3 and DCL data sets. Without the wavelength correction, thickness maps derived using DCL data are about 1 µm thicker than
Figure 8: TV3 test data (shaded squares with error bars) and model fringe amplitude (black circles). Data are average values for three frames at 910.00 nm, with error bars showing the standard deviation of normalized signal level among the three frames. Data are shown for column 3000 of Chip 1. Model amplitudes are based on the TV3-derived thickness solution shown in Fig. 5.

Fringe data for a single wavelength is compared to model fringe amplitude in Figure 8. The error bars in the data show frame-to-frame variation among three frames, all illuminated at the same nominal wavelength. The model fringe amplitude data provide an excellent fit within these error bars, but highlight a problem with the fringing solution: that intrinsic noise in the test data limit the accuracy of the retrieved fringe model parameters.

6. The Wavelength Error

Thickness maps retrieved using TV3-only data and DCL-only data differed by about 1 µm. In Wong and Malumuth (2011), this discrepancy was shown to be partly due to an error in the wavelength calibration of the DCL data. After this correction, the difference between thickness maps based on the two test data sets was reduced to less than 0.3 µm (Figure 7).

Measurement of the wavelength error was done in two steps. First, fringe model output for the entire spectral range was calculated using the TV3-derived thickness map. Then, for each monochromatic frame of DCL data, an optimal wavelength was found, such that the residuals were minimized between the test data and the model output. Figure 9 shows this optimal wavelength, plotted as a function of the commanded wavelength in the DCL test data set. A strong systematic relationship is evident.

The ratio of optimal to commanded wavelength is plotted in Figure 10. The mean ratio between 700 and 1000 nm is 0.972 ± 0.003.

The distribution of data points in the figure is well understood in terms of order errors. Order errors result because fringe amplitude, as a function of either wavelength or detection layer thickness, is oscillatory. Fringe amplitude (and thus residuals between data and models) therefore can be very similar for nearby values of wavelength that differ by one order. When adjacent orders of fringing produce very similar residual minima, small amounts of noise in the data can cause automated minimization methods to select one residual minimum over the other in a random fashion. This explains the behavior in Figure 10 near the commanded DCL wavelength of 840 nm, where frame-by-frame values of the optimal/commanded wavelength ratio (black points) are distributed above and below the
Figure 9: Optimum (y-axis) vs. commanded (x-axis) wavelengths for Chip 2 DCL data frames.

Figure 10: Optimum scale factor vs. commanded wavelength for all the Chip 2 DCL data. Red line shows the mean scale factor of 0.972, with a standard deviation of 0.003.
mean ratio of 0.972 (red line), but do not lie exactly at 0.972. A single wavelength scale factor may be an oversimplification of the problem, because at this wavelength, the optimal wavelength actually corresponds to a local maximum in the fit residuals; two adjacent minima are randomly selected from frame to frame here. Over the entire wavelength range, however, a single wavelength scale factor provides a good fit, because frame-by-frame values oscillate around the mean value of 0.972. Section 7 (and Wong and Malumuth, 2011) discusses possibilities for further refining corrections that may make it possible to combine the two data sets to find a common solution.

7. Future Work

Improvements to the understanding and correction of fringing effects in the WFC3/UVIS detector can be made in four areas: understanding and correcting discrepancies between the TV3 and DCL data sets, using flat field data as input to the fringing model, using on-orbit cluster observations to evaluate various fringe model solutions, and developing fringe-flat software to allow users to correct their data for fringing effects.

Although it is demonstrated above that the wavelength correction to the DCL data considerably improves the agreement between the TV3-derived and DCL-derived thickness maps, three clues show that the correction is not perfect:

• Fringe amplitude, but not phase, is better fit by the DCL data (at wavelengths shorter than about 850 nm) without the wavelength correction (Wong and Malumuth 2011).

• The mean difference between retrieved thicknesses is non-zero (Figure 7).

• There is a gradient in the absolute thickness difference map, increasing from lower right to upper left in both chips (Figure 7).

Despite the ∼70% better agreement between thickness maps after the wavelength correction, the first clue suggests that no wavelength correction should be applied. It is important to recognize that fringe amplitude is also affected by the secondary parameters (Malumuth et al. 2003b; Wong 2010a). Parameter set maps derived from TV3 data are much closer to uniformly distributed across the six secondary parameter set values, whereas the DCL parameter set maps are dominated by one of the six possible values, again suggesting that the wavelength correction is inconsistent with the fringe amplitudes in the data. The second clue indicates that more work must be done before the two data sets can be combined, but provides little obvious insight.

The third clue makes an interesting parallel with the pattern seen in low-frequency flat field corrections derived from on-orbit data (Rajan 2010, Sabbi et al. 2011), which also show diagonal gradients across both chips. The orientation of Chip 2 was reversed in the DCL tests (Wong and Malumuth 2011), so if the gradient were due to a problem with the DCL data, then it should have a reversed orientation between the two chips. Because the orientation of the gradient is the same in Chip 1 and Chip 2, the problem must lie with the TV3 tests instead. This problem could have a common origin with the gradient in the low-frequency flat fields, and may be related to window ghosts from the TV3 illumination source. If so, it may be possible to apply the low-frequency flat correction, or window ghost correction, to the TV3 data before solving for detector thickness (and secondary parameters). This approach could provide new results, including an improved wavelength correction for the DCL data.

The entire motivation for including both DCL and TV3 data is to cover a larger wavelength range, and thus to enable fringing correction for shorter-wavelength filters such as F656N. If this is not possible, the TV3 flat field data themselves could be used instead. Monochromatic data are preferable, because they create larger fringe amplitudes (and
thus better signal to noise ratios), but a trade for wavelength coverage may make it worthwhile to consider including flat field data as inputs to the fringing model. Unfortunately, this approach is hampered by the low amplitude of fringing at short wavelengths.

Two Cycle 17 calibration proposals (IDs 11922 and 12091) are being executed to observe the Omega Centauri star cluster in narrow-band red filters affected by fringing. These observations will provide the ultimate test of various fringe corrections, as produced by the Malumuth et al. (2003b) model using parameter sets derived from TV3 and DCL data, or additional parameter sets created after further refinements to the analysis.

Ultimately, the parameter set that best corrects the on-orbit fringe test data will be used to generate fringe flat fields, which can correct science data. Because fringing is a strong function of wavelength, users will need to provide SEDs to the model in order to get fringe flats appropriate for the observations. This includes observations of narrow-band red sources observed through wide filters, because such sources will also be affected by fringing. Modeling fringing in WFC3/UVIS is still at an early stage, with the best correction not yet defined, so work on software to create fringe flats has not yet begun.

8. Conclusions

As measured in ground flat field data, fringing affects narrow-band red filters at the 0.5–16% level. Although pipeline flat fields may correct for most of the photometric error associated with fringing, sources with SEDs significantly different from the calibration source will not be properly corrected. Two monochromatic data sets (DCL and TV3) were used to retrieve detection layer thicknesses, which can be used to model fringing for any desired SED.

Although thickness solutions for the two data sets disagree, the solutions are more consistent after a correction is made to the wavelength of illumination in the DCL data set by decreasing the wavelength of illumination by a factor of 0.972 with respect to the commanded wavelength. This correction reduces the difference in retrieved detection layer thickness from about 1 µm to about 0.28 µm. However, a gradient in the retrieved thickness difference between the two data sets, diagonally across each chip, implies that further corrections must be made to the TV3 data before retrieving detection layer thickness maps. The gradient resembles the gradient in the low-frequency flat field correction from on-orbit calibration data (Sabbi et al. 2011), and may result from window ghosts created by the illumination source. If further analysis is unable to reconcile the two calibration data sets, then it may be useful to include flat field data as a supplementary source of input data to the model, because the TV3 tests did not extend to short enough wavelengths to reliably correct filters at wavelengths less than 850 nm. The use of ground flat data as input data for the fringing model is limited by the degree to which the spectrum of the calibration source is known.

On-orbit cluster observations currently being collected will be the ultimate test of various fringe model parameter sets. Once the optimal parameter set is determined, fringe flats can be generated from the model for any arbitrary SED. Fringe modeling is the only way to correct science data in cases where the SED differs significantly from the calibration source, such as line emission data, or data with strong spectral slopes across filter bandpasses.

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WFC3 IR ”Blobs”, IR Sky Flats and the measured IR background levels

N. Pirzkal

Space Telescope Science Institute, Baltimore, MD, 21210

Abstract. The near infrared background, as seen using the WFC3 IR channel, varies by several orders of magnitude. Ever since WFC3 was installed on HST, we have been monitoring and examining incoming observations of sparse fields. This monitoring has enabled us to identify the WFC3 IR ”Blobs” (Pirzkal et al., ISR 2010-06). It has also allowed us, for a few broad band filters, to assemble IR delta sky flats. We have also been able to measure the observed IR background levels in many of the WFC3 IR filters. Here, we show how we have used WFC3 observations to construct the delta sky flats as well as how we detected and created maps of IR ”Blobs”. We finally show how the measurements of the IR background vary and how they correlate with HST pointing and Zodiacal light levels.

1. Introduction

WFC3 was installed on board of HST during SM4 in May 2009. Following its installation, some of the images obtained using its IR channel have been affected by a small number of blemishes that appear to have 10–15% lower count rates than the surrounding areas. These ‘Blobs’ were not seen or detected during ground testing and have been progressively appearing in IR images obtained after July 2009. They are most visible in images containing a large uniform object, or alternatively images with high background levels. Over the last few months, we have systematically examined all available WFC3 IR data to detect and monitor the Blobs. We have been able to determine their likely cause, and their physical properties, such as positions, sizes, and depths. We have also been able to determine the rate at which new Blobs appear, and the stability of a Blob once it has appeared. Blobs and their properties are discussed in Section 3. The method we used to monitor WFC3 for IR Blobs consists of combining as many masked WFC3 images as possible. Thus, in the course of this project, we were able to also assemble high signal to noise sky flats in several medium and broad bands WFC3 filters. We discuss the properties of these sky flats in Section 4. Since we had to estimate the median IR sky background of individual WFC3 IR images before combining them to generate sky flats and look for IR blobs, we are in a position to report on the direct measurements of the IR sky background as seen using WFC3. The WFC3 IR background levels are discussed in Section 5.

2. Assembling the data

All of the data discussed here were calibrated by the STScI archive pipeline. They were thus calibrated using the current best calibration products such as flat fields and darks. We first identified a series of WFC3 Proposals observing fields that were neither too crowded nor contained large extended objects and with integration times larger than 300 seconds. The latter requirement was imposed to insure that the IR background would generate at least a few hundreds counts in these images. We first ran Sextractor to locate sources aggressively in each image. The Sextractor object segmentation map was used to create an object mask...
which was then grown (to further mask out the faint edges of sources). We additionally masked pixels likely to be affected by IR persistence by examining all WFC3 IR images taken in the preceding 24hr period and by masking out any pixel that were filled with more than 30,000 e-. A typical science image and its mask are shown in Figure 1.

3. IR Blobs

The first instance of a Blob was detected in data taken in July 2009, shortly after launch and the installation of WFC3 on HST. A round, darker area was readily visible in the pipeline processed FLT files of a few GO datasets. These were not detected in any of the associated and contemporary bias, dark or flat-field calibration files. Further examination of all available WFC3 IR data showed that this artifact likely appeared between July 19, 2009 and August 6, 2009. Another Blob was then observed to have appeared some time between August 6, 2009 and August 8, 2009. In both instances, once a Blob appeared, it remained detectable in all subsequent observations as long as these were deep enough to contain a significant number of IR background photons. In practical terms, this means that Blobs are visible in long exposures, with more than 300 seconds, broad band filter images. In no cases, and this remains true to this day, has a Blob appeared and then either disappeared, changed its appearance or physically moved in a significant way. As we discuss below, the sizes and position of Blobs have not being changing. The one notable exception, data where the Blobs appeared to have shifted, allowed us to infer that Blobs were not physically located on the WFC3 IR detector itself but were instead caused by the Channel Select Mechanism (CSM). The CSM mirror diverts incoming light into the WFC3 IR channel. The CSM moves the mirror out of the way when the UVIS side of WFC3 is used. As long as the CSM settles in the exact same commanded location, the Blobs appear at the same physical location on the detector. When the CSM is positioned differently, the Blobs are seen to shift in the bottom-left to top-right direction. In most cases, the CSM is always commanded to be at the same position and the Blobs hence appear always at the same location in the final images. Figure 2 shows a few Blobs, easily visible in a deep science observations, marked using red circles.

As discussed more thoroughly in the WFC3 ISR 2010-06, Blobs are best detected on images with count rates of at least a few 100's of e-/pixel. There are very few fields that have been observed by WFC3 that are sparse enough to provide a uniform illumination of the detector. We instead detected and monitored Blobs by combining many deep WFC3 IR observations. While some of the brighter Blobs are visible in deep single images, as shown in Figure 2, many more are visible in the monthly combined images that we produced, as we demonstrate in Figure 3 where we show F160W sky flat that we generated for the month of February 2010. Such images have allowed us to both detect and monitor IR Blobs on a monthly basis.

The physical properties of these Blobs are described in details in ISR 2010-06. To summarize, we identified 19 bright Blobs in the WFC3 field of view and this number appears to have now stabilized and is no longer increasing, even though the number of Blobs was initially seen to increase quickly immediately after launch. The central region of the larger Blobs absorb as much as 15–20% of the incoming light. Their median radius is on the order of $\approx 5$ pixel and shows a small color variation between the F125W and F160W filters.

A full map of these Blobs is currently incorporated in the Data Quality (DQ) array of all FLT files obtained from the STScI MAST archive. A bit value of 512 is used to mark the location of these Blobs. The current DQ mask is shown in Figure 3. Currently, only $\approx 0.4\%$ of WFC3 pixels are affected, less than a quarter of all pixels marked in the DQ map (1.2%). The effect of the blobs can be mitigated by taken judiciously dithered images. Proper dithering strategies for WFC3 are discussed in details in ISR 2010-09.
4. Delta sky flats

The Blob monitoring effort has required us to combine many WFC3 images in order to generate deep images of the background, against which the effect of the Blobs could be detected, measured and monitored, and this using different WFC3 broad band filters.

Now that WFC3 has been in full operation for several months, we are in a position to combine a significant number of WFC3 science observations (e.g. Figure 1) and produce deep, high signal to noise delta sky flats. The delta sky flats that we discuss here are our estimate of the additional flat fielding correction that is required to make a uniform sky background appear flat in pipeline calibrated data. These were generated using data obtained from the MAST archive and were hence already flat-fielded using the current pipeline flats (Obtained during ground testing of WFC3 and prior to launch). As shown in Figure 4, we currently see as much as 3% residuals in the flat fielding of WFC3 IR data that are currently being flat fielded using the current, ground based, pipeline flats. Delta sky flats are now being assembled for several filters (F098M, F105W, F125W, and F160W). Preliminary analysis of delta sky flats obtained using different broad band filters show that the same pattern exists in all delta sky flats. We also measure very little variation from one filter to the next, as shown in Figure 4 where we show the fractional difference between the F125W and F160W data. As shown there, the delta sky flats agree to within 0.5%. We expect delta sky flats to be made available to the community in the Fall of 2010, together with an ISR describing in details how they were created and their relative merits.

5. The IR background

As described above, the creation of IR delta sky flats, initially undertaken to look for IR Blobs, has required us to measure the background level of individual IR image. This was necessitated by the fact that the IR background, as seen using WFC3, varies greatly as a function of position in the ecliptic plane and also as a function of how far the bright Earth limb is with respect to the line of sight. During this exercise, we thus ended up measuring, individually, the IR sky levels over a long period of time and a wide range of HST pointings, and using several filters.

As shown in Figure 5, the first result of this is that we were able to verify, using a large number of observations, that the background estimates from the ETC are close to reality. While we in fact see a large variation of the background levels (by a factor of 2–3), as illustrated in Figure 6, the range of observed values match well the range of background levels returned by the ETC for low, average, and high zodiacal light levels. In principle, as more data are obtained using the WFC3 IR channel, we will be able to map out the IR background level in several filters over a wide range of Helio ecliptic latitudes and longitudes. Figure 7 shows a preliminary plot showing the measured background levels at ecliptic latitudes of 20, 40, 50, and 80 degrees and as a function of helio ecliptic longitude. Note how a slightly lower background is measured in the anti Sun direction (180 degree) at increasing ecliptic latitudes, and how, for a given latitude, the background is observed to vary by a factor of $\approx 2$ (we have attempted to exclude any observations taken too close to the Bright Earth Limb, which would cause the background to be another factor of several higher). This work is ongoing as more WFC3 data are been accumulated.

6. Conclusion

The appearance of IR Blobs in WFC3 IR images has prompted us to continuously monitor incoming WFC3 IR data. This process has required us to regularly measure the background sky level in WFC3 IR observations and to assemble deep IR sky flats using several broad and medium band filters. We showed here how the IR Blobs location is now known and
how their number has stabilized. The IR Blobs affect a relatively small number of pixels in the field of view and can be currently dealt with with proper dithering strategies. The sky flats we assembled demonstrated that the current pipeline flat-fielding, performed using flat-fields obtained during ground testing of WFC3 and prior to launch, leaves flat-fielding residuals that are on the order of ±3%. Deeper, higher signal to noise sky flat-fields continue to be assembled as more WFC3 data is accumulated. Finally, we have shown here how the IR background level is seen to vary significantly and is highly correlated with the pointing of HST in the Ecliptic plane. Current data however confirm that the ETC estimates are in reasonable agreement with observations. In the future, we expect to be able to provide better estimates of the IR background as a continuous function of Helio ecliptic longitude and latitude, as we continue to measure the background in more IR observations.
Figure 2: A deep WFC3 IR science exposure with Blobs marked using red circles.
Figure 3: Left: A WFC3 F160W IR delta sky flat created using observations from February 2010. While some of the brighter Blobs are visible in deep single images, as shown in Figure 2, many more are visible in monthly delta sky flats generated by combining between 50 and 100 individual exposures. Right: The current Data Quality map, showing the location of the Blobs in the WFC3 field of view. While appearing large visually, IR blobs affect less than $\approx 0.4\%$ of all of the WFC3 pixels and only slightly increase the number of otherwise flagged as bad pixels.
Figure 4: Top Row: delta sky flats generated using several hundred WFC3 images. F125W is shown on the left and F160W on the right. In both of these delta sky flats, a $\pm 3\%$ variation can be seen. We see very little difference between the F125W and the F160W delta flats, as shown on the bottom left panel. There, we show the fractional difference between the two delta sky flats shown above. The F125W and F160W appear to be within 0.5\% of one another over most of the field.
Figure 5: IR background levels as measured using WFC3, as a function of filter used. The vertical blue bars show the full range of IR background measured in each filter. The blue dot shown the median background value in each filter. The red, green, and yellow dots show the predicted values using the current Exposure Time Calculator (ETC), in cases of low, average, and high zodiacal light levels, respectively.
Figure 6: Two typical WFC3 F160W observations, shown using the same color stretch. The image on the left, has a high, uniform background of 0.7 electrons/pixel. The image on the right has a much lower background level of 0.37 electrons/pixels. The difference in background level is caused by scattered Zodiacal light. The image on the left was taken while pointing at a field with a Helio ecliptic longitude that was less than 50 degrees. The image on the right was taken at a Helio ecliptic longitude of 255 degrees, and hence away from both the Sun and any back scattered light originating in the anti-Sun direction.
Figure 7: The IR background levels, as measured from F160W observations, as a function of helio galactic longitude. This is our first, preliminary map, of the IR zodical light in this filter, shown at different ecliptic latitudes (blue: 20 degrees, green: 40 degrees, red: 60 degrees, light blue: 80 degrees).
HST/WFC3 In-Orbit Grism Performance

H. Kuntschner, M. Kümmel, and J. R. Walsh

Space Telescope European Coordinating Facility, European Southern Observatory, Karl-Schwarzschild-Str. 2, 85748 Garching, Germany

H. Bushouse

Space Telescope Science Institute, 3700 San Martin Drive, Baltimore, MD 21218, USA

Abstract.

The Wide Field Camera 3 (WFC3) is fitted with three grisms for slitless spectroscopy. In the UVIS channel there is one grism, G280, for the near-UV to visible range (200 - 400nm; 1.4nm/pix). The IR channel has two grisms: G102 for the shorter (800-1150nm; 2.45nm/pix) and G141 for the longer (1100-1700nm; 4.65nm/pix) NIR wavelengths. Using SMOV and Cycle 17 calibration data taken on WR stars, planetary nebulae and flux standard stars, we have assessed the performance of the grisms. We have measured the field-dependent trace locations and dispersion solutions and determined the throughputs. The trace and wavelength solutions for the IR grisms were found to be linear functions, varying smoothly across the field-of-view. The UVIS grism exhibits a highly bent trace and significantly non-linear dispersion solutions. The maximum throughputs for the G102 and G141 grisms, including the telescope optics, are 41% at 1100 nm and 48% at 1450 nm, respectively. Limiting magnitudes at S/N=5 and a 1h exposure are $J_{AB}=22.6$ and $H_{AB}=22.9$ for the G102 and G141 grisms, respectively. The results are published as sensitivity and configuration files that can be used with the dedicated extraction software aXe to reduce WFC3 slitless data.

1. Introduction

The Wide Field Camera 3 (WFC3) is fitted with three grisms for slitless spectroscopy. In the UVIS channel there is one grism, G280, for the near-UV to visible range (200 - 400nm). The NIR channel has two grisms, G102 and G141, for the shorter (800 - 1150nm) and longer NIR wavelengths (1100-1700nm), respectively.

There are several noteworthy areas where slitless spectroscopy differs from slit spectroscopy, and an overview can be found in Walsh et al. (2010, this volume). Fig. 1 presents a visual impression of grism spectra for point sources in the three WFC3 grism modes. Prior to launch, calibrations of the trace and wavelength solutions as function of target position within the field-of-view (hereafter FoV) and also throughput measurements were determined in three thermal vacuum (TV) ground calibration campaigns. The final pre-flight calibration results are summarized in Kuntschner et al. (2008a,b; 2009a).

In order to ease the process of WFC3 grism data reduction, a dedicated, semi-automatic spectral extraction software package, called aXe (Kümmel et al. 2009; see also Kümmel et al. 2010, this volume) is available to extract, flat-field, wavelength- and flux-calibrate spectra. The spectral trace and dispersion solutions are a function of source position within the FoV (see also Secs. 2.1. & 2.2.). These 2-dimensional variations can be well characterized and the resulting reference and calibration files are used in the extraction software aXe.
Figure 1: **Top:** Appearance of the UVIS G280 grism spectral orders on the detector with a F300X direct image (circled) superimposed to illustrate the relative positions. The stronger 1st order is to the left and the 0th order is in the center. Higher spectral orders are strongly overlapping and occupy the full extent of the detector. The image shows the full extent of the detector in the x-axis (4096 pixels) and about 500 pixels in the y-axis. **Middle:** G102 grism observation of the flux standard star GD153 (program 11552) with a F098M direct image (circled) superimposed to illustrate the relative positions. Spectral orders 0, +1, and +2 can be seen on the image. The image shows the full extent of the detector in the x-axis (1014 pixels) and about 200 pixels in the y-axis. **Bottom:** G141 grism observation of the flux standard star GD153 (program 11552) with a F140W direct image (circled) superimposed to illustrate the relative positions. Spectral orders 0, +1, +2, and +3 can be seen in the image. The image shows the full extent of the detector in the x-axis (1014 pixels) and about 200 pixels in the y-axis.

This paper presents updated in-orbit calibrations for the grisms derived from SMOV and selected Cycle 17 calibration observations. An early analysis of the data was described in Kuntschner et al. (2009b,c). Using data taken on WR stars, planetary nebulae and flux standard stars, we have assessed the performance of the grisms. We have measured the field-dependent trace locations and dispersion solutions and determined the throughputs for the near-IR grisms. All data shown in this paper were processed with HST pipeline software *calwf3* version 2.0 and reference files available as of May 20, 2010.

2. The near-IR grisms

2.1. Trace calibration

In order to establish a good in-orbit trace (the location of the center of gravity in the spatial direction) calibration, we utilize the calibration observations of the wavelength standard planetary nebulae (PN) HB12. For the trace calibration, we are not interested in the primary target itself but in the other point sources around the primary target. This field (Galactic latitude: −02.85 degrees) comprises a relatively good compromise between field coverage and object density and thus avoiding too much spectral overlap between different sources in the FoV (see Fig. 2). The ST-ECF aXe software package for the reduction of slitless spectroscopy data treats the spectral traces and wavelength solutions as defined with respect to the position of the source in the direct image. The centroids of all sources above a given threshold in the F098M images \((X_{\text{ref}}, Y_{\text{ref}})\) were determined with SEExtractor (Bertin & Arnouts 1996). These positions are assumed not to change between the observations of each direct image – grism image pair. The spectra of all sources were traced as a function of \(\Delta X = X - X_{\text{ref}}\) in the detector X-direction by measuring the centroids of 7-10 pixel wide bins, using custom-written IDL programs. We found the traces of all orders (excluding the
Figure 2: Image of a G102 grism exposure (iab907j6qflt.fits). The primary target HB12 is located in the centre. Most of the other prominent spectra are first orders of different (stellar) objects in the FoV that were used for the trace calibration.

The zeroth order) to be well fit by straight lines with standard deviations of 0.1 pixels. The trace definitions are of the form

\[ (Y - Y_{\text{ref}}) = DYDX_{0} + DYDX_{1} \cdot \Delta X, \]  

where DYDX_{0} and DYDX_{1} are field dependent and given in the usual format used by the ST-ECF aXe reduction package, e.g. DYDX_{1} = a0 + a1*X_{ref} + a2*Y_{ref} + E (see also the aXe manual\(^1\) for more details).

The final, field-dependent trace solution for the +1\(^{st}\) order of the near-IR grisms was derived from 137 (G102) and 162 (G141) different traces covering semi-uniformly the FoV. A graphical representation of the measured offsets and slopes for the G102 grism as a function of X_{ref} and Y_{ref} positions is given in Fig. 3. The offset shows a marked trend with Y_{ref} position while the slope shows a more complex dependence on X_{ref} and Y_{ref}. The offset is well represented with a field dependent fit using only a linear function (rms = 0.07 pixel), while the slope required a fit with quadratic terms (rms = 0.0004). The in-orbit calibrations show reasonably good agreement with the ground calibrations obtained in TV2 and TV3 with differences of ~0.3 pixel in the offset and good agreement for the slope.

For the -1\(^{st}\), +2\(^{nd}\) and +3\(^{rd}\) orders, a similar procedure was carried out, however, due to the reduced field coverage for these orders the field dependent fits were more restricted and the trace prediction accuracy is reduced. For the 0\(^{th}\) order we adopt the trace determined in the ground calibrations. The same analysis was carried out for the G141 grism with similar results. Note, that the 0\(^{th}\) order is slightly dispersed (by the prism on which the grooves are ruled) and thus not a perfect image of the source.

\(^1\)http://www.stecf.org/software/slitless_software/axe/
2.2. Wavelength solutions

In order to establish the in-orbit wavelength calibration, the Galactic PN Vy2-2 was observed as part of the Cycle 17 calibration program (Proposal 11937) over 9 different field positions.

For wavelength calibration one would ideally want point sources with clearly identifiable emission lines over the entire wavelength range and at the resolution of the grism. While bright, nearby PNs do show a reasonably good distribution of emission lines, they are slightly extended at the spatial resolution of the WFC3 IR camera. The PNs Vy2-2 and HB12 have already been used for HST/NICMOS grism calibrations (Pirzkal et al. 2009) and they offer a reasonable balance between brightness, number of emission lines and wavelength coverage, hence we opted for using those targets in the first wavelength calibrations of the WFC3 IR grisms.

For the G102 grism a total of nine emission features and for the G141 a total of seven emission features could be identified and used to establish the wavelength solution (see Tab. 1). Prior to determining the wavelength solution, the radial velocity of -71 km/s for Vy2-2 and an air to vacuum conversion was applied to the tabulated wavelength of the emission lines. The ground calibration efforts suggested that a linear wavelength solution is a good representation of the true wavelength solution for these grisms (Kuntschner et al. 2008a,b). Therefore, we established for each spatial position a linear wavelength solution (see Fig. 4) where fits give a typical rms scatter of 0.2-0.3 pixel. The resulting field-dependent wavelength solution for G102 (see also Fig. 5) shows typical errors of 5 Å for the zeropoint and 0.04 Å/pixel for the dispersion (∼24.5 Å/pixel). For G141 the typical errors are 8 Å for the zeropoint and 0.06 Å/pixel for the dispersion (∼46.5 Å/pixel).
Sometimes the emission lines are affected by e.g. cosmic ray hits, hot/dead pixels etc. and thus good emission line centroids could not be established. Therefore we used a 3-sigma rejection iteration for each linear fit to remove outliers. The He I line at 1083.03 nm is saturated in the data. However, we removed the saturated pixels from the fit and used the line wings to establish a fit to the line. There is marginal evidence in the in-orbit data for a non-linear wavelength solution, however, the deviations appear to be < 0.5 pixel over the full wavelength range. More wavelength calibration observations, including the use of true point sources (e.g. extra-galactic PN, Be stars), are needed to confirm any non-linear terms in the dispersion solution.

Table 1: Rest air wavelength of emission features fitted to Vy2-2

<table>
<thead>
<tr>
<th>Name</th>
<th>Wavelength (Å)</th>
</tr>
</thead>
<tbody>
<tr>
<td>[Ar III]</td>
<td>7751.1</td>
</tr>
<tr>
<td>O I</td>
<td>8446.4</td>
</tr>
<tr>
<td>H P11</td>
<td>8862.8</td>
</tr>
<tr>
<td>[S III]</td>
<td>9068.6</td>
</tr>
<tr>
<td>H P9</td>
<td>9229.0</td>
</tr>
<tr>
<td>[S III]</td>
<td>9530.6</td>
</tr>
<tr>
<td>H P7</td>
<td>10049.4</td>
</tr>
<tr>
<td>[S II]</td>
<td>10321.9</td>
</tr>
<tr>
<td>He I</td>
<td>10830.3</td>
</tr>
<tr>
<td>He I</td>
<td>11969.1</td>
</tr>
<tr>
<td>He I</td>
<td>12784.9</td>
</tr>
<tr>
<td>H B13</td>
<td>16109.3</td>
</tr>
<tr>
<td>H B12</td>
<td>16407.2</td>
</tr>
<tr>
<td>H B11</td>
<td>16806.5</td>
</tr>
<tr>
<td>H B10</td>
<td>17362.1</td>
</tr>
</tbody>
</table>

Notes: Using a ground based spectrum (Hora et al. 1999) of Vy2-2 the wavelengths used in the fits were adjusted: The rest wavelength of the [S II] emission line is 10328.9, however, the ground based spectrum suggests that at the spectral resolution of the G102 grism the effective peak wavelength is reduced by ∼7 Å. For the G141 grism the ground based spectrum suggests that at the spectral resolution of the G141 grism the effective peak wavelength requires adjustments of -10.4 and +34.6 Å, for the He I emission lines at 11969.1 and 12784.9 Å.

2.3. Throughput measurements

Using the trace and wavelength solutions described in the previous section, the spectra of the flux standard star GD153 observed in Proposal 11552 were extracted. Note, that we make use of our own "3D" flat-field calibrations for the G102 and G141 grisms, which significantly improve the quality of the spectra (Kuntschner et al. 2009b,c). There were four, slightly dithered, exposures for each of the near-IR grisms taken near the central position of the FoV. A further two exposures cover the top left and bottom right corners of the FoV. For the flux calibration we combine the +1st order of the four central exposures into a single spectrum with the help of the spectral drizzle option in the aXe software (aXedrizzle). This spectrum is used to establish the flux calibration for the near-IR grisms. The spectrum was converted to units of [e−/Å/sec] and divided by a smoothed version of
Figure 4: Linear fit for the wavelength solution of the G102 grism in the +1\textsuperscript{st} order. The example shows a fit to one of the central pointings ($X_{\text{ref}} = 342.84$, $Y_{\text{ref}} = 553.73$). The top plot shows the fit whereas the bottom panel shows the deviations from the linear fit; the dashed lines indicate ±0.25 pixel deviation from the fit.

Figure 5: Example of the field-dependent wavelength solution for the G102 1\textsuperscript{st} order spectra is shown as a function of $X_{\text{ref}}$ and $Y_{\text{ref}}$ position (diamond symbols). The final field-dependent wavelength solution is shown as the grid of black dots. The zeropoint shows a marked trend with $X_{\text{ref}}$ whereas the slope is mostly dependent on $Y_{\text{ref}}$ position.
the model spectrum taken from the HST CALSPEC library\(^2\) (gd153_mod_007.fits). Median smoothing in a sliding 7 pixel window provides the final sensitivity function. The error is evaluated by taking the standard deviation of the four individual sensitivity curves from the central pointings at each wavelength bin and imposing a minimum error of 1%. Significant differences (e.g. >5\%) between individual observations of the standard star are only seen when the throughput of the system drops at the blue and red edges. For the G102 grism this occurs for wavelengths below 775 nm and above 1160 nm. Sensitivity curves were also derived for the 0\(^{\text{th}}\), +2\(^{\text{nd}}\), +3\(^{\text{rd}}\) and -1\(^{\text{st}}\) orders. While these orders are typically not used to extract science spectra, they are used in aXe to evaluate potential cross-contamination of other +1\(^{\text{st}}\) order science spectra.

Fig. 6 shows the total instrument throughput including telescope and detector for the WFC3 near-IR grisms G102 and G141. The peak throughput for G102 is 41\% at 1100 nm, while it is > 10\% over the wavelength range 805 to 1150 nm. The G141 grism shows a peak throughput of 48\% at 1450 nm while it is > 10\% from 1080 to 1690 nm. An example of science data extracted from WFC3 grism observations, with the help of the aXe software and calibration files discussed in this paper, is given in Kümmel et al. (2010, this volume).

3. The UV G280 grism

The analysis of the calibration observations for the UVIS G280 grism is still ongoing at time of the writing of this contribution. This grism was not designed for WFC3 but is a WF/PC1 spare. In contrast to the near-IR grisms, there is an offset of about 175 pixels in the y-direction between the direct image and the spectra (see Fig. 1), the zeroth-order is relatively bright (i.e. the star-like feature near the center of the G280 image in Fig. 1) due to a lower grating efficiency and clear substrate, and there is significant curvature of the spectra at the blue ends of the first orders (nearest the zeroth order; see also Fig. 1).

\(^2\)http://www.stsci.edu/hst/observatory/cdbs/calspec.html
The amplitude of the trace curvature was determined from in orbit data of the wavelength calibration star WR14 observed as part of the Cycle 17 calibration program 11935. For a fiducial, central position the trace amplitude is about \( \sim 25 \) pixels in the detector y-direction (see Fig. 7) and the trace is well fit with a 5\(^{th}\) order polynomial. However, the present data does not allow a field-dependent calibration of the trace or wavelength for this grism.

In order to establish a wavelength solution for the UVIS G280 grism at the fiducial, central position we identified the wavelengths of several emission features in the IUE spectrum of WR14. Wavelengths of emission features redwards of 320 nm were taken from an optical spectrum of WR14 (Torres-Dodgen & Massey 1988). Fig. 8 shows a 4\(^{th}\) order polynomial fit to 17 emission features identified in WR14 where a robust estimate of the standard deviation yields a value of 2.5 \( \AA \). The mean dispersion at 2200 \( \AA \) is 12.2 \( \AA \)/pixel, varying from \( \sim 10.8 \) \( \AA \)/pixel at 1850 \( \AA \) to \( \sim 14.4 \) \( \AA \)/pixel at 4000 \( \AA \).

A comparison of the IUE spectrum with an extracted and calibrated UVIS G280 spectrum of WR14 is given in Fig. 9. The overall agreement is satisfactory, however, there are indications of a wavelength calibration error in the G280 spectrum for the bluest wavelengths.

4. Conclusions

This paper presented calibrations of the WFC3 grisms based on SMOV data and Cycle 17 calibration programs. We have established in-orbit, source position dependent calibrations of the trace and wavelength solutions for the near-IR grisms. Both, the local trace and
Figure 8: Dispersion solution for the UVIS G280 grism as derived from the calibration star WR14. A 4th order polynomial fit (blue solid line) to 17 emission features in the spectrum of this star is shown. A robust estimate of the standard deviation yields a value of 2.5 Å.

Figure 9: An IUE spectrum of the star WR14 is shown as black solid line. The calibrated UVIS G280 spectrum as extracted from the Cycle 17 calibration observations (program 11935; ibbr01zmq_flt.fits) is shown in red. When spectral resolution differences between the observations are taken into account, the overall flux level of the UVIS G280 spectrum appears \( \sim 10\% \) higher compared to IUE for this preliminary flux calibration of the UVIS grism channel.
wavelength solution can be approximated with a linear function. The mean dispersion of the G102 grism is 24.5 Å/pixel, varying from 23.5 to 25.0 Å/pixel across the FoV. The mean dispersion of the G141 grism is 46.5 Å/pixel, varying from 45.0 to 47.5 Å/pixel across the FoV. The trace and wavelength solutions for the UVIS G280 grism are currently under study. Preliminary investigations show a highly bent trace and non linear wavelength solutions. The mean dispersion is 12.2 Å/pixel at 2200 Å.

For the near-IR grisms we establish flux calibrations for the $+1^{st}$ order. The total throughput of HST and the WFC3 G102 grism peaks at 1100 nm with 41% and is above 10% between 805 and 1080 nm. The G141 grism peaks at 1450 nm with 48% and is above 10% from 1080 to 1690 nm.

References

WFC3/IR Channel Behavior: Dark Current, Bad Pixels, and Count Non-Linearity

Bryan Hilbert and the STScI WFC3 Team

Space Telescope Science Institute, Baltimore, MD 21218

Abstract.
Using data taken during Servicing Mission Observatory Verification (SMOV) and Cycle 17, we have characterized many aspects of the on-orbit behavior of the IR Channel. We find the mean dark current in the IR channel to be 0.042 e−/s/pixel. We have also recently finished the creation of a bad pixel mask for the IR detector, which contains a list of pixels with non-nominal behavior that should be ignored in WFC3/IR data analyses. An update to the non-linearity correction file will be produced soon. Analysis of cycle 17 non-linearity calibration data is on-going.

1. Introduction

With the installation of WFC3 into HST during Servicing Mission 4 in May 2009, the primary task of the WFC3 team has been to ensure a high quality calibration of the data produced by the instrument. We present here some basic results of the behavior of the IR channel. This includes dark current and signal non-linearity behavior, as well as a description of the updated bad pixel mask. Full details of each of these behaviors can be found in the referenced Instrument Science Reports (ISRs). We summarize here the important highlights of each investigation.

2. Dark Current

Unlike the dark current behavior observed in CCD-derived data, the timing pattern used to collect data with the WFC3/IR detector has a large impact on the dark current accumulation rate. Analyses of ground testing data have shown that dark current signal does not increase linearly with time, but instead is initially zero (or even negative) followed by a regime where the signal does increase linearly with time. (Hilbert and McCullough, 2009) Figure 1 shows the mean dark current versus time for several sample sequences as measured in Cycle 17, where this pattern is apparent. These observations are very similar to those obtained during ground testing.

Given this behavior, dark current in WFC3/IR data must be subtracted from science data on a read-by-read basis, using a dark current file produced using the same sample sequence as the science data. Calculating an image of the mean dark rate and scaling that image by the exposure time of each read would result in an over- or under-subtraction of the dark current signal, as implied by the non-linear dark signal in Figure 1. Figure 2 is a diagram showing this dark current removal strategy.

In preparation for dark current characterization and calibration, we have collected a large quantity of WFC3/IR dark current data during (Servicing Mission Observatory Verification) SMOV and Cycle 17. Ramps were acquired at all supported full-frame and subarray sizes for all sample sequences. For each array size/sample sequence combination, we calculated a pixel-by-pixel sigma-clipped mean dark current ramp. These ramps have all either been delivered to the calibration database (in the case of the full-frame ramps, on
Figure 1: Mean dark current signal versus time for three different sample sequences. Note that dark current behavior in the early reads of a ramp depends on the sample sequence used to collect the data.

Figure 2: Diagram showing the dark current subtraction strategy for the WFC3/IR channel.

April 8, 2010), or will be delivered shortly (subarray observations). These updated dark current reference files represent a significant (4-8X) increase in signal to noise ratio over the previous dark current files in the calibration pipeline. Each new dark current reference file is a mean ramp created from 16 to 63 individual ramps, depending on sample sequence. The old dark current ramps were each the mean of only 3 ground-based dark current ramps.
For any observers with science goals where an accurate removal of dark current is critical, we recommend reprocessing your data with these new dark current files.

To give a general idea of the dark current behavior, we provide some basic statistics on the highest signal to noise mean dark current ramp below. Figure 4 shows a histogram and the cumulative distribution of dark current rates for all pixels. The peak dark current rate is 0.042 e-/sec. Only 0.7% of the pixels have a dark current rate above the hot pixel limit of 0.40 e-/sec. Figure 3 shows an image of the final read of the highest signal-to-noise dark current ramp, in order to give an idea of the relative dark current levels across the detector. We have also monitored the hot pixel population during SMOV and Cycle 17 and found it to be stable. We see no increase in the total number of hot pixels since the beginning of SMOV. The hot pixel population is also unchanged after monthly UVIS channel anneals, during which the IR detector temperature increases by 35°C.

Figure 3: Image showing the final read of a high signal-to-noise dark current ramp. Histogram equalization stretch from 0 to 0.4 e⁻/second.

3. Bad Pixel Table

During the various ground testing campaigns for the WFC3 IR channel, efforts were made to identify and flag any pixels on the detector which were bad in ways that made them scientifically useless. (Hilbert et al., 2003 and Hilbert, 2007) With WFC3 successfully installed in HST, we wished to update this list of bad pixels to reflect any changes since ground testing. During SMOV and the early stages of Cycle 17 a large amount of calibration data were obtained, allowing for an on-orbit bad pixel search. This was accomplished using a combination of dark current and internal flat field observations. Results of this bad pixel table update are detailed in Hilbert and Bushouse (2010).

Using the flat field and dark current observations, we searched for three types of bad pixels. These included: “dead”, “unstable”, and “bad in the zeroth read”. Our goal was to produce a table that lists pixels which are permanently bad. For example, a dead pixel, which exhibits little or no sensitivity to illumination, will likely always remain in that state. By focusing on permanently bad pixels, our aim is to have a table that changes little over time and can be applied to any WFC3/IR observation. Pixels which are bad in potentially
more transitory ways, such as those with higher than average dark current, are flagged in other types of reference files, such as dark current reference files.

The populations of bad pixels mentioned below were all combined into a single table which was uploaded into the calibration database on April 12, 2010. This new table replaced that produced from ground testing data, and should be used for the identification of bad pixels in all on-orbit data. Table 1 shows the flag values associated with each type of bad pixel. Final values in the bad pixel table are calculated using bit-wise addition. For example, a pixel found to be both bad in the zeroth read (8) and unstable (32), then that pixel will have a value of 40 in the final bad pixel table.

Figure 5 shows an image of the new bad pixel table, where all pixels flagged as bad in any of the ways described below appear white.

<table>
<thead>
<tr>
<th>Flag Value</th>
<th>Bad Pixel Type</th>
<th>Pixels Affected</th>
</tr>
</thead>
<tbody>
<tr>
<td>4</td>
<td>Dead</td>
<td>3,910</td>
</tr>
<tr>
<td>8</td>
<td>Bad Zeroth Read</td>
<td>4,990</td>
</tr>
<tr>
<td>32</td>
<td>Unstable</td>
<td>10,885</td>
</tr>
<tr>
<td>512</td>
<td>Affected by Blob</td>
<td>4,534</td>
</tr>
</tbody>
</table>

Table 1: Bad pixel flag values and descriptions, along with the number of science pixels flagged as each type in the new bad pixel table.

3.1. Bad Zeroth Read Pixels

Bad zeroth read pixels exhibit anomalous signal in the zeroth read of a data ramp (the read immediately following the detector reset at the beginning of an IR exposure sequence), usually due to being shorted or unbonded (Hilbert et al., 2003). We used the zeroth reads
of raw (uncalibrated) dark current ramps for this search. We first created a sigma-clipped mean zeroth read image. We then created a histogram of the values in this mean zeroth read for each quadrant of the detector. A Gaussian was fit to each histogram, and pixels falling more than 3\(\sigma\) from the mean value were flagged as bad. In order to catch any pixels which may have been inconsistent from ramp-to-ramp, we also repeated the above process on each input ramp individually. Any pixels which were more than 3\(\sigma\) from the mean in some ramps but not others were caught with this additional step. With this two-step method, we identified 4,990 pixels (0.5% of the detector’s light-sensitive pixels) which are bad in the zeroth read. Figure 6 shows an example of a histogram and gaussian fit. Pixels identified as bad in the zeroth read were given a value of 8 in the bad pixel table.

3.2. Dead Pixels

These are pixels with very low quantum efficiency and measure little or no signal when illuminated. As with the bad zeroth read pixels, we used a two step process to find dead pixels. First, we created a mean flat field image, using the final read from each ramp in a large set of flat field files. Any pixels exhibiting zero or negative signal in the mean flat field were flagged as dead. Next, we looked for pixels with very low, but not necessarily zero, quantum efficiency. For this, we moved pixel-by-pixel across the detector. For each pixel, we calculated the sigma-clipped mean value in the surrounding 50 by 50 pixel box. If the pixel had a signal less than 30% of the local mean signal, we flagged it as dead. By using this strategy, we hoped to compensate for the non-uniformity of the flat field illumination level. By comparing the signal from a pixel to that of its neighbors, we avoided marking a pixel as bad simply because it was located on a portion of the detector with a lower illumination level compared to other areas. Finally, we manually marked the pixels comprising the “death star” as dead. This feature (seen as the largest circular feature along the bottom edge of Figure 5) is a collection of poorly performing pixels resulting from a manufacturing defect. Combing the results from these three searches, we found 3,910 dead pixels (0.4% of the detector’s light-sensitive pixels), which were marked with a 4 in the bad pixel table.
3.3. Unstable Pixels

Unstable pixels were first observed on the WFC3/IR detector during ground testing. Pixels within this population display variable or unrepeatable signal measurements across a set of nominally identical ramps. The physical reason behind this behavior is unknown. A basic census of unstable pixels was taken using flat field ramps during ground testing. With WFC3 on orbit, the unstable pixel search was repeated using a larger dataset and a more thorough search.

Unstable pixels display a wide range of behaviors. Some unstable pixels appear stable and repeatable in almost all of a data set, but will then measure appreciably different signal values in just one or two ramps. Other unstable pixels display signal values that vary wildly from ramp to ramp in all observations of a data set. Figures 7 and 8 respectively, show examples of these two types of unstable pixels. Further examples can be found in the appendix of Hilbert and Bushouse (2010). Investigations across a wide range of IR data as part of this study also revealed some pixels which were unstable in a set of dark current ramps but stable in a set of flat field ramps, and others with the opposite behavior.

Due to the complexities of this behavior, we used multiple search methods and datasets to identify unstable pixels. By performing the search on a set of flat field ramps, and then repeating the search on a set of dark current ramps, we hoped to arrive at an accurate map of unstable pixels across the IR detector.

The first search method involved comparing the measured signal rate in a mean flat field image versus those in each of the individual flat field images in a dataset for a given pixel. If a pixel showed a signal rate in any individual image that was more than N-σ from the mean value, the pixel was marked as unstable. Figure 7 shows an example of an unstable pixel found with this search method. For our final unstable pixel search, we settled on a value of 5 for N. Using this limit, we identified 2,223 unstable pixels using the flat field dataset, and 6,001 unstable pixels using the dark current dataset. Since this
Figure 7: Example of an unstable pixel. This plot shows the signal measured up the ramp for a set of nominally identical flat field ramps. In this case, the pixel’s response was consistent for all but three ramps in the data set.

method relies on the uncertainty value in the mean flat field, it worked best at identifying pixels which were largely repeatable from ramp to ramp, but displayed anomalous signal in a small number of ramps.

The second search method was designed to identify unstable pixels which had much more variable signal rates from ramp to ramp. These pixels produced large uncertainty values in the mean flat field and therefore were potentially able to slip through the $N \sigma$ search. In this case, we searched for pixels with a measured signal rate that varied by more than a certain percentage of the mean measured rate. Figure 8 shows an example of a pixel falling into this category. Details of the calculation can be found in Hilbert, 2010. After some trial and error, we set our threshold to be a 0.93% variation relative to the mean signal rate for the flat field dataset, and a 61.5% variation for the dark current data. Using these values, we found 3,795 unstable pixels in the flat field data, and 3,318 in the dark current ramps.

Combining the populations of unstable pixels from these various searches, we find a total of 10,885 unstable pixels (1.06% of all science pixels) on the IR detector. These pixels are flagged with a value of 32 in the bad pixel table.

### 3.4. Blobs

The final type of bad pixel added to the bad pixel mask was pixels which were affected by blobs. WFC3 ISR 2010-06 (Pirzkal, Viana & Rajan 2010) provides details on these blobs. Essentially, these blobs are collections of pixels with decreased throughput relative to other pixels. This is most likely due to particulates resting on the Channel Select Mechanism (CSM). For the purposes of the bad pixel table, we took a list of blob positions and diameters, provided by Pirzkal, and mark the blob-affected pixels with the value of 512 in the bad pixel table. In all, 4,534 pixels (0.44% of all science pixels) are flagged as affected by blobs.
Figure 8: Another example of an unstable pixel, where the measured signal is highly variable.

4. Non-Linearity Behavior

The measured signal level from HgCdTe detectors, such as that in the WFC3/IR channel, is a non-linear function of the number of incident photons. Unlike that in CCDs, this non-linearity effect is significant down to signal levels which are a relatively low fraction of the full well.

Using flat field ramps as well as observations of external targets, we have devised a non-linearity correction algorithm for IR channel data. The correction currently in place was derived using flat field data from ground testing. On-going analysis using on-orbit flat field ramps as well as observations of 47Tuc will produce an updated correction to be put in place shortly. Here we describe the basic method used to calculate the non-linearity correction. Further details on this method can be found in Hilbert (2008).

In order to calculate the magnitude of the non-linearity at various signal levels, we began by creating a mean flat field ramp. Figure 9 shows a plot of the measured signal versus time for one pixel of this mean flat field in the series of connected black diamonds. The red line shows a best-fit line created using only reads with less than 4,500 DN, and then extended out to the remaining reads. Note that the measured signal appears linear at low signal levels, but begins deviating from this line at a signal level well below the marked saturation level. We define the signal level at which the measured signal is below the best-fit signal by 5% as the saturation level for that pixel. Our goal is to remove the non-linearity effects for signals below this saturation level.

By taking the difference between the best-fit line and the measured signal in each read, and dividing the difference by the calculated best-fit signal, we are able to calculate the percentage of the measured signals’ non-linearity at each signal. The black diamonds in Figure 10 show this calculated non-linearity versus signal level. The red curve in Figure 10 shows the best-fit 3rd order polynomial to the black diamonds. The best-fit coefficients of this red curve can be used to remove non-linearity effects from all measured signals in this pixel using Equation 1.
\[ s_{\text{corr}} = s_{\text{meas}} \times (1 + A + B \times s_{\text{meas}} + C \times s_{\text{meas}}^2 + D \times s_{\text{meas}}^3) \] (1)

In this equation, A through D are the best-fit polynomial coefficients, while \( s_{\text{meas}} \) is the measured signal, and \( s_{\text{corr}} \) is the corrected signal, with non-linearity effects removed. During the calculation of these coefficients when using the ground testing data, the limited amount of
data implied large uncertainties in the best-fit polynomial coefficients. We therefore decided to calculate the mean coefficients across each quadrant of the detector, and use these when correcting signals for all pixels in each quadrant. The blue line in Figure 10 shows the curve produced by the mean quadrant 2 coefficients. In this case, the difference between the best-fit polynomial for the individual pixel and that for the quadrant mean is small. For other pixels, this difference is more significant. Part of our on-going analysis using the on-orbit flat field data is an investigation into whether the pixel-by-pixel or quadrant-averaged correction provides a better non-linearity correction. We are also investigating whether the non-linearity behavior is independent of the source of illumination, by comparing results obtained from flat field data to those from observations of point sources. A forthcoming ISR will detail those results.

5. Conclusions

Data from cycle 17 have been used to evaluate the behavior of the WFC3/IR channel, and produce appropriate calibration files. Calibration files derived from on-orbit data represent a significant improvement over the previous versions of these files, which were derived from ground tests. New dark current subtraction files and a new bad pixel table were uploaded to the calibration database system in early April of 2010. Any data retrieved from the archive after this time will have had these new files applied. If an accurate dark current subtraction or bad pixel flagging is important to your science goals, we recommend re-retrieving your data from the archive.

We will continue to collect dark ramps for all sample sequences in Cycle 18 and will provide periodic updates to the dark current reference files. Similarly, the bad pixel population will continue to be monitored during Cycle 18, and any necessary changes to the bad pixel table will be made.

The non-linearity correction reference file will be updated soon, based upon results from the on-going analysis.

6. Observation Planning and Data Reduction Tips

In order to mitigate the effects of bad pixels (as well as other effects), we strongly recommend that observers dither their observations. By making multiple observations of a field while moving targets around on the detector, observers will decrease the chance of having all observations of an individual target contaminated by a bad pixel. Dithered individual observations can then be combined by software such as Multidrizzle (Fruchter et al. 2009), where observations affected by bad pixels are ignored, resulting in a combined image where effects of bad pixels are minimized.

The current version of the cosmic ray rejection table (u6a17488rLcrr.fits) causes calwf3 to calculate a signal rate for all pixels on the detector, regardless of whether they are flagged as bad or not. Due to this behavior, it is important that observers consider their science goals, decide which bad pixel types are not acceptable to use in their analyses, and mask the appropriate pixels.

References

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Persistence in the WFC3 IR Detector

Knox S. Long, Sylvia M Bagett, Susana Deustua, and Adam Riess

Space Telescope Science Institute, Baltimore, MD 21218

Abstract. As is the case for most if not all modern IR arrays, bright sources observed with the IR detector in WFC3 leave faint residual images in subsequent exposures. This image persistence has been observed not only in dithered exposures by one observer of a single target within an orbit but also and more often in exposures of a different target by another observer in subsequent orbits. The amount of image persistence in the WFC3 IR detector is a function of the degree of photo-generated charge saturation of a pixel and time since the pixel was (over)exposed. The persistence decays roughly as a power law with time, and is typically $0.3 \text{ e s}^{-1}$ for a pixel that was highly saturated 1000 s earlier. Here, we show examples of persistence which have been observed and characterize the effect. We also describe ways for observers to find the pixels that are likely to have been affected by persistence, and to mitigate the effects of persistence when planning observations and reducing their data.

1. Introduction

Persistence is a residual image observed in most types of IR arrays after they have been exposed to signals that fill individual photodiodes, or pixels, to a substantial fraction of the full-well. Persistence has been observed on HST’s first IR instrument, NICMOS (Daou & Skinner 1997), which like the IR detector incorporated into WFC3 is a HgCdTe device. Persistence has also been observed in images taken with the IR detector on WFC3. The time scale for the persistence to decay in the HgCdTe detector on WFC3 is sufficiently long that persistence can be observed not only within an orbit as the source is dithered but also on subsequent orbits. This means that the persistence signature on the array can on occasion reflect an observation by another observer of a completely different region of the sky. The characteristics of the persistence in orbit are similar to those measured in thermal vacuum tests (McCullough & Deustua 2010).

Persistence is understood to arise from traps in the material that comprises the active area of the detector, especially regions of the photodiode that become accessible to free electrons and holes when the depletion region changes location due to photo-generated charge (Smith et al. 2008). Persistence primarily reflects the total number of photo-generated electrons accumulated from the beginning of an exposure to the end; thus persistence is seen from bright sources exposed for a short period of time or less bright sources exposed for a longer period of time. Other factors, such as the amount of time spent in a specific charge-state, also affect the amount of persistence in some HgCdTe devices, but have not yet been investigated systematically in on-orbit observations with the IR detector incorporated into WFC3.

In the remainder of this report, we describe what has been learned about persistence in the WFC3 detector since observations began shortly after it was installed into HST, our initial attempts to identify persistence in on-orbit data and to develop a protocol to remove persistence in subsequent images.
The nature of persistence in the WFC3 array

An example of persistence as seen in the WFC/IR array is illustrated in Figure 1. The frame being shown is the first image of a sequence of 1300 s exposures of a high latitude field observed with the F160W filter. The image is displayed in histogram equalization mode that emphasizes faint features and shows very clear signatures of at least two dither sequences from observations in earlier orbits. These observations contained a number of very bright stars that produced heavily saturated images. Since persistence is a function of pixel exposure the previous dither pattern is imprinted on the array several hours later. The brightest persistence in this image is about 0.09 e s$^{-1}$, and corresponds to about 120 electrons.

A plot of persistence (measured in e s$^{-1}$) as a function of earlier illumination (measured in e) is shown in Figure 2a. In this case, the observation that caused the persistence contained a single extremely bright star and the persistence is seen in the pixels in the array that are included in the wings of the point spread function (PSF) of the star. The persistence image was obtained about 14 minutes after the original exposure. Persistence in the WFC3 array is minimal in regions of the detector where the exposure, measured in electrons, is less than 50,000 e, rises rapidly as the exposure grows to 100,000 e, and rises only slowly for values that are larger than this.$^1$ The turnover at 100,000 e reflects the full-well depth of the pixels in the WFC3 array.

The rate at which persistence decays, as measured in a calibration program in which a sequence of darks were obtained after first illuminating the array with the Tungsten flat-field lamp, is shown in Figure 2b. The various curves are for lamp exposures ranging from 50,000 to 1,500,000 e. At times greater than 100 s, the persistence decays roughly as a power law with a slope of about -2. The decay rate appears to vary somewhat as a function of illumination, but the decay is not a strong function of the illumination.

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$^1$In describing exposures of greater than the well-depth, we mean the product of the instantaneous rate and the exposure time.
Figure 2: a) The persistence seen in a dark exposure about 14 minutes after an observation involving multiple exposures of an extremely bright star. Most of the pixels shown are due to the wings of the PSF of the bright star. Every tenth pixel is plotted in green to make the dispersion in individual pixels a little clearer. b) The rate of decay of persistence measured in es$^{-1}$ as a function of time in seconds following Tungsten lamp flat-field exposures.

3. Finding persistence in WRC3 images

As indicated in Figure 1, one simple way to inspect images for evidence of persistence, especially persistence from observations taken at a different position on the sky, is to display the first image of a visit in histogram equalization mode. Another very good way to find persistence is to construct difference images of exposures taken at different dither positions in your observation. A convenient way to do this, if your own images are well-dithered, is to use MultiDrizzle (Fruchter et al. 2009) to produce individually drizzled, distortion-corrected images from each exposure (single_sci.fits files) projected to the same coordinate and pixel scale. Subtracting any two images that you have dithered will reveal the persistence since the signature of the field you were observing will be largely removed by this procedure, but due to the offset, the persistence will not. An example of this is shown in Figure 3. In this case observations of a high latitude field were preceded, unbeknownst to the observer, by a calibration observation of the outskirts of 47 Tuc.

Self-induced persistence, that is persistence occurring within a single visit, is often harder to detect than persistence from earlier visits involving other astronomical targets. This is because many observers use dither patterns involving steps that are small compared to the wings of the PSF of a bright star. This is a good strategy from the point of persistence since it confines persistence to small regions of the detector that are usually not part of the primary science of an observer. However, if an observer uses larger steps, as one might to mosaic a larger field or to step over macroscopic features on the detector (such as the “Death Star” or IR blobs), then persistence can be an issue. In this case, a simple way to check for persistence is to simply step through images displayed in image coordinates and

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2If one suspects persistence induced by earlier observations or indeed if one wishes simply to check whether persistence is a possibility, one can obtain a time history of WFC3 IR exposures using a search form located at http://archive.stsci.edu/hst/history_search.html (If you copy and paste this link, make sure the ‘’ appears in your browser.)
Persistence in the WFC3 IR Detector

Figure 3: a) An image that looks fairly normal until one actually checks for persistence b) The difference of the first and last single sci.fits files from this observation. The two images were dithered on the sky and so the persistence, which is tied to a pixel instead, shows up as a positive and negative signals which are separated by the size of the dither.

look for features that gradually fade as you step through the images beginning of course with the image that contained a bright star. An example of this is shown in Figure 4.

Figure 4: A sequence of images obtained with WFC3/IR when a bright star was observed and the following exposures involved large dither steps from the initial position.

4. Mitigating the effects of persistence

From the point of view of planning observations, the best way to mitigate persistence is to avoid it. The best strategies for this are to use as short a series of exposures as is possible, consistent (usually) with the desire to avoid time gaps for storing the data. This will minimize the number of pixels that are any image that are saturated or close to saturation. This is especially important if one’s science involves crowded star fields or very bright diffuse objects. Obviously, one should also be sure to make sure that dithers do not bring the areas on the sky that contain the prime science into regions of the detector that have been previously exposed to bright stars. Observers need to be especially careful in situations where they are looking for faint targets in fields crowded with moderately bright stars. An exposure of an 18th magnitude star with WFC3/IR using the broadband filters, e. g. F110W that lasts about 150 s will saturate the central pixel to an effective depth of 100,000 e. Estimates of which stars will saturate in WFC3 are available using the Bright Object Protection in APT.

At first glance one might suggest that a good strategy to reduce the effects of persistence would be to order the exposures such that the “thinnest” exposures occur first, however usually, observers balance exposure times so that all of the images are comparably deep.
In many instances, one cannot avoid persistence in images. The STScI does attempt to screen observations for the targets and exposures that will cause the worst persistence and to prevent IR exposures taken too close in time to these observations, but to do so with all IR observations would reduce overall scheduling efficiency of HST, and in most cases the regions of the IR detector that are affected by persistence in a subsequent observation are small, comparable to or less than the areas of the detector that are marked as being less than optimal data quality for other reasons.

It is also possible to at least partially remove the effects of persistence from images if one has access to the earlier images. Tools to correct for persistence are currently under development at STSCI; our prototype software assumes that persistence $P(x,y,t)$, measured in $\text{e}\cdot\text{s}^{-1}$, can be modeled as

$$P(t) = P_{1000} \left( \frac{t}{1000\text{s}} \right)^{-\gamma} \left( \frac{1}{\text{e}^{\frac{\delta z}{z_o}} + 1} \right)$$  \hspace{1cm} (1)$$

where $P_{1000}$ is the maximum persistence at a fiducial time of 1000 s and characteristic exposure level $z_o$ measured in electrons. Here $\delta z$ describes the scale of exposure over which the persistence rises from a minimal to maximal value. We measure $t$ from the end of a previous exposure to the mid-point of the current exposure. (This is not a perfect description of the persistence as it does not track the slow rise in persistence as the image becomes extremely heavily saturated, but at present pixel to pixel variations in persistence is likely a larger problem.) If, as is often the case, there are multiple earlier images that potentially contribute to persistence in the current exposure, we assume that only the image contributing the maximum persistence matters. Based on this model, we calculate a persistence image and subtract it from the current fit file. The persistence image for our first example (Figure 1) is shown in Figure 5, along with the persistence subtracted image (again displayed in histogram equalization mode). In this case, we have set $z_o$ to be 80,000 e and $\delta z$ to be 20,000 electrons and $P_{1000}$ to be 0.3 $\text{e}\cdot\text{s}^{-1}$. Most of the persistence has been subtracted. Whereas persistence peaked at 0.09 $\text{e}\cdot\text{s}^{-1}$ previously, it now peaks at less than 10% of this, so about 90% has been removed.

Figure 5: a) An estimate of the persistence affecting the image shown in Figure 1, using the model described here. The image is plotted on a linear scale. b) The image shown in Figure 1 now with the persistence removed. The image is displayed in the same histogram equalization mode as Figure 1.
5. Summary and Future Work

Persistence is a residual image seen in most if not all HgCdTe detectors and is seen in the IR detector on WFC3. In particular, the WFC3 IR detectors exhibit persistence whenever the photo-generated charge is a substantial fraction of the full-well. This persistence can be observed in exposures within a single visit or in exposures which follow observations (by others) of completely different star fields by several hours. It can be both scientifically and cosmetically deleterious, but can be reduced by the averaging process that takes place when dithered images are “drizzled” and/or when models based on the time history of exposures are used to subtract the persistence from images.

We are continuing to investigate persistence in the IR, including quantifying pixel by pixel variations in persistence and attempting to determine how to model effects due to the length of time a pixel is near full well. We do not yet have a complete model for persistence but are working with individual observers on removing persistence from their images. We are considering the possibility of producing persistence images on a routine basis for observers, but in the meantime observers who find significant persistence in their images are encouraged to contact helpstsci.edu, and we will, on a case by case basis, work with them to see if persistence can be removed from their images. The latest information on image persistence will be posted on http://www.stsci.edu/hst/wfc3/ins_performance/persistence/.

References

WFC3 Image Calibration and Reduction Software

Howard A. Bushouse

*Space Telescope Science Institute, Baltimore, MD 21218*

**Abstract.** Standard WFC3 image processing consists of the calwf3 task, which removes instrumental signatures from the images, and multidrizzle, which corrects images for geometric distortion and combines dithered sets of exposures. In addition, the aXe software package is used off-line to perform spectral extraction and calibration of WFC3 grism data. We present an overview of the standard WFC3 pipeline processing and calibration, including the status of calwf3 and multidrizzle. Reasons for reprocessing data outside the pipeline environment are also discussed.

1. **WFC3 Pipeline Overview**

   The pipeline data reduction and calibration system at STScI uses several steps to process WFC3 exposures when they arrive on the ground. A brief outline of those steps is listed here.

   - **Generic Conversion.** Raw telemetry data from all the HST instruments goes through this process, which reformats the data into FITS files containing the raw detector readouts and a host of information about the exposure and the instrument stored in header keywords. In addition, this process also queries the STScI Calibration Data Base System (CDBS) to determine the appropriate calibration reference files (biases, darks, flats, etc.) to be used later when calibrating the data. The names of the selected reference files are stored in header keywords in the raw FITS files.

   - **calwf3.** The calwf3 task is run for each raw or set of associated raw images to remove instrumental signatures. This includes the usual types calibrations, such as bias and dark subtraction, non-linearity correction, and flat fielding. For IR exposures it also includes up-the-ramp fitting to produce a final image from the multiple non-destructive reads of the IR detector. UVIS channel exposures obtained with the CR-SPLIT option are also combined by calwf3, but the resulting combined and CR-cleaned image is now essentially made obsolete by subsequent multidrizzle processing.

   - **multidrizzle.** The multidrizzle task drizzles calibrated images to remove geometric distortion, combine dithered exposures, and remove cosmic-ray (CR) hits. For dithered exposures the resulting combined image is also cleaned of bad pixels and other cosmetic defects.

   - **Archiving.** The raw telemetry files are stored in the Multimission Archive at STScI (MAST) and metadata gleaned from the calibrated images is stored in the archive catalog.

   - **OTFR.** Image retrieval requests submitted to the MAST archive use the On-The-Fly-Reprocessing (OTFR) system to completely reprocess the data before delivering it to the requestor. This process starts from the raw telemetry data stored in the archive and reapplies all generic conversion, calwf3, and multidrizzle processing using the latest software and calibration reference files.
• **aXe.** aXe is an off-line software package that is used exclusively for extracting and calibrating spectra from grism exposures. aXe processing is not applied within the pipeline. See K"ummel et al. (2009), K"ummel et al. (2010), and Kunthacker et al. (2010) for more information about aXe and its application to WFC3 grism images.

2. **Using MAST For Reprocessing**

Because WFC3 is still a relatively new science instrument on HST, the detailed instrumental calibrations are still evolving as we continue to learn more about the behavior of the instrument and continue to refine the calibration data. The simplest way for GO’s to have the latest calibrations applied to their data is to put in a retrieval request to MAST, which will result in a complete reprocessing of the data before being delivered, as mentioned in the previous section. This is also usually the best and only way to have the latest versions of the processing software applied to the data, because the software used in the STScI pipeline is updated more frequently than what is made available publically in the IRAF/STSDAS package.

WFC3 users are encouraged to check the “Late Breaking News” section of the WFC3 web site (http://www.stsci.edu/hst/wfc3) for information about new calibration reference files and updates to the calwf3 and multidrizzle software. Specific information about software updates can be found via the “Pipeline” link. The WFC3 team also periodically publishes WFC3 Space Telescope Analysis Newsletters (STANs), which contain information about calibration and software issues.

Users can determine which version of calwf3 was applied to their data by checking the value of the “CALVER” keyword in the header of their FITS files. Similarly, the “HISTORY” keywords contain a record of which version of multidrizzle, and associated sub-tasks, was used to produce the images.

3. **Recent CALWF3 Updates**

As discussed in section 1, calwf3 is the software task that applies instrumental calibrations to all WFC3 images. At the time of this writing, the latest version of calwf3 is v2.1, which was released to the public in June 2010 in the STSDAS package release v3.12. See the “Pipeline” link on the WFC3 web site for detailed information about all of the changes contained in this and previous versions.

Some of the more important changes that have occurred in the last few months include the following. First, a mistake in the way the IR flat fielding and gain conversion (FLATCORR step) were being applied resulted in DC offsets between the four amplifier quadrants in calibrated IR images. This problem was fixed in calwf3 v1.8. Secondly, the IR up-the-ramp fitting and CR rejection process (CRCORR step) was incorrectly computing the estimated uncertainties for the fits to each pixel, which resulted in severely underestimated ERR array values in the output IR flt.fits files. This problem was fixed in calwf3 v2.0.

Finally, an updated version of the IR cosmic-ray rejection parameters table (CRREJTAB) was delivered in June 2010, which changed the behavior of the IR up-the-ramp fitting process (CRCORR step). Previously, when the CRCORR process encountered a pixel that was flagged as bad in all of the readouts for the exposure, it would not fit the pixel ramp data and instead simply enter a value of zero for that pixel in the output flt.fits image. This happens for pixels that have a “static” problem condition, such as a dead, hot, or photometrically unstable pixel, in which case all the readouts for that pixel are flagged as bad. This differs from pixels for which only some of the readouts are bad, such as those that saturate part way through the exposure or contain a CR hit in one or more readouts.
The updated version of the CRREJTAB that was put into use in June 2010 changes this behavior so that the CRCORR process still tries to fit the ramp data for pixels with all readouts flagged, coming up with a “best effort” value for the slope of the ramp. Such pixels are therefore no longer zeroed out in the calibrated image. All of the data quality (DQ) flags associated with those pixels are recorded in the DQ array of the output .flt.fits file, so users should be careful to check the DQ values to decide which pixels are safe to use in their analysis.

An example of the effects of this change is shown in Figure 1. The left panel of this figure shows a portion of a calibrated IR image that was produced before the change and therefore has bad pixels zeroed out. The right panel shows the same image produced with the new processing, which no longer zeroes out the bad pixels. Notice that a few pixels still have anomalous values, which are pixels that simply have very erratic behavior.

4. Recent MultiDrizzle Updates

WFC3 pipeline processing did not include the use of multidrizzle until early February 2010, when initial WFC3 geometric distortion solutions became available and the software itself had been verified to be working properly for WFC3 images. A few issues have been discovered and fixed since that time.

First, after having gained some experience with the application of multidrizzle to WFC3 IR images, an updated version of the IR multidrizzle parameters table (MDRIZTAB) was delivered in April 2010. This reference table sets the various multidrizzle processing parameters for use in the STScI pipeline. There were several changes to the table. The values of the “drizsep_bits” and “final_bits” parameters were modified so that IR image pixels flagged with DQ=512, which indicates that the pixel is affected by an IR “blob” (see Pirzkal, Viana & Rajan 2010 and Pirzkal 2010) are considered to be good, while pixels with all other DQ values are considered bad and hence rejected. The CR detection threshold was increased to reduce the number of false detections and detected CRs are no longer flagged in the DQ arrays of the input .flt.fits images. Finally, sky subtraction was turned on for all exposures using wide- and medium-band filters and the IR grisms.
The second important change involved an update to both the WFC3 generic conversion process and multidrizzle, both of which occurred in April 2010. The generic conversion process was updated to compute the velocity aberration factor for all WFC3 exposures, as is done for other HST instruments. This value is stored in the “VAFCTOR” header keyword of WFC3 FITS files and specifies the overall image scale factor that should be applied to correct for the expansion or contraction that results from the orbital motion of HST. Multidrizzle was also updated to make use of this keyword value when processing WFC3 images and include the appropriate scale factor when removing geometrical distortions from the images. Users who require precise relative astrometry or image registration should make sure to use reprocessed data products that have this correction applied.

5. Rerunning CALWF3 and Multidrizzle

Occasionally it may be necessary for a user to rerun calwf3 or multidrizzle on their data to achieve the optimum results for their science program. Rerunning calwf3 may be done, for example, when you want to apply an alpha release of a new calibration file that is available for download from the WFC3 web site, but not yet installed for use in the OTFR system. You might also want to use a custom version of a particular reference file, such as a modified bad pixel table to flag features unique to your data, or use different CR or bad pixel rejection parameters in the IR ramp fitting via the CRREJTAB table.

The procedure to do this is relatively straightforward. Once you have downloaded WFC3 reference files from the web and possibly made modifications to them, you simply need to edit the values of the header keywords in your raw data files that contain the names of these reference files, and then run calwf3 on the raw files. Details of this procedure can be found in the WFC3 Data Handbook, which is available at the WFC3 web site.

Multidrizzle processing must be closely tailored to different types of imaging situations, which can not be done easily in the pipeline environment. The multidrizzle processing parameters that are used in the pipeline are therefore a “one size fits all” compromise. This makes it very important for users to rerun multidrizzle themselves on their data in order to achieve the best results. Typical parameters that may need optimizing include the final output image pixel size and corresponding “pixfrac” setting, and different values of the “drizsep\_bits” and “final\_bits” parameters to accept and reject different families of flagged pixels. Perhaps most important is the use of user-supplied alignment information, which adjusts the alignment of the individual images being combined. This image alignment information can often be generated through the use of the “tweakshfits” task available in PyRAF. See the Multidrizzle Handbook for detailed information, the WFC3-specific multidrizzle information on the WFC3 web site, and Mutchler (2010).

Table 1 lists the WFC3 data quality (DQ) values that are used to flag various problem conditions with pixels in UVIS and IR images. Note that there are some unique conditions for the two imaging channels. When rerunning multidrizzle on your data it is important to consider which types of flagged pixels may be included in the combined image and which should be rejected. For example, permanently dead or bad pixels (dq=4) and saturated pixels (dq=256) should probably always be rejected. Some other types of flagged IR pixels, however, may be OK to include for some types of data. If you are mainly interested in bright sources in the images, then it’s probably OK to include pixels flagged as deviant in the IR zero read (dq=8) and possibly even photometrically unstable (dq=32), because the errors associated with these conditions are likely to be negligible compared to the source signal. For faint sources, on the other hand, these types of problem pixels should be rejected.

The choice of which types of flagged pixels to reject also depends on the dither pattern used for the exposures. If the dither steps are large enough to step over large features like the IR “blobs”, then you may reject pixels flagged with dq=512. The regions of rejected pixels will be filled in with good data from adjacent images. If the dither steps are smaller...
Table 1: WFC3 Data Quality Flags

<table>
<thead>
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<th>DQ Value</th>
<th>UVIS channel</th>
<th>IR channel</th>
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</thead>
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<tr>
<td>1</td>
<td>Reed-Solomon decoding error (on HST)</td>
<td>Reed-Solomon decoding error</td>
</tr>
<tr>
<td>2</td>
<td>Missing data packet (during downlink)</td>
<td>Missing data packet</td>
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<tr>
<td>4</td>
<td>Dead/Bad pixel</td>
<td>Dead/Bad pixel</td>
</tr>
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<td>8</td>
<td>(currently unassigned)</td>
<td>Deviant zero-read bias value</td>
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<td>16</td>
<td>Hot pixel</td>
<td>Hot pixel</td>
</tr>
<tr>
<td>32</td>
<td>CTE tail (not yet implemented)</td>
<td>Photometrically unstable</td>
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<tr>
<td>64</td>
<td>Warm pixel</td>
<td>(currently unassigned)</td>
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<td>Bad bias value in overscan</td>
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<td>512</td>
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<td>Bad flat field value (blobs)</td>
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<tr>
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<td>(currently unassigned)</td>
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<td>Ghost/Crosstalk (not yet in use)</td>
<td>Ghost/Crosstalk (not yet in use)</td>
</tr>
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</table>

than the blob size, however, then those regions can not be filled in and therefore should be rejected. Figure 2 shows an example of filling in the IR blobs via large dither steps. The left panel in the figure shows an individual calibrated image (flt file) from the set of exposures, in which you can see the effect of one of the blobs. The resulting drizzled image (right panel) has the region of the blob filled in with good data from adjacent exposures.

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Figure 2: On the left is a section of a single IR image showing a region of decreased signal due to an IR “blob”. The image on the right is the result of multidrizzling a set of these images that used dither pattern steps larger than the size of the blobs, thus filling in the region of the blob with good data from other images.
First On-orbit Measurements of the WFC3-IR Count-rate Non-Linearity

Adam Riess

Space Telescope Science Institute, Baltimore, MD 21218

Abstract.

Previous HgCdTe detectors on HST have suffered from a count-rate dependent non-linearity, motivating an investigation of a similar effect on the WFC3-IR detector. An initial measurement of this effect was made by comparing the photometry of star clusters observed over a wide dynamic range and at overlapping wavelengths in WFC3-IR and NICMOS and/or ACS-WFC. Utilizing a color term to account for differences in the observed bandpasses, we find a significant detection of a non-linearity in WFC3-IR photometry which is in the same direction but a few times smaller than that of NICMOS. From 235 stars in 47Tuc observed with WFC3-IR in F110W and F160W and in similar bandpasses in NICMOS Camera 2, we measure a non-linearity of WFC3-IR of 0.011+/-0.0023 and 0.010+/-0.0025 mag per dex, respectively, over a range of 10 magnitudes (4 dex). An independent measurement utilizes 1390 stars in NGC 3603 observed with ACS-WFC F850LP and WFC3-IR F098M and yields a very similar result, 0.010 +/- 0.0033 mag/dex. The consistency of this measurement from two different comparison detectors of different technology indicates this result is robust. The impact of this non-linearity is that photometry of faint (i.e., sky dominated) sources calibrated with WFC3-IR zeropoints will appear 0.04 +/-0.01 mag too faint.

1Johns Hopkins University, Baltimore, MD 21218
ACS after SM4: New Life for an Old Workhorse

David Golimowski\(^1\), Jay Anderson\(^1\), Amber Armstrong\(^1\), Steve Arslanian\(^2\), Luigi Bedin\(^1\), Ralph Bohlin\(^1\), Kevin Boyce\(^2\), George Chapman\(^3\), Edward Cheng\(^3\), Marco Chiaberge\(^1\), Colin Cox\(^1\), Tyler Desjardins\(^1\), Darryl Dye\(^2\), Tracy Ellis\(^1\), Brian Ferguson\(^1\), Andrew Fruchter\(^1\), Norman Grogin\(^1\), Pey Lian Lim\(^1\), Markus Loose\(^4,5\), Ray Lucas\(^1\), Olivia Lupie\(^2\), Jennifer Mack\(^1\), Aparna Maybhate\(^1\), Kathleen Mil\(^1\), Max Mutchler\(^1\), Raphael Ricardo\(^5\), Barbara Scott\(^2\), Beverly Serrano\(^2\), Marco Sirianni\(^6\), Linda Smith\(^1\), Anatoly A. Suchkov\(^7\), Augustyn Waczyński\(^2\), Alan Welty\(^1\), Thomas Wheeler\(^1\), and Erin Wilson\(^2\)

Abstract. The ACS CCD Electronics Box Replacement (CEB-R) installed during SM4 features a Teledyne SIDEAR ASIC that permits optimization of the WFC via adjustment of CCD clock voltages, bias voltages, and pixel transmission timing. An on-orbit campaign to optimize the performance of the WFC was undertaken at the start of the SMOV period. Initial tests with pre-SM4 default voltages and timing patterns showed that WFC’s performance matches or exceeds its pre-failure levels, notwithstanding the expected increases in dark current and hot pixels and the decline in charge-transfer efficiency due to prolonged exposure to HST’s radiation environment. One WFC CCD exhibited anomalous behavior when operated with nondefault settings of its reset drain voltage. Consequently, the optimization campaign was truncated after two iterations, and ACS science operations commenced with the pre-SM4 default configuration. Several artifacts attributed to the CEB-R appear in post-SM4 WFC images: large-scale but stable bias gradients, low-level but temporally variable horizontal stripes, a signal-dependent bias shift, and amplifier crosstalk. STScI has developed algorithms for the correction or mitigation of these electronic artifacts as well as for the restoration of images affected by continuously degrading CTE. Standalone correction packages are now or will soon be publicly available. These packages will be incorporated into the calacs package of the OPUS data pipeline by September 2011.

1. Introduction

The Advanced Camera for Surveys (ACS) suffered component failures in its Side 1 and Side 2 Low Voltage Power Supplies (LVPS) in June 2006 and January 2007, respectively.

\(^1\)Space Telescope Science Institute, 3700 San Martin Drive, Baltimore, MD 21218
\(^2\)NASA’s Goddard Space Flight Center, 8800 Greenbelt Road, Greenbelt, MD 20771
\(^3\)Conceptual Analytics LLC, 8209 Woburn Abbey Rd, Glenn Dale, MD 20769
\(^4\)Markury Scientific Inc, 518 Oakhampton Street, Thousand Oaks, CA 91361
\(^5\)Teledyne Scientific & Imaging LLC, 1049 Camino Dos Rios, Thousand Oaks, CA 91360
\(^6\)European Space Research and Technology Centre, Keplerlaan 1, Postbus 299, 200 AG Noordwijk, The Netherlands
\(^7\)Department of Physics and Astronomy, Johns Hopkins University, 3400 North Charles Street, Baltimore, MD 21218
that afterwards prevented operation of the Wide Field Channel (WFC) and High Resolution Channel (HRC). Both channels were revived after the Side 1 failure by switching the power and control to the Side 2 electronics. The subsequent fatal short circuit in the Side 2 LVPS or Auxiliary Power Box (APB) blew a fuse in the Power Distribution Unit (PDU) and damaged the hold bus. The Solar Blind Channel (SBC) was unaffected by these failures and remained available for scientific use throughout this problematic period.

The WFC was restored to operation after the successful installation of the WFC CCD Electronics Box Replacement (CEB-R) and the Low Voltage Power Supply Replacement (LVPS-R) during Servicing Mission 4 (SM4). Figure 1 is a schematic diagram of the ACS repair (ACS-R) configuration. The LVPS-R receives power from the Side 1 PDU via the Power Intercept Element and then distributes the +5, −15, +15, and +35 V supplies to the CEB-R and the HRC CEB via the Power Output Element. Unfortunately, additional damage to the HRC power harness during the Side 2 failure prevented the recovery of the HRC during the ACS-R mission. Consequently, the HRC remains unavailable for scientific use.

The CEB-R is equipped with an Application-Specific Integrated Circuit (ASIC) manufactured by Teledyne Scientific & Imaging, LLC, that provides a System for Image Digitization, Enhancement, Control, and Retrieval (SIDECAR). The SIDECAR ASIC generates the bias and clock voltages needed to operate the WFC CCDs. It is fully programmable and its flight assembly code can be modified on-orbit to adjust the CCD voltages if desired. The video signals from WFC’s four output amplifiers can be sensed by an internal oscilloscope mode and sent directly to the ASIC. The CEB-R is also equipped with a Field Programmable Gate Array (FPGA) that controls the timing and packaging of 16-bit pixel data from the CEB-R to the Main Electronics Box (MEB). Figure 2 is a schematic diagram of the CEB-R and its interfaces with the WFC CCDs and MEB. Together, the SIDECAR ASIC and the FPGA allow on-orbit optimization of WFC performance through concurrent and iterative adjustments of the CCDs’ operating voltages and timing patterns.

Although ground tests verified the functionality of the CEB-R with a flight-spare WFC unit, the intrinsically variable performance of the SITe CCDs manufactured for ACS precluded a ground-based determination of the optimal voltage and timing settings for the...
on-orbit WFC CCDs. Ground tests did show that the external preamplifiers attached to the CCDs display non-ideal, transient settling behavior. This behavior was not tested in isolation, so its effect on the CEB-R’s signal processing circuitry could not be assessed prior to SM4. Furthermore, ground tests revealed that the read noise depends on the timing of the data transmission from the CEB-R to the MEB relative to the timing of the pixel integration and analog-to-digital (A/D) conversion. The CEB-R’s ability to adjust the timing pattern of the data transmission provides a potential means of noise mitigation that was not available with the original CEB.

To account for the different characteristics of the ground and flight detectors, and to achieve the best possible performance of the newly repaired WFC, the ACS instrument teams at Goddard Space Flight Center (GSFC) and STScI devised an ACS-R Optimization Campaign (OC) that would be conducted early in the Servicing Mission Observatory Verification (SMOV) period. The OC originally comprised seven distinct tests of CCD performance that were conducted wholly or partially over a maximum of eight iterations spanning approximately three weeks. The number and order of iterations and the planned duration of the OC changed as events and scheduling constraints evolved. For reasons described in §2.2, the OC as executed in May and June 2009 was truncated after two iterations. The rationale and strategy of the OC are described at length in Technical Instrument
2. The ACS-R Optimization Campaign

2.1. Test Plan and Timeline

The principal goal of the ACS-R Optimization Campaign was to converge upon a configuration of the CEB-R that yields WFC performance that matches or exceeds the performance achieved with the original CEB. This performance was assessed using the previously supported A/D gain settings of 1.0 $e^-$/DN and 2.0 $e^-$/DN, and with two types of correlated double-sampling (CDS) of the video output from the external preamplifiers. The CEB-R offers both the Clamp-and-Sample (C&S) and Dual-Slope Integrator (DSI) types of CDS, whereas the original CEB offered only C&S. DSI generally provides lower read noise than C&S, so DSI was the presumed default type of CDS for post-SM4 ACS operations.

The metrics by which the WFC performance was judged are read noise, linearity, average full-well depth, charge transfer efficiency (CTE), and amplifier cross-talk. Each iteration of the OC contained a diverse set of bias, dark, and internal flat-field exposures obtained with different A/D gains, CDS types, and clock rates that collectively formed a general performance test at the initial voltage and timing settings established for that iteration. Each iteration also contained a test to verify that the internal ASIC signals were properly phased with the incoming video signal. The iterations also contained various combinations of five specific tests (pre-amp settling, serial clock feed-through, bias voltage optimization, clock voltage optimization, and science data transmission), which were designed to explore the effects of successive and/or simultaneous adjustments of a small number of voltage or timing settings. Only two iterations contained all five specific tests; the others featured only one or two of the tests. Figure 3 shows the test plan and timeline as intended for a maximal 8-iteration OC and the actual STS-125 launch date of 11 May 2009. A detailed description of the test plan and timeline is given by Golimowski et al. (2010).

2.2. Results of the Optimization Campaign

Iteration 1 of the OC was executed on 28–29 May 2009. All general and specific tests were performed using the default voltages and timing patterns used at the time of the Side 2 failure in January 2007. The operating temperature of the CCDs was also set to the pre-failure value of $-81^\circ$ C. Iteration 1 thus provided an extensive set of baseline calibration images that could be compared directly with images obtained with the original CEB. Table 1 lists the performance characteristics of the WFC as measured in January 2007 and May 2009, and compares them against the predicted post-SM4 performance based on ground tests of the CEB-R and continued exposure to HST’s radiation environment. Also listed are the
values for each characteristic that, if realized, would be problematic for science conducted with the ACS WFC and that would require special pre- or post-exposure processing to mitigate their effects (Gilliland et al. 2008).

Table 1 shows that the measured post-SM4 performance of the WFC after Iteration 1 matches or exceeds the predicted performance, except in the cases of dark current and CTE. We attribute these exceptions to ill-constrained projections of radiation effects, as discussed in §3. The significantly lower read noise obtained with the DSI confirmed the pre-SM4 expectation, and justified the adoption of the DSI as the default CDS mode for all post-SM4 science exposures. The electronic components of the DSI do, however, produce artifacts in the WFC bias frames that are not present when using the C&S mode (see §4). Fortunately, these artifacts are either sufficiently stable or predictable to allow their removal during routine data reduction and processing.

Figure 4 shows the photon transfer curve (Janesick 2001) obtained for WFC quadrant D after Iteration 1. The curve shows excellent linearity of response up to the nominal full well depth of $\sim 80000 \ e^{-}$ and is representative of the performance of all four WFC quadrants. The average A/D conversion factor among all four quadrants is $2.036 \pm 0.082 \ e^{-}/DN$ for the commanded default setting of ATODGAIN=2.0.

Although the general performance tests conducted after Iterations 1 and 2 of the OC showed that the functionality of the WFC was fully restored by the CEB-R, one of the specific tests revealed an anomaly that prevented further testing during the OC. In the bias voltage optimization test (labeled “D” in Figure 3), the reset drain voltage ($V_{OD}$) of each CCD was offset by $\pm 1 \ V$ to explore the local minimum read noise and the local maximum full well depth. Previous ground tests showed that these offsets lay within the safe operational range of the flight-spare WFC unit. However, when the on-orbit value of $V_{OD}$ was offset by $-1 \ V$, one of the CCDs (WFC2) exhibited anomalous behavior analogous to charge injection. Proper functionality of WFC2 was restored when $V_{OD}$ was reset to its default value. This behavior had not been seen in earlier tests of flight- or engineering-grade WFC CCDs, and it was not reproducible in subsequent tests with the flight-spare WFC unit. Because the default performance of WFC already met the requirements for post-SM4 science operations, the OC was truncated after discovery of the $V_{OD}$ anomaly to avoid any risk of compromising subsequent WFC science observations.

### 3. Effects of Prolonged Radiation Exposure

The dark current, CTE, and number of hot pixels in ACS’s CCDs deteriorate steadily over time because of continuous exposure to electrons and ions in HST’s trapped radiation environment. The growth of the hot pixel population has been slowed by annealing the CCDs monthly to $\sim 20^\circ \ C$. The anneals restore a significant fraction of new hot pixels to normal operation, but they do not heal the defects in normal pixels that are responsible
## Table 1: WFC performance summary

<table>
<thead>
<tr>
<th>Characteristic</th>
<th>Pre-SM4 (measured)</th>
<th>Post-SM4a (predicted)</th>
<th>Problematica (measured)</th>
<th>Post-SM4 (measured)</th>
</tr>
</thead>
<tbody>
<tr>
<td>Read noise</td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>(e−; ATODGAIN=2.0)</td>
<td>5.5</td>
<td>4.0</td>
<td>10</td>
<td>4.0–5.7</td>
</tr>
<tr>
<td>Dark current</td>
<td>10.7</td>
<td>15</td>
<td>100</td>
<td>20–25</td>
</tr>
<tr>
<td>(e−/pix/hr)</td>
<td>0.68</td>
<td>1.1</td>
<td>1.5</td>
<td>1.1</td>
</tr>
<tr>
<td>Hot pixels (%)</td>
<td>&lt; 0.1</td>
<td>&lt; 0.1</td>
<td>∼0.5</td>
<td>&lt; 0.2</td>
</tr>
<tr>
<td>Full well depth (e−)</td>
<td>84000</td>
<td>84000</td>
<td>40000</td>
<td>&gt; 80000</td>
</tr>
<tr>
<td>Non-linearity (%)</td>
<td>0.02</td>
<td>&lt; 0.1</td>
<td>&lt; 0.2</td>
<td>0.02–0.30</td>
</tr>
<tr>
<td>CTE (EPER; 1620 e−)</td>
<td>0.999949</td>
<td>0.999921</td>
<td>0.9999</td>
<td>0.99989</td>
</tr>
<tr>
<td>Crosstalk (50 ke− source)</td>
<td>4 × 10−5</td>
<td>4 × 10−5</td>
<td>4 × 10−4</td>
<td>(5 ± 4) × 10−5</td>
</tr>
<tr>
<td>Bias shift (%)</td>
<td>0.02</td>
<td>&lt; 0.1</td>
<td>&lt; 0.2</td>
<td>0.02–0.30</td>
</tr>
</tbody>
</table>

*Predicted and problematic values from Gilliland et al. (2008). Dark current and CTE predictions are based on limited data since the change in operating temperature in July 2006.

*Signal-dependent shift of pixel bias due to changing video offset in high-pass AC filter. The effect can be corrected by a software algorithm (see §6).
Figure 4: Photon transfer curve for WFC quadrant D, showing excellent CCD linearity and a nominal full well depth. The curve typifies the performance of all four quadrants.

for increasing the mean dark current. In an attempt to improve science efficiency, STScI decreased the duration of the monthly anneals from 12 hr to 6 hr in late-2004. This change had the undesired effect of increasing the rates of hot pixel and dark current growth. To counter these trends and to improve CTE, the operating temperature of the WFC CCDs was lowered from $-77^\circ$ C to $-81^\circ$ C after the switch to ACS’s Side 2 electronics in July 2006 (Sirianni et al. 2006). The combination of lower CCD temperature and 6 hr anneal periods remained in effect until the Side 2 failure in January 2007. After SM4, the CCD temperature was maintained at $-81^\circ$ C, but the anneal period was restored to its original 12 hr duration.

Table 1 shows that the number of WFC hot pixels measured immediately after SM4 matched exactly the predicted values, despite ACS’s complex anneal and temperature history. On the other hand, the measured dark current and CTE were significantly worse than the predicted values. This contradiction suggests that the predictions suffer large uncertainty because of the changing anneal durations and CCD temperatures, and/or that HST’s radiation environment has become more intense since the Side 2 failure in January 2007. There is evidence for both conditions. Figure 5 shows the mean dark current of one WFC CCD measured from SM3B in March 2002 to SM4 in May 2009. Given the relatively few dark current measurements after the temperature change in July 2006 and the large intrinsic scatter of those measurements, the discrepancy between the predicted and measured post-SM4 dark current is hardly surprising. However, Massey (2010) has recently suggested...
that the density of charged particles in HST’s low-Earth orbit may have increased after the solar cycle reached its maximum in 2007, so the dark-current discrepancy may have an external physical cause.

The post-SM4 predictions and measurements of CTE also suffer from ambiguity and uncertainty. Table 1 indicates that the actual post-SM4 CTE, measured via the Extended Pixel Edge Response test (EPER; Janesick 2001) with a flat-field illumination of 1620 $e^-$, is worse than both the predicted (0.999921) and problematic (0.9999) values given by Gilliland et al. (2008). However, the accuracy of the EPER technique depends on several conditions, including the uniformity of the illumination (which governs the accuracy of the determination of the intrinsic pixel signal) and the spatial sampling of the deferred charge trail. Neither of these conditions is ideal for on-orbit testing of WFC, especially as the CTE is now sufficiently poor that much of the deferred charge extends beyond the 20 rows of virtual overscan. Fortunately, CTE degradation can also be measured from the relative changes in deferred charge trails from hot pixels over time. Figure 6 tracks the magnitude and shape of the trails for hot pixels of varying intensity since 2002. The trends are shown for both vertical (parallel) and horizontal (serial) charge transfers. The CTE steadily worsened from December 2002 to June 2006, after which the CCD temperature was lowered to $-81^\circ$ C. This temperature change effectively restored the vertical CTE (the dominant component) to its December 2004 level. The predicted post-SM4 values (solid black points) are highly uncertain because they are based on only two measurements obtained in mid- and late-2006. Nevertheless, the actual post-SM4 measurements of hot pixel trails (open squares and triangles) indicate that both the vertical and horizontal CTE fall in the “good corners” of the $1\sigma$ error boxes surrounding their predicted values. Thus, the hot pixel test – which is arguably more indicative than the EPER technique – shows that the post-SM4 CTE measurements are completely consistent with the pre-SM4 trends extrapolated forward 2.5 years to May 2009.

Consistency of predictions aside, ACS’s CTE has degraded to a degree that corrective action must be taken to achieve accurate photometric, astrometric, and morphological analyses of science images. STScI has developed two techniques for correcting systematic
Figure 6: Evolution of WFC horizontal (serial) and vertical (parallel) CTE as a function of time in orbit. The plots track the amounts of deferred charge from 1000 $e^-$ hot pixels (ordinates) and the decay profiles of the deferred-charge trails from all hot pixels (abscissae) at discrete epochs (colored points). The solid black points represent the predicted post-SM4 values extrapolated from the two measurements (July and December 2006) obtained after the CCD temperature was lowered to $-81^\circ$ C. The open squares and triangles represent the measured post-SM4 values with and without removal of the horizontal bias stripe artifacts, respectively (see §4).

errors from poor CTE, one aimed at point-source aperture photometry (Chiaberge et al. 2009) and the other at pixel-based image restoration (Anderson & Bedin 2010; Anderson 2010). Both techniques have been tested against “truth images” obtained shortly after the installation of ACS in March 2002 (at which time the CCDs had pristine CTE), and they yield mutually consistent results for signal levels above $\sim 20 e^-$. Refinement of the pixel-based CTE correction at lower signal levels is underway. The algorithm will be publicly released first as a standalone STSDAS package by December 2010 and then as part of the calacs package in the OPUS data pipeline by September 2011.

4. Post-SM4 Bias and Dark Frames

Figure 7a shows the median of 34 WFC bias frames obtained between monthly anneal cycles in mid-2009. Besides the familiar blocked columns and discontinuous gains and biases across the quadrant boundaries, there are two new electronic artifacts imposed by the CEB-R. The more prominent artifact is a bias gradient across each CCD quadrant that starts at the nominal bias level near the readout amplifiers (i.e., the corners of the full bias image in Figure 7a) and rises by $\sim 10$ DN at the quadrant boundaries. This gradient is the
combined effect of (1) a slow drift in the offset of video output signal caused by the high-pass filter in the CCD’s external preamplifier, and (2) slightly different gains between the up and down stages of the DSI caused by imperfectly matched feedback resistors associated with each stage. The bias gradient is very stable between frames clocked with the same timing pattern, so it can be removed straightforwardly through subtraction of a standard “superbias” reference image. However, the gradient strongly depends on the timing pattern of the CCD readout, so distinct superbias reference files must be created to calibrate each of the full-frame and subarray readout modes supported for WFC science observations.

A second artifact of the CEB-R seen in Figure 7a is a low-amplitude (±1 DN) vertical modulation of the bias level that imposes uniform horizontal stripes over the entire image. These stripes are caused by 1/f noise in the bias offset imposed upon the DSI output signal by the SIDECAR ASIC after each parallel transfer. Although the noise per frame imposed by the stripes (σ ≈ 0.75 e−) is small compared with the amplifier read noise (4–5 e−), the constancy of the stripes along each row and the 1/f vertical power spectrum present a highly correlated structure that may interfere with the analysis of astronomical images near the noise floor. Because the bias stripes are not constant between successive frames, they cannot be removed though subtraction of a superbias reference frame. STScI has developed an algorithm for mitigating the bias stripes that is particularly effective for sparsely filled scenes. This algorithm is now available as a standalone STSDAS task, and will be incorporated into the calacs package of the OPUS data pipeline in September 2011. A detailed description of the stripe removal algorithm is presented in this volume by Grogin et al. (2010).

Figure 7b shows the median of 24 long-exposure (~ 1000 s) dark frames obtained between the same anneal cycles that yielded the superbias image in Figure 7a. The bias gradient in this “superdark” has been effectively removed by subtraction of the superbias image. The average dark current increases uniformly from the serial register to the opposite edge of the CCD because of the increasing deferred charge from hot pixels and cosmic ray events due to degraded vertical CTE. The large-scale horizontal structure in the superdark is a combination of the residual 1/f noise in the bias offset and the intrinsic variation in the CCDs’ dark current. Consequently, the bias stripe removal algorithm of Grogin et al. (2010)
cannot be applied to dark frames. Nevertheless, a superdark generated biweekly from 24 individual dark frames has sufficiently small residual stripe noise to accurately calibrate nearly all ACS science images.

5. Amplifier Crosstalk

Because the four WFC quadrants are read out concurrently, electronic crosstalk between the output amplifiers can occur. An imaged source in one quadrant may appear as a faint, mirror-symmetric ghost image in the other quadrants. The ghost image is often negative, so bright sources in the “offending” quadrant will appear as dark depressions on the “victim” quadrants. Table 1 shows that the post-SM4 crosstalk between all pairs of amplifiers is consistent with pre-SM4 levels for ATODGAIN=2.0 and offending signals of $\sim 50000$ e$^-$. However, pre-SM4 WFC images exhibited significant crosstalk from offending signals of a few hundred e$^-$ (Figure 8a). Similar exposures obtained after SM4 showed no evidence of this low-level crosstalk at ATODGAIN=2.0 (Figure 8b).

Analysis of hot pixels in long-exposure dark frames obtained with ATODGAIN=1.4 show that the crosstalk is comparable to that measured with ATODGAIN=2.0. On the other hand, analysis of ATODGAIN=1.0 dark frames shows that crosstalk ghosts from saturated offending sources are as much as 15–20 times stronger than their ATODGAIN=2.0 counterparts and more than an order of magnitude stronger than their pre-SM4 ATODGAIN=1.0 counterparts. Consequently, ACS users are advised against using ATODGAIN=1.0 unless their science goals demand it. Note that STScI only provides calibration reference files for ATODGAIN=2.0, so ACS users who desire other gain settings must include a calibration plan as part of their science program.

An algorithm for correcting pixels affected by crosstalk has been developed by STScI and is currently available as a standalone IDL procedure. STScI plans to incorporate
this algorithm in the \textit{calacs} package of the standard OPUS data pipeline by September 2011. Users should contact the STScI Help Desk for more information about this crosstalk correction algorithm.

6. Signal-Dependent Bias Shift

The idiosyncracies of the external preamplifiers and the DSI that cause the bias gradients described in §4 also conspire to produce small bias shifts that depend on pixel brightness and that decay over many serial clock cycles. The high-pass AC filter in the preamplifier causes the video offset to vary proportionally with pixel signal. If several bright pixels are successively encountered during read out, then the average video offset of the AC filter rises; if several dark pixels are successively read out, then the average video offset drops. This changing video offset is sensed by the DSI, which subtracts the integrated analog reference signal for each pixel from the integrated analog pixel signal. Because the resistors that govern the gains of the two integration phases are imperfectly matched, the drift in the video signal offset between the two phases is amplified proportionally with the pixel signal.

The resultant effect of these electronic idiosyncracies is a pixel-to-pixel bias shift that is 0.02–0.3\% of the average signal from the preceding pixels in the serial clocking sequence. This range reflects the variation of the effect among the four WFC quadrants; the effect is stable within each quadrant. Moreover, the duration of the bias shift depends on the time constant of the AC filter associated with each quadrant’s preamplifier. For most science programs, the bias shift has a negligible impact on image photometry or morphology, but programs that involve high-precision astrometry or high-contrast scenes (e.g., faint nebulae surrounding bright stars) may be adversely affected. Fortunately, the electronic behavior of the bias shift is well-understood and can be removed during image processing with an analytical algorithm developed by Markus Loose. Preliminary tests of the uncalibrated algorithm show that the bias shift is effectively removed to within 1 DN for all pixel signals. STScI will refine the calibration of the bias-shift correction during Cycle 18 and will incorporate it into \textit{calacs} and the OPUS data pipeline by September 2011.

7. Summary

The ACS CCD Electronics Box Replacement (CEB-R) installed during Servicing Mission 4 (SM4) has fully restored the functionality of ACS’s Wide Field Channel (WFC). Unfortunately, the High Resolution Channel was not recovered during SM4. The Solar Blind Channel was not affected by the WFC failure and has remained operational throughout ACS’s on-orbit service.

A campaign to optimize the performance of the WFC under the command of the CEB-R showed that the WFC’s performance matches or exceeds its pre-failure levels, notwithstanding the expected increases in dark current and hot pixels and the decline in charge transfer efficiency (CTE) due to continued exposure to \textit{HST}’s radiation environment. Because of an electronic anomaly associated with the WFC2 CCD’s reset drain voltage (\(V_{\text{OD}}\)), the optimization campaign was truncated, and WFC has since been operated with the same voltages and timing patterns used before the failure of the original CEB in January 2007.

Several artifacts attributed to the CEB-R appear in post-SM4 WFC images, but all can be effectively removed or mitigated to enable uncompromised scientific analysis. Bias gradients of ~ 10 DN are present in all four WFC quadrants, but they are stable and can be removed completely by subtraction of a “superbias” reference image during normal image processing. Low-amplitude (±1 DN) but temporally variable stripes appear along the rows of all WFC images. These stripes contribute less than 1 e\(^{-}\) to the electronic noise budget, but their variability and 1/\(f\) power spectrum prevents their complete removal during standard image processing. The local bias level also has a small dependence on the
average signal from the preceding pixels in the serial clocking sequence. Finally, the amplifier
cross-talk at high signal levels ($\sim 50000$ $e^-$) is substantially higher than pre-SM4 levels for
ATODGAIN=1.0, but it is comparable to pre-SM4 levels for the default ATODGAIN=2.0.
At signal levels of a few hundred $e^-$, the significant cross-talk seen in pre-SM4 images is no
longer present in post-SM4 images.

STScI has developed algorithms for the correction or mitigation of the new electronic
artifacts and the restoration of images affected by continuously degrading CTE. Standalone
correction packages are now publicly available for bias-stripe mitigation and cross-talk cor-
rection; other standalone packages for bias-shift and pixel-to-pixel CTE correction will be
available by December 2010. All these corrective algorithms will be incorporated into the
\textit{calacs} package of the OPUS data pipeline by September 2011. Users should contact the
STScI Help Desk for more information about any of this corrective software.

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An Empirical Pixel-Based CTE Correction for ACS/WFC

Jay Anderson

Space Telescope Science Institute

Abstract. This presentation summarizes a paper that has been recently published in PASP, Anderson & Bedin (2010). The paper describes our pixel-based approach to correcting ACS data for imperfect CTE (charge-transfer efficiency). We developed the approach by characterizing the size and profiles of trails behind warm pixels in dark exposures. We found an algorithm that simulates the way imperfect CTE impacts the readout process. To correct images for imperfect CTE, we use a forward-modeling procedure to determine the likely original distribution of charge, given the distribution that was read out. We applied this CTE-reconstruction algorithm to science images and found that the fluxes, positions and shapes of stars were restored to high fidelity. The ACS team is currently working to make this correction available to the public; they are also running tests to determine whether and how best to implement it in the pipeline.

1. Introduction

The harsh radiation environment of space takes a toll on CCD detectors. It is well known that when high-energy protons, neutrons, and ions impact the detector they generate cosmic-ray pocks and trails. Particularly high-energy events not only displace electrons, but they can even displace silicon atoms, creating a defect in the CCD lattice. These defects cause two problems. Depending on where the defect is, it can create a short-circuit in the detector and generate a steady stream of electrons in a pixel. This is known as a “warm” or “hot” pixel — pixels with unusually high dark current.

Defects can also trap electrons during the readout process. The trapped charge ends up getting disassociated from its original pixel and eventually gets released into another pixel. The result is that the original image gets smeared out in the direction away from the readout amplifier. This is known as imperfect CTE (charge-transfer efficiency), or alternatively as CTI (charge-transfer inefficiency). Figure 1 shows a typical short exposure taken with ACS after SM4. The streaks emanating downwards from the bright objects correspond to charge that should have been read out with the objects, but got delayed because of the readout process.

One way to think of CTE losses is to imagine little gremlins in the CCD (see Figure 2). They like to grab and stare at bright, shiny electrons as they are shuffled past. After a little while, they get bored and put the electrons back on the conveyor belt, only to grab others later. The result is that some fraction of the charge ends up in upstream pixels.

This can cause problems for image analysis. If we are measuring flux through an aperture, some of the flux from the source may end up getting delayed so much that it ends up in a pixel that is outside of the aperture. This causes stars to appear to be fainter than they really are. Over the years, the Institute has come up with a great many empirical corrections to account for these losses (see Cawley et al. 2001, Biretta et al 2005, Riess & Mack 2004, Chiaberge et al. 2009), but these corrections are complicated by the fact that they depend sensitively on the size of the aperture, the flux of the source, the date of observation, the level of the background, etc. The delayed charge also causes severe...
problems for astrometry and shape-based types of image analysis, such as weak lensing (see Kozhurina-Platais et al. 2007).

For all these reasons, it has long been a desire to come up with CTE corrections that can correct the smearing in more than just a parametrized way. Ideally, we would like to correct CTE at the pixel level.

2. Previous CTE Efforts

Given how many projects would benefit from such a pixel-based correction, there have been may attempts over the years to come up with an algorithm to restore images to their pre-smeared state. Bristow & Alexov (2002) and Bristow (2003a and 2003b) and Goudfrooij (2006) made several attempts to treat CTE at the pixel level in STIS. The models produced nice images, but were not able to restore all the flux, and as such, the current stat-of-the-art is still to use empirical parametric corrections. Furthermore, the corrections took so long that it was not practical to use them for large numbers of images. Riess (2000) made similar attempts with WFPC2 and found that he was able to restore galaxy shapes, but again the procedure took prohibitive amounts of time.

Recently, Massey et al. (2010) developed a CTE correction based on a study of the warm pixels (WPs) in the COSMOS science data. They used lab-data to model the profile shape and empirically determined the amplitude of the tails as a function of the warm-pixel intensity. Their correction did a nice job removing the trails in their images. Unfortunately, since their images all had backgrounds of $50 \, e^-$, they were unable to probe the impact of CTE losses when the background is lower. Also, Massey et al. did not have access to true fluxes and positions for the objects they corrected, so they were unable to determine whether their restoration of the flux in the observable trails really did restore all the flux that was lost. Nevertheless, their paper was a great step forward. It demonstrated that...
a pixel-based correction could produce a clear aesthetic success in a reasonable amount of
time, with the hope of a quantitative success with more effort.

3. Warm Pixels in the Darks

With this in mind, we decided to extend the Massey et al. warm-pixel study to the darks. The
darks should have the lowest possible background and should allow a more comprehensive
study of the trails. Furthermore, a great many dark exposures can be stacked together
to remove any read-noise-related limitations. A small portion of this dark stack is shown
in the right panel of Figure 3, with a warm pixel identified. The trails are striking in this
high signal-to-noise stack.

We first identified the warm pixels as being those pixels that were brighter than their
nearby neighbors in 120 out of 168 dark exposures. We then extracted their trails by taking
their upstream pixels and subtracting a local “sky” value, taken from the set of ±5 pixels
in the same row. The left panel of Figure 3 provides a schematic of our set-up.

In Figure 4, we plot the average trail upstream of the brightest warm pixels (those with
about 10,000 DN$_2$, or 20,000 $e^-$). We have isolated the WPs that are farthest from the the
amplifier in an effort to highlight the CTE signal. These warm pixels tend to have about
100 DN$_2$ in their first upstream pixel, which is about 1% of the WP intensity. The next
pixel up tends to have about 50 DN$_2$. Subsequent pixels have fewer and fewer counts, but
there is appreciable flux in the pixels all the way out to 60 pixels.

The red curve is an attempt to fit this trend with a dual-Gaussian (as has been suggested
by some laboratory research). A dual-Gaussian is able to match the inner parts and outer
parts, but not the middle part. It may be that three Gaussians could represent this curve,
but in this work we will simply adopt an empirical profile for the trail.

Figure 5 shows the profiles for several different WP-intensity levels. The lower-right
panel is the same we examined in Figure 4, and the dotted blue curve shows this trend in
Figure 3: The schematic on the left shows how we measured the profile upstream of each warm pixel. The WP is identified (as seen in the figure to the right), and we measure the flux in the Y-upstream pixels above a local sky (found from the surrounding 10 pixels in each row). The hook to the left seen in the WPs is from X-CTE, or serial-CTE, and the differences ($XU_{n} - XD_{n}$) can be used to measure that profile.

Figure 4: The schematic on the left shows the vertical profile of a bright WP. We see 1% of the WP height in the first upstream pixel, 0.4% in the second, etc. The plot on the right shows the entire trail profile. The red curve is an attempt to fit this with a dual Gaussian.
Figure 5: The profiles for four different intensities of warm pixels. Lower right shows the bright ones studied in Figure 4. The lower left shows WPs with an intensity of 1000 DN\(^2\). The upper right panel shows the profile behind WPs of \(\sim\)100 DN\(^2\). Finally, the upper left panel shows the profile behind faint WPs with about 10 DN\(^2\). The blue curve traces the lower-right profile exactly and is repeated in the other panels.

The ACS/WFC was manufactured to have a “mini-channel” or “notch” that would allow the lowest levels of charge to be transferred through the detector without significant CTE losses. Unfortunately, there is no empirical evidence for such a channel. Electron packets with even very small amounts of charge suffer extremely large CTE losses. This is consistent with previous photometric investigations, which have clearly shown that fainter stars suffer much more from CTE than brighter stars. ACS technical lead Golimowski confirms that it is likely that when the mini-channel doping was attempted, it was not stable and ended up mixing in with the rest of the silicon.
4. A Readout Model

Before we can correct images for CTE losses, we must first construct a model that tells us how charge-transfer issues affect the pixels in the first place. We based our model on the trails observed in the warm pixels.

The model has two parameters. The first parameter is \( \phi(q) \), which tells us the chance that a given electron will be trapped during its journey from the top of the chip at \( j = 2048 \) to the serial register at \( j = 0000 \). It goes from about 80% for the first electron in a packet to about 28% for the 100th electron to about 3.8% for the 10,000th electron. The idea is that there are traps within the pixels that can access only certain electrons. Some traps will grab the first electron in a packet, others will grab the 10th, etc. A physical explanation for this can be seen in the simulations in Hardy & Deen (1998). Electron packets containing different numbers of electrons have different cross-sectional areas as they are shuffled through the pixel. The smaller the cross-section of the packet, the fewer traps it will encounter. This is why smaller electron packets see fewer traps total, even though they experience more traps per electron.

The second parameter is \( \psi(q,n) \), and it tells us how long it will take for a trapped electron to be released. Just reading off from the panels in Figure 5, we see that \( \psi \) appears to be a function of both \( q \) and \( n \), meaning that the trail profile appears to depend on the charge level. This would imply that traps that affect different marginal electrons may have different release times. One could imagine that the different traps could be at different distances from the electrodes, and therefore might experience different electric fields during the cycling of the potential that shuffles charge along. It could also be the case that the traps all have the same release profile, but the profile we observe is impacted by some 2nd-order CTE (the charges in the trails are themselves subject to CTE smearing).

Since it would be impossible to determine which traps lie in which pixels, our model assumes that all pixels have the same fractional number of traps that can affect electrons at all levels. We will keep track of which traps in each pixel are filled, empty or partially filled (in the process of being released).

Our aim was to construct a readout model that would reproduce the observed trails from the (presumably) delta-function warm pixels. But before we could construct this model, we needed to answer one final question. When an electron encounters a trap that it has access to, does it always get trapped, or does it only sometimes get trapped? In other words, if there are two consecutive pixels of 100 DN in a column, will the first pixel fill in all the traps that the second could see, or will the first only have some chance of filling the traps such that the second could come through and fill some more of the same traps? In the second case, there would be a trap-filling probability that we would have to determine somehow. This would be difficult to constrain. It would be much easier to assume the first case where traps are deterministically filled completely every time.

To determine which of these assumptions is the case, we examined the cosmic rays in the dark exposures to see whether a CR (cosmic-ray) event with two relatively equal bright pixels had a trail that reflected the brightest pixel, or whether the trail reflected the sum of the pixels. Figure 6 shows that the trails behind CRs are consistent with the simple deterministic-filling case: the trail we see is only as bright as that of the brightest pixel. All traps get filled instantly.

5. Optimizing the Parameters of the Readout Model

Once we had a model for what the readout-process does to the pixel values, we used iterative forward modeling to find what “original” distribution of pixel values could be pushed through the readout process to result in the observed distribution. We used an algorithm
Figure 6: To determine how shadowing may effect trap-filling, we performed an experiment. We looked at the trails behind CRs and compared them against the pixels behind WPs. We found that a CR that had a two consecutive pixels of $\sim 2500$ DN$^2$ or $\sim 5000$ DN$^2$ intensity, showed a trail similar to that seen behind a single WP of the same intensity. A CR that had two pixels that summed up to the WP intensity showed considerably smaller trails.

It was not sufficient to simply read off the parameters for our model from the curves in Figure 5. We started with these values, but found that the 2nd-order CTE effects (the tails themselves suffer imperfect CTE) required us to iterate several times on the form of $\phi(q)$ and $\psi(q, n)$. We ended up parametrizing these functions by tabulating their values at $q=10$, $q=100$, $q=1000$, and $q=10,000$ and used log interpolation in between. Similarly, $\psi$ was tabulated only for $n=1, 2, 3, 5, 7, 10, 15, 20, 25, 30, 40, 50, 60$. We found no evidence for trails that extended beyond $n=60$.

Figure 7 shows the results of our final model. The black curves show the trails behind the warm pixels in the original dark stack, and the yellow curves the trails in the corrected stack. It is clear that our algorithm has removed the average trails.

Figure 8 shows a portion of the dark exposure before and after correction. It is clear that our correction has restored most of the trails to delta-function warm pixels. Not all of the trails are perfectly restored; some are under-corrected and others are over-corrected, but on average the correction is good.

We ran the correction on a 30 s raw science image and show the results in Figure 9. It is clear that the correction does a nice job removing the trails from the stars and warm pixels. It is worth noting that the algorithm doesn’t always do a great job on the cosmic rays, since sometimes the CRs hit the detector during the readout, and as such would not have participated in as much charge shuffling as their pixel location would suggest. There is nothing that can be done for these CRs: they will be blemishes with or without this correction. But it is extremely encouraging that the correction works so well on the rest of the science-image pixels.

Many corrections in the past have managed to pass this aesthetic test, though those routines typically took a prohibitive amount of time to correct the images. Our corrections
Figure 7: The profiles for WP trails in the original (black) and corrected (yellow) images. The profiles are shown for eight logarithmically spaced bins to show that the trends are corrected in between the node-points of the solution.
Figure 8: (Left) The stack of the dark exposures before correction. (Right) the stack of the dark exposures after correction.

here take about 10 minutes on a typical 3 GHz workstation. The real test, though, is how well photometry, astrometry, and shape are preserved. We will examine these in the next section.

6. The Rubber Meets the Road

We note that the algorithm as previously described was developed solely to restore the trailed WPs to delta functions, so we have not optimized anything thus far for the stellar case. The fact that the algorithm restores the flux in the visible trails to the apparent source does not necessarily imply that the algorithm will restore photometry, astrometry, and shape in a quantitative sense. It is possible that the trails are longer than we can perceive in the image (i.e., longer than 60 pixels in this context). If this is the case, then there will be flux lost from the object that we will not restore with our algorithm.

In the PASP article (AB10), we perform several tests to show that the algorithm does in fact restore flux, position, and shape. We do not have room here to show all the results, but will summarize them here.

6.1. The Standard CTE Tests

The standard way CTE is evaluated in the ACS/WFC detector takes advantage of the fact that the readout of the top chip is done in a different direction on the sky from the readout in the bottom chip. This allows us to take an image of the sky, shift the field by one chip height in $y$, then take another image. Stars close to the readout register in the first exposure will be far from the register in the second, and vice versa. Comparing the fluxes of stars in this way ends up doubling the CTE signature.

The left panels of Figure 10 show a comparison of two pairs of 30s images taken in June 2009, soon after SM4. Each dot represents a star. Its horizontal location corresponds
to its $y$ location in one chip. For its vertical position, we plot the difference between its
photometry in the two offset exposures. The green line reports the slope, which is 6% for
the brightest stars ($25,000 \, e^{-}$ total counts) to over 40% for the faintest stars shown ($250 \, e^{-}$
total). (The actual CTE suffered from an individual image is half this.) The right panels
show the same photometry for the same stars, but as measured in the corrected images. It
is clear that the remaining trend is essentially zero everywhere.

6.2. Summary of other tests run

We perform some additional photometric and astrometric tests in AB10. In particular, we
compare some short 10 s and 100 s images taken of the 47 Tuc calibration field against 1200
s images taken immediately afterwards. The 10s images had a background of about $2 \, e^{-}$,
the 100 s images had a background of about $15 \, e^{-}$, and the 1200s images had a background
of about $150 \, e^{-}$. We find that before correction the photometry loss can be as large as 0.5
magnitude for stars with low flux in the short (low-background) exposures. After correction,
the residuals are reduced considerably, with any remaining errors in the correction being
less than about 20% of the correction itself (see Figures 14 and 15 in AB10).

We also compare the positions of stars as measured in the corrected and uncorrected
short-deep pairs of images. We find that the uncorrected positions show CTE-related offsets
of up to 0.3 pixel for the faint stars in the short exposures relative to the deep exposures.
The residuals for the corrected short-deep position comparisons are close to zero (also in
Figures 14 and 15 in AB10).

In Figure 11 we compare the net photometric corrections that we obtained with the
parametrized corrections that Chiaberge et al. (2010) would have suggested for the uncor-
rected stars. The trends agree nicely for the two backgrounds explored.

An additional test we ran concerned the shape of stars. We took a short-deep pair of
images of the 47 Tuc calibration field. We identified known stars near the gap (and therefore
far from the readout amplifier) and examined their shapes as a function of brightness in four
different exposures: the corrected deep, the uncorrected deep, the corrected short, and the
uncorrected short. This test is documented in Figure 18 of AB10. Figure 12 here illustrates
one of the specific cases examined.
Figure 10: Results of the standard chip-shift CTE test for before the correction and after the correction. Each symbol represents one star measured in the two different exposures. Its horizontal location here represents its \( y \) location in the first exposure. Its vertical location represents the difference in apparent flux for the stars in the two images. The green line is a fit to the slope (which is reported at the top, in mag/4096 pixels). The slope on the left represents \( 2 \times \) the standard CTE trends, since both frames suffer CTE losses. The listed slopes on the right show that the trends are largely removed.
Figure 11: Comparison of our net correction with the parametrized correction from Chiaberge et al (2010) for ultra short and short exposures. The Chiaberge corrections correspond to a 3-pixel-radius aperture and the sky background as listed. The open circles come from taking the difference between the photometry from the listed short exposure and a deep 1200s exposure. The black dots correspond to the same difference after our correction has been applied. The error bar shows the scatter about the average.

Figure 12: In each panel, the red line corresponds to the vertical profile of the PSF, as measured from bright stars in corrected deep (339s) images. The left panel shows the profiles of several stars in the corrected short (30s) exposures. The stars selected here have about 75 $e^{-}$ in their centermost pixel, and about 300 $e^{-}$ total. The corrected stars appear to follow the PSF shape nicely. The middle panel shows the profile for the uncorrected stars. It is clear they lose flux from the leading edge and gain flux on the trailing edge. The right panel shows the correction explicitly.
6.3. Future plans

More details of this algorithm can be found in AB10. The Institute is currently working to make this correction available to the public via a stand-alone PYRAF task. It is also evaluating how this new understanding and treatment of CTE might lead to improvements in the pipeline, including how best to deal with the dark exposures. One of the limitations of the current effort is that the darks that were used were all about the same length: 1000 s. There are so many WPs that the faint WPs (\(\sim 10 \text{DN}_2\)) are beginning to crowd each other, making it difficult to explore CTE issues at the lowest electron-packet levels. To remedy this, a new set of shorter darks has been taken and will help us explore CTE at the lower charge levels.

Finally we note that AB10 also explores CTE smearing in the serial (x) direction and finds that it is appreciable, but the tails are so short that it should have no significant effect on photometry. Astrometry, however, will be affected at the 0.01-pixel level. AB10 also discusses ways to prevent read-noise from getting amplified by the CTE correction.

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Red Leak Effects in Observations of Solar System Objects with ACS/SBC

Paul D. Feldman  
Department of Physics and Astronomy, Johns Hopkins University, 3400 North Charles Street, Baltimore, MD 21218

Harold A. Weaver  
Space Department, Johns Hopkins University Applied Physics Laboratory, 11100 Johns Hopkins Road, Laurel, MD 20723

Joachim Saur  
Institute of Geophysics and Meteorology, University of Cologne, Cologne, Germany

Melissa A. McGrath  
NASA Marshall Space Flight Center, Huntsville, AL 35824

Abstract.

Following the failure of STIS in August 2004, attempts to obtain ultraviolet spectroscopy and photometry of solar system objects shifted to the Solar Blind Channel (SBC) of the ACS. These included spatially resolved spectroscopy using the PR130L prism of comet 9P/Tempel 1 (at the time of Deep Impact) and Europa, and imaging and photometry of the asteroid 21 Lutetia, one of the targets of ESA’s Rosetta mission. Initial estimates of long wavelength (“red”) contamination of the data due to impurities in the FUV MAMA detector suggested that these observations were feasible. Subsequent analyses produced better sampled, more reliable response curves that showed the long wavelength response to be much worse than expected. The data from the Europa and Lutetia observations confirm this conclusion and provide additional quantitative measurements of the magnitude of the “red leak.”

1. Comet 9P/Tempel 1

At the time of the STIS failure, a program of ACS/HRC imaging and STIS FUV spectroscopy had been approved for cycles 13 and 14, spanning the times of approach and impact, on July 4, 2005, of the Deep Impact spacecraft into comet 9P/Tempel 1. A key goal of this program was to determine the CO abundance in this Jupiter-family comet, particularly after the impact when the excavated material from several meters below the surface would harbor the more volatile species such as CO and CO$_2$. With the data on the throughput of the ACS/SBC available at the time, it appeared possible to obtain FUV spectra of the comet using the PR130L prism without serious contamination from long wavelength photons scattered from the solid grains in the ejecta plume (see Fig. 2 of Boffi et al., 2008, for the throughput curve then found in Synphot). Moreover, the throughput of the prism was twice that of the STIS G140L grating and the CO spectrum was well known so that the lack of a slit while observing an extended target was not considered to be a difficult obstacle. In retrospect it was very difficult to untangle the emissions, and only the
imaging data using the F140LP filter were used in the subsequent analysis (Feldman et al., 2006).

2. Europa

A similar rationale was used to justify a program to study the electromagnetic interaction of the Jovian field with the tenuous atmosphere of Europa in June 2008 using SBC prism spectroscopy. Previous observations of the Galilean satellites had been made with STIS using the 2″ wide slit (the maximum satellite diameters were all < 2″) so that monochromatic images of the emissions of O1 at 1304 and 1356 Å were well separated in the dispersed images. Shortly before these observations were made, the report of Boffi et al. (2008) appeared, pointing out that on-orbit observations made with the PR110L prism revealed a much larger sensitivity to optical and near-infrared light of the MAMA detector than previously thought. That report proposed a revised throughput curve that also applied to the PR130L prism and the long-pass filters.

![Figure 1: Spectral images taken with the PR130L prism of Europa (left) and the solar analogue star 16 Cyg B (right). The positions of the disk images of Europa centered at the wavelengths of the atomic oxygen emissions are indicated.](image)

The dispersed Europa image taken with the PR130L prism is shown in the left panel of Figure 1. For comparison, a similar image of the solar analogue star 16 Cyg B is shown in the right panel. The pile-up of long wavelength photons is apparent in both images. Spectral traces along the dispersion axis are shown in Figure 2. In the case of Europa the data are summed over the full disk which reduces the effective spectral resolution. Nevertheless, the decrease in the albedo of Europa below 2000 Å is clearly seen. From the spectrum of 16 Cyg B we find that only 15% of the recorded photons are from wavelengths below 2000 Å. Saur et al. (2010) have used this spectrum, multiplied by a model albedo, and convolved with the full Europan disk, to determine a background for quantitatively isolating the oxygen emission features assumed to originate from Europa’s atmosphere. They also derive a long wavelength throughput curve that is qualitatively in accord with the revised Synphot curves of Boffi et al. (2008).
3. Asteroid (21) Lutetia

In anticipation of the flyby of Lutetia on July 10, 2010, by ESA’s Rosetta spacecraft en route to a rendezvous with comet 67P/Churyumov-Gerasimenko in 2014, HST observations were made in November 2008 with the goal of measuring the asteroid’s albedo from the far ultraviolet to the near infrared. With the ACS visible light cameras not operational, WFPC2 was used in five wide band filters ranging from F218W to F606W. For the far ultraviolet, the SBC was used with the F140LP and F165LP filters so as to produce differential photometry from the difference images in the two filters. Once again the “red leak” dominated the observed images and the analysis of Weaver et al. (2010) found that less than 10% of the detected photons were at wavelengths below 1895 Å or 1975 Å for the F140LP and F165LP filters, respectively. In order to consistently match the observed count rate in both filters and also in the difference image, Weaver et al. (2010) found that the system throughput for \( \lambda > 2000 \) Å had to be increased by a factor of 2.5 from the current Synphot values. The validity of this conclusion will be tested by analysis of the far ultraviolet spectra of Lutetia obtained by the Alice spectrograph on Rosetta during the Lutetia fly-by. These data are now in hand, but at the time of this writing were still being analyzed. The magnitude of the “red leak” thus remains somewhat uncertain, and may be affected by detector temperature and ageing, as suggested by Boffi et al. (2008).

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References

Application of the SIDECAR ASIC as the Detector Controller for ACS and the JWST Near-IR Instruments

Markus Loose
Markury Scientific

Abstract.

The SIDECAR ASIC is a fully integrated controller for high-performance optical and infrared detectors. It combines all functions on a single microchip, including output signal amplification and A/D conversion, bias voltage generation, clock generation, and housekeeping telemetry. The SIDECAR ASIC has been implemented into the CCD electronics box for ACS (during the SM4 repair of the instrument), and it is built into all of the JWST near-infrared instruments. The presentation will give a brief overview of the core capabilities and features of the SIDECAR ASIC, and then focus on the performance and operational aspects relevant to HST/ACS and JWST. Challenges with respect to 1/f noise on the bias voltages and overall strategies for further noise reduction will be discussed. In particular with respect to JWST, flexible data acquisition modes combined with elaborate post-processing have shown to provide significant noise improvements over the baseline approach. In this context, the full in-system programmability (even on-orbit) provides a valuable tool to adapt the operation and configuration of the SIDECAR ASIC to changing conditions.
The Astrometric Context of HST in 2010

William van Altena

Yale University, New Haven, CT

Abstract.

For the past 20 years, the HST has occupied a unique niche in astrometry, that of high-resolution imaging and interferometry-based astrometry. As a consequence of its small field of view, HST astrometry has been primarily oriented towards the observation of carefully selected targets rather than surveys searching for new classes of objects. Ground-based astrometric, photometric and radial velocity surveys have played crucial roles in characterizing classes of objects and identifying astrophysically-interesting targets worthy of observation by the HST. The symbiotic nature of HST and ground-based astrometry is illustrated by surveys of visual and spectroscopic binaries that lead to the identification of binaries in critical parts of their orbits where one or a few HST Fine-Guidance Sensor observations of their angular separation can lead to a definitive mass determination. Photometric surveys have identified numerous Cepheid variables in the Milky Way that might be used to calibrate the cosmic distance scale, however their distances generally remain estimated only from their spectro-photometric characteristics. HST-FGS observations of carefully selected and relatively nearby Cepheids have led to a solid independent calibration of the cosmic distance scale. Ground-based observations of galactic globular clusters have determined the physical makeup of those clusters and characterized the evolutionary tracks of stars with differing masses and metallicities, but only HST imaging astrometry in conjunction with radial velocity observations have made it possible to determine the dynamical state of the clusters. The above examples highlight only a few contributions of HST astrometry to the critical calibration of astronomical objects, while in this talk I will try to place HST in the context of how we are using astrometry to understand our local universe.
Multi-Wavelength Geometric Distortion Solution for WFC3/UVIS and IR

V. Kozhurina-Platais\textsuperscript{1}, C. R. Cox\textsuperscript{1}, L. Petro\textsuperscript{1}, M. Dulude\textsuperscript{1}, J. Mack\textsuperscript{1}

\textit{Space Telescope Science Institute, Baltimore, MD, 21218}

Abstract. The standard astrometric catalog based on ACS/WFC observations of the globular cluster \(\omega\) Cen field has been used to examine the geometric distortion of WFC3/UVIS and IR as a function of wavelength. Multiple observations of this field taken with large dither patterns and a large range of the HST roll-angles were exposed through 10 UVIS and 5 IR band-passes. A 4th order polynomial model was used to derive the geometric distortion coefficients relative to the distortion free coordinates of our astrometric field in the UVIS and IR channels. The main results of this calibrations are: 1) geometric distortion can be successfully corrected at the 0.1 pixels (4 and 10 mas) precision level in UVIS and IR channels, respectively; 2) the non-perpendicularity of coordinate axes (skew) can be used to assess the scale change from filter to filter, time, and HST roll-angle; 3) the coefficients of the geometric distortion in the UVIS and IR channels are used in STScI Multidrizzle software in order to correct WFC3 images for distortion.

1. Introduction

A new instrument, \textit{Wide Field Camera 3} (WFC3), a fourth generation imaging instrument of \textit{HST}, was installed on \textit{HST} during Servicing Mission 4 in May 2009. The Servicing Mission Orbital Verification (SMOV) observations (proposal 11444, and 11445, PI–L. Dressel) with the WFC3/UVIS and IR channels were used to derive the geometric distortion only for one filter, F606W and F160W, in UVIS and IR, respectively. A preliminary optical ray-tracing model demonstrated that the geometric distortion in WFC3/UVIS and IR cameras is severe, on the order of \(\sim 7\%\) across the detector. This distortion in the form of displacement of celestial sources from their true positions on the sky can reach up to about 120 pixels or \(\sim 5''\) in UVIS and 35 pixels, or \(\sim 4''\) in IR. The knowledge of accurate geometric distortion is important not only for deriving accurate positions, parallaxes and proper motions of the scientifically interesting objects but also to rectify the WFC3 images. The Multidrizzle software (Koekemoer 2002, ), currently installed in the STScI on-the-fly pipeline (OTFR), requires accurate distortion correction in order to combine dithered WFC3 images (UVIS and IR), to enhance spatial resolution, and to deepen the detection limit. If the geometric distortion correction implemented in Multidrizzle is not accurate enough, then the WFC3 combined frames can produce blurred images and distort the under-sampled Point Spread Function (PSF). Any significant uncertainty in the geometric distortion correction is detrimental to alignment of WFC3 images with MultiDrizzle and to mitigating the effect of an under-sampled PSF.

The goal of this multi-wavelength astrometric calibration was to derive an accurate geometric distortion in 10 UVIS and 5 IR filters to a precision level of 0.1 pixels (or 4 and 10 mas respectively), which is sufficient to combine dithered and mosaic \textit{WFC3} UVIS and IR images using the STSDAS Multidrizzle software.
Here, we present the analysis and results of geometry distortion calibration, now sufficiently accurate to use in Multidrizzle with observations of the new HST camera WFC3 in UVIS and IR channels.

2. Observations and Reductions

The standard astrometric catalog in the vicinity of globular cluster ω Cen was used for a multi-wavelength geometric distortion calibration of UVIS and IR. The tangent–plane projection type positions (further referred to as the U,V rectangular coordinate system) of stars in the standard astrometric catalog are globally accurate to ∼0.02 ACS/WFC pixels (Anderson & van der Marel, 2010). The globular cluster ω Cen was observed with UVIS and IR detectors near the center of the standard astrometric catalog with large dither patterns and at different HST roll-angles (CAL–11911, PI–E. Sabbi and CAL–11928, PI–V. Kozhurina-Platais). The observations were taken through 10 UVIS filters and 5 IR filters. The supplemental calibration program (CAL–12094, PI–L.Petro) also observed ω Cen with UVIS through F606W filter at different HST roll-angles from –95° to –135° at a step of –10°.

The reductions and analysis of multi-wavelength geometric distortion for UVIS and IR observations are similar to the reductions used in the SMOV UVIS and IR geometric distortion calibration and described in detail by Kozhurina-Platais et al., (2009a, 2009b). At the time of our analysis, there were no high precision tools available such as an effectivePSF library as has been derived for the ACS/WFC camera (Anderson, 2002). Therefore, IRAF task – DAOPHOT/PHOT, which includes a Gaussian fit to the PSF centroid and simultaneous aperture photometry was used to obtain the X & Y positions of stars on each of the UVIS CCD chips for a total of 200 UVIS images. Nevertheless, the IR ePSF library was available (Anderson, 2010) at the time of IR analysis, Thus, the ePSF fitting technique was used to obtain an accurate and high-precision X & Y positions of stars on each of the IR images for total of 90 IR images.

3. Analysis

**Polynomial Solutions.** Similar to the SMOV calibration (Kozhurina-Platais et al., (2009a, 2009b)), a fourth–order polynomial was used to derive the coefficients of geometric distortion for UVIS and IR, namely:

\[
U = A_1 + A_2 X + A_3 Y + A_4 X^2 + A_5 Y^2 + A_6 X Y + A_7 X^3 + ... + A_{15} Y^4 \tag{1}
\]
\[
V = B_1 + B_2 X + B_3 Y + B_4 X^2 + B_5 X Y + B_6 Y^2 + B_7 Y^3 + ... + B_{15} Y^4 \tag{2}
\]

where U and V are the tangent–plane positions in our astrometric standard catalog, and X,Y are the measured pixel positions in the observed UVIS or IR frame. About 40,000 stars were detected in each of the UVIS CCD chip and about 6,000 stars were detected in IR frames from ω Cen observations – a sufficient number to model properly the geometric distortion with a high-order polynomial. The RMS of these solutions were about 6 mas in both UVIS and IR. It is interesting to note, that the RMS of the solution is a factor of three larger than the RMS of 2 mas from the Large Magellanic Cloud (LMC) solution, used in SMOV calibration. A simple explanation for such a high RMS is unaccounted intrinsic dispersion of the proper motions of ω Cen. According to Dinescu et al., (1999), the absolute proper motion of ω Cen are μ_αcosδ=–4.9 mas and μ_δ=–3.5 mas per year. The epoch difference between WFC3/UVIS & IR observations and the standard catalog is at about 4 years. However, a significant contribution to our solutions is the substantial internal velocity dispersion in proper motion of ω Cen. As reported by Anderson & van der
Marel (2010) the internal velocity dispersion in the proper motion of ω Cen is at the level of 0.9 mas per year. Approximately, it scales up proportionally with the epoch difference and thus, in four year contributes to the RMS as much as ∼4 mas or ∼0.1 UVIS pixels.

For each of the 10 observed UVIS filters and for each of the 5 observed IR filters, unique polynomial coefficients were obtained which accurately represent the geometric distortion in UVIS and IR filters.

**Multi–Wavelength Distortion.** The various terms in the general polynomial characterize separate components of the geometric distortion. For example, in the X solution: $A_7$ through $A_{21}$ terms in the general polynomial are classical cubic- and fifth-order distortion terms: the 3rd–order indicate the pin–cushion type of distortion; and the 5th-order terms would indicate the presence of barrel–type of distortion. Specifically to our solution, the 4th-order terms (the terms $A_7$ through $A_{15}$, Eq.1,2) indicate the distortion between pin–cushion and barrel type; while the 2nd order terms, $A_4$ and $A_5$, are the plate tilt terms (van de Kamp, 1967); the linear terms, $A_2$ and $B_3$, are most significant and represent the relative plate scale in X and Y. In order to characterize the filter dependency distortion, the linear coefficients $A_2$ and $B_3$ from F606W UVIS solutions (or F160W IR solution) were chosen as a reference scale. Then, the same linear coefficients from the solution in the other filters were transformed into the F606W linear coefficients, namely:

$$X_{\text{scale}} = (X_{\text{F606W}} - X_{\text{Filter}}) \times S$$
$$Y_{\text{scale}} = (Y_{\text{F606W}} - Y_{\text{Filter}}) \times S$$

where $X_{\text{F606W}}$ and $Y_{\text{F606W}}$ are linear terms $A_2$ and $B_3$ from the UVIS F606W solution (or IR F160W solution) and $X_{\text{Filter}}$ and $Y_{\text{Filter}}$ are linear terms from the appropriate filter solutions and $S$ is the scale which provides the size of the effect in pixels at the far edge of UVIS detector ($S=4000$ pixels) or IR detector ($S=1000$ pixels).

Figures 1 and 2 show the maximum displacement in UVIS1 and UVIS2 CCD chips due to the scale change $w.r.t.$ F606W filter. As seen from these figures, there is a clear correlation in the amount of geometric distortion as a function of wavelength. For example, the difference in the scale in the F225W filter is higher by ∼0.03% than in F606W. However, the scale difference in in F390W filter is ∼0.01% lower than in F606W.

![Figure 1: UVIS1 relative scale displacement $w.r.t.$ F606W filter.](image-url)
The main reason for the variation in plate scale vs. wavelength is variations in the thickness of a filter itself, which is especially significant in the UV filters. Similar filters dependency distortion in WFPC2 filters was reported by Trauger et al. (1995), the WFPC2 wavelength-dependent geometric distortion were computed by analyzing the results of ray tracing. The coefficients were presented as quadratic interpolation function of the relative index $n$ of $\text{MgF}_2$ field-flatten lenses. According to Kozhurina-Platais et al. (2003), the scale difference in F300W filter is 0.5% higher than in F555W, but in F814W is 0.3% smaller than in F555W.

Figure 3 shows the maximum displacement due to the scale change w.r.t. F160W filter in the IR detector. As seen from Fig.3, the IR scale from filter to filter is quite stable and small (with the exception of F098M). It is a clear indication that the amount of geometric distortion from filter to filter remains nearly constant. Therefore, it is safe to adopt the distortion coefficients from the F160W filter for the other uncalibrated IR filters.

Skew parameter. The main parameters, used to quantify the linear part of geometry distortion, are the terms in the general linear transformation between two coordinate systems. As defined in Eq.1 and Eq.2, $U$ and $V$ represent the orthogonal reference coordinate
system, which is free of any systematics in the positions. The $X$, $Y$ coordinate system is the measured frame after applying the distortion correction. Then the linear transformation between these two systems can be expressed as:

$$U = A_1 + A_2 X + A_3 Y$$  \hspace{1cm} (5)$$
$$V = B_1 + B_2 X + B_3 Y$$  \hspace{1cm} (6)$$

This $3 \times 2$–parameters transformation is a simple linear translation between two coordinates systems (Taff 1980, Kiselev 1989). The rotation angle between these two systems then can be defined as:

$$\tan(\theta) = \frac{A_2}{B_3}$$  \hspace{1cm} (7)$$

The skew term, is the amount of non-orthogonality between the two principal X & Y axes and can be derived from the following ratio of linear terms:

$$\tan(\gamma) = -\frac{A_2 A_3 + B_2 B_3}{A_2 B_3 - B_2 A_3}$$  \hspace{1cm} (8)$$

The skew term is one of the principal geometric distortion parameters, which validates the distortion model. The skew will be constrained accurately in the distortion model if the HST observations are taken with different HST roll–angle, but the skew will not be constrained in the distortion model if the observations are taken with the same roll–angle. Also, as it was noticed and discussed by Anderson (2007), the variation of the ACS/WFC linear geometric distortion terms changed monotonically since ACS was installed in 2002.

The skew term, defined above, is the parameter used here for characterization of UVIS geometric distortion. Therefore, we tested the following assumptions:

1. the adopted geometric distortion for the WFC3/UVIS is free of skew;
2. the adopted geometric distortion for the WFC3/UVIS is time–independent.

In order to validate the derived UVIS geometric distortion and possible detect some residual systematics, we used the observations of $\omega$ Cen with different HST roll–angles. The supplemental UVIS calibration program (CAL-12094, PI–L.Petro) was designed so that $\omega$ Cen would be observed through the F606W filter with different roll–angles from $-95^\circ$ to $-135^\circ$ at the step of $-10^\circ$.

Table 1 lists all observations of $\omega$ Cen with different roll–angles (and $\text{POSTARGS} = 0:0$) available to the date (CAL–11452, CAL–11911 & CAL–12094) which were used for characterization of the UVIS geometric distortion.
Table 1: The list of ω Cen Observations.

<table>
<thead>
<tr>
<th>Image ID</th>
<th>Proposal ID</th>
<th>Date-Obs (yy:mm:dd)</th>
<th>WFC3 UVIS Filters</th>
<th>Exp. Time (sec)</th>
<th>Orientat (°)</th>
</tr>
</thead>
<tbody>
<tr>
<td>iaby02x7q</td>
<td>11452</td>
<td>2009-07-15</td>
<td>F606W</td>
<td>35.0</td>
<td>-28.30</td>
</tr>
<tr>
<td>ibc301qrq</td>
<td>11911</td>
<td>2010-01-14</td>
<td>F606W</td>
<td>40.0</td>
<td>149.76</td>
</tr>
<tr>
<td>ibe801nq</td>
<td>12094</td>
<td>2010-04-25</td>
<td>F606W</td>
<td>40.0</td>
<td>-95.25</td>
</tr>
<tr>
<td>ibe802nq</td>
<td>12094</td>
<td>2010-04-25</td>
<td>F606W</td>
<td>40.0</td>
<td>-105.25</td>
</tr>
<tr>
<td>ibe803nq</td>
<td>12094</td>
<td>2010-04-25</td>
<td>F606W</td>
<td>40.0</td>
<td>-135.25</td>
</tr>
<tr>
<td>ibe804ntq</td>
<td>12094</td>
<td>2010-04-25</td>
<td>F606W</td>
<td>40.0</td>
<td>-135.25</td>
</tr>
<tr>
<td>ibe805nvq</td>
<td>12094</td>
<td>2010-04-25</td>
<td>F606W</td>
<td>40.0</td>
<td>-115.25</td>
</tr>
<tr>
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<td>12094</td>
<td>2010-04-25</td>
<td>F606W</td>
<td>40.0</td>
<td>-95.25</td>
</tr>
<tr>
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<td>2010-04-25</td>
<td>F606W</td>
<td>40.0</td>
<td>-105.25</td>
</tr>
<tr>
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<td>2010-04-25</td>
<td>F606W</td>
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<td>-125.25</td>
</tr>
<tr>
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<td>2010-04-25</td>
<td>F606W</td>
<td>40.0</td>
<td>-125.25</td>
</tr>
<tr>
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<td>2010-04-29</td>
<td>F606W</td>
<td>40.0</td>
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</tr>
<tr>
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<td>11911</td>
<td>2010-07-04</td>
<td>F606W</td>
<td>40.0</td>
<td>-35.24</td>
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</table>

The so-called *flt.fits* flat-fielded images, listed above were run through the Multidrizzle software (Koekemoer et al. 2002) in order to correct for the geometric distortion. Multidrizzle software used the new derived multi-filters UVIS geometric distortion coefficients in the form of IDCTAB (Instrument Distortion Correction Table, Hack & Cox (2001)). The X,Y positions from the drizzled images were obtained using the IRAF/DAOPHOT/PHOT task, which includes a Gaussian fit to the PSF and simultaneous aperture photometry. The standard astrometric catalog of ω Cen as reference frame (U,V in Eq. 5–6) and the measured positions (X,Y in Eq.5–6) from UVIS drizzled images were used to calculate the skew term (Eq.8) from the linear transformation.

Figure 4 shows the skew as a function of time over one year of the WFC3/UVIS observations. Figure 5 shows skew as a function of time over 3 consecutive HST orbits (short period of time ~ 5 hours). As seen in Fig. 4 the variation of skew appear to be closer to zero, over one year of WFC3 on board. However, Fig.5 with skew over three consecutive HST orbits of observations shows the linear dependency of a skew on the location of HST on orbit.

One of the factors in the scale change is the velocity aberration, which across the UVIS detector can vary as much as 5 parts in 100,000 (Cox & Gilliland, 2002). The velocity
aberration scale correction factor to the image is available in the header of the science fits file and is used by Multidrizzle. Thus, the skew parameter calculated from from drizzled images are velocity aberration free. Another factor in the scale change is the telescope breathing, which takes place on an orbital time scale and causes an effective change of focal length in the optics. Thus, the linear dependency of scale change on a orbital time scale seen in Fig.5 is the so–called thermal breathing. The temporal variations in the scale change due to the thermal breathing from image to image is at the level of \( \sim 0.02\% \).

As discussed by Anderson (2007), the variation of the skew term for the ACS/WFC has changed monotonically since ACS was installed in 2002. The size of this change is clearly noticeable over 5 years at about 15\,mas off from the original 2002-based distortion solution. This amount of change in distortion is significant for accurate mapping of ACS images with Multidrizzle.

From one year of observations, the WFC3/UVIS geometric distortion is stable within the linear and periodic deviations at about \( \pm 3'' \) due to the thermal breathing. Scaling by 4000 provides the size of the effect of maximum displacement at the far edge of drizzled images at the level of 0.05 UVIS pixels. From one year of WFC3 UVIS observations, there is no evidence of a secular change of the WFC3/UVIS scale.

One of the goals of WFC3 astrometric calibration is to monitor the variation of UVIS and IR distortion over time with a specially designed program so that the standard astrometric field in the vicinity of \( \omega \) Cen will be observed several times per year with a nominal HST roll-angle with the steps of \( \pm 10^\circ \).

**Lithographic pattern.** After applying the best–fitting polynomial, the residuals of X & Y positions between the WFC3/UVIS observed \( \omega \) Cen frame and the astrometric standard catalog are essentially flat, i.e. all large–scale residuals are successfully removed. Nevertheless, there are a noticeable fine–scale systematic pattern in the residuals from our best–fitting polynomial solutions. Figure 6 shows a 2–D post–solution XY residual map between an observed \( \omega \) Cen frame and the astrometric standard catalog. These systematics residuals are typically 0.15 pixels in amplitude and depends on positions of CCD chip.

A complicated and correlated residual structure, placed in fixed–size box seen in Fig.6 is due to the photo–lithography pattern with a size of \( \sim 700\times1000 \) pixels, imprinted onto the detector itself during the manufacturing process. The boundaries of the mask used for photo–lithography are evident. The discontinuities between the vectors coincide with the lithographic boundaries (red lines) as measured in the flat-field images. This very specific complicated structure of residuals shows a typical amplitude of \( \sim 0.1 \) pixels can not be removed by a polynomial model. It would take a polynomial of higher (up to 15th) order to remove these fine-scale variations. The simple way to remove the fine–scale variations is to...
Figure 6: 2-D XY residual map between the UVIS positions corrected for the geometric distortion and the standard astrometric catalog. The top panel shows the XY residuals for the WFC3/UVIS1 CCD chip and the bottom panel shows the XY residuals for WFC3/UVIS2 CCD chip. The boundaries of the photo-lithography mask with size of $\sim 700 \times 1000$ pixels are over-plotted by red boxes. The largest vector is $\sim 0.15$ pixels, magnified by 2500. The units are WFC3/UVIS pixels.

model the fine-scale structure with look-up table, which can be linearly interpolated at any point in the image. This look-up table would be analogous to ACS/WFC and ACS/HRC (Anderson & King, 2002) as a reference file DGEO in Multidrizzle.

4. Conclusion

Here, we present the results of the WFC3 UVIS and IR multi-wavelength geometric distortion calibration. For each of the 10 observed UVIS filters and for each of the 5 observed IR filters, unique polynomial coefficients were obtained, which accurately represent the geometric distortion in UVIS and IR filters. The derived geometric distortion coefficients in the form of IDCTAB can be successfully used in STSDAS/Multidrizzle software for: 1) stacking of WFC3/UVIS and IR images with different dither pattern; 2) rejecting the CRs; 3) enhance the spatial resolution; 4) deepen the detection limit. The geometric distortion
can be successfully corrected at the 0.1 pixels or 4 and 8 mas precision level in UVIS and IR images, respectively.

Summarizing the multi-wavelength geometric distortion for WFC3/UVIS and IR, we conclude that linear scale in the adopted distortion model varies as a function of filters, due to the filters thickness itself. Although, the precision and accuracy of the geometric distortion depends on how accurate and precise is the centering technique for measurement of X and Y positions of under–sample PSF. Because of that, the effective PSF library for each calibrated UVIS filters and PSF fitting technique, developed by Anderson, will allow us to better characterize the WFC3/UVIS multi-wavelength geometric distortion. The next steps in the improvement of UVIS geometric distortion are: 1) implementation of ePSF library and PSF fitting technique for UVIS observations; 2) implementation of the look-up table (DGEQ file) in Multidrizzle software to remove the fine–scale variations due to photolithography patterns in the WFC3/UVIS images. These two implementations will improve the precision level in the UVIS distortion down to 1 mas. However, the precision level in IR distortion mainly depends on how accurate and precise is the centering technique for measurement of X and Y positions of severely under-sample PSF on drizzled IR images.

Summarizing the assessment of the linear part of the WFC3/UVIS geometric distortion, we conclude that for over one year of the WFC3 on HST board, the WFC3/UVIS geometric distortion is time–independent and stable within the linear and periodic deviations at the level of ±0.05 pixels at the far edge of UVIS drizzled images due to the orbital breathing. We will continue to monitor the variation of the WFC3/UVIS and IR distortion over time.

Acknowledgments. We thank Jay Anderson for sharing with his IR effectivePSF library and fitting code. V.K-P. especially appreciates the lengthy discussions with Imants Platais and Andrea Bellini for very valuable comments, which improved some aspects of the WFC3 geometric distortion and the text of this paper.

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Fine Guidance Sensors: Calibration Status

E. Nelan
Space Telescope Science Institute, Baltimore MD, 21218

B. McArthur
University of Texas at Austin, Austin, TX, 78712

Abstract. Servicing Mission Four (SM4) include the installation of a refurbished FGS, designated FGS2r2. The commissioning and calibration of this FGS for guide duty is outlined in this paper. Prior to SM4, the FGS1r interference fringes were re-optimized for the first time since 1999 by an adjustment of the instrument’s articulated mirror assembly. While this adjustment improved the FGS1r angular resolution when operated as a science instrument in Transfer mode, it also changed the FGS1r geometric distortion by a sufficient amount to render the instrument inoperable as a astrometric science instrument. An interim calibration was performed that has restored the FGS to nearly its previously well calibrated configuration, but additional on orbit calibration observations may be required to fully restore the FGS1r astrometric performance.

1. Introduction

The HST pointing control system (PCS) relies upon the Fine Guidance Sensors (FGSs) to provide line of sight measurements of the direction to guide stars to fine point and stabilize the observatory to better than 7 mas, allowing the science instrument to achieve exquisite image quality. The ability to accurately measure a guide star’s position also enables the FGS to serve as a very accurate astrometric science instrument. This paper has two parts. First we discuss the commissioning and partial calibration of FGS2r2 following its insertion into HST during Servicing Mission 4 (SM4). Second we discuss the re-optimization of the FGS1r interference fringes (or S-curves) prior to SM4, and the resultant changes in its geometric distortions, and the re-calibration that was needed to restore its performance as an astrometric science instrument. Background information on the design, operations, and performance of the FGS as an HST guider and as a science instrument can be found in the FGS Instrument Handbook (Nelan, E., et al., 2010).

2. FGS2r2

During SM4, on EVA5 FGS2r was replaced by the refurbished FGS2r2. This was the third time a new FGS was installed in HST. (FGS1r was installed in SM2. FGS2r had been installed in HST during SM3A, but the intensities of several LEDs on its star selector shafts were declining to the point that guide star acquisition failures were expected to become increasingly more prevalent in the future.) The first step to commission an FGS is the optimization of its interference fringes via adjustment of the instrument’s articulated mirror assembly, or AMA (figure 1) to align the FGS interferometric optical elements with the optical axis of the HST, thereby mitigating the otherwise degrading affect of the spherically aberrated primary mirror on FGS performance (figure 2). This involves observations of a
point source (single star) with the FGS in Transfer mode to measure the interference fringes at several (typically 5) widely separated locations in the FGS field of view. These data are used to inform the adjustment of the AMA. Four iterations of the observation-adjustment cycle were planned and executed to achieve the optimal interferometric performance of FGS2r2.

Following the AMA interferometric optimization, the plate scale of FGS2r2 and its alignment relative to FGS1r and FGS3 was measured. FGS2r2 observed seven stars distributed across its field of view in Position mode, while FGS1r and FGS3 acquired guide stars. This was repeated for a total of six orbits, with FGS2r2 observing the same seven stars while FGS1r and FGS3 acquired a different guide star pair for each orbit. The coordinates of the FGS2r2 target stars and the FGS1r & FGS3 guide stars are known to an accuracy of about 30 milli-arcseconds from the US Naval Observatory CCD Astrometric Catalog (UCAC-III). This calibration enabled subsequent use of FGS2r2 for guiding HST. However, the instrument’s geometric distortion was not calibration on-orbit since it was expected (based upon experience with FGS1r and FGS2r) that its early on-orbit evolution (out gassing of water vapor from its graphite epoxy composites) would quickly render early calibrations invalid; STScI decided to defer this calibration until such time that FGS2r2 demonstrated adequate stability (and if needed, after a re-optimization of the S-curves by an AMA adjustment). In the interim, the FGS2r2 design prescription for the geometric distortion has been used. To avoid imprinting FGS2r2 calibration errors on to the HST science products while its geometric distortion is not accurately calibrated, FGS2r2 is not used as the “dominant guider” if possible, rather it is used only to stabilize the roll angle of the observatory.

Since completing its early commissioning in the summer of 2009, FGS2r2 continues to successfully guide HST. Monitoring programs are tracking the relative changes in the FGS2r2 geometric distortion and the evolution of its interferometric fringes. Surprisingly, the changes have been much smaller than expected when compared to the evolution of FGS1r and FGS2r in their first years on orbit (figure 3). This suggests that a second AMA re-optimization will not be needed and the five orbit geometric calibration test can proceed.

Figure 1: FGS Optical Train with the AMA (fold flat mirror 3)
Figure 2: Aligning FGS Interferometric Elements with Telescope Optics mitigates the impact of HST’s spherically aberrated primary mirror on FGS performance. Left, the optical axis of the telescope (orange line) is not aligned with the shear line (blue line) of the Koesters prism; the spherically aberrated wavefront produces additional aberrations (coma & astigmatism) that result in a degraded interferometric response (S-curve). Right, after adjustment of the AMA, the FGS is properly aligned w.r.t the telescope and optimal performance is achieved.

Figure 3: FGS2r2 S-curve Stability During First Year On Orbit observed with “full aperture” (F583W) element. This is remarkable stability compared to the observed changes in FGS1r and FGS2r during their first years on orbit. Note, the change in the vertical “bias” along the X-axis may be due to small field stop migration in the X-channel. This is being monitored over time.
3. **FGS1r**

FGS1r has been used as a science instrument for high angular resolution studies of binary star systems, and for high precision astrometry to determine the distance to astrophysically interesting objects, or to measure the reflex of a star’s hosting planets. FGS1r was installed in HST during SM2 in 1997, and was the first FGS outfitted with the AMA to mitigate the deleterious affects of spherical aberration on its interferometric performance. After two years on orbit it was found to have stabilized sufficiently to warrant a re-optimization of its S-curves via an AMA adjustment (the interference fringes had degraded significantly, but had stabilized). In early 2000 FGS1r was re-optimized and fully calibrated to serve as a high precision astrometric and ultra-high angular resolution science instrument. But alas, the evolution of FGS1r continued for another five years, but in such a way to impact primarily the interferometric response along its y-axis; however, after 2005, no further evolution is evident (figure 4). This prompted a request by the STScI FGS team to restore the instrument to its 2000 performance via an AMA adjustment. This request was approved and the adjustment was carried out on January 22, 2009 (figure 5).

Movement of the AMA also causes a shift (of about 5") of the affected FGS relative to the other apertures in the HST focal plane. It was also expected that the AMA move would affect the optical distortion across the FGS FOV by an amount that is significant relative to the desired accuracy of the astrometric measurements being carried out by ongoing General Observer programs (but this change would be insignificant for the guide function). To search for possible changes in the distortion, and to acquire the data needed to restore the knowledge of the alignment of FGS1r relative to the other HST apertures, a field of stars with UCAC-III coordinates was observed in four HST orbits before the AMA move, and four orbits after the move. Comparison of the before/after data strongly indicated that the FGS1r optical distortion had changed, with errors approaching 10 mas when the pre-AMA moved distortion coefficients are applied to post-AMA move data (figure 6). Moreover, the distortion change appears to have a radial structure.

To restore the FGS1r geometric distortion calibration, a special 6-orbit calibration test was carried out in March 2010 to observe stars in the M35 open cluster that have relative positions known to an accuracy of 1 mas from previous FGS1r (and FGS3) calibrations.
Figure 5: FGS1r Y-axis S-curve after AMA adjustment on Jan 22, 2009. The X-axis interference fringe was not changed.

Figure 6: Change in FGS1r Geometric Distortion after AMA move. The 10 mas residuals reflect the error introduced into the calibration. Note the apparent radial structure in the residuals.

Figure 7: March 22, 2010 FSG1r Geometric Distortion Calibration using M35. The telescope roll angle was constrained to be within +/- 25 deg of nominal roll for that date.
Figure 8: Preferred method for calibrating the FGS1r Geometric Distortion. Here the target field M35 is in the anti-Sun direction, for which the HST roll angle is unrestricted.

Figure 9: Updated geometric distortion calibration derived from the March 22, 2010 observations. The residuals from comparison of the pre and post AMA-adjustment observations yield 3 mas residuals, which is larger than the goal of 1.7 mas.

Figure 10: Residuals from the comparison of pre and post AMA-adjustment using the update geometric distortion calibration and a radial term added to the plate overlay algorithm, resulting in 2 mas residuals. However, star fields less “rich” than this may lack sufficient degrees-of-freedom for the radial term application. Such astrometry programs may be degraded if the geometric distortion calibration is not further improved by additional observations of the M35 cluster over a wider range of HST orientations.
The observations executed with HST restricted to +/25 degrees from nominal roll (figure 7), rather than the preferred time when M35 is near the anti-Sun direction and the HST roll is unrestricted (figure 8). Nonetheless the observations were successful in restoring the quality of the FGS1r geometric distortion calibration to nearly the level it had been prior to the AMA move. However, this apparently requires the use of additional (radial) terms in the plate overlay model that can be applied only when observing fields with more than about ten stars (these radial terms are not part of the geometric distortion calibration, rather they are semi-ad hoc enhancements to the plate overlay model that can not be universally applied to sparse plates). Further investigations are needed to determine if the March 22, 2009 observations need to be augmented by additional observations when M35 can be observed at roll angles that differ from the March 2010 observations by about 180 degrees to maximize the apparent distortion across the FGS FOV.

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References

A New Technique for Measuring Absolute Proper Motions with HST: Using Background Galaxies as Positional References

S. Tony Sohn, Jay Anderson, & Roeland P. van der Marel

Abstract. Over the past few years, it has been demonstrated that HST is well suited for high accuracy astrometric sciences. For example, HST’s imagers have been used to measure relative proper motions of stars for deriving the internal dynamics of globular clusters. Absolute proper motions have also been measured in several studies but so far, most of them relied on a single reference object, a quasi-stellar object (QSO) within the field. In this work, we present a new way of measuring proper motions of stars. Instead of relying on background QSOs, we developed a technique that uses numerous background galaxies as positional references. We discuss this technique in detail and demonstrate how well we are able to measure proper motions of stars. Our new method can in principle be applied to any existing multi-epoch HST data and opens up a possibility to measure accurate transverse motions for galaxies out to a few Mpc.

1. Introduction

Thanks to the excellent resolving power of HST, high precision astrometric science is now a well-established field. For example, HST’s imaging cameras have been used to measure the relative motions of stars for studying the internal dynamics of Galactic globular clusters (e.g., McLaughlin et al. 2006; Anderson & van der Marel 2010).

The current onboard imagers of HST (e.g., ACS and WFC3) are capable of measuring relative positions of stellar sources to better than 0.5 mas. However, the absolute astrometric accuracy of HST itself is limited by the quality of the Guide Star Catalog (∼0.2 arcsec for GSC2.3). Nevertheless, it is possible to measure very accurate absolute proper motions (PMs) of stars by measuring their displacement over time relative to stationary background references sources in the same field.

What are considered good background reference sources? Conventionally, quasi-stellar objects (QSOs) have served as excellent reference sources due to their star-like appearance and luminous nature (see e.g., Piatek et al. 2002; Kallivayalil et al. 2006). However, using QSOs has the disadvantage that they must be spectroscopically identified in advance, and that the number density of known bright QSOs is limited (but rapidly increasing with new surveys). For the small field of view of HST imagers, these limits require the observer to deliberately image a QSO in the field in at least two different epochs to measure the absolute PM of stars around them. Background galaxies, on the other hand, are abundant and on average homogeneously distributed over the sky. A glance at any HST image shows that there are many galaxies in its background. In addition, thanks to the exquisite spatial resolution, HST can distinguish galaxies from stars down to a level where galaxies become numerous. It is clear that if the positions of these background galaxies can be reliably measured, they can serve as excellent local positional references and stars could then be measured directly against them. Examples of such approach are demonstrated by Bedin et al. (2005) and Kalirai et al. (2007) on measuring absolute PMs of two stellar clusters, but
the galaxy samples used as reference objects were limited to the ones that are very compact and point-like.

In this article, we introduce a new technique for measuring absolute PMs of stars in the field using background galaxies as reference points. The core of our technique lies in determining positions of galaxies using a template-fitting method. This allows us to measure not only compact galaxies but also extended ones with great precision. We discuss the details of our measurements and provide estimates on the uncertainties.

2. Measuring Absolute Proper Motions

Measuring absolute PMs involves measuring the displacement between the position for a target object at one epoch and its position at another epoch with respect to stationary reference objects. In this section, we discuss the details on how we obtain PM measurements. We assume that we are working on a deep and well-dithered set of first epoch data that will be used to (1) construct the reference image and to (2) build templates of the background galaxies, and a less deep set of second epoch data taken with the same imager as the first epoch. Exposure times are assumed to be nearly identical for all individual images including the second epoch ones. The majority of stars found in this field would belong to a co-moving stellar system (e.g., halo stars of M31 or Local Group dwarf galaxies) and our goal will be to measure the bulk motion of those stars over time.
2.1. Overview

Our technique starts with the construction of a median-combined master frame based on all individual exposures of a field. This co-addition is similar to MultiDrizzle but better optimized for astrometry. Since the observations were well-dithered, we supersample the image by a factor of 2 to enhance the spatial resolution of our master image. The supersampling ensures that aliasing effects are minimized when building the galaxy templates via interpolation. An automatic detection scheme is then run on the master frame to identify and classify sources into stars and galaxies. For the stars in the field, we adopt their positions determined from fitting library PSFs and use those to derive transformation relations that tie each individual exposure to the master frame. The transformations are derived for both first and second epochs. Because most of the stars in the field will have moved towards the same direction on the sky in the second epoch, the way we have set up the transformations will allow us to measure the displacement of galaxies with respect to a fixed stellar moving reference frame. The bulk reflex motion of galaxies will then directly translate into the negative motion of stars. In the course of deriving such transformations between the master and individual frames, we adopt the available distortion solutions (e.g., Anderson & King 2004) and use six-parameter linear transformations to correct for any time-dependent linear skew variations (Anderson 2007).

For each individual galaxy, we create a template from the master frame that takes into account the galaxy morphology, the PSF, and the pixel binning. This template is used to measure the accurate position of galaxies (which will be discussed below in detail) in each exposure. Once we derive the positions of galaxies in each frame, we take the average of those positions for each epoch [i.e., \((\bar{X}, \bar{Y})_1\) for epoch 1 and \((\bar{X}, \bar{Y})_2\) for epoch 2], calculate the difference between the positions, and divide them by the time baseline \((\Delta T)\) to calculate the proper motions in each X and Y. The final measured reflex displacement of each individual galaxy implies a PM of the stars in the field equal to

\[
\mu(X, Y) = \frac{-(\bar{X}, \bar{Y})_2 - (\bar{X}, \bar{Y})_1}{\Delta T}.
\] (1)

In principle, one can use the position of galaxies for deriving the transformations that relate individual frames to the master frame and directly measure the stellar motion as has been done by Mahmud & Anderson (2008). However, positions of stars are much reliably measured than those of galaxies, and to reduce the uncertainties, the optimal way to define the transformations are by using the numerous stars in the field. We show the diagram illustrating the overall process discussed above in Fig 1.

2.2. Measuring Positions of Galaxies: GSF-Fitting Method

The master frame we created is used to construct a template for each galaxy that gives us an idea on how the galaxy’s flux is expected to be distributed among the pixels in the pixel grid of each individual exposure as a function of our assumed center for that galaxy. This template is then fit to the observed pixels to determine the position for the galaxy, just as an effective point spread function (ePSF) is used to derive the position of stars (Anderson & King 2000). Since our galaxy templates are similar to the concept of ePSF, they will be denoted by galaxy spread functions (GSFs).

An important question one might ask at this point would be what exactly the definition of a position of galaxy is. For stars, positions are relatively well defined since they are point sources by nature. However, for galaxies, the definition of position is arbitrary – a galaxy may be intrinsically asymmetric, have less pronounced peak than stars, or even have multiple peaks. Fortunately, the definition of position is not crucial because of the following reason. Our galaxy templates can be built around any point of reference as long as there is flux that belongs to the galaxy. For the sake of minimizing the uncertainties associated with determining the positions (described below), the best point of reference would of course be...
where it provides the highest signal-to-noise ratio. Hence, we adopted the galaxy’s brightest pixel in the master image as its point of reference. This provides us with a handle for each galaxy. Consequently, when we measure a position of a galaxy in an image, we are simply finding the location of this handle in each individual exposure.

Once we have positions of the handles in the master frame, we locate the predicted positions of them in each individual exposure using the transformation relations derived in the earlier stage. Since the center of the template corresponds to the handle for the galaxy, we find the position for the galaxy in the exposure by finding the location within the exposure’s pixel grid that corresponds to the center of the template. To do this, we take the 5 × 5 pixels centered on the galaxy’s brightest pixel in the exposure. We then resample the master frame about this center and construct the galaxy templates. The template is moved right, left, up, and down in order to better fit the exposure in increments of 1/100 pixel. At each trial offset, we evaluate the template at the location of each of the 25 pixels (using bicubic interpolation to interpolate the finely sampled template) and compare it with the observed pixel values. This gives us a χ²-type quality-of-fit estimate for each trial location. We find the minimum of this surface and associate the location of the template center in the individual image with the location of the handle in the master frame. Fig 2 illustrates the process of the GSF-fitting method.
2.3. Uncertainties of the Proper Motion Measurements

The random error of PM ($\mu$) in Equation (1) can be expressed as

$$\Delta \mu = \frac{1}{\Delta T} \sqrt{\frac{\sigma_r^2}{M_1} + \frac{\sigma_i^2}{M_2}},$$

(2)

where $\sigma$ is the RMS scatter and $M$ is the number of exposures per epoch. Because the stars have far better positional accuracies than the galaxies, the RMS scatter $\sigma$ in Equation (2) will be dominated by the uncertainties in our measurements for the galaxy positions. If we assume that the number of exposures in the first epoch is considerably higher than those of the second epoch (e.g., $M_1 = 100$ and $M_2 = 2$; then, $M_1 >> M_2$), the final PM error is simplified as $\Delta \mu = \sigma/\Delta T \sqrt{M}$. The typical RMS positional accuracy $\sigma$ per galaxy and exposure will be similar for the two epochs since both epochs are observed with the same detector. We may then derive $\sigma$ from the repeated measurements of the epoch which has the higher number of exposures (the first epoch in our test case).

In order to estimate the positional uncertainties of our GSF-fitting method with real HST data, we have carried out simulations on a test data set that consists of 108 ACS/WFC images of the “M31 Spheroid” field (Brown et al. 2003) taken with the F606W filter in Dec 2002 and Jan 2003. The master frame was created as described in Section 2.1. To simu-
late individual images, we first replaced the pixel values of the master frame with random deviates drawn from a Poisson distribution with the mean equal to the original counts, and then used the linear transformation relations and distortion solutions to project the noise-added master image onto exposure planes (this latter process is called “blotting back” in MultiDrizzle). Positions of galaxies were derived on each simulated exposure using our GSF-fitting method. We carried out 1,000 simulations and measured the RMS scatter of the distributions for difference in the average and individual positions in $X (\Delta X)$ and $Y (\Delta Y)$.

In Fig 3, we show examples of our results for two different cases: circular galaxies (left panel) and highly-elongated ones (right panel; mostly edge-on spirals). In each panel, the third column shows how consistently the GSF-fitting method found the positions of galaxies among independent simulations, i.e., a narrower scatter indicates that the measured positions are precise whereas a wider scatter implies that the positions are uncertain. The RMS of the distribution in $X$ and $Y$ are displayed in the fourth columns. For the circular shaped galaxies the scatter plots generally show isotropic distributions. Our results indicate that even for galaxies that appear very faint such as the two faintest galaxies in Fig 3, the one-dimensional RMS is $\leq 0.1$ pix ($\sim 5$ mas). In order to compare our GSF-fitting method against classical approaches in finding positions, we have repeated the simulations this time using various centroiding methods (e.g., Gaussian-fitting and derivative search methods) for measuring the positions of galaxies. For nearly all galaxies, the GSF-fitting method provided lower uncertainties. The two faintest galaxies in the left panel of Fig 3, for example, were measured with RMS larger than $\sim 0.3$ pix using the centroiding methods. For the highly-elongated galaxies, we find that the scatter plots show elongated distribution about the average position. Consequently, the RMS is small on one side and significantly larger on the other. This demonstrates that our GSF-fitting method allows even the edge-on galaxies to be used as positional references for stars located along the perpendicular direction of the disk.

### 2.4. GSF-Fitting Method: Application to Real Data

We finally demonstrate how well the GSF-fitting works on real data. For each galaxy in the four-panel examples of Fig 4, we show an extraction from the individual image, the GSF that gives the minimum residual (see Section 2.2 above), and the actual residual image. We find that for most of the galaxies, the residual images look very clean albeit with some cosmic rays. However, for some of the galaxies, the residual images show under- or over-subtracted features. The most likely cause of these cases is the cosmic rays affecting the inner $5 \times 5$ pixel region of the individual images. We examine the residual images one after another to reject the positions of the galaxies affected by the cosmic rays.

### 3. Conclusions

We have shown that our GSF-fitting technique for measuring positions of galaxies works very well. We conclude this proceeding by listing the benefits of using background galaxies as positional references:

- Abundance of sources provides an $\sqrt{N}$ advantage in averaging.
- For any given frame, background galaxies are generally found all over the place. By using these objects, we are able to sample the entire detector, thereby reducing the chance to introduce systematic errors caused by local effects.
- Background galaxies are virtually found anywhere in the sky. There are some instances where galaxy groups or clusters are found in the background of a target field, and one will benefit further by using those galaxies in the field.
Figure 4: Four-panel examples showing how well the GSF-fitting method works on real data. For each panel, the first column shows galaxies in each individual image, the second column shows the best-fit GSF, and the third column shows the difference between the first and second columns.

- Finally, for the reasons stated above, there are many potential scientific applications for the Local Group and beyond.

References

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Cross-matching the Hubble Legacy Archive

Nathan Cole and Tamás Budavári

Dept. Physics & Astronomy, Johns Hopkins University, Baltimore, MD 21218

Steve Lubow

Space Telescope Science Institute, Baltimore, MD 21218

Abstract. The Hubble Legacy Archive (HLA) is composed of hundreds of thousands of images taken by a number of different instruments all across the sky. A large number of these images are of overlapping fields and therefore contain many of the same objects. Through the use of a probabilistic cross-matching technique, it is possible to use these overlapping fields to determine a linkage between similar objects. This linkage can then be used to improve astrometry, as well as provide a means for connecting images taken via different instruments. We present a complete description of this algorithm including the cross-matching technique and solution for correcting the astrometry based upon a single 3D rotation. We then test this algorithm upon two different sets of overlapping Advanced Camera for Surveys (ACS) data, one with a common filter and one with differing filters, to great success. In both the single filter and different filter cases, a cross-match was performed and an astrometric correction possible in which the relative astrometry was improved from the 0.12” inherent to the ACS images in HLA to 0.0034” and 0.0020”, respectively. These values are well into the subpixel level.

1. Introduction

Ever since its deployment in 1990, the Hubble Space Telescope (HST) has been stunning all those who see the images that it takes on its many instruments. Both scientists and laymen alike have marveled at the discoveries that the HST has made possible. Having survived 5 servicing missions and gone well beyond its initial projected lifespan and into its 20th year of service, the HST continues to marvel and amaze.

In order to maximize the amount of science that can be done with data from the HST, as well as preserve that data, the Hubble Legacy Archive (HLA) was created. The HLA consists of a database maintaining all of the data products available for analysis and a web interface providing enhanced versions of those data products through advanced browsing capabilities. Through the HLA there is easy access to all of the multitudes of data that has been taken from the HST’s instruments and all of the analyses of that data.

The HST’s long service history and large number of instruments has caused a huge amount of data to be generated. This data is mostly in the form of images, most notably the Advanced Camera for Surveys (ACS) and Wide-Field Planetary Camera 2 (WFPC2), that are then processed to create source lists of all of the objects seen within a given image. Many of these images overlap, either through design or coincidence, providing repeat viewings of the same objects. These repeat views mean, however, that there could be two, three, or even hundreds of entries within the HLA for a single object simply because it appears in multiple images.
Figure 1: A small patch of sky with overlapping images in the same filter. Here we have zoomed in enough to discern the small, \( \sim 0.1'' \), angular separations. Image HST\_10503\_13\_ACS\_WFC\_F606W is represented by the cross (+), while image HST\_10503\_18\_ACS\_WFC\_F606W is represented by the x (X).

Through the use of a cross-identification algorithm it is possible to analyze these overlapping images for objects that are common to them. Thus, linkages can be created to connect all instances of the same object within all images it is found in, as well as providing a link between instruments. In this way, one will be able to determine all information about a single object including: how many images it occurs in, what the analysis says about each one of those occurrences, which instruments it was viewed by, etc.

Furthermore, once objects have been cross-matched, it is possible to utilize those matches to improve the astrometry of those images. Using those objects that have been found in multiple fields, it is possible to calculate a transformation for each image which would correct the astrometry of that field. In this manner, the cross-identified objects can be used to significantly improve the relative astrometry of all overlapping HST fields, and therefore create new source lists within the HLA which maintain these new positions and linkages to the matched objects.

In the following sections, we develop a algorithm to perform the cross-identification of objects between images providing a linkage of common objects between images. We then extend the algorithm to calculate the image transformations necessary to correct the relative astrometry between images. Combining these two pieces, two tests are then performed to demonstrate the effectiveness of the developed algorithm.
2. Algorithm

In this section, we will discuss the formulation of the algorithm that was developed to cross-identify objects in the HLA and use those identifications to improve the relative astrometry of those fields. While this method was developed for use with the HLA, there is nothing to tie it specifically to that set of data. Any set of data in which there exists positions and errors for those positions would be a candidate for use with this method. Section 2.1. will discuss the specifics of the cross-identification method while section 2.2. will discuss and define the method for which the astrometry is corrected.

2.1. Cross-matching

To perform the act of cross-identifying sources between images, we have chosen a proper probabilistic method which is assured to give us all matches and probabilities. To do this we employ the Bayesian method described in Budavári & Szalay (2008). In the usual limit of high precision and small separation, the Bayes factor in the two observation case is

$$B = \frac{2}{\sigma_1^2 + \sigma_2^2} \exp \left[ -\frac{\psi^2}{2(\sigma_1^2 + \sigma_2^2)} \right], \text{ where}$$

$$\psi$$ is the angle between source positions and $$\sigma_i$$ is the astrometric precision.
After having calculated the Bayes factor for all matches, the posteriors (P) can be found by relating it to the priors ($P_0$):

$$P = \frac{BP_0}{1 + BP_0},$$  \hspace{1cm} (2)

$$P_0 = \frac{N^*}{N_1N_2},$$  \hspace{1cm} (3)

in the two-way case with $N^*$ representing the number of matched sources, which we do not know. However, we can calculate it from self-consistent arguments.

With the value of $N^*$, the probabilities for each match can be calculated. These probabilities can then be used to determine a subset of “good” matches while discarding those which are deemed erroneous. This subset can then be used to determine the astrometric corrections as seen in section 2.2.

### 2.2. Astrometric Correction

Given a set of catalogs with absolute astrometry that is not as precise as their internal accuracy, we want to solve for the transformation of the individual catalogs such that their relative registration is optimal. Assuming the astrometry is good enough for an initial crossmatch, let $c^{(\beta)}$ denote the best guess position of an association $\beta$. Choosing a single field, partially correct the positions $r^{(\beta)}$ towards $c^{(\beta)}$ and repeat this procedure considering a different field every time.

It is advantageous to perform these corrections within a spherical coordinate system instead of using a tangent plane projection. In this manner, we are able to describe the
rotation and translation of the tangent plane with a single 3D rotation, \( T \). Also, this makes it much easier to cross-register many fields. The problem with this approach is that stretching, should it be necessary, of the field is not easily implementable.

In general, we need to solve the minimization problem

\[
\min_T \left\{ \frac{1}{2} \sum_{\beta} \frac{1}{\sigma_{\beta}^2} \left( c^{(\beta)} - T r^{(\beta)} \right)^2 \right\}
\]

where \( \sigma_{\beta}^2 \) is the astrometric error squared and the constraint is

\[
T' T = I
\]

The problem is analytically tractable if we assume that the rotations are infinitesimal. This is a good approximation, for the area of sky covered in a single field is very small and the astrometry is good to start. Since infinitesimal rotations are commutable and can be described using the cross-product operator, the minimization becomes unconstrained but still quadratic

\[
\min_\omega \left\{ \frac{1}{2} \sum_{\beta} \frac{1}{\sigma_{\beta}^2} \left[ c^{(\beta)} - r^{(\beta)} + \omega \times r^{(\beta)} \right]^2 \right\}.
\]

Let \( o^{(\beta)} = c^{(\beta)} - r^{(\beta)} \) and be the offset vector. Therefore,

\[
\min_\omega \left\{ \frac{1}{2} \sum_{\beta} \frac{1}{\sigma_{\beta}^2} \left[ o^{(\beta)} - \omega \times r^{(\beta)} \right]^2 \right\}
\]
Figure 5: A small patch of sky with overlapping images in different filters. Here we have zoomed in enough to discern the small, $\sim 0.1''$, angular separations. Image HST 10503.13 ACS WFC F606W is represented by the cross (+), while image HST 10503.18 ACS WFC F814W is represented by the x (X).

or if we use components and the Levi-Civita symbol:

$$\min_{\omega} \left\{ \frac{1}{2} \sum_{ij} \frac{1}{\sigma_\beta^2} \left( o_i^{(\beta)} - \sum_{jk} \varepsilon_{ijk} \omega_j r_j^{(\beta)} \right)^2 \right\}.$$  

(8)

Taking the partial derivatives $\partial/\partial \omega_j$, we get the linear equation

$$\sum_j A_{lj} \omega_j = b_l, \text{ where}$$  

(9)

$$b_l = \sum_{im} \varepsilon_{ilm} \sum_{\beta} o_i^{(\beta)} r_m^{(\beta)} / \sigma_\beta^2.$$  

(10)

$$A_{lj} = \sum_{ikm} \varepsilon_{ijk} \varepsilon_{ilm} \sum_{\beta} r_k^{(\beta)} r_m^{(\beta)} / \sigma_\beta^2.$$  

(11)

Thus allowing the correction of any fields that have been cross-matched.

This method allows for very fast calculation of the transformation. It requires only the summation of the vector, $b$, the summation of the 3x3 matrix, $A$, and then simply solving for the transformation vector itself. It is also very stable.
Figure 6: The same patch of sky as Figure 5 at convergence of the algorithm. The relative astrometry has been noticably improved after processing.

2.3. Method

Using the algorithms developed in sections 2.1. and 2.2., it is now possible to combine these into a method for iteratively correcting the astrometry for overlapping fields. The procedure begins by selecting two or more overlapping fields with a common intersection. The cross-matching algorithm is then applied to determine all probable matches and their probabilities. A cut is then made on probability, currently at 0.95, to create a subset of matches. The best guess position, \( c \), is then calculated as a weighted average for all members of this trimmed list.

The astrometric correction step begins by choosing a single field, and determining the transformation \( \omega \) as seen in section 2.2.. After \( \omega \) has been calculated, an updated source list for the field is created by applying the transformation to the entire source list:

\[
r' = r - r \times \omega
\]

The astrometric correction is done symmetrically, therefore, the next field is chosen and the transformation for that field calculated and applied similarly. This is repeated until all fields have been corrected. Once all fields have an updated source list, convergence is checked by calculating the change in position \( r' - r \) for every source. If the maximum change in position is larger than the preset threshold, \( 1 \times 10^{-17} \), the entire process is repeated until this condition is met.
Figure 7: The residuals, in arcseconds, at each step of processing for images HST_10503_13ACS_WFC_F606W and HST_10503_18ACS_WFC_F814W. Rapid convergence is seen over 8 iterations.

3. Data

The following tests were performed using the HLA DR4 database released in March 2010. SExtractor (Bertin & Arnouts 1998) and DAOPhot (Stetson 1987) source lists both exist within the database; however, we have chosen to use the SExtractor sources. The SExtractor sources were chosen because SExtractor provides error information that is not seen in DAOPhot which is crucial to the developed algorithm. The SExtractor source positions were taken for all sources with a FLAGS value of 0, thus only point sources were examined for the following tests.

4. Results

A set of fields were chosen for a two-way matching from the ACS WFC dataset, HST_10503_13ACS_WFC_F606W and HST_10503_18ACS_WFC_F606W, to determine the algorithms’ effectiveness. These fields were chosen due to having a high degree of intersection as well as being from different exposures using the same filter. The large intersection allows for a high number of sources to be used in the determination of the transformations, while the exposures using the same filter should reduce the number of drop out sources that are visible within one filter and not another. A small patch of intersection from these fields can be seen in Figure 1 in which the source positions are plotted as X’s (HST_10503_18ACS_WFC_F606W) and +’s (HST_10503_13ACS_WFC_F606W). The plotted sources are those determined to be part of a “good” match that made the probability
The difference in the positions of these sources are clearly visible and, while small, they are certainly not negligible.

The algorithm was then allowed to run to convergence on these two fields with a relaxation constant of the Golden Ratio (≈ 1.6). After seven iterations, the algorithm converged, to the final positions seen in Figure 2. Clearly the relative astrometry has been improved with the matched sources aligning in almost all cases. The residuals, angular separation of matches between the two fields, can be seen in Figure 3. Prior to any correction, the peak separation between fields can be seen to be approximately 0.12". However, as the algorithm iterates, the separation can be seen to rapidly decrease until convergence. Figure 4 is a zoom of Figure 3 and shows more detail at smaller separations. At convergence, the peak angular separation is 0.0034". This is a correction of three orders of magnitude and well into the subpixel range.

A second set of fields were chosen for a second two-way matching test from the ACS WFC dataset. These fields were HST_10503_13_ACS_WFC_F606W and HST_10503_18_ACS_WFC_F814W. These fields again have a large intersection; however, they are taken using different filters. In this manner, we will be able to see how much the algorithm will be affected by drop outs between filters. Figure 5 shows the initial positions of matches within a small piece with the two fields with HST_10503_13_ACS_WFC_F606W represented by ‘+’s and HST_10503_18_ACS_WFC_F814W represented by ‘X’ s. Again, there is a considerable separation between the two fields.

The algorithm was, again, allowed to run to convergence on these two fields with a relaxation constant of the Golden Ratio. This run took eight iterations to converge to the final positions seen in Figure 6. Substantial correction and alignment of the fields is again
seen. The residuals of this run, can be seen in Figure 7. Initially, the peak separation is approximately $0.12''$ with considerable improvement over each of the eight iterations. A zoom of Figure 7 can be seen in Figure 8 for more detail. At convergence of this run, the peak separation has been reduced to $0.0020''$. Again, a correction of three orders of magnitude and into the subpixel regime. As can be seen here, the algorithm is clearly unaffected by the drop out potential between filters.

5. Conclusions

We have developed and demonstrated the effectiveness of an algorithm to cross-identify sources from overlapping HST fields and improve their astrometry. Through two examples, we have shown the ability to match and correct the astrometry of fields from the inherent separation between ACS fields of approximately $0.12''$ by three orders of magnitude and into the subpixel range. Specifically, we have corrected overlapping fields in the same filter to a separation of $0.0034''$ and overlapping fields in differing filters to a separation of $0.0020''$.

After these successful tests, we are continuing development of this algorithm to increase the speed of the cross-identification step and provide for n-way matching of fields on the visit level. Once these improvements have been completed, a complete cross-matching and astrometric correction is to be done upon the entire HLA. Specifically, for all overlapping fields within and between ACS and WFPC2 fields. This will allow for not only an astrometric correction for all fields but a linkage of common sources to be created between both images and instruments. In regards to ACS and WFPC2, this will provide easy access to both the ACS and WFPC2 filter bands providing an extended amount of data about a specific source visible to both instruments.

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Some Thoughts on Cross-calibration in the Mid-Infrared

Sean J. Carey

Spitzer Science Center, IPAC, Caltech, Pasadena, CA, 91125

Abstract. I provide some examples of the cross-calibration between the instruments aboard the Spitzer Space Telescope. Current cross-calibration accuracy between the Spitzer instruments is better than 5% and consistent with the calibration uncertainties of the instruments and the current uncertainties in the absolute calibration of infrared zero-points. Utility of cross-calibration extends beyond reconciling measured fluxes; examples are given of how cross-calibration can be used to identify/mitigate instrumental signatures. The Spitzer cross-calibration efforts are placed in context by a brief review of the current state of absolute calibration in the infrared. The current accuracy of calibration is limited by knowledge of the fundamental zero points to about 2%. Rapid progress in the field is being made which suggests that the 1% absolute photometric accuracy requirement of some future dark energy probes can be met.

1. Introduction

Cross-calibration, or comparison to the absolute photometric accuracy of other observatories/instruments is fundamental to placing data from one instrument in scientific context with existing data. Note that for the remainder of this contribution, I will use instrument as the generic term for a fundamental quanta of device that conducts observations that need to be calibrated. In addition to verifying that observations from different instruments can be placed on the same physical scale, cross-calibration is often used during development and data reduction to identify other calibration issues which can be missed due to the resource-limited nature of most space-based astronomical instruments. As a general rule, the best way to improve and understand the calibration of any individual data set is to use it for science which necessarily means comparing the data to other existing data sets for most (but not all applications). Section 3 provides specific examples of cross-calibration between the various instruments aboard the Spitzer Space Telescope (Werner et al. 2004). These examples span the gamut of uses from refining the absolute calibration, to correcting for instrumental signatures to sanity checks for science purposes.

Cross calibration is related to but is often a separate activity from the absolute calibration of a instrument. Absolute calibration is the tying of the observed (spectro)photometry of an instrument to a known physical standard. Ideally, there would be no need for cross-calibration if all instruments were perfectly absolutely calibrated using the same or consistent fundamental standards. However, this is particularly not the case in the infrared as different authors use different standards or different values for the same standards. As a result, the differences arising in the cross-calibration of instruments can be attributed to uncertainties/errors in the instrument calibrations and differences in the assumed fundamental calibrator used in the calibration of each instrument. Section 2 provides a summary of the current state of absolute calibration in the infrared to place the cross calibration examples discussed in this paper in better context.

Fortunately, for many science goals, a perfect absolute calibration is not essential. In fact, for exoplanet transit observations, arguably one of the more exciting applications of
space-based infrared observations from Spitzer, the relative calibration is also unimportant as long as the calibration remains stable throughout the observations. Implicit in this statement is the assumption that the parameters of the host star (effective temperature and radius) are extremely well-known (e.g. van Belle & van Braun 2009), but this detailed characterization is well beyond the scope of most calibration campaigns. It is worth mentioning that as the precision of transiting planet observations improves the knowledge of the fundamental properties of the host stars will be the limiting factor in measuring planet radii and determining their temperatures.

In many cases, as long as the instruments are consistently calibrated then colors and flux ratios will be correct. Derived quantities such as temperature, hardness of incident radiation field or extinction can be well determined from data with a good relative calibration. If the uncertainties in absolute calibration are a few percent or so, which is currently the state of the best infrared calibration, then quantities such as derived luminosity and distance will have small errors that are acceptable for most science applications.

However, absolute calibration is significant for one major line of inquiry, using supernovae as a probe of dark energy requires absolute photometric accuracies of order 1%. The use of supernovae as probes of dark energy is one of the goals of the WFIRST mission, a primary recommendation of the 2010 Decadal Survey of Astronomy and Astrophysics. Kim et al. (2004) provide a discussion of the effect of zero point shifts on the derivation of cosmological parameters using supernovae as standard candles. Currently, the required absolute calibration is not met in the near- to mid-infrared, but current work on absolute calibration is promising and suggests that this goal could be met in the near future.

In addition to the discussion on absolute calibration in Section 2 and the cross-calibration examples in Section 3, I contemplate a suggestion to improve the absolute and cross-calibration of future space-based instruments in Section 4 and provide some final comments in Section 5.

2. Infrared Absolute Calibration

The absolute calibration of modern infrared instruments is complicated by the high sensitivity of the instruments which precludes direct observation of the fundamental standards. The calibration is typically done by observing a set of standards which are well spectral typed and have excellent photometry at shorter (optical) wavelengths. The absolute flux scaling of the standards is done by transferring the standard to a fundamental calibration through optical and near-infrared photometry. A variety of standards are used: A dwarfs as they are relatively featureless in the infrared, K giants as they are bright in the infrared and solar analogs as there exist good models based on the solar spectrum. White dwarfs have been used extensively in the near-infrared calibration of HST with internal consistencies of 1% (e.g. Bohlin et al. 2001). KIII and solar analog models have uncertainties of order 1-2% at the relevant wavelengths. Figure 1 displays the difference in the Cohen et al. (2003) and Engelke et al. (2006) templates for the KIII star, NPM1p68.0422, used as a calibrator for IRAC (Fazio et al. 2004). The differences in the templates are profound at the several percent level and attest to the current difficulties in mid-infrared calibration. As instruments become increasingly more sensitive and can no longer access the brighter calibrators, solar analogs and white dwarfs are becoming more important and will be key components of the calibration methodology for JWST.

Unfortunately, the fundamental standard, Vega, is problematic. At wavelengths of 12 $\mu$m and longer, Vega has a pronounced infrared excess (Aumann et al. 1984) due to circumstellar material. An inner, hotter circumstellar disk (Ciradi et al. 2001) further complicates analysis in the mid-infrared. For shorter wavelengths, it has been shown that Vega is a rapid rotator (e.g. Gray 1988) and is not well represented by models of A0V type stars. Coupled with the uncertainties in the templates/models of the primary standards, the
uncertainties in using Vega as a fundamental calibrator result in the best current infrared calibrations having uncertainties of order 3%. This is reflected by the current state of the literature as different workers in the field can arrive at calibrations that differ by 2% starting with the same data and apparently (at least to this author) reducing and analyzing the data in equally valid ways. This is clearly the case in comparing the work of Rieke et al. (2008) and Price et al. (2004), both of whom heavily weight the MSX data to determine the flux of Vega at around 10 $\mu$m. Engelke et al. (2010) have recently attempted to improve the situation by developing a fundamental zero point spectrum based on 109 Vir in the optical and Sirius in the infrared. The situation is even worse at longer wavelengths as there are few stars that can be observed and many of the calibrators used such as planets or asteroids are inherently variable.

For Spitzer, each of the three science instruments used a different methodology to develop their primary calibration standards. The choice of calibration method was left to the teams building the instruments and adopted by the Spitzer Science Center during standard operations. For the 3-8 $\mu$m cameras comprising the IRAC instrument, a network of four AV and seven KIII standards were developed using absolutely calibrated spectral templates (Cohen et al. 2003). These templates are rooted in the fundamental calibrations associated with MSX and the methodology described in Price et al. (2004). During operations, it was noted that the KIII standards and AV standards produced calibrations that were discrepant by 7.3%, 6.5%, 3.6% and 2.1% for the 3.6, 4.5, 5.8 and 8.0 $\mu$m channels, respectively. Reach et al. (2005) adopted the AV star calibration as it was more likely that uncertainties in molecular absorption features in the KIII spectra were complicating the templates. This hypothesis was confirmed through IRTF observations of some of the IRAC primary calibrators (Figure 2). In the final processing of IRAC data, updated templates (see Figure 1) for the KIII stars are being used. With the new templates, the discrepancy between the calibration factor determined using the KIII stars and AV stars is less than 1% in all
channels. The 24, 70 and 160 \(\mu\)m photometers (and 70 \(\mu\)m spectrometry) comprising the MIPS (Rieke et al. 2004) instrument used a combination of AV stars at 24 \(\mu\)m (Engelbracht et al. 2007) and 70 \(\mu\)m (Gordon et al. 2007, Lu et al. 2008 for the spectroscopic mode) and asteroids and red extragalactic sources at 160 \(\mu\)m (Stansberry et al. 2007). I will concentrate on the 24 \(\mu\)m calibration in the remainder of the discussion as that is the MIPS channel that overlaps with other well-calibrated datasets from \textit{Spitzer} and \textit{MSX}. The 24 \(\mu\)m calibration uses \([K] - [24] = 0\) for AV stars and a photospheric model of Vega for the zero point flux density at 24 \(\mu\)m (Rieke et al. 2008). The IRS spectrographs (Houck et al. 2004) are calibrated using MARCS stellar models (Decin et al. 2004) of one primary standard, HR 7341 (K1III spectral type) and several secondary standards.

3. **Specific examples of cross-calibration**

My case studies in the utility of cross-calibration are all drawn from personal experience and deal with the three instruments, IRAC, MIPS and IRS aboard \textit{Spitzer}. The examples span the spectrum of uses of cross-calibration from the traditional comparison to diagnosis of issues in the calibration of a specific instrument than cannot be well-handled without the use of data from another instrument. In all cases, the cross-calibration is between instruments that have overlapping bandpasses/spectral responses. Cross-calibration can also be done between instruments without overlapping responses; however, the assumed spectral energy distribution of the calibration sources can become a larger source of error.
for cross-calibration without overlapping responses. For photometers, if the instruments are perfectly, relatively calibrated then

\[ \frac{F_A \times S(\nu_B)}{C_A} = \frac{F_B \times S(\nu_A)}{C_B} \]  

(1)

where \( F \) is the measured flux density of instrument A or B, \( C \) is the appropriate color correction for the given instrument and calibration source spectrum and \( S(\nu) \) is the source flux density at the effective frequency of the given instrument. Typically, color-corrections are the order of a few percent for well-behaved calibration sources and for instruments calibrated using the isophotal assumption. The ratio, \( S(\nu_B)/S(\nu_A) \), is will typically differ from 1 by \( \delta \nu/\nu \) for stellar calibrators and similar instrument bandpasses. For extended sources, the variation in effective flux density can be quite large (up to 50%) depending on the spectral energy distribution of the source.

In addition to the examples in Sections 3.1-3.6, ongoing programs exist to cross-calibrate the Spitzer instruments with WISE, IRAC and MIPS with HST and JWST (Gordon & Bohlin, this proceeding) and warm IRAC observations are planned to cross-calibrate with DIRBE. The DIRBE cross-calibration observations will include at least one map of a DIRBE calibrator using IRAC for the entire DIRBE beam (42 arcminutes) to assess how the additional flux in the DIRBE beam due to extended sources and field stars affects the comparison.

3.1. Spitzer Instrument cross-calibration

While there is no cross-calibration requirement for Spitzer, the Spitzer Science Center conducted several cross-calibration experiments. The experiments consisted of observations of calibrators for the spectrometer, IRS, by all three instruments as the photometers could achieve high signal-to-noise data with very short observations of the IRS standards. Gizis et al. (2010) will discuss these experiments in great detail, but the salient points are summarized below.

Between IRAC and IRS, observations of three IRS calibrators, HR 7341, HR 2194 (A0V), and HR 6606 (G9III) were compared by integrating the IRS SL1 and SL2 spectra with the IRAC 8 \( \mu \)m response function. For all three sources the photometry agrees to 0.3%. The uncertainty in absolute calibration of IRAC is 3% and IRS has a uncertainty of 5%. Considering that the calibration methodology of the two instruments is entirely different this agreement is surprisingly good. The comparison of IRS to MIPS also uses overlapping bandpasses. In this case, IRS LL observations of HR 2194 and HR 6348 (K1III) are compared to the MIPS 24 \( \mu \)m passband. IRS is 2.2\%±1.0\% fainter than MIPS at 24 \( \mu \)m. The calibration uncertainty of MIPS is 4%, and while the IRS/MIPS offset is measurable, it is comfortably within the instrumental calibration error budgets. The disagreement is in accord with the good agreement of IRS with IRAC and the results of Rieke et al. (2008) which suggest that IRAC is 1.5% lower than MIPS. The difference between IRAC and MIPS is attributed to the differences in determination of the flux density of Vega at 10 \( \mu \)m between Rieke and MSX (Price et al. 2004) as the IRAC templates developed by Cohen are based on the MSX results.

3.2. IRAC extended source

Comparison between different instruments can often help characterize instrumental signatures/artifacts that would be difficult to diagnosis without external information. One example is the pronounced internal scattering present in the Si:As detectors used by IRAC in the 5.8 and 8.0 \( \mu \)m passbands. Approximately 30% of the light incident on an IRAC Si:As detector pixel is scattered more or less uniformly throughout the remainder of the array. Cohen et al. (2007) independently confirmed this result. As IRAC is calibrated using point sources and a small photometric aperture (12.2 arcsecond radius), this scattered light
Figure 3: Extended source photometry comparisons between IRAC 8.0 µm and MSX 8.3 µm for three HII regions as a function of aperture size in arcseconds. The black symbols and line are the ratio of the uncorrected photometry. The green symbols/line are the data after correcting for the internal scattering and the blue symbols/line are the ratios after correcting for the different bandpasses and effective frequencies of the instruments.

is not considered in the calibration. Instead, it only becomes apparent in the photometry of extended sources where a portion of the scattered light from some pixels in the aperture is scattered back into the other portions of the source aperture. To quantify this effect, the IRAC instrument support team conducted measurements of HII regions and bright reflection nebula previously imaged by MSX at 8.3 µm and with good quality ISO SWS spectra to provide SED information. Figure 3 displays the ratio of aperture photometry for three HII regions observed by IRAC and MSX. Where substructures were present, multiple aperture sizes were used to sample the different spatial scales. After applying a correction of 30% for the scattered light (called aperture correction in Figure 3) and correcting for the different passbands/responses as shown in Equation 1, the MSX and IRAC results are in very good agreement.

3.3. Warm IRAC / Cryogenic IRAC

Cross-calibration is also useful for the same instrument if the operating parameters change significantly. One such example is the operation of the IRAC InSb detector arrays at 28.7 K after the exhaustion of cryogen ending the prime mission for Spitzer. As part of the warm recalibration, the photometry of the primary IRAC calibrators was compared to the cryogenic values. Not surprisingly, the flux conversions between data numbers and
flux density changed as the detectors operated at significantly higher temperatures (during cryogenic operations the detectors were at 15 K) and at a lower applied bias setting (500 mV compared to 750 mV) for the 3.6 µm array. However, we were also able to improve the calibration of the cryogenic data with new information on the nature of the intra-pixel gain variations exhibited in these detectors. In warm operations, the variations are 2× and 4× more significant at 3.6 and 4.5 µm, respectively. Figure 4 displays an intra-pixel gain map made from dedicated observations during the checkout phase of the warm mission. The gain variations for each pixel are parameterized by a two-dimensional Gaussian in x and y offset of the centroid of a star from the center of a pixel.

As the significance of the intra-pixel gain variation was not understood during the initial characterization phase of the cryogenic mission, no dedicated experiment was conducted to determine the nature of the effect. In the Reach et al. (2005) calibration, a one-dimensional parameterization was used which left a residual scatter of ∼1%. With the additional information from the warm characterization, the final cryogenic processing was able to use the same 2d Gaussian parameterization and significantly reduce the scatter in the resulting photometry.
3.4. MIPSGAL point source

The MIPSGAL legacy survey (Carey et al. 2009) mapped the inner Galactic plane with MIPS at 24 and 70 μm. It is complementary to the GLIMPSE survey (Churchwell et al. 2009) which covered the same region of the Milky Way with IRAC. To extract science from the combined data sets, the relative accuracy of the photometry of the data products needed to be assessed. As part of the preparation of the MIPSGAL point source catalog, we compared the measured [8] − [24] color for a sample of KIII and AV stars imaged by both surveys. Figure 6 displays the difference in observed color from the predicted color. Several of the stars (about 15%) in the sample had significantly redder color than predicted which is indicative of circumstellar disks. After removing the disk candidates (color differences less than -0.2 magnitude), the average color difference is +0.03 indicating that the IRAC 8 μm photometry is 3% too bright compared to the 24 μm photometry which is contrary to the results of Section 2 (although still within the uncertainties) and has also been noted by the SAGE legacy team in their data of the Large Magellanic Cloud. This result is inconsistent with the cross-calibration of Section 3.1 and may be possibly due to the different methods used to determine the photometry.

3.5. MIPSGAL extended source

As part of the MIPSGAL processing, a more exact droop correction was applied in the presence of saturated data. This and other pipeline enhancements specific to the bright, diffuse backgrounds of the Galactic plane are discussed in Mizuno et al. (2008). To verify that these enhancements did not adversely effect the data, we smoothed the 24 μm image data to the same resolution as the MSX 21.3 μm data and compared the surface brightness on a per-resolution element basis. Assuming that underlying emission had a spectrum resembling a massive star-forming region such as M16, we color corrected each data set and compared the data to a ν^{-1.5} spectrum. The MSX and MIPS data are in excellent agreement up to about 1700 MJy sr^{-1} at which point the MIPS detectors saturate.
3.6. MIPSGAL/HiGal

At wavelengths longer than 24 μm, absolute calibration is extremely challenging. As an independent check of the photometry of the Herschel Galactic plane survey (HiGal; Molinari et al. 2010), the HiGal team compared the 70 μm PACS maps to the 70 μm MIPSGAL maps of the same regions. The result for one of the 2° × 2° tiles is shown in Figure 8. The agreement is quite good considering the degree of striping still present in the MIPS 70 μm maps and the preliminary state of the PACS calibration.

4. A Modest Calibration Proposal

With an increasing number of observatories planned for L2, a useful resource for those instruments would be an NIST referenced, external calibrator similar in spirit to the aluminum reference spheres used by MSX (Price et al. 2004). The idea would be to fly a reference source which is absolutely calibrated to better than 1%. Unlike the MSX reference spheres which were passive in nature, the calibrator should be active and able to adjust its temperature and output luminosity. The calibrator should have an internal integrating sphere with some ability to adjust the spectral signature either through filters at the output window of the cavity or perhaps a Fabry-Perot to output a pure spectral frequency. By adjusting the output, the calibrator would serve instruments from the UV to the submillimeter. The calibrator should provide telemetry on power, temperature and location. The power would be supplied via solar panels probably requiring the calibration source to be either on a boom or tethered. Consumables would consist of propellant for station keeping at L2. The calibration source would probably not require its own launch vehicle; it is likely that it could be launched along with a larger mission that had some space in the fairing.
My back of the envelope estimate for cost is $\approx 50$ million which is likely to be $< 1\%$ of the cost of missions sent to L2.

5. Concluding Remarks

While the current state of the absolute calibration in the infrared is not optimal, there has been much recent progress and a good reason to be optimistic that the calibration will achieve the better than $1\%$ requirement needed for WFIRST. An improved absolute infrared calibration will also greatly facilitate science to be done with JWST although it is not needed to achieve any of the major science goals. There has been excellent recent progress in attempting to move away from Vega as a primary calibrator (Engelke et al. 2010), improve stellar models (Castelli & Kurucz 2004; Decin & Eriksson 2007), and unify the calibration scheme through well-planned cross-calibration experiments such as the HST/Spitzer work of Bohlin & Gordon (this proceeding, also Bohlin et al. 2010). The ACCESS mission (Kaiser, this proceeding), a rocket borne, NIST standard calibrated mission will help provide a framework for securing the photometry of stellar standards to an absolute scale. Ongoing development of calibration sources for instruments such as SNDICE (Schahmaneche, this proceeding) and ground based absolute calibration campaigns such as described by Zimmer et al. (2010) will continue to improve the situation. The calibration mission sketched in Section 4 would permit absolute flux measurements to be made limited only by the signal-to-noise of the science observation and would obviate the need for cross-calibration.

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Some Thoughts on Cross-calibration in the Mid-Infrared

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James Webb Space Telescope Calibration

John C. Mather

NASA Goddard Space Flight Center, Greenbelt, MD 20771

Abstract. The James Webb Space Telescope (JWST) is the planned successor to the magnificent Hubble Space Telescope and the smaller but remarkably powerful Spitzer Space Telescope. It will extend the Hubble and Spitzer science in many areas, ranging from the first stars and galaxies, to the current formation of stars and planets, and the evolution of planetary systems to conditions capable of supporting life. The JWST is a NASA-led project in partnership with the European and Canadian space agencies. The deployable cooled 6.5 meter telescope will cover the wavelength range from 0.6 to 28 µm with imaging and spectroscopy. With diffraction-limited image quality at 2 µm and and zodiacal background-limited camera sensitivity for λ < 10 µm, the JWST will be the most powerful space observatory yet constructed. To enable the huge telescope to fit into the rocket fairing, it is very carefully folded up for launch. It has a primary mirror with 18 segments, each one able to be positioned with 6 degrees of freedom and a radius of curvature adjustment. While it is quite well protected from thermal variations, it is nevertheless expected that the JWST primary mirror may be readjusted on the order of every two weeks. This design enables a primary mirror larger than the rocket fairing, but also leads to very interesting calibration issues. In the years since JWST was conceived, the potential scientific benefits of greatly improved calibration and stability have become apparent. Now the challenge is to find ways to achieve those improvements with hardware that has already been designed. In this paper, I outline the basic issues and some strategies to pursue.

1. Introduction

Even before the launch of HST, a conference was held in 1989 at the STScI to consider possible successors, and the proceedings are very informative (Bély et al. 1989). The Next Generation Space Telescope (NGST) was NASA’s response to recommendations of the report “HST and Beyond” (Dressler et al. 1996), which called for an IR-optimized telescope with an aperture of 4 m or more, along with equipment to study Earth-like planets around Sun-like stars. The NGST project was initiated in October 1995 with small studies that showed the amazing potential benefits of a large radiatively-cooled telescope operating far from the Earth. When NASA administrator Dan Goldin told the American Astronomical Society that NASA would build the recommended telescope but even larger, he received a standing ovation. The NGST was described along with the scientific motivations in the Black Book (Stockman et al. 1997). In 2000, the National Academy’s Decadal Survey ranked NGST as top priority among large space missions. In 2002, The NGST was named for NASA’s second Administrator James E. Webb to honor his leadership of NASA as we prepared the Apollo mission. For further details on the JWST history, please consult the web site: http://www.stsci.edu/jwst/overview/history/index.html.

More recent documentation includes the science team’s report (Gardner et al. 2006), and the Arizona conference proceedings (Thronson et al. 2009). Many additional documents are available on line at http://www.jwst.nasa.gov and http://www.stsci.edu/jwst/, and in
white papers (Google “JWST white paper”). Some early technical papers are at STScI (http://www.stsci.edu/jwst/externaldocs/technicalreports), and Casertano et al. (2001) gives an initial view of JWST calibration. See also SPIE articles (Google JWST NGST calibration SPIE). After more than two decades of calibrating HST, there are many good tools available that can be augmented and updated for JWST.

2. Instruments

The JWST carries four instruments and a fine guidance sensor sharing the focal plane. Some calibration, such as dark current measurements, will be done in parallel to science observations. The Near Infrared Camera (NIRCam) covers wavelengths from 0.6 to 5 µm in two bands separated by a dichroic filter. It has two modules providing side-by-side 2.2 arcmin square fields of view; the short wavelength channels are Nyquist-sampled at 2 µm, and the long wavelength channels at 4 µm. NIRCam is provided by the University of Arizona (M. Rieke, PI) with support from Lockheed Martin. NIRCam is also used for wavefront sensing, using thin lenses and diffraction gratings in the filter wheels, and using a special pupil-imaging lens. Detectors are 2048² pixel arrays of HgCdTe from Teledyne, with different cutoff wavelengths for the short and long wavelength channels. NIRCam provides coronagraphic spots using a thin wedge in the filter wheel.

The Near Infrared Spectrometer (NIRSpec) provides spectroscopy over the same wavelength range, using gratings and grisms to obtain spectral resolutions of 100, 1000, and 3000. NIRSpec uses a microshutter array located at an image plane to select up to 100 objects for simultaneous observation in a field of view of 9.7 square arcmin. Over 250,000 shutters are individually addressable by ground command. In addition, NIRSpec provides fixed slits and an integral field capability using image-slicing mirrors. The NIRSpec is provided by ESA with support from Astrium; the Project Scientist is P. Jakobsen. NASA provides the microshutter array (S. H. Moseley, PI) and the detector arrays (B. Rauscher, PI). These detectors are also from Teledyne.

The Mid IR Instrument (MIRI) provides imaging, integral-field spectroscopy, and coronagraphic capability from 5 to 28 µm over a field of view of 1.4 × 1.9 arcmin. The optical system is provided by a European consortium of many institutes, led by G. Wright of the UKATC, and the detector assembly is provided by the Jet Propulsion Laboratory using Raytheon Si:As detectors. G. Rieke (Arizona) co-leads the team with G. Wright. MIRI requires active cooling to about 7 K for detector operation.

The Fine Guidance Sensor (FGS) and the Tunable Filter Imager (TFI) are provided by the Canadian Space Agency. The FGS provides two modules with adjacent 2.3 arcmin square fields of view and generates position information for guide stars in its field of view. The TFI includes a tunable Fabry-Perot interferometer with a spectral resolution of about 100, covering the range from 1.6 to 4.9 µm.

The instruments are housed within the Integrated Science Instrument Module (ISIM). All of the instruments are in final preparation and are expected to arrive at Goddard Space Flight Center by October 2011 for integration into the ISIM.

3. Telescope

The JWST telescope uses a 3-mirror anastigmat design, to provide a much larger diffraction-limited field of view than a Cassegrain telescope. Its 6.5 m aperture produces a final f/20 image that is diffraction-limited at 2 µm after deployment and adjustment on orbit. The primary mirror is made with 18 hexagonal beryllium segments, lightweighted by removing 92% of the material to leave a triangular rib structure on the back of each piece, with the remaining material of order 2 mm thick. At the time of this writing, all of the primary mirror segments have been polished at room temperature, all have been measured cold, and
one has been polished to final form and coated with gold. All the mirror segments, the secondary and tertiary mirror, and the fine steering mirror will be completed by September 2011.

4. Observatory and Orbit

The observatory combines the telescope, the ISIM, the spacecraft bus, and a giant sunshield to enable passive cooling of the telescope and instruments to about 40 K. The sunshield, roughly the size of a singles tennis court, uses 5 layers of metallized plastic to obtain a Sun Protection Factor of $>10^6$. The selected shape allows observation for lines of sight between 85° and 135° from the Sun, so that 35% of the sky is accessible at any one time. Telemetry to the Earth is handled by K_A band transceivers and a small parabolic dish, to relay 464 Gbit/day of data. Onboard command processors allow event-driven observations, following a sequence rather than absolute times, to improve efficiency of operations. Primary pointing is provided by applying torque to large momentum wheels in the spacecraft bus. The spacecraft bus also carries the compressors to operate a helium pulse-tube cooler for the MIRI.

The selected orbit is a large loop centered on the Sun-Earth Lagrange point L_2, about $1.5 \times 10^6$ km from Earth. This point, first found by Euler in 1750, moves around the Sun once a year with the Earth. The JWST avoids the shadow of the Earth. JWST will near its desired orbit about 2 months after launch, about the same time that the telescope and instruments reach equilibrium temperature. The orbit is unstable, so small jets are fired every few weeks to maintain it. The jets are also used to remove accumulated angular momentum (due to solar radiation pressure torques) from the wheels. Fuel is provided for 10 years of scientific observations.

5. Wavefront Sensing and Control

The JWST is deployed and focused after launch. There are many steps leading to final alignment, and all have been demonstrated through simulation and using a 1/6 scale telescope model. The final fine adjustment uses the same mathematics as were used to repair the HST. Stellar images are taken in focus and out of focus, and least-squares-fit algorithms compute the required mirror adjustments. The fitting algorithms use the fact that the electric field of an image is analytic in three dimensions, so that phase errors in the focused image are related to amplitude changes in the out-of-focus images. The NIRCam provides the needed capabilities for this fine adjustment, but the earlier stages require the use of data from the other instruments as well, to ensure that large-scale (low-order) aberrations such as field curvature and field tilt are within specifications.

Periodic adjustment of the mirrors will be needed because the temperatures change slowly with time, with time constants of weeks to respond to changes of orientation relative to the Sun. In addition, as the orbit is 1.7% eccentric around the Sun, there is a 7% (peak-to-peak) variation of the incident solar radiation through the year. It is conceivable that experience will show us how to adjust the mirrors better and better, or that we could decide to optimize the focus for one instrument or another, according to the scientific objectives. Calculations suggest that optimizing the focus for just one point in the NIRCam field of view could give diffraction-limited imaging at 1 μm near that point, but this optimization could make the imaging significantly worse for the other fields of view.
6. Calibration Challenges

The JWST requirements for calibration accuracy were set several years ago. Photometric accuracy for NIRCam, MIRI, and TFI were specified at 5%, and spectroscopic flux accuracy was set at 10% for NIRSpec and 15% for MIRI. Much work has already been completed, generating error budgets to enable the construction of the needed tools and procedures. Calibration algorithms and data are to be provided by the instrument teams, and the STScI is to convert them into tools that run routinely on the flight data.

Since the JWST was conceived in 1995, the importance of calibration has dramatically increased. The discovery of the accelerating universe was based on a discrepancy of about 20% in the brightness of distant supernovae of type Ia. While this may seem to be a large effect, the proof requires comparison of extremely faint and redshifted objects with bright local standards at different wavelengths, and many types of non-ideal instrument behaviors must be calibrated or bounded. Also, the observation of transiting Earth-like planets around Sun-like stars requires exceptional (10 parts per million) photometric stability over the period of a day or so. The Kepler mission has achieved such stability, but it was custom-designed with that requirement foremost. The HST was designed without such a requirement, but has nevertheless been used to observe exoplanet transits and detect various atmospheric constituents.

A paper by Cohen (1998) says that there is a network of standard stars with \( \pm 3\% \) accuracy, and outlines the problems to be overcome. The ACCESS sub-orbital mission will set up several NIST-traceable bright standard stars with \( \pm 1\% \) accuracy (Kaiser et al. 2010.) There are some effects that can not be measured on the ground, others that can not be measured in flight, and some things that ought to be measured and usually are not. Papers by the Spitzer MIPS team (e.g. Engelbracht et al. 2007) give calibration factors based on standard stars, but with differences between approaches that are of order 1-2\%. Can JWST achieve this or better photometry? That depends on addressing many challenges, such as:

- Faint target stars and galaxies. As JWST can observe fainter objects than any other telescope, by definition there is no experience in finding or utilizing a network of standards down to the 1 nJy level.

- Residual images. Tests show that the JWST detectors do have residual images after observations of bright objects, but they are much better behaved than many earlier IR detectors. Nevertheless the effects might be important after slews, or after intentional observation of bright objects. We do not know whether the residual images would have any effect in using dithered observations, where each dither pointing will still have a small after-image from the previous one.

- Detector nonlinearity. Our detectors store charges on capacitive pixels until they are read out, and each pixel may have its own linearity correction. These effects are substantial but presumably measurable through study of the time-dependence of the signals as we sample up the ramps.

- Detector reciprocity failure. This is a different effect, one in which the rate of arrival of the photons matters, instead of the total number in the pixel well. It can limit the ability to compare faint and bright objects, and it can make the response to a point source depend on the background light from zodiacal dust or stray light.

- Detector pixel response functions. While the JWST detectors are well-behaved there are detectable photometric effects based on where exactly a star is located relative to the pixel boundaries. These effects, if significant, might also differ across the face of each detector, or differ with wavelength.
• Detector sampling sequence. It is not unlikely that the calibration factors are detectably different if we change the timing of the sampling of the detectors; hence, the smaller the number of choices, the less work is involved in calibrating them.

• Cosmic ray removal. As we send back multiple samples of the detector signals through each integration period, we have the ability to recognize cosmic rays in the data and remove them. But we do not know whether the corrected data will have the same quality as those from pixels that are not hit by cosmic rays. There are possible nonlinear effects caused by the cosmic rays themselves, that might act as after-images and decay over long periods of time.

• Detector degradation with time. Our detectors are exposed to much higher doses of cosmic rays than those used in low Earth-orbiting telescopes. On the other hand, those in the Spitzer Space Telescope have held up well. Unlike CCDs, which suffer changes in charge transfer efficiency, all the JWST detectors operate with an amplifier for every pixel. Hence, cosmic ray damage is expected to be localized in the form of “hot pixels”.

• Point spread function variation in time. We have an official specification: “SR-13: Encircled Energy Stability. The total encircled energy of an image of a point source over the FOV of the Near-Infrared Camera within a circle of 0.08 arc-second radius and at a wavelength of 2 micrometers shall be stable to better than 2.0% over a period of 24 hours without intervention by ground command.” Proving that we will meet this specification has been very challenging, and it is possible we will have to loosen this specification to match what can be built and promised. On the other hand, the worst case variations probably occur very infrequently, and typical performance may be much better than this specification. Temperature variations or other effects may change the telescope image quality over week time scales, and periodic adjustments to optimize the image quality may divide the calibration data files into many subsets. The image quality also varies strongly across the field of view of each instrument, and with wavelength. While this part of the problem may be well modeled by physical optics, we must certainly account for it, and users must be aware of it.

• Multi-slit spectroscopy with microshutters. No microshutter instrument has been built before NIRSpec, so we have no practical experience with the challenge of hundreds to thousands of slit functions for every observation. We are also not able to place each target in the middle of a slit, so we must learn to cope with this by observing strategies (dithering, etc.), modeling, or inclusion of slit-loss errors in the analysis.

• Spectral response functions. All the filters in the instruments have been measured before they are integrated into the instruments, but some may change with age, launch vibration, or cryo-cycling. The gratings and grisms are supposed to be well understood, but we will have to know how close the spectral responses are to the calculated ones. Spectral stray light (spectral purity) might be an issue with our complex optical systems.

• Vibrations. The JWST spacecraft bus contains momentum wheels and refrigerator compressors that run all the time, along with the telemetry antenna, which is adjusted occasionally, and the jets, which fire infrequently but cause significant disruption of observations. The wheel vibrations change with time as the wheels change speed, and at some speeds the vibrations will be amplified by mechanical resonances in the spacecraft or telescope. Although the effects on the point spread function are required and predicted to be small most of the time, they may be detectable sometimes. These vibrations may limit the ability of the JWST to do transit spectroscopy on exoplanets.
• Fine (fast) steering mirror. This small flat mirror, located at the image of the primary mirror formed by the tertiary mirror, is the core of the pointing control system. It responds immediately to the error signals from the fine guidance sensor. Its range of motion is small, but there is one important effect: differential image distortion. The mirror servo acts to keep the guide star on its required location in the FGS field of view, but all other parts of the telescope field suffer tiny residual motions. While the specifications show that this effect must be extremely small, there may be observations where the effect can be detected. There is another effect as well: noise from the fine guidance sensor is partially tracked by the fine steering mirror, and hence the effect on the PSF depends on the brightness and color of the guide star.

• Steep off-axis optics. The focal surfaces of the instruments are significantly curved and tilted, the plate scale (arcsec/pixel) is not constant across the fields of view, and circular images become ellipses. While this is purely geometrical optics and well understood in principle, we can anticipate that some studies of weak lensing and cosmic shear might still suffer from errors in this area.

• Optical contamination. We do not expect significant changes of the dust coverage or condensed volatile materials during the JWST mission, because the sources are all supposed to be on the warm side of the sunshield, with no way to reach the telescope or instruments. Nevertheless we will need to be alert to changes. Micrometeoroid pits on the primary and secondary mirrors will increase gradually with time but are predicted to cover only 0.1% of the primary mirror after 10 years.

• Stray light and zodiacal light. These sources of light are not supposed to influence the point source photometry, but they do influence the noise and if the detectors are nonlinear, or have reciprocity failure, they might also affect the calculated source brightness. We need to look for this possibility by comparing photometric fields taken under different stray and zodiacal light levels. We also have to be alert to structure in the stray light, due to some undesired path like the “rogue path” in which unfocused light from the sky goes past the baffles all the way to the detectors.

7. Astrometry

Much work has already gone into plans for astrometric calibration for JWST. Good astrometry is essential to placing spectrograph slits or coronagraph spots in the right place. Astrometric reference fields have already been calibrated by HST and will be ready for use by JWST. The stability of the astrometric calibration for JWST can be readily checked, and there are no known factors expected to make it change much. There is significant potential for scientific discovery based on JWST astrometry: follow-up of microlensing planet observations may be able to measure image deflection as well as amplification, proper motion studies of newly discovered brown dwarfs, etc.

8. Summary and Conclusions

We are preparing to calibrate the JWST, both photometrically and astrometrically, but there are many interesting effects due to the extreme sensitivity (ability to observe sources that are far fainter than any calibration standards), and the potentially variable PSF due to changes in the temperature of the segmented primary mirror.
9. Acknowledgments

The JWST project has enjoyed continuous support from NASA Headquarters and from ESA and CSA, bringing the needed technologies from concept to reality in an orderly fashion, and enabling an extraordinary advance in observational astronomy. The STScI has been a partner in the project from the very beginning. I especially thank Ed Weiler of NASA Headquarters for his leadership in initiating the mission and guiding it to its present stage of maturity, with flight hardware being built and delivered.

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Status of the James Webb Space Telescope Integrated Science Instrument Module System


Goddard Space Flight Center, Greenbelt, MD 20771

Abstract.

The Integrated Science Instrument Module (ISIM) of the James Webb Space Telescope (JWST) is discussed from a systems perspective with emphasis on development status and advanced technology aspects. The ISIM is one of three elements that comprise the JWST space vehicle and is the science instrument payload of the JWST. The major subsystems of this flight element and their build status are described.

1. Introduction

The James Webb Space Telescope (Figure 1) is under development by NASA for launch during 2014 with major contributions from the European and Canadian Space Agencies. The JWST mission is designed to enable a wide range of science investigations across four broad themes: (1) observation of the first luminous objects after the Big Bang, (2) the evolution of galaxies, (3) the birth of stars and planetary systems, and (4) the formation of planets and the origins of life [1],[2],[3].

Integrated Science Instrument Module (ISIM) is the science payload of the JWST [4],[5],[6]. Along with the telescope and spacecraft, the ISIM is one of three elements that comprise the JWST space vehicle. At 1.4 mT, it makes up approximately 20% of both the observatory mass and cost to launch. In order to maximize efficiency with respect to mass, power, and nonrecurring engineering cost, the four JWST science instruments are not designed as traditional stand-alone systems; rather, the ISIM provides nine key subsystems that are shared by each of the science instruments. These are the: (1) optical metering structure, (2) ISIM Electronics Compartment (IEC), (3) harness radiator, (4) ISIM Command and Data Handling System (ICDH), (5) ISIM Remote Services Unit (IRSU), (6) electrical harness system, (7) thermal control system, (8) flight software system, and (9) on-board script system. In the following sections, we briefly describe the science instruments and provide detail on their shared support systems.

2. Science Instrumentation

2.1. The Near-Infrared Camera (NIRCam)

Near-infrared imagery is provided by the NIRCam instrument [7] & [8] shown in Figure 2. This camera provides high angular resolution wide-field imagery over the 0.6 – 5 micron spectrum. Table 1 summarizes the key elements of NIRCam’s capabilities, which remain aimed at providing efficient surveying in two filters simultaneously with a 2x2 arc minute field. The detector pixel scale is chosen to optimally sample the telescope point spread function across this wavelength range by use of a dichroic beam splitter. Two identical
optical modules image adjacent fields of approximately 4 square arcminutes to provide full redundancy for telescope wavefront sensing. The detector arrays for each wavelength regime in NIRCam have sensitivity matched to the needs of the particular channel – the short wavelength channel uses HgCdTe with a 2.5 $\mu$m cut-off and the long wavelength arm uses HgCdTe with a 5.1 $\mu$m cut-off. Using two types of detector material relaxes out-of-band blocking requirements for NIRCam’s bandpass filters. Occulting coronagraphy, yielding a rejection ratio of $\sim 10^4$, is provided in both long and short wavelength channels. All focal plane arrays support high cadence sub-array exposures to provide a high dynamic range capability for exoplanet transit observations. The instrument is cooled to $\sim 40$K using passive radiators.

The NIRCam Engineering Test Unit (ETU) and the associated instrument-level test program has been completed (Figure 2) and the ETU has been delivered to GSFC for fit-check integration and other testing with the flight ISIM. The NIRCam also serves as the wave front sensor for figure control of the JWST primary mirror and includes pupil plane optics (dispersed Hartman sensors, a grism, and defocus lenses) and deployable pupil imaging optics for this purpose. The NIRCam is being developed by the University of Arizona with Lockheed Martin Advanced Technology Center.

2.2. The Mid-Infrared Instrument (MIRI)

Imagery and spectroscopy over the 5 – 29 micron spectrum is provided by the MIRI instrument [9]. This instrument (Figures 5 and 6) provides broad-band imagery, low (1%) spectral resolution long slit spectroscopy, and medium ($\sim 10^3$) spectral resolution integral field spectroscopy (Table 2). The imaging mode includes both occulting and quadrant phase mask coronagraphy. The latter type enables very small inner working angle observations of stellar debris disks and exoplanet systems. When used in combination with the NIRSpec instrument, an optimally sampled integral field spectrum covering the whole 0.6 – 29 micron JWST wavelength range can be obtained at medium spectral resolution.
Figure 2: The NIRCam optics train layout

Figure 3: The NIRCam refractive optics design employs advanced high stability cryogenic lens mounts
Figure 4: The NIRCam engineering test unit

Table 1: Key NIRCam Features

<table>
<thead>
<tr>
<th>Feature</th>
<th>Description</th>
</tr>
</thead>
<tbody>
<tr>
<td>Wavelength Range</td>
<td>0.6 to 5.0 µm</td>
</tr>
<tr>
<td>Optical Design</td>
<td>Refractive with dichroic channels: 0.6-2.4 µm and 2.4-5.0 µm</td>
</tr>
<tr>
<td>Pixel Scales</td>
<td>0.0317 arcsec/pixel (short wavelength)</td>
</tr>
<tr>
<td></td>
<td>0.0648 arcsec/pixel (long wavelength)</td>
</tr>
<tr>
<td>Field of View</td>
<td>4 square arc minutes</td>
</tr>
<tr>
<td>Spectral Resolution</td>
<td>Multi-filters, λ/Δλ ~ 4, 10, 100</td>
</tr>
<tr>
<td>Coronagraphy</td>
<td>Focal plane and pupil masks available for both wavelength ranges</td>
</tr>
<tr>
<td>Bench</td>
<td>Bonded Be</td>
</tr>
<tr>
<td>Filter/Pupil Wheels</td>
<td>DC Direct Drive w/ position feedback</td>
</tr>
<tr>
<td>Pick-off mirrors</td>
<td>Adjustable in focus</td>
</tr>
<tr>
<td>Detectors</td>
<td>4Kx4K using 2.5 µm cutoff HgCdTe,</td>
</tr>
<tr>
<td></td>
<td>2Kx2K using 5.0 µm cutoff HgCdTe</td>
</tr>
<tr>
<td>Redundancy</td>
<td>Fully redundant opto-mechanical design w/ cross-strapped electronics</td>
</tr>
</tbody>
</table>
Figure 5: MIRI

Figure 6: The MIRI verification model
The MIRI development is a partnership between NASA and the European Space Agency. NASA/JPL is providing a cryo-cooler system to meet the 7K operating temperature requirements of this instrument, a 3 Mega-pixel Si:As detector system, and overall instrument flight software. The MIRI optical assembly, mechanisms, and instrument-level integration is provided by a European consortium led by the UK Advanced Technology Center and includes the Netherlands, France, Germany, Sweden, Belgium, Spain, Denmark, Ireland, and Switzerland. The MIRI has completed its verification model test program and is currently in flight model integration.

<table>
<thead>
<tr>
<th>Mode</th>
<th>Wavelengths µm</th>
<th>Spectral Resolution</th>
<th>Field of View</th>
<th>Comments</th>
</tr>
</thead>
<tbody>
<tr>
<td>Imaging</td>
<td>5.6,7,7.10,11.3,12.8</td>
<td>R=5</td>
<td>1.9x1.4 arcmin</td>
<td>Discrete filters</td>
</tr>
<tr>
<td></td>
<td>15,18,21 &amp; 25.5</td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>Coronography</td>
<td>10.65,11.3,16, and 24</td>
<td>R=10</td>
<td>25x25 arcsec</td>
<td>Lyot &amp; phase mask</td>
</tr>
<tr>
<td>Spectroscopy</td>
<td>5-11</td>
<td>R=100</td>
<td>5x0.2 arcsec</td>
<td>Long slit</td>
</tr>
<tr>
<td>Spectroscopy</td>
<td>5-28.5</td>
<td>R≈ 2000</td>
<td>3.5x3.5 to 7x7 arcsec</td>
<td>Integral field unit</td>
</tr>
</tbody>
</table>

2.3. The Near-Infrared Spectrograph (NIRSpec)

NIRSpec is the first multi-object spectrograph to be flown in space. It is designed to obtain near-IR spectra of 100 or more astronomical sources simultaneously at a spectral resolution of R≈ 100 over the 0.6 - 5 micron wavelength range, or at a spectral resolution of R≈ 1000 over 1 - 5 microns. The R≈ 100 mode employs a single prism as its dispersive element and is intended for measuring the redshifts and continuum spectra of faint galaxies. The R≈ 1000 mode utilizes diffraction gratings to cover the 1 - 5 micron spectral region, and is primarily intended for detailed follow-up observations using conventional nebular emission lines as astrophysical diagnostics. Lastly, three R≈ 3000 gratings, also covering 1 - 5 micron, will allow kinematic studies of individual galaxies to be carried out in single object or integral field mode.

The NIRSpec design employs all reflective optics, with most of its optical and structural elements manufactured out of an advanced ceramic material (SiC - 100). The optical chain (Figure 7) has three main components. The fore-optics re-image and magnify the focal plane image of the JWST telescope onto the slit selection mechanism. The collimator converts the light emerging from each slit into a parallel beam and projects it onto the grating wheel which carries six (flat) reflective gratings, a (dual pass reflective) prism and a (flat) mirror for target acquisition. The camera finally focuses the dispersed collimated light coming off the grating onto the detector array. A filter wheel located at an internal pupil of the fore-optics carries order separation filters for the diffraction gratings and also serves as the instrument shutter.

The NIRSpec multi-object aperture control (Figure 7) is a programmable Micro-Shutter Array (MSA) provided by the Goddard Space Flight Center (GSFC). The MSA is made up of four sub-arrays, each comprising 365 x 171 individually programmable shutters. The open area of each shutter is 200 mas wide and 460 mas long. The active area of the whole MSA spans a field of view measuring 3.4 x 3.6 arc-minutes on the sky. In addition to the programmable micro shutters, the MSA also carries several fixed slits that can be used for high contrast observations of single objects at any of the three spectral resolutions, and an integral field unit for use with three R≈ 3000 gratings.
The NIRSpec detector system, also provided by GSFC, consists of two close-butted 2k x 2k Rockwell HgCdTe detector arrays. In order to optimize the detector noise limited sensitivity of NIRSpec, the detector samples the spectra at a relatively coarse 100 mas per pixel.

NIRSpec is being built by the European Space Technology Center with EADS Astrium.

2.4. The Fine Guidance Sensor (FGS) and Tunable Filter Instrument (TFI)

In order to provide sensing for the observatory’s 7 mas line-of-sight stability control, an image-based fine guidance sensor is located in the ISIM with the science instruments. The optical assembly of the FGS instrument (Figure 8) consists of two fully redundant modules. These modules image two adjacent fields of 2.3 arcminutes on a side, with 68 mas pixels. This scale will enable the required guiding precision of 4.9 mas at 16 Hz, with a faint limit that gives 95% guide star probability in the emptiest part of the sky. The two redundant detector systems are fed by a single set of relay optics, and operate over a wavelength range of 0.6 to 5.0 microns.

The TFI shares an optical bench with the FGS. It consists of a low order Fabry-Perot tunable filter that images a similar field of view (2.2’) with a 65 mas pixel scale and operates at a resolving power of approximately 100 over the 1.5 to 5.0 µm spectral range with a non-functional region between 2.6 and 3.1 µm. The TFI optics image a single order of interference onto the focal plane array such that the scanning etalon produces a spatially resolved spectra of all objects in the field of view. Coronagraph spots at the edge of the field provide central light suppression of 10^{-4} at 1 arc-sec radius.

3. Science Support Systems

3.1. Optical Metering Structure System

The ISIM Optical Metering Structure (Structure) is a state-of-the-art bonded composite structure, and is the first large (2.5m length) cryogenic optical metering structure to be
built and tested for JWST (Figure 9). The Structure houses the Science Instruments (SI), provides structural support for ground testing and launch, and maintains the correct position of the science instruments with respect to each other and the Optical Telescope Element (OTE) to a precision of approximately 80 microns upon cryogenic cycling and launch loading. The cryo-cycling portion of this allocation has been test verified to 25 microns. It is a three-dimensional frame of 75 mm cross-section square tubes made of carbon-fiber/cyanate-ester composite material, bonded together with gussets made of composite and clips made of Invar 36. Large node fittings made of Invar 36 are used in a few select locations, most notably at the interfaces to the OTE and to Ground Support Equipment (GSE). The SIs are supported by saddle fittings made of Invar 36 bonded to the tubes, and attached via Titanium Interface Plates that are individually customized to provide precise alignment of the Instruments with respect to each other and to the Telescope. Composite material was selected for high specific stiffness and low Coefficient of Thermal Expansion (CTE). Invar 36 was selected for low CTE, but at the cost of low strength and high density. Titanium was selected for high specific strength. The tube topology was driven by OTE and SI accommodation, specifically the need for stable attachment to the OTE near the primary mirror, stable attachment of the SIs near the light coming from OTE, and the limits of physical space available within the overall JWST architecture.

The key design challenges were to support the SIs in precise alignment, maintain alignment stability through 270 degrees of temperature with cryogenic cycling, survive the large loads applied during rocket launch, survive exposure to extreme cryogenic temperatures (27K), and distort during cool-down from room temperature to cryogenic temperatures less than 0.5mm. In addition, critical JWST optical performance predictions are made using mathematical models, including a math model of the ISIM Structure. Accurate correlation of the Structure thermal distortion math model to measured thermal distortion test results was required and has been achieved on the flight structure.

To address these challenges, the project executed a comprehensive "building block" development program to measure cryogenic cool-down distortion, room temperature strength, and cryogenic temperature strength of the constituent materials and structural elements at the following successively higher levels of assembly: 1) composite lamina/laminates, adhesives, and metals, 2) composite tubes, 3) bonded tube/gusset/clip joints and bonded tube/fitting joints, 4) a representative bonded frame, and finally 5) the flight Structure. All
of the non-flight development work needed to validate the flight design and correlate the Structure thermal distortion math model concluded in 2008, the Flight Structure build was completed in 2009, and cryo-distortion testing of the Flight Structure at GSFC completed in May 2010 (Figure 9). Photogrammetric measurements with > 0.025 mm accuracy were made at room temperature and cryogenic operational temperature, demonstrating that Structure cool-down distortions are well within the required 0.5mm limit. This is a major risk reduction milestone, and a major success for the JWST program.

3.2. ISIM Electronics Compartment (IEC)

The ISIM Electronics Compartment (IEC) is one of the primary engineering challenges of the observatory. It accommodates warm (300K) instrument electronics that must reside close to the cryogenic (40K) science instruments due to cable length restrictions. As a result, it must be located on the cryogenic side of the sunshield (Figure 10). The eleven electronics boxes in the IEC dissipate a total of 230W which must be rejected to space such that the heat leak (both radiative and conducted by structural interfaces and electrical harnesses) to the surrounding cryogenic system is controlled to approximately 350 mW representing roughly 1/3 of the total ISIM cryogenic heat load to the Observatory.

Control of the IEC power dissipation is critical to the operation of the JWST. Precise control of heat leaks is essential to achieving thermal balance of the telescope and instruments as well as avoidance of infrared stray light. The IEC power is rejected to space via a flat panel radiator that is located behind an array of composite cylindrical mirrors (baffles) to create a beam pattern that avoids impingement on adjacent cryogenic surfaces (Figure 11).

The IEC protective enclosure (Figure 12) utilizes composite technology. The "Shell’s" composite construction is robust and can properly house the electronics boxes and yet strategically emit the heat so it does not interfere with the science equipment. The shell is primarily made of fiberglass and Nomex® honeycomb core. The baffles incorporate a unique gold vapor deposition process over its fiberglass structure to give it a low emissivity surface.
Figure 10: The IEC and its location in the JWST

Figure 11: The IEC radiators must emit with a 20 deg beam pattern
The radiator panels are constructed of expanded aluminum honeycomb core sandwiched by two thin face-sheets. The majority of the components for the IEC are fabricated with integration and test underway (Figure 13).

### 3.3. Harness Radiator System

The harness radiator component of the ISIM system was a late addition to the architecture (Figures 14 and 15). It was added after the IEC was implemented and is intended to capture and radiate to space the parasitic heat flowing in electrical harnesses attached to the electronics within the IEC. These electronics are standard designs and operate in the -20°C to 40°C range. This warm set of electronics is in close proximity (~2 m) to the cryogenic instruments. As a consequence, it is necessary to intercept the heat from these short thermal transition harnesses and radiate it to space with a dedicated radiator system.

A detailed set of requirements for the Harness Radiator (HR) have been developed and are being implemented. Among these are its overall mass, the launch and testing environments, and mechanical interface loads. Three of the principle performance requirements will be described here.

The most critical performance requirement for the HR is the allowable heat load to the instruments and telescope due to the harnesses. This heat load is used to establish the heat intercept needs of the Harness Radiator. The total allocated heat load to the ISIM cryogenic section is 460 mW. Of this, an allocation of 150 mW has been given to the electrical harness that are managed by the HR. The HR is therefore required to reduce the heat load below this 150 mW value. The harness radiator must also limit the total heat load to the Optical Telescope Element, and the Backplane Support Fixture (BSF) to less than 30 mW. At present, the HR design is predicted to reduce the heat load to the instruments to less than 95 mW.

A second key performance requirement is the need to accommodate the relative motion of the ISIM, the telescope (BSF), and the IEC (Figure 15). Each of these components are mounted on kinematic (or iso-static) mounts. So motion between these components will be seen both during launch, and during the system cool-down to reach on orbit operating temperatures. The requirement for this displacement is to accommodate 18 mm of relative motion between the HR and the BSF.

The third key performance requirement is to shield the electrical harnesses from micro-meteor (MM) damage. To achieve the best thermal performance of the HR and the harnesses, the HR design routes the harnesses external to the set of enclosures in which the instruments are mounted. This routing provides the best view to space for the harnesses but also exposes the harnesses to potential damage from micro-metors. An extensive study
Figure 13: Full scale mock-up of the IEC showing electronics boxes integrated to the interior surface of the IEC flat panel radiator (top left), the flight IEC enclosure (top right), and the flight baffle mirror assemblies (bottom).

Figure 14: The harness radiator assembly
of the behavior, performance, and potential damage mechanisms of exposed harnesses was conducted by the NASA Engineering Safety Council (NESC). A set of study guidelines and questions regarding micro-meteor damage and performance were developed by the ISIM engineering team and provided to the NESC for evaluation. The NESC team reviewed past mission harness performance, conducted analysis of the MM environment, determined the damage potential of the harnesses based on these studies, and conducted hypervelocity MM tests of representative electrical harnesses. The recommendation of the NESC was that harnesses routed external to the ISIM enclosure should be shielded by 1 to 2 mm of titanium in order to protect the harnesses and meet the overall probability of success rates imposed on the harnesses. This recommendation has been implemented.

The HR accommodates two basic types of electrical harnesses. These are the instrument detector interface harnesses, and the instruments engineering (lamps, mechanisms, heaters, temperature monitoring) interface harnesses. The detector interface harnesses are flat laminated cables composed of fine gauge wires (36 AWG and smaller). The engineering interface harnesses are also fine gauge wires (26 AWG and smaller) but are of a more standard wire harness construction using kapton and Teflon insulated harness bundles.

To improve the ability to draw heat from the harnesses and radiate it to space, the HR has implemented four distinct radiator panels each mounted to a common backbone support structure. As the harnesses cross over each of these panels, a harness clamp has been implemented which compresses the harnesses to a controlled compression value. Testing has been done to show the harnesses survive and properly function while compressed at clamp pressures up to 200 psi. Because there is differential contraction of the harnesses as well as the different panels of the Harness Radiator, strain relief loops/hoops are designed into the harness sections between the harness clamps.

The Harness Radiator has been designed to allow harnesses for the four science instruments to be separately integrated in any order. This has been done to improve overall
3.4. Command and Data Handling and Remote Services Subsystems

Figure 16 shows the ISIM Command and Data Handling (IC&DH) and Remote Services Unit (IRSU) in context of the overall ISIM electrical system. The IC&DH/IRSU hardware and software together provide four major functions: (1) Coordinate collection of science image data from the NIRCam, NIRSpec MIRI, and FGS instruments in support of science objectives; (2) Perform multi-accum & lossless compression algorithms on science image data to achieve data volume reduction needs for on-board storage and subsequent transmission to the ground; (3) Communicate with the Spacecraft Command and Telemetry Processor (CTP) to transfer data to the Solid State Recorder (SSR) and prepare for subsequent science operations; and (4) Provide electrical interfaces to the Thermal Control Sub-system.

The IC&DH hardware consisting of 7 separate assemblies is shown in Figure 17 along with the general science image data collection, processing and storage flow. The IRSU consists of 3 assemblies and is shown in Figure 18.

The IC&DH/IRSU communicate with the Instruments over Spacewire Point-to-Point interfaces operating at link rates between 9.4 Mbps to 60 Mbps. The IRSU contains a Multi – SpaceWire Concentrator Card (MSC) that interfaces via SpaceWire to the instruments and concentrates and routes the packets to the IC&DH subsystem. The internal SpaceWire communications within the IC&DH occur at a link rate of 80 Mbps.

The IC&DH provides the centralized ISIM Command and Data Handling functions for the NIRCam, NIRSpec, MIRI, and FGS instruments and unifies the Image collection system integration and testing workarounds by allowing one instrument to be removed, with its harness, while another instrument can continue testing.
with a common Spacewire point to point topology. The JWST IC&DH is unique in that it centralizes the C&DH functions for NIRCam, NIRSpec, MIRI, and FGS away from the detector electronics region in order to manage stringent thermal requirements while maintaining the ability to receive image data at peak detector readout rates. Multiple SpaceWire 4-port Routers within the IC&DH and IRSU provide high speed extraction of image data from each of the instruments and data routing to each of the Focal Plane Array Processors (FPAP). The FPAPs are unique in that they are dynamically configurable and assignable to any one of the 18 total NIRCam, NIRSpec, MIRI, and FGS instrument Sensor Chip Assemblies.

The IRSU also provides several electrical functions for the Thermal Control Sub-system. IRSU provides power to the heaters used to perform contamination control and to operate the Instrument Trim heaters. Also, IRSU interfaces to temperature sensors within ISIM and will perform readout of those sensors, relaying that data to the S/C for downlink. The IC&DH/IRSU interfaces to each of the Instruments support unique and independent science operations while providing a standard protocol across the interface ultimately resulting in integration and test risk reduction during ISIM Integration and Test.

Finally, the IC&DH/IRSU has been architected to meet high data throughput and processing requirements while minimizing overall lifecycle costs and risks. Both ICDH and IRSU Flight Units have been fabricated and are undergoing final environmental tests prior to delivery.

3.5. Cryogenic Thermal Control System

The cryogenic portion of the ISIM lies on the anti-sun side of the large sunshields, permanently shaded from the sun, the warm spacecraft bus and solar panels. When operating, the detectors and optics of the three Near-infrared Instruments must be cooled to the range of approximately 36 to 40K and the Mid-Infrared Instrument (MIRI) must be cooled to approximately 6K. In order to support long integration observation of faint objects, the thermal environment of the instruments must also be kept stable to within 0.1K and 0.25K over 10,000 seconds, and 24 hours respectively. Additionally, thermal isolation of the science instruments from each other enhances their stability and maximizes science data return. To accomplish these key requirements, the entire instrument suite is passively controlled to this temperature range via direct instrument coupling to dedicated radiators. Individual thermally isolated radiators are provided to cool the NIRCam, FGS, and NIRSpec instruments, the MIRI electrical harnesses and the bases of the MIRI support struts. The instruments
are connected to their dedicated radiators through a complex series of segmented high purity aluminum heat straps designed to minimize the thermal gradient between instrument and radiator. These flexible aluminum straps are supported in composite channels, which are mounted to the ISIM structure and the surrounding BSF through composite flexures. Figure 19 shows a photo of three proto-flight strap segments with support flexures, prior to vibration testing at the vendor.

A two stage mechanical cooler maintains the MIRI at its operating temperature of approximately 6K, as this temperature is too cold to be achieved using only passive techniques. Tests performed at the MIR instrument provider, combined with thermal analyses performed at the GSFC, recently indicated that the cooler’s 6K cold stage would have had insufficient heat lift margin when the instrument is immersed in ISIM’s 40K thermal environment. As a result, the ISIM is also providing a thermal radiative shield around the MIR instrument, cooled through direct contact with the cooler’s second stage heat exchanger operating at approximately 20K. The resulting trade of heat loads between the cooler’s second and third stages raises the performance margin on the 6K stage to acceptable levels. Measures taken to assure the ISIM compartment reaches its designated operating temperatures include assignment of milliwatt-level heat load allocations to each of the instrument optical assemblies, and establishment of interface temperature requirements with the Observatory. Additionally, extensive efforts have been devoted to the design and detailed analysis of the Harness Radiator (see Section 3.3), a multi-stage radiator designed to reject most parasitic heat from the instrument harnesses to space as they enter the cold ISIM from the ambient temperature IEC (see Section 3.2).
The carbon-based composite ISIM structure is a potential source of entrapped water. A contamination avoidance cool-down procedure has been devised to avoid contamination of the instrument detectors and sensitive optics during the instrument cool-down following launch: contamination control heaters on the instruments will be activated to keep them warmer than 165K until the entire ISIM structure cools to below 140K, entrapping the moisture within it or on non-sensitive surfaces. Once the instruments reach operational temperatures, it would take multiple faults to cause significant contamination onto the instruments due to release of entrapped water. A procedure has therefore been developed to decontaminate the instruments should this occur; the same heaters will be used to warm the instruments alone to 165K, before the structure has time to warm to 140K. Special control algorithms have been devised to accomplish both the contamination avoidance and decontamination.

Small "trim" heaters provide make-up heat (if needed) to each SI to assure minimum operational temperatures are maintained during science operation. Temperature monitoring throughout the cryogenic ISIM is provided by 93 calibrated CERNOX sensors on the ISIM structure, instruments, radiators, harness radiator, and other surfaces to verify that thermal requirements are met, and to be able to monitor ISIM health and safety at all times, including during launch and at other times when science instrument telemetry is unavailable. Figure 20 is a schematic of the ISIM and instrument thermal control scheme.

3.6. Flight Software System

The ISIM Flight Software (IFSW) is composed of a set of independent tasks that interoperate to process transactions originating from both internal and external sources. The tasks are contained in an operating system environment that is adapted to the mission-specific hardware components with a set of custom device drivers shown in Figure 21 as OS Services. The tasks can be classified into three categories: (1) Command and Telemetry Services, (2) Applications that operate the Science Instruments (SIs) and support science operations, and (3) System Support Services that manage the task environment and external system interfaces.

These tasks have all been developed using IBM® Rational® Rose Real-Time® (RRT), an object-oriented development environment that features visual design of tasks as state machines and definition of inter-task messaging as "ports" with "protocols", and which utilizes conventions established in the Unified Modeling Language™ (UML™) Transitions between states are defined graphically, and behavior code can be associated with transitions or with entry to or exit from a state. Transitions are triggered by specified events such as expiration of a timer or receipt of a signal at a port. The framework code is automatically generated and combined with the behavior code by RRT, which greatly reduces the amount of code that must be written manually. The same development environment was provided to the Science Instrument software development teams to promote ease of integration when their applications are delivered to the ISIM Flight Software Lab at Goddard.

Goddard has defined and implemented, using RRT, a set of standard task interface ports to which every task conforms that allows it to send command packets, receive telemetry packets, receive internal table values for task configuration, receive system time information, perform file operations, use centrally managed timers, and respond to check-in requests to verify that each task is still running. At startup time, each task binds these ports and registers key information that the system needs to provide those services. To date, at least two versions of each SI Application have been integrated with only few and minor problems that were for the most part straightforward to identify and remedy.

To illustrate how the tasks interoperate to provide the desired instrument operation functions, we will describe a few transactions tracing the path taken by successive messages the transaction comprises. Transactions typically originate with the receipt of a command, the expiration of a timer, or the arrival of a telemetry packet. All commands entering the
Figure 20: ISIM thermal control schematic
system are messages, sent by a task from its command port, that are queued for service by the Command Manager Task which sends the command to the task indicated by the Application ID in the command packet. A command from the ground will be received by the spacecraft communications system and transferred to the ISIM by the spacecraft Command and Telemetry Processor (CTP). The IFSW CTP Interface Task, which handles the exchange of all commands and telemetry between ISIM and the spacecraft, places the command into a message sent out its command port to the Command Manager Task.

The primary command source in normal operations is the Script Processor Task (SP), which runs scripts written in JavaScript upon receiving a command to do so. The script execution is performed by a JavaScript engine running as separate task that supports multiple concurrent JavaScripts running independently of each other. A set of extensions to the JavaScript language have been implemented that provide the interface to SP, which in turn can access ISIM FSW services through the standard task interface ports. Also, to provide communication between independently running JavaScripts, there are extensions that can set and retrieve the values of shared parameters.

A collection of JavaScripts, stored as ASCII files, make up the Operations Scripts System, discussed in Section 3.9, which provides the capability for automatic operations (see Figure 22). A JavaScript can send a command by communicating to SP, which sends the command packet to the Command Manager. If the command that originated from a JavaScript is an SI function, such as to move a Grating Wheel to a certain position, the command would be routed to the Application Task for that SI. That SI Application Task may generate many commands to the SI hardware to complete the requested operation. These hardware commands are sent via the Command Manager to the bus interface task, either 1553 or SpaceWire, that connects with the SI component being commanded.

The same SI Application Task can send a command to the SI hardware, via the Bus Interface Task, requesting status information to verify that the previous command was properly executed. This hardware status information will be received by the Bus Interface Task and formatted into a Telemetry Packet that is sent out its telemetry port to the Telemetry Manager, which routes the packet to any task that has subscribed to receive packets identified with that Application ID. When the SI Application Task receives that packet, it can decide whether the command succeeded, the operation is complete, or an error has occurred. Every task issues a “housekeeping” telemetry packet at regular intervals providing general

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**Figure 21: ISIM Flight Software Layered Task Architecture**

<table>
<thead>
<tr>
<th>Applications</th>
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</thead>
<tbody>
<tr>
<td>Command and Telemetry Services</td>
</tr>
<tr>
<td>System Support Services</td>
</tr>
<tr>
<td>OS Services</td>
</tr>
<tr>
<td>Physical Layer</td>
</tr>
</tbody>
</table>

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status information that can be viewed in real time during a ground contact, or recorded on the Solid-State Recorder and analyzed later on the ground. Also, every task can issue event messages to report errors or successful operations.

The Telemetry Manager Task filters telemetry bound for the SSR and real-time link based on parameters stored in on-board filter tables because the channels that carry telemetry data to the spacecraft CTP have limited capacity. In general, the SI Application Tasks will sustain a higher communication rate through their instrument hardware interface to monitor and control their associated instrument hardware than the rate of commands and telemetry over the user interface through which the SI System is operated, typically by JavaScripts via the Script Processor.

The Data Acquisition Function is implemented as several tasks that collaborate to marshal science image data through the system to the SSR, and for providing selected image data to other tasks to support boresight guiding and target acquisition. Image data from science observations is never processed by the ISIM Processor. To perform a science data exposure, the user first configures the Data Acquisition Task specifying the entire sequence of images that the exposure is expected to generate. Data Acquisition uses this information to generate a set of frame processing instructions, which it loads into one of the 18 channels provided by the three Focal Plane Array Processor (FPAP) cards. If multiple detectors will be exposed simultaneously, a separated channel is configured for each detector. Once configured, the Data Acquisition Task and the FPAP channels wait for the arrival of science data. Then, the desired SI hardware detector subsystem is configured by sending commands to the SI’s Application Task, which in turn sends commands to configure the hardware. When the exposure is commanded to start, data flows into the designated FPAP channel and each frame is processed in accordance with the associated frame processing instruction, which includes information about when to output an image to the SSR. All the image processing is done by the hardware channel so very little CPU time is consumed to
capture and record science data. The Data Acquisition Task responds to interrupts and creates header and trailer records for the file sent to the SSR and then triggers the various elements in the science data path to orchestrate the movement of data from the FPAP channels to the SSR. Data Acquisition can also be commanded to retrieve image data subwindows and either write them to files or send them in packets so that the locations of stars can be determined. The FGS System uses this data to send guidance information to the spacecraft CTP to stabilize the telescope by fixing on a guide start. JavaScripts programmed to locate a target star and move it into the desired instrument aperture also use image data.

The remaining software elements perform various monitoring and maintenance functions to keep the software running smoothly, detect problems with the system operation, and restart the software or the processor if uncorrectable problems occur. The Telemetry Monitoring task checks telemetry values against limits stored in a table, and if a value stays outside of the specified limits for a specified length of time, a stored command sequence can be initiated by the Stored Commanding task to take a predefined corrective action. The Memory Scrub task systematically reads all memory locations to detect memory errors caused by radiation using the Error Detection and Correction circuitry in the CPU hardware. The Health and Safety task periodically sends a check-in request to every other task and if a task fails to check in, a software reset is initiated. The Executive Services task can be commanded to send very detailed diagnostic information to ground to assist with analysis of anomalies and the Memory Manager task can update portions of memory to install patches in software to correct minor errors that are discovered.

3.7. Operations Scripts System

The Operations Scripts System (OSS) is a collection of JavaScripts (Figure 23) that execute planned science observations and supporting activities, which are specified as a set of text files that are uploaded to the ISIM file store after ISIM initialization. The JavaScripts remain on board to carry out requests uplinked to the spacecraft from the ground nominally every 10 days. These request files are organized around the concept of a visit, which refers to a group of associated science, calibration or engineering activities. A science visit corresponds to a single telescope pointing and includes all the activities that are associated with a single guide star, beginning with the spacecraft slew, and including the guide star acquisition, target acquisition, SI configuration, science exposures, and any small telescope movements needed to fulfill the science objectives.

An Observation Plan, maintained as a JavaScript object, consists of a list of Visit Records. Each Visit Record contains the acceptable execution time window and points to a
Visit File that contains lists of activities organized as groups of activity sequences. All the sequences in a group are executed in parallel, and all sequences in a group must complete before advancing to the next group. Groups are processed serially. Observation Plan segment files are uploaded and appended to the on-board Observation Plan JavaScript object. Visits are executed in the order they appear in the Observation Plan. After a visit is executed, that Visit Record is deleted from the Observation Plan and the corresponding Visit File is deleted as well. The Activity Sequences within a Visit File consist of activity statements that are executed in order. Each activity statement refers to an activity JavaScript and contains parameters that are passed to the JavaScript. Every type of activity for each SI has a corresponding JavaScript to perform that activity, and parameters in the calling statement are used to control settings that vary from one operation to the next, such as the image window size or which filter wheel to use.

Performing SI operations involves sending commands to the appropriate SI Application task and requesting telemetry back to confirm successful execution of the command. To facilitate the construction of these commands, OSS contains a dictionary file of all the commands that can be sent that provides the format for each command and its associated parameters. There is a similar dictionary file for all the telemetry values that can be requested that specifies the telemetry packet ID and offset where that parameter can be found in the packet. There are interface JavaScripts that provide the access to this information to the JavaScripts that are performing these command and telemetry operations.

The reading, interpretation and execution of the Observation Plan and the Visit Files are conducted by a set of JavaScripts, collectively called the Observation Plan Executive (OPE), that run under the control of the Script Processor Task (SP). SP can run up to ten independent JavaScript threads each in a separate "context". This allows the OPE to simultaneously manage the Observation Plan and the Visit File execution, to manage multiple concurrent visit file activities, to interact with the spacecraft to coordinate reaction wheel momentum unloading, and to respond to requests from the ground to start, stop or modify the on-board Observation Plan. SP provides the interface between the JavaScripts and the ISIM FSW services described in Section 3.8, which include sending and receiving commands and telemetry and reading and deleting files. A command from the ground to SP to run the top-level OPE script starts the automated observing process, which continues as long as there are visits to process, until a command is sent from the ground to stop or suspend automatic observing, or until an error occurs that requires ground attention.

SP carefully monitors and regulates resources for JavaScript execution within each context including memory allocation and release, CPU execution time slicing, and limiting the frequency that telemetry packets are issued. Limiting CPU execution time is necessary because some scripts, such as the ones that perform target acquisition, can be very CPU intensive and can run for a long time. Limiting how long each context can use the CPU ensures that tasks with real-time deadlines always have enough time to perform their functions.

The fact that the OSS is written in JavaScript and stored on-board as text files is significant because this gives the operations personnel greater visibility, control and flexibility over the telescope operations. As they learn the ramifications and subtleties of operating the instruments, they can modify the JavaScripts and, after thorough testing in a ground facility, they can simply replace an onboard file to make the change. The JavaScripts are written using the terminology of the user’s domain of commands and telemetry. This is in contrast to the ISIM FSW tasks, which involve real-time state machine logic, intricate hardware interaction, intertask message routing, a hierarchy of task priorities, and complex operating system interfaces. The OSS is in the user domain and is consistent with the users’ perspective, needs and objectives. The ISIM FSW is in the system domain and provides the user with access to the functions and features of the system without the need for detailed knowledge of the underlying implementation.
Another significant feature of the OSS, a feature that pervades the end-to-end science operations philosophy of JWST, is the notion of event-driven operation. In the JSWT context, event-driven operation means simply that the sequence of operations is specified, and that the next step is performed when the previous step completes. Earlier generations of spacecraft relied on an uplinked command sequence stored in “command memory” with every command specified down to the bit with an associated time tag. The on-board computer would simply read each command, and would, at the designated time, send the command to the intended piece of hardware. This meant that every detail of spacecraft and science operations had to be modeled on the ground so that accurate calculations of the command execution times could be performed and that worst-case times for every operation had to be assumed. If an operation failed, for example, because a guide star was not acquired, operations stopped. At the next real-time contact, the ground staff would discover the problem and engage in a flurry of activity to create a new command load, which would be uplinked at a subsequent real-time contact. As flight processors became faster and radiation-hardened memory increased in capacity, on-board automation evolved as well. With event-driven operations, if a guide star acquisition fails, that visit is skipped and the OPE simply advances to the next visit, which greatly increases the observing efficiency of the telescope and greatly simplifies the ground system as well.

4. Optical Integration and Test

The major subsystems that make up the ISIM Element are separately verified before delivery for Element-level integration and test (I&T), during which the Element is verified to meet its requirements after assembly: The SI developer verifies that the SI meets its optical performance and optomechanical alignment requirements as a stand-alone unit, assuming nominal Optical Telescope Element (OTE) performance. The ISIM structure is verified to meet its mechanical alignment requirements, and it is optically tested after integration to verify that the whole ISIM meets its performance requirements after integration of the instruments with their supporting systems (Section 3) and exposure to simulated launch and operational environments.

4.1. Integration and alignment

For ISIM Element development, the JWST optomechanical coordinate system for the science payload, the “vehicle coordinate system” (VCS), is represented on interface tooling [10],[11]. The SIs are optically aligned to their mechanical interface to the ISIM structure using this tooling (Figure 24a). The ISIM Element is aligned to its interface to the OTE using another precision fixture, the ISIM Test Platform (ITP) (Figure 24b and 25a).

The alignment fixture used for SI alignment and verification to the ISIM structure is called the ambient science instrument mechanical interface fixture (ASMIF). It is a metal platform about the size of the SI footprint on the ISIM structure. When the SI is mounted to its ASMIF, the ASMIF supports the full SI load. The type of metal and dimensions vary from SI-to-SI, because each interface is mechanically different and each SI has a different mechanical design. The ASMIF has a flight-like, precision representation of the SI-ISIM interface (i.e., tapped holes, precision pin holes and slots for clocking pins). The ASMIF is only used at ambient. The SI developer relates the cryogenic alignment of the SI to this ambient fixture. The ASMIF is thus a "master gauge" for SI-ISIM alignment.

The ITP is the alignment fixture used for ISIM alignment and verification relative to its interface to the OTE. The ITP is an invar platform that holds the ISIM structure and SIs and, thus, has a footprint of ~2.2 x 2.2 m and contains OTE-like interface features. The ITP is a bolted and pinned, invar structure that is stress relieved for optimum stability and repeatability at ambient and cryogenic conditions. The difference between the ITP and
Figure 24: a. The ETU NIRSpec instrument (middle, left) during post-shipment metrology operations. b. Foreground: The ISIM Structure (left) undergoing coordinate system metrology using LIDAR and theodolite instruments. Background: The ITP is being readied for cryogenic testing.

Figure 25: a. The ITP surrogate telescope backplane, showing integration of the MATF fixture in preparation for cryogenic metrology. b. CAD rendering of the Flight ISIM structure and cryogenic alignment performance verification via PG in the SES chamber.
a nominal telescope backplane is an important factor in ground alignment adjustment prior to delivery of ISIM to OTE for final integration.

The SI interface on the ISIM structure is expected to change very little (i.e., $\sim 0.6$ mm in translations) as a result of temperature change from ambient to the cryogenic operating temperature and all other ground-to-orbit effects (e.g., moisture desorption, g-release uncertainty). The various sources of ground-to-orbit change have been identified and estimated based on extensive modeling validated by component and assembly test articles. The levels of change are small compared to ISIM top-level performance requirements. We are therefore able to treat excursions from nominal in the SI-ISIM interface as error budget terms.

Commercial laser tracker, LIDAR, and theodolite instruments are used for ambient metrology and envelope measurement[12]. Photogrammetry (PG) is used for cryogenic metrology of the ITP and ISIM structure (Figure 25b)[13]. Cryogenic characterization of the ITP was completed in early 2010. The ITP dimensional change from ambient to cryogenic temperatures is in good agreement with model expectations. The structure cryogenic alignment was measured in spring 2010, with good results. The custom, cryogenic PG system performed better than its required measurement uncertainty.

4.2. Optical testing and verification

Before integrating the ISIM with the OTE, the ISIM is verified to meet its requirements relative to its interface. Furthermore, the performance of the SIs is calibrated at the ISIM-level in order to support later OTE-level ground testing and on-orbit, Observatory-level commissioning. This battery of optical test and verification is accomplished using a full-field, cryogenic optical stimulus, the OTE SIMulator (OSIM) developed by GSFC and Ball Aerospace (Figure 26) [14].

For the optical verification or “end-to-end” test of the ISIM element, the OSIM is used to simulate a star field generated by a nominal OTE, to within limits of alignment uncertainty, fabrication of the OSIM optics, design residual, etc. In order to meet OSIM-to-nominal-telescope-interface alignment requirements, OSIM must calibrate its output relative to this mechanical interface on the ITP. This calibration is performed during a series of optical tests prior to testing ISIM with OSIM.

During the calibration of OSIM, PG measures the location and orientation of the Master Alignment Target Fixture (MATH) (Figure 25a). The MATH is an invar plate that is pinned to the ITP, and allows OSIM to align its optical output to the mechanical OTE-ISIM interface. The MATH hosts an array of PG targets on its surface that are used for measuring MATH six degree-of-freedom alignment relative to the ITP. The MATH is also equipped with an array of five optomechanical targets that are not visible to the PG system, but are designed to provide an optomechanical reference for OSIM’s in situ alignment. The optical verification of the ISIM Element using OSIM is divided into multiple subtest categories: SI basic functionality and internal calibration; focus; pupil shear; boresight and field of view; vignetting and stray light; wavefront error; co-boresight stability; wavefront sensing. The first category of tests is performed to establish the functionality of the SI mechanisms, detectors, calibration lamps, etc. — this battery of checks will performed periodically to trend behavior after certain I&T evolutions and environmental exposure. The tests related to optical alignment (i.e., focus, pupil shear, boresight) are associated with ISIM Element requirements for SI alignment as an assembly of instruments. SI alignment relative to each other and, as a suite, relative to the ISIM-OTE interface is verified via this test, with significant analysis to account for ground effects. The co-boresight stability test is a thermal over-drive test, where the relative alignment of the SIs is monitored as a function of temperature. The SI wavefront error (i.e., image quality) is measured using phase retrieval techniques. The as-built wavefront error is calibrated for future reference during verification activities at the Observatory-level and on-orbit alignment of the telescope. Also in support
Figure 26: a. Ray trace schematic of the OSIM optical layout. b. CAD image of the ISIM Element and cryogenic performance verification using OSIM in the SES chamber.
of on-orbit telescope assembly and commissioning, the ability to perform wavefront sensing-related measurements is also verified.

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Design Aspects of the JWST Calibration Pipeline Software

R. Jedrzejewski

STScI, Baltimore MD 21218

Abstract. Some considerations relating to the design of the JWST Calibration pipelines are presented. Much of the code for calibrating the detectors can be shared among the instruments, and the fact that several observing modes are common to more than one instrument means that it makes sense to design the pipelines to calibrate observing modes rather than instruments.

1. Introduction – The HST Calibration Pipelines

One successful aspect of the Hubble Space Telescope mission has been the development of automatic calibration pipelines. These pipelines apply a sequence of calibration steps to HST datasets with the purpose of removing the instrumental effects that would otherwise compromise the quantitative interpretation of the data. Each calibration step applies an algorithm that is designed to reverse the effect of a phenomenon that degrades the data; for example, not every pixel in most area detectors has the same sensitivity, so the flatfield calibration step corrects for this by multiplying each pixel by a factor (the flatfield) that has been measured to equalize the sensitivities of each pixel.

The HST calibration pipelines were designed to use reference files to hold the quantitative information to be used for calibration. So, in the example given above, the flatfield is stored in a flatfield reference file that is identified by a header keyword. There are also header keywords that state whether a particular calibration step is to be performed or omitted.

The quality of calibration achieved in the automatic calibration pipeline is sufficient for a significant fraction of GOs and archival researchers. The existence of these pipelines means that calibrated data of good and uniform quality is available for every exposure taken by HST, and researchers can mine these data without having to become an expert on the intricacies of calibration and data analysis.

Users are not limited to the default calibrations provided by STScI when the data are retrieved from the HST archive. The HST calibration pipelines are distributed as part of the STSDAS data analysis software package, and users can run these pipelines themselves. They can choose which steps to perform and which steps to omit, and if they have calibration reference files that are superior to the default STScI reference files, they can use them instead just by changing the value of a header keyword.

2. History of the HST Calibration Pipelines

The first set of calibration pipelines were for the initial instrument set on HST (WFPC, FOC, FOS, GHRS and HSP). They were all programs written in the IRAF spp (subset pre-processor) language, reading and writing data in GEIS (Generic Edited Information Set) format. The use of spp mandated integration with IRAF, while GEIS was a data format used only for HST images that was also not platform-independent (i.e. you cannot just copy GEIS files from one platform to another and expect them to work).
The introduction of STIS and NICMOS in 1997 allowed some of these limitations to be fixed. Both the STIS and NICMOS pipelines were written in C, although they used the IRAF image i/o libraries through a layer (called hstio) that provided a consistent C interface to the applications. They also read and write data in FITS format, which is generally regarded as the ‘universal’ astronomical data format and is platform-independent. However, there was still a considerable tie to IRAF through the use of the image i/o libraries.

While the calibration pipelines for each instrument incorporated some overall systems engineering that provided several common design elements, the pipeline programs themselves were written completely independently with no sharing of application code. At the time of the installation of STIS and NICMOS, this was not unreasonable, since the new C-based pipelines for STIS and NICMOS shared little functionality, and the re-using spp code from the first-generation instruments was no advantage over writing the code in C, even if the functionality was very similar.

The Advanced Camera for Surveys (ACS) was installed in 2002, and development of the ACS calibration pipeline closely followed the STIS imaging pipeline, since the detector in the High Resolution Channel of ACS was the same type as the CCD detector in STIS. At this point, the calibration pipelines for each instrument had been following the independent ‘stovepipe’ development strategy, and the ACS pipeline was developed in the same way. In fact, the STIS imaging pipeline was used as a starting point for the ACS calibration pipeline, with the code morphing to meet the specific needs of ACS.

Similarly, when the calibration pipelines for WFC3 and COS were written, the stovepipe model was continued. The WFC3 pipeline was developed in C by starting from the ACS pipeline in the UVIS channel, and from the NICMOS pipeline in the IR channel. The COS pipeline was developed in Python, and so did not inherit much legacy code from the earlier spectrographs.

3. HST Pipeline Decisions

Some of the design decisions that went into the HST Pipelines were good.

3.1. Separation of code from data

A conscious effort was made to separate things that change often (data obtained from calibration programs) from those that don’t (algorithms, housekeeping, error checking). The data are stored in calibration reference files that are catalogued in the Calibration Data Base System, and updated as new measurements are obtained. Even some of the parameters to be used in the calibration pipelines are stored in reference files; for example, the many parameters used by Multidrizzle in the ACS pipeline are stored in a reference file that is read at run-time, and the optimum parameters can thus be changed without making any changes to the code.

3.2. Re-use of code from earlier instruments

Since much of the required functionality for the later pipelines had already been written for the earlier ones, it made very good sense to re-use as much of that code as was practical.

3.3. Regression test suite and robust error handling

Each delivery of the calibration pipelines to operations was accompanied by regression testing using a large and varied collection of datasets that exercise most of the pipeline functionality. In this way it was assured that datasets would be able to pass through the pipeline without encountering fatal errors, and that data products could be created for all types of observations.

However, some of the design decisions worked out less well...
3.4. Code duplication

This is probably the biggest issue. There are whole classes of functionality that involve duplicated and slightly altered code. For example, the directories in the calacs/acsccd and the calwf3/wf3ccd directories are strikingly familiar:

<table>
<thead>
<tr>
<th>calacs/acsccd</th>
<th>calwf3/wf3ccd</th>
</tr>
</thead>
<tbody>
<tr>
<td>acscd.c</td>
<td>blevdrift.c</td>
</tr>
<tr>
<td>blevdrift.c</td>
<td>doatod.c</td>
</tr>
<tr>
<td>blevfits.c</td>
<td>dobias.c</td>
</tr>
<tr>
<td>doatod.c</td>
<td>doblev.c</td>
</tr>
<tr>
<td>dobias.c</td>
<td>doccd.c</td>
</tr>
<tr>
<td>doblev.c</td>
<td>dofflash.c</td>
</tr>
<tr>
<td>doccd.c</td>
<td>findblev.c</td>
</tr>
<tr>
<td>dofflash.c</td>
<td>findover.c</td>
</tr>
<tr>
<td>findblev.c</td>
<td>getccdsw.c</td>
</tr>
<tr>
<td>findover.c</td>
<td>getflags.c</td>
</tr>
<tr>
<td>getacsflag.c</td>
<td>mainccd.c</td>
</tr>
<tr>
<td>getccdsw.c</td>
<td>mainccd.o</td>
</tr>
<tr>
<td>mainccd.c</td>
<td>mkpkg</td>
</tr>
<tr>
<td>mainccd.o</td>
<td>wf3ccd.c</td>
</tr>
</tbody>
</table>

These are completely independent files that are maintained separately, yet they fulfil very similar functions and share much code that is either identical or else differs only in whitespace.

3.5. Stovepipe Design

The code duplication in the previous subsection is largely a result of the stovepipe designs of the calibration pipelines. The duplication of the ccd calibration steps results from the compartmentalization of the instrument pipelines: each pipeline is a completely independent entity whose functionality is all contained within the directory structure of that instrument. Some of the lower-level library code is shared, but the main algorithmic code is written and maintained separately for each instrument.

This design was a perfectly natural result of the instruments being added over a long period of time. Each time a new instrument was added, the new calibration pipeline for that instrument was written in the most economical way, usually by taking the part of earlier pipelines that are most similar to the needed functionality and using them as a starting point for the new pipelines. In this way, the new pipeline could be up and running in a very short time, but with the disadvantage of a maintenance headache when updating the common code sections.

3.6. Different Data Product Formats

All HST data products are FITS format, but there the resemblance ends. The data products of each instrument are slightly different. STIS data is packaged with separate associated exposures in one file. NICMOS data has associated exposures in separate files. ACS and WFC3 data are like NICMOS data, except exposures using more than one detector are packages in one file, whereas contemporaneous NICMOS exposures using more than one camera are packaged in separate files. WFPC2 data packages all 4 detectors in 1 file, but they have different orientations on the sky because the readout amplifiers are in different quadrants.

This heterogeneity can make working with HST data confusing, to say the least.
4. JWST Pipelines

The JWST pipelines can take advantage of several factors.

4.1. Only One Generation of Instruments

JWST will be located at L2, and was not designed to be serviced. The instrumental complement at launch will last for the whole mission. So there will be no new instrument pipelines to add down the road. This means that the calibration pipelines for the science instruments can benefit from a more integrated approach, involving designing in elements that are common to several instruments, not keeping each instrument pipeline separate.

4.2. Only Two Types of Detectors

JWST will have 5 instruments capable of producing science data: NIRCAM, NIRSpec, MIRI, FGS-TFI and FGS-Guider. Of these, NIRCAM, NIRSpec, FGS-TFI and FGS-Guider will use the same Hawaii-2RG detector types. MIRI will use three of the same Si:As detectors. So many of the detector calibration steps will be identical algorithmically for the near-infrared instruments, and also for each of the MIRI detectors.

4.3. Identical Packaging for Science Data

The decision was made to put each individual exposure in its own data file. In JWST language, an exposure is a series of one or more integrations, and an integration is a sequence of consecutive non-destructive readouts of a detector. Some of these readouts may be averaged on-board the observatory to reduce the readout noise and lower the data volume, at the expense of increasing the susceptibility to cosmic rays. So unlike HST, the data from each detector goes in a separate file, even when the detectors are in the same instrument and were collecting data at the same time. The advantage of this is that each exposure can be treated identically when detector calibrations are being performed, but there is the disadvantage that it is necessary to develop the capability to keep track of which exposure goes with which when combining data from different detectors, as might occur when exposures are combined to produce a mosaic.

4.4. Common Science Modes

Even though the JWST science instruments have very different science goals, there is a significant degree of commonality in the science capabilities. The science modes of each of the instruments are detailed below:

NIRCam

- Imaging
- Coronagraphy
- Grism Spectroscopy

NIRSpec

- Multi-Object Spectroscopy
- Slit Spectroscopy
- Integral Field Spectroscopy
- Imaging

MIRI

- Imaging
• Integral Field Spectroscopy
• Slit Spectroscopy
• Coronagraphy

FGS-TFI

• Imaging
• Coronagraphy
• Non-redundant Mask Imaging

This commonality means that a significant amount of code can be shared among the instruments. Once the detector calibrations have been performed, many of the instrumental calibrations are purely mode-specific, rather than instrument-specific.

This raises an important question: why not develop the JWST calibration pipelines by MODE rather than by instrument? The algorithmic code for performing the calibration steps tends to depend much more on the observing mode rather than on the instrument. The differences between instruments are handled more by the reference files and some of the control code rather than the algorithmic code, so it makes much more sense to break up the JWST calibration pipelines by observing mode rather than by instrument.

5. Towards a Different Type of Calibration Pipeline(s)

Looking at all of the calibration steps that will be necessary to calibrate, for example, JWST NIRCam data, we find that the simplest breakdown is:

• Detector Calibration

• Mode-specific Calibration

The mode-specific calibration can be further decomposed:

• Detector Calibration

• Mode-specific calibration
  – Imaging Calibration
  – Coronagraphy Calibration
  – Spectroscopy Calibration

Doing the same exercise for MIRI and NIRSpec, one would find exactly the same decomposition:

• Detector Calibration

• Mode-specific Calibration
  – Imaging Calibration
  – Coronagraphy Calibration
  – Spectroscopy Calibration
Since each of the instrument calibrations would look almost exactly the same, it makes more sense to design the pipeline like this, without reference to the instruments at this high level. Now the pipeline might look something like what is shown in Figure 1: The detector calibration (calDetector) can be either a near-IR detector calibration (calNearIRDetector), or a mid-IR detector calibration (calMidIRDetector). Similarly, the instrument/mode calibration (calInstrument) can be either an imaging calibration (calImaging), a spectroscopic calibration (calSpectroscopy), or a coronagraphic calibration (calCoronagraphy). The spectroscopic calibration may be even further decomposed to either a slit spectroscopy calibration (calSlitSpectroscopy), a multi-object spectroscopy calibration (calMultiObjectSpectroscopy) or an integral-field spectroscopy calibration (calIFUSpectroscopy).

It is important to stress that this simplification will not be at the expense of performing the best calibration algorithms. It is quite likely that some of the calibration steps can be performed using identical algorithms on all the instruments. For example, we might expect that a nonlinearity correction will involve applying a low-order polynomial to the ramp, and this algorithm will apply to both near-infrared detectors and the mid-infrared detectors in MIRI. However, other calibration steps will be quite different for each instrument. The slit spectroscopy calibrations will almost certainly be quite different for MIRI and NIRSpec, and yet they may share some of the primitive operations so there will still be plenty of opportunity for code sharing.

6. Summing Up

For a variety of reasons, it makes sense to depart from the practice of coding each of the JWST instrument calibration pipelines separately, as was done for HST. We can make use of the fact that there are only two different types of detector on JWST to simplify the detector calibrations, and use the common science modes in the JWST instruments to craft pipelines around these modes rather than around the instruments. But the overall driver is making sure that the BEST algorithm is used in each situation, not the most convenient one.
Acknowledgments. Much of the driver from this came from Massimo Robberto’s 'CalWebb' concept.
Abstract. Planning for the JWST absolute flux calibration has begun. The science goal is to be able to predict the absolute flux of stars for all JWST instruments to sub-1% accuracy. A set of A0V, solar, and white dwarfs stars have been picked as a preliminary calibration sample. These stars have been observed with a combination of the Spitzer and Hubble Space Telescopes. The mapping of these stars to the sensitivities of the JWST instruments is shown along with the comparison of the Hubble and Spitzer flux calibrations. The Hubble calibration is ~1% lower than the official IRAC calibration of Reach et al. (2005) and 1.5% lower than the proposed calibration of Rieke et al. (2008).

1. Introduction

The planning for the JWST absolute flux calibration is briefly described in this paper. The science goal for JWST is to predict the calibrator star fluxes to sub-1% for all JWST instruments, in order to minimize this component of the 2% total error budget for the imaging modes of the JWST instruments. To achieve the 1% accuracy, three different types of stars and multiple stars in each type are chosen. A0V, solar, and white dwarf stars have been picked given the relative ease of modeling their spectral energy distributions (SEDs). A preliminary sample of these stars consists of 4 white dwarfs, 4 solar-type stars, and 6 A0V stars. These stars have been observed with a combination of the Spitzer and Hubble Space Telescopes. How these stars map to the sensitivities of the different JWST instruments is shown. In addition, the comparison of the Hubble and Spitzer calibrations is summarized. Most of the material presented verbally at the Workshop is published or in press. For more details on the required flux ranges for each JWST instrument and how these calibrators map to them, see the JWST report (Gordon, Bohlin et al. 2009). The Spitzer-IRAC vs. Hubble cross calibration is in Bohlin et al. (2011).

2. Preliminary JWST Calibrator Sample

The full details of the preliminary JWST calibration sample can be found in Gordon et al. (2009). The goal of the calibration sample is to provide a set of stars for each instrument capability (e.g., band/spectral order). The preliminary sample provides at least 1 star for each instrument capability but is clearly not enough stars given that the modeled IR SED for each star will have some statistical uncertainty. Nevertheless, the preliminary sample provides a starting point for the development of the JWST flux calibration.

The predicted flux distributions of the preliminary sample are shown along with the range of sensitivities for NIRCam imaging and MIRI imaging and coronagraphy are shown in Fig. 1. For the other instruments and capabilities, see Gordon et al. (2009). As can be
Figure 1: The NIRCam and MIRI sensitivity limits and the spectra of the proposed primary calibrators are shown with blue for WDs, green for A types, and red for G stars. The sensitivity in each filter is given by the vertical bar where the min, middle, and max marks give the min full frame, max full frame, and max subarray sensitivities, respectively. A Kurucz model for Vega that matches the observed flux below 1 \( \mu \text{m} \) has been included to illustrate that MIRI can observe very bright targets at the longest wavelengths.

seen, not all of the stars are observable in all the modes, but all stars can be observed in some modes. This work demonstrates that a larger sample is required to fully cover the JWST wavelength and dynamic ranges. The immediate goal is expanding the sample to allow for at least 5 stars per type (AV, solar, and white dwarf/hot stars) in each instrument filter/grating mode. The expanded sample will be reported in a future JWST report.

3. **Comparison of HST and Spitzer-IRAC Flux Calibrations**

The goal is to establish absolute flux standard with a 1% accuracy over the 0.8–30 \( \mu \text{m} \) range. Even though the goal has not yet been achieved, current agreement among the IRAC flux calibrations is better than 2.5%. HST fluxes are based on the modeled SEDs of three pure hydrogen WDs: GD153, GD71, and G191B2B (Bohlin 2007, Bohlin et al. 2001). Fluxes for sets of A and G stars are precisely measured by HST on this WD flux scale; and models are fit and used to predict the mid-IR flux distributions. In order preclude systematic errors, our strategy for calibrating JWST uses three different stellar types: WDs, A, and G stars in equal proportion. If all three types yield the same JWST instrumental sensitivities, then a precision at the level of agreement can be claimed.

An example of the quality of fit of Castelli & Kurucz (2004, hereafter CK04) models to the HST G star flux distributions is in Figure 2 per Bohlin (2010). While Bohlin & Cohen (2008) found larger residuals among A star models, Bohlin (2010) found agreement for G stars between the best fitting CK04 and MARCS models (Gustafsson et al. 2008) to 1% in broad continuum bands as shown in Figure 3. Better grids of models are needed, especially for A stars.
Figure 2: The ratio of the HST fluxes to the best fitting CK04 model SED at $R = 100$ (small circles) and in the broad bands of Table 1 (large circles). The denominators are interpolated from the CK04 model grid using the IDL routine model_int.pro. The stellar parameters for the CK04 models are in Bohlin (2010). The models are normalized to the HST fluxes over the 0.6–0.9 $\mu$m range. Before binning to $R = 100$, the models are smoothed to the resolution of the observations.
Figure 3: Ratios of the best–fit MARCS models to the best fitting CK04 models at R=100 (small open circles) for the seven HST standard stars. The large thick open circles are the ratios in the broad bands that are used to derive the best fits. The two sets of best–fit models agree to ∼1% in the continuum, which is a measure of the uncertainty in the adopted CK04 model SEDs beyond the long wavelength 2.5 µm limit of the NICMOS measurements. The broad peaks near 4.6 µm are caused by slight differences in the strength of the CO fundamental band, which is the most prominent solar feature in the mid–IR. The MARCS models do not extend longward of 20 µm; but the flatness of the ratio from 10–20 µm attests to the validity of the CK04 grid from 20–30 µm.
Figure 4: Ratios of calibration constants computed from the HST based fluxes to the IRAC values of Re05. WDs are in the top panel, solar analogs are in the center panel, and A stars are at the bottom. One sigma error bars are shown on the points that differ from unity by more than 3σ. The key to the individual symbols is at the right side of each panel. The scales are the same for all three stellar types, but there is an offset of 0.05 in the top panel.
Figure 4 shows the level of agreement between the calibration of IRAC by Reach et al. 2005, hereafter Re05) and by the sets of proposed JWST standards with HST heritage. The ratios for G and A stars average about 1% lower than Re05 and ~1.5% lower than Rieke et al. (2008). Most of the HST values agree with their average within the error bars. The worst exception is for G191B2B with its 4σ deviation, which suggest that there is some problem with the NLTE model or the IRAC photometry. To help reduce the uncertainties, a larger set of stars with HST heritage is being assembled.

References

Bohlin, R. C. 2010, AJ, 139, 1515
Candidate Calibrators for the In-Orbit Spectrophotometric Calibration of the MIRI Medium Resolution Spectrograph Onboard the James Webb Space Telescope

E. K. J. Bauwens¹, T. Nakos², L. Decin¹,³, J. A. D. L. Blommaert¹, G. Rieke⁴, C. Waelkens¹, P. Degroote¹, C. Engelbracht⁴, M. Groenewegen⁵

Instituut voor sterrenkunde, KULeuven, Celestijnenlaan 200D, B-3001 Leuven, Belgium
Department of Physics and Astronomy, University of Ghent, Krijgslaan 281 S9, WE05, B-9000 Ghent, Belgium
Sterrenkundig Instituut Anton Pannekoek, University of Amsterdam, Kruislaan 403, NL-1098 Amsterdam, The Netherlands
Steward Observatory, University of Arizona, 933 North Cherry Avenue, Tucson, AZ 85721, USA
Royal Observatory of Belgium, Ringlaan 3, B-1180, Belgium

Abstract. The James Webb Space Telescope (JWST), due to launch in 2014, will provide an unprecedented wealth of information in the near and mid-infrared wavelengths, thanks to its high-sensitivity instruments and its 6.5 m primary mirror, the largest ever launched into space. MIRI, the Mid-InfraRed Instrument, will be the only instrument (onboard JWST) with a spectrograph operating in the 5 – 28 µm wavelength range. We present a list of 32 stellar sources that have been selected as candidates for the in-orbit calibration of the Medium Resolution Spectrometer (MRS) of MIRI. We describe the selection criteria, the modelling of the stellar atmospheres, and also present some preliminary results.

1. Introduction

We present the current status of an effort to generate a list of stars for the in-orbit calibration of the Medium Resolution Spectrograph (MRS) of MIRI, the Mid-IR Instrument onboard the James Webb Space Telescope. By comparing the predicted theoretical spectra (computed using stellar atmosphere models) to the spectra obtained with MIRI observations, it will be possible to compute the Relative Spectral Response Function (RSRF), that characterises the wavelength-dependent response of the spectrometer.

To reach a calibration accuracy of a few percent, a careful selection of calibrators is essential. The final accuracy depends critically on both (a) observational issues (e.g., the non-uniformity of the data used for the selection) and (b) uncertainties in the theoretical spectra fitted to the data. In the following sections, we describe (a) the selection criteria used to select candidate calibration stars; (b) the observational data sets associated with these candidates, and (c) some results about the theoretical spectra.

2. JWST & MIRI

The James Webb Space Telescope is a project led by NASA, with major contributions from the European and Canadian Space Agency (ESA, CSA, respectively). JWST is due to
Table 1: Properties of the MIRI integral field units. The region of the sky corresponding to a spatial sample is set by the slice width in the across slice direction, and by the pixel field of view in the along slice direction.

<table>
<thead>
<tr>
<th>$\lambda$ $\mu$m</th>
<th>Slices</th>
<th>Spatial sample dimens.</th>
<th>FoV</th>
<th>R</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td></td>
<td>Across ($''$)</td>
<td>Along ($''$)</td>
<td>Across ($''$)</td>
</tr>
<tr>
<td>5.0 – 7.7</td>
<td>21</td>
<td>0.18</td>
<td>0.20</td>
<td>3.7</td>
</tr>
<tr>
<td>7.7 – 11.9</td>
<td>17</td>
<td>0.28</td>
<td>0.20</td>
<td>4.5</td>
</tr>
<tr>
<td>11.9 – 18.3</td>
<td>16</td>
<td>0.39</td>
<td>0.25</td>
<td>6.1</td>
</tr>
<tr>
<td>18.3 – 28.3</td>
<td>12</td>
<td>0.64</td>
<td>0.27</td>
<td>7.9</td>
</tr>
</tbody>
</table>

launch in 2014. The minimum life of the mission is five years, although there is a possibility of extension up to ten years. A detailed description of the space observatory and its science cornerstones can be found in Gardner et al. (2006).

MIRI (Wright et al. 2004) is a European-American collaboration and is unique, in many aspects. First of all, it will be the only instrument operating in the mid-IR wavelength range (5 – 28 $\mu$m). Second, it is also the only JWST instrument that can perform imaging, spectroscopy and coronagraphy. MIRI consists of three identical 1024 $\times$ 1024 Si:As sensor chip assemblies (SCA), of 25 $\mu$m pixel size, that will operate at 7 K. As the spacecraft will be passively cooled down to $\approx 35$ K, the 7 K temperature will be achieved by means of a mechanical cooler. The MIRI SCAs and cryo-cooler are provided by Jet Propulsion Laboratory, while MIRI’s optical system is provided by a European Consortium (EC) that consists of 28 institutes from 10 European countries.

Two of the MIRI SCAs will be used for integral field spectroscopy, via four image slicers. The two integral field units (IFUs) will provide spectral and spatial information of the targeted objects with a resolution of $R \approx 2000 – 3700$ over the $5 \mu$m < $\lambda$ < 27 $\mu$m wavelength range. The IFUs provide four simultaneous and concentric fields of view, that increase in size with wavelength. One of the two IFUs will cover the short-wavelength range (5 – 11.9 $\mu$m) while the second one will cover the long-wavelength range (11.9 – 28.3 $\mu$m).

The main properties of the MIRI medium resolution spectrometer (MRS) are presented in Table 1.

3. Selection criteria

Quantitative use of the MRS ultimately depends on observations of well-understood calibration stars. The selection of the candidates has been based on a combination of (a) requirements (imposed by the instrument specifications), (b) strict criteria that have to be respected by all stars and, (c) a “wish-list”, i.e. conditions that we would like at least some of the calibrators to fulfill. In the following sections we give a detailed description of the steps followed for the selection of the calibration candidates.

3.1. Spectral Type

To avoid systematic errors introduced by features associated to a specific spectral type of stellar sources, the calibrator list should include sources spanning a large spectral type range. According to Decin et al. (2007), a combination of A dwarfs, cool giants (G9III to K5III) and solar analogs provides an ideal data set, as the modelling of the stellar atmospheres of stars of the above spectral types is well understood. For an elaborate discussion on this topic we refer to Decin & Eriksson (2007), and we may summarise as follows:


- For the A dwarfs, the possible presence of a debris disk will result in a flux excess in the MIRI wavelength range. Spitzer MIPS 24\(\mu\)m (Rieke et al. 2004) photometric data should allow checking for IR excess and rejecting objects where one is present. Concerning the modelling of the stellar spectra, the main uncertainties for the A dwarfs arise from the hydrogen line predictions. However, as ISO spectra have demonstrated, there are no hydrogen lines present in the MIRI wavelength range.

- The late type giants have strong molecular absorption bands in the MIRI wavelength range. These features can be modelled with sufficient accuracy, provided that stellar parameters like effective temperature, gravity and chemical abundances, are accurately known. A mid-IR flux excess might also be introduced due to a circumstellar shell, which can also be identified thanks to MIPS 24\(\mu\)m photometric data.

- For the solar analogs, both atomic and molecular lines are present. However, high resolution solar spectra places constraints that can lead to accurate models of their atmospheres.

### 3.2. Other criteria

Some additional requirements for considering a stellar source as a good spectrophotometric calibrator are:

- the star is photometrically stable in time
- its 10\(\mu\)m flux density is in the range 4 – 500 mJy (this is a MIRI MRS requirement).

Additional factors that were taken under consideration for the selection, are

- the Right Ascension and Declination of the sources, as it is essential to be able to observe at least a few calibrators at any time, during the mission,
- the flux density of some sources to be close to the faint limit of the requirement range (\(\approx 4\) mJy), so that a cross-check calibration with the MIRI imager can be performed for some of the calibrators in our list\(^1\), and
- MIPS 24\(\mu\)m data should either exist at the Spitzer archive, or should have been possible to obtain, to check for a possible IR excess.

### 4. The list of candidate calibrators

A preliminary list of 79 MIRI-MRS candidate calibrators was presented in Decin et al. (2007), containing

- calibration sources from Spitzer’s IRAC (Reach et al. 2005), MIPS 24\(\mu\)m (Engelbracht et al. 2007) and IRS (Houck et al. 2004) instruments
- ISO’s ISOCAM (Blommaert 1998)
- an object list presented in Cohen et al. (2003), originating from the Landolt standard stars (Landolt 1992) and the Carter-Meadows catalog of faint IR stars (Carter & Meadows 1995), and
- solar analogs, suggested by G. Rieke (priv. comm.).

\(^1\)According to the latest specifications, it will be possible to observe with the MIRI imager even the brightest of the sources in the list, in sub-array mode.
The above list was finally reduced to 32 sources with MIPS 24 \( \mu \text{m} \) photometry, which is essential to reject sources with a circumstellar dust shell or a disk. The current list of candidate calibrators, containing seven solar analogs, seven late-type giants and 18 A dwarfs, is presented in Table 2. The term “candidate” has the meaning that the present list can still be reduced. In the subsequent sections we shall explain how the refinement of the list will be achieved.

Table 2: Candidate calibrators for the MIRI MRS, with spectral type, position and flux at 10 \( \mu \text{m} \), in mJy. These fluxes are estimates made by extrapolating observations at different wavelengths, using an appropriate Kurucz model (Vega, Arcturus and the Sun for the A, K and G-type stars, respectively).

<table>
<thead>
<tr>
<th>nr</th>
<th>Identifier</th>
<th>Sp. Type</th>
<th>RA (J2000)</th>
<th>Dec (J2000)</th>
<th>( F_{10} ) (mJy)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>HD 001644</td>
<td>K0</td>
<td>00 20 43.6</td>
<td>+03 26 58.0</td>
<td>376.0</td>
</tr>
<tr>
<td>2</td>
<td>HD 002811</td>
<td>A3V</td>
<td>00 31 18.5</td>
<td>−43 36 23.0</td>
<td>57.0</td>
</tr>
<tr>
<td>3</td>
<td>HD 014943</td>
<td>A5V</td>
<td>02 22 54.7</td>
<td>−51 05 31.7</td>
<td>175.9</td>
</tr>
<tr>
<td>4</td>
<td>HD 015646</td>
<td>A0V</td>
<td>02 28 04.4</td>
<td>−64 17 59.0</td>
<td>113.8</td>
</tr>
<tr>
<td>5</td>
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<td>A2V</td>
<td>02 44 10.6</td>
<td>−52 34 14.0</td>
<td>172.4</td>
</tr>
<tr>
<td>6</td>
<td>HD 021981</td>
<td>A1V</td>
<td>03 30 37.0</td>
<td>−47 22 30.5</td>
<td>204.0</td>
</tr>
<tr>
<td>7</td>
<td>HD 037962</td>
<td>G4V</td>
<td>05 40 52.0</td>
<td>−31 21 40.0</td>
<td>120.0</td>
</tr>
<tr>
<td>8</td>
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<td>G1V</td>
<td>05 48 20.1</td>
<td>−24 27 50.0</td>
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</tr>
<tr>
<td>9</td>
<td>HD 040335</td>
<td>A0</td>
<td>05 58 13.5</td>
<td>+01 51 23.0</td>
<td>57.0</td>
</tr>
<tr>
<td>10</td>
<td>HD 042525</td>
<td>A0V</td>
<td>06 06 09.4</td>
<td>−66 02 23.0</td>
<td>169.4</td>
</tr>
<tr>
<td>11</td>
<td>HD 046819</td>
<td>K0III</td>
<td>06 30 22.7</td>
<td>−66 01 23.0</td>
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</tr>
<tr>
<td>12</td>
<td>HD 057336</td>
<td>A0IV</td>
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<td>−79 25 55.0</td>
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</tr>
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</table>

5. Available spectroscopic and photometric data

For the candidate MRS calibrators, ground-based observations are essential to (a) test the reliability of the candidates and (b) constrain the stellar parameters, needed for comput-
The in-orbit calibration of the MIRI MRS, onboard JWST

ing the theoretical spectra of the calibrators. More specifically, the optical and infrared observations can be used in the following way:

- **UBVRI** photometric data, to constrain the effective temperature, $T_{\text{eff}}$.
- NIR photometric data, in particular $K$-band, to constrain the effective temperature in the case of late type giants.
- Optical and/or NIR spectroscopy, to set constraints on the effective temperature, $T_{\text{eff}}$, the surface gravity, $\log g$, and the abundances.
- Parallaxes (if possible), to derive the distance and hence to estimate the dust reddening $E(B - V)$ and the surface gravity.

To minimize the impact of the observational uncertainties to the total calibration error budget, the data should be obtained (and reduced) in a consistent and homogeneous way. As a consequence, the number of instruments involved in the data acquisition should be reduced as much as possible. For the purposes of the MIRI MRS calibration, data have either been collected thanks to telescope time allocated to our project, or have been retrieved from astronomical databases:

- The $U, B$ and $V$ band photometry where obtained using the Swiss Euler telescope, at La Silla, Chile and the Flemish Mercator telescope, at La Palma, Spain (which is identical to the Euler telescope). Three instruments were involved in the data acquisition. The first one is the P7 photometer. As it was installed in the past on both telescopes, it has been used to observe targets covering a wide declination range. Unfortunately P7 was decommissioned before all stars in our list were observed. Data for the remaining stars were obtained either with the MEROPE CCD camera on the Mercator telescope, or the C2 CCD camera on the Euler telescope.

- $R$ and $I$ band observations were obtained using the MEROPE and C2 cameras. However, due to technical issues, very limited information can be extracted from the data. For most objects these data will therefore not be used in the modelling.

- The 2MASS catalogue provides $J$, $H$ and $K$-band photometry for all but one star (SA 112-595). The quality flags for the results were checked to make sure only high quality data was used. This resulted in excluding the $J$, $H$ and $K$-band photometry for HD 159048, the $J$ and $H$-band photometry for HD 113314 and HD 175510, and the $H$-band photometry for HD 196724.

- Recently, Ishihara et al. (2010) published the AKARI results for the IRC point-source catalogue, which includes photometry at 9 and 18 $\mu$m. These data are particularly useful to verify if stars in our list show any flux excess, in the MIRI wavelength range, with respect to the models. For 27 of the candidate calibrators there is 9 $\mu$m photometry available (there is no information for HD 040335, SA 103-526, SA 112-595, SA 112-275 and [RMC2005]KF09T1). 18 $\mu$m photometry is available for just four sources, namely HD 046819, HD 113314, HD 159048 and HD 196724.

- In addition to the AKARI data, IRAC photometry (covering the wavelength range 3.6 – 8.0 $\mu$m) is also available for 17 of our targets, but the data still have to be reduced. Finally, MIPS data at 24 $\mu$m are being used for all stars to check whether the calibrator candidates show an IR excess. These data were either retrieved from the Spitzer archive, or were obtained based on observing proposals submitted by us.
6. Models, SED fitting and preliminary results

The data mentioned above were used as input for a $\chi^2$ minimisation procedure, that selects the best-fit Spectral Energy Distributions (SED) among 1615 models, to fit the observed data points. These SEDs, taken from the literature, originate from eight different codes used to model stellar atmospheres. The 1615-SED grid spans an effective temperature range from $2400 - 55000$ K and a surface gravity from $-1$ up to $9$ (in log $g$ units).

As the MRS candidate calibrators are A dwarfs, solar analogs and late type giants, the best-fit model is chosen from a subset of 556 SEDs, which originate from three different stellar atmosphere codes. More specifically,

- 409 SEDs come from the Kurucz (1993) models, that span the effective temperature range $3500 - 50000$ and $0 \leq \log g \leq 5.0$. The $T_{\text{eff}}$ step size is $250$ K between $3000$ K–$10000$ K and $500$ K between $10000$ K and $13000$ K (we do not expect higher temperatures for the candidate calibrators, given their spectral type).
Figure 3: Grid of SEDs used to compute the stellar properties of the calibrators (see Section 6, for more details). The Kurucz SEDs are from the Kurucz (1993) library, the MARCS models from Gustafsson et al. (2008) and the MARCS2 from Decin (priv. com.). The evolutionary tracks are from Schaller et al. (1992).

- 28 SEDs come from the Gustafsson (2008) models, with 4000 K ≤ T_{\text{eff}} ≤ 6000 K and 0 ≤ log g ≤ 5.0. The step size is 100 K between 2500 – 4000 K and 250 K for temperatures higher than 4000 K.

- 119 SEDs are an improved version of the Gustafsson (2008) models, developed by members from our team (namely L. Decin, which we shall call MARCS2), with 3000 K ≤ T_{\text{eff}} ≤ 5500 K and 0 ≤ log g ≤ 5.0. The log g step-size is 0.5 for all models, and we assume solar abundances for all cases. Although the available models have been computed at a resolution of ΔT_{\text{eff}} = 250 K and Δ log g = 0.5, it is possible to interpolate between the models, to reach a more precise effective temperature and surface gravity estimation. Our goal is to derive the best fitting values for T_{\text{eff}} and log g at a resolution of 100 K and 0.25 respectively. The grid of models used for the SED fitting is presented in Fig. 3.

Additional information derived from the χ² minimization is the E(B – V) reddening factor and the calibrator’s angular diameter. It should also be noted that in order to check whether the star shows a possible flux excess in the MIRI wavelength range, the photometric points above 5 µm are not included in the fitting procedure. An example of the resulting best fitting model for HD 037962 is shown in Fig. 4, where the photometric points longer than 5 µm are simply over-plotted.

For a subset of 17 MRS candidate calibrators, for which we required the objects to have available P7 photometry, we derived the stellar parameters T_{\text{eff}}, log g, reddening, E(B – V) and angular diameter, θ, using the SED fitting procedure described above. The results are presented in Table 3.
Figure 4: Example of SED fitting results for HD 037962, based on six photometric points (three coming from the Geneva catalog \((UBV)\) and three from the 2MASS archive \((JHK)\)). Longer wavelengths do not participate in the fit, to check whether the calibrator experiences an infrared excess. The photometric errors are also plotted but they are smaller than the symbols, except for the 24 \(\mu\)m measurements.

7. Future steps

Future steps for the further characterisation of the calibrators candidates include (a) the reduction of the IRAC photometric data, (b) the SED fitting for the remainder of the sources, and (c) the analysis of the spectroscopic data. From the obtained high resolution optical spectra, the \(T_{\text{eff}}\), \(\log g\) and microturbulent velocity \(\xi_t\) will be derived, by measuring the equivalent width of the Fe lines. Through detailed modelling of the spectral lines of other elements it will be possible to refine the chemical abundances of the calibrators. From the above process it will be possible to generate spectra that better represent their stellar parameters.

We currently posses near-IR spectra for seven targets, obtained with the SOFI spectrograph on the ESO NTT telescope at La Silla observatory, Chile. These spectra cover the 0.9 – 2.5 \(\mu\)m wavelength range, at a resolution between 600 and 2200. There are also Spitzer IRS spectra available for 17 of the objects in our list. These infrared spectra will be used to compare to the theoretical spectra we will derive for the candidate calibrators, to check whether there are any discrepancies. After rejecting cases with disagreements between models and observations, a final set of 15 to 20 calibrators will be retained from the list of candidate calibrators.

8. Summary

The list of 32 candidate calibration stars presented in this work, contains well known sources of three spectral classes (A dwarfs, solar analogs and late type giants). An extended set of data, photometric as well as spectroscopic and both in the optical and IR wavelength ranges has been obtained for these sources.

For a subset of the sources the first estimates for the stellar parameters based on the photometric data are presented here. In the coming months this analysis will be done for the remaining sources, as well as the analysis of available spectroscopic data. Based on
Table 3: List of the sources for which preliminary stellar parameters have been determined. This subset originates from Table 2 and includes sources with P7 photometry. The accuracy for the \( T_{\text{eff}} \) and the \( \log g \) is 100 K and 0.25 respectively, unless indicated otherwise.

<table>
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<tr>
<th>nr</th>
<th>Identifier</th>
<th>Sp. Type</th>
<th>( T_{\text{eff}} ) [K]</th>
<th>( \log g )</th>
<th>E(B-V)</th>
<th>( \theta ) [mas]</th>
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<td>0.168 ± 0.009</td>
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<td>A2V</td>
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<td>4.00</td>
<td>0.000</td>
<td>0.222 ± 0.012</td>
</tr>
<tr>
<td>5</td>
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<td>0.279 ± 0.014</td>
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<tr>
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<td>2.00</td>
<td>0.000</td>
<td>0.098 ± 0.005</td>
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</table>

\textsuperscript{a}Not in agreement with the \( T_{\text{eff}} = 9230 \) K found by Wright et al. (2003).

\textsuperscript{b}Not in agreement with the \( T_{\text{eff}} = 9520 \) K found by Wright et al. (2003).

\textsuperscript{c}The accuracy is 250 K.

\textsuperscript{d}Not in agreement with the \( T_{\text{eff}} = 11 000 \) K found by Wright et al. (2003).

\textsuperscript{e}Not in agreement with the \( T_{\text{eff}} = 9520 \) K found by Wright et al. (2003).

the agreement or discrepancies between the models and the observations for the candidate calibrators, 15 to 20 calibrators will be selected from this list to be used as the fiducial calibrators for the MIRI MRS. These stars will be used to calibrate the MRS based on both continuum and spectral line measurements, at a level of \( \approx 2\% \).

References

JWST Operations and the Phase I and II Process

Tracy L. Beck

*Space Telescope Science Institute, Baltimore, MD 21218*

**Abstract.** The JWST operations and Phase I and Phase II process will build upon our knowledge on the current system in use for HST. The primary observing overheads associated with JWST observations, both direct and indirect, are summarized. While some key operations constraints for JWST may cause deviations from the HST model for proposal planning, the overall interface to JWST planning will use the APT and will appear similar to the HST interface. The requirement is to have a proposal planning model similar to HST, where proposals submitted to the TAC must have at least the minimum amount of information necessary for assessment of the strength of the science. However, a goal of the JWST planning process is to have the submitted Phase I proposal in executable form, and as complete as possible for many programs. JWST will have significant constraints on the spacecraft pointing and orient, so it is beneficial for the planning process to have these scheduling constraints on programs defined as early as possible. The guide field of JWST is also much smaller than the HST guide field, so searches for available guide stars for JWST science programs must be done at the Phase I deadline. The long range observing plan for each JWST cycle will be generated initially from the TAC accepted programs at the Phase I deadline, and the LRP will be refined after the Phase II deadline when all scheduling constraints are defined.

1. JWST Operations and Observing Overheads

The James Webb Space Telescope (JWST) is a 6.5 meter segmented mirror space telescope designed to investigate astronomical targets at infrared wavelengths (∼1.0-28.0 μm) with unprecedented sensitivity. JWST will operate at the Sun-Earth Lagrangian 2 (L2) point where it can take advantage of very dark infrared viewing conditions.

The operation of JWST will require several activities that will subtract from available community observing time on the telescope. JWST has a segmented primary mirror, so regular wavefront sensing observations will be needed every ∼2 days to monitor the telescope point spread function (PSF) shape. Wavefront control maneuvers will also be required to correct the PSF shape with a frequency of approximately ∼2 weeks. The solar wind will apply a torque to JWST, and momentum management tasks will be executed every ∼2 days to correct for this. To maintain the proper orbit at the L2 point, JWST will have orbital “station keeping” maneuvers executed with an approximately 3 week frequency. These and similar operations activities designed to maintain the observatory performance and location are collectively referred to as indirect observing overheads.

JWST slews 90 degrees in 1 hour, hence the slew time from target to target could be significant. The observing plan will be generated to minimize slew times between astronomical targets to maximize the observing efficiency. The slew tax assigned to science observations will likely be based on statistical models for the average slew distances (slew overhead is TBD). The overheads associated with a given science program include: 1) the (average) time needed to slew to a target, 2) guide star acquisition, 3) target acquisition (when necessary), 4) the total elapsed exposure time, 5) the time for instrument setup and
mechanism movements, 6) overheads for dithers, 7) other observing taxes that might be levied (for example, for timing or scheduling constraints). These overheads and observing taxes are associated with the observing program and not the observatory. Hence, these are direct overheads that must be taken into account within the observing time requested in the science proposal.

2. The JWST Proposal Planning Process

As for HST, the planning of science programs for JWST will be done using the Astronomer’s Proposal Tool (APT) software for both the Phase I (project assessment) and Phase II (project execution) deadlines. The interface for planning JWST programs will likely look familiar to experienced HST users. For the initial proposal process, the high level JWST mission requirement is for the Phase I submission to capture all program information that is necessary for the time allocation committee (TAC) to assess the scientific strength of a project. However, a goal of the JWST proposal process is to have the program definition as complete as possible at the Phase I deadline (which may not be feasible for large projects, TBD). The JWST guiding field is much smaller than for JWST, so guide star availability must be verified in the Phase I feasibility check. For easier interaction with programs, the JWST Phase II process will use the same file syntax and the same upload as the Phase I submission. JWST will have limited space and volume constraints on the data recorders, so there are presently no plans to include science observations with a different instrument in parallel with the prime science (limited parallel calibrations only are planned).

In the APT, the available JWST instrument modes map to specific observing templates that will be used to define an observing program (Fullerton et al. 2007). Table 1 presents the available instrument capabilities and observing modes, and the corresponding APT templates. Prior to launch of JWST, future releases of the HST APT will have model example JWST templates included for testing the interface. At the present time, the plan is that unsupported or uncalibrated instrument modes, beyond those listed in Table 1, will not be available for program definition within APT. General Observers (GOs) will fill out the instrument templates in Phase I to describe their proposed observations. The goal is to have a flexible planning tool that streamlines the Phase I process for most programs, with only small changes needed at Phase II, if required. Of course, there are some instrument modes that will not fit into this model. For example, the NIRSpec multi-object spectroscopy (MOS) mode will require a “pre-image” of the science field for accurate target coordinates (for early cycles, HST WFC3 or ACS images may be used). So, NIRSpec MOS observations with no prior imaging will likely require a multi-step Phase II process and sparser program definition at the Phase I deadline.

JWST will have strict pointing limitations, and as a result the spacecraft orient is going to be constrained for all targets. When defining a science program, GOs will be allowed to chose an available instrument aperture position angle (PA), but not the actual spacecraft orient (Sparks et al. 2009). Moreover, the selection of a specific instrument aperture PAs will be treated as a scheduling constraint. Observing a target at a specific PA will translate to a very small observing window, making the visit scheduling more difficult. Observations of point sources or symmetrical targets with imaging, IFU or long-slit spectroscopy may be executed at any PA for maximum scheduling flexibility, while observations with the NIRSpec MOS mode will need a fixed position angle in order to specify the visit. There will be three opportunities for GOs so specify an aperture PA: 1) in the Phase I (as a scheduling constraint on the observation), 2) at the Phase II deadline, a GO can specify the exact PA for the observation after getting an allowed PA range from the initial long range plan (LRP) and 3), observing PAs are assigned by the telescope schedulers for flexible observations that do not have any constraints on the observing angle.
For streamlining the JWST program definition process, proposers will be allowed and encouraged to submit fully executable proposals at the Phase I deadline. In order to best define a program at Phase I, several scheduling constraints might be required. Scheduling constraints permitted at the Phase I deadline include:

1) Target coordinates - will be needed to constrain the scheduling windows and estimate allowed aperture PAs.

2) Aperture PA constraint - at the Phase I, GOs can select a given aperture PA from a list of available positions based on the target coordinates.

3) Guide star availability - checks for available guide stars for a given target position are necessary at the Phase I deadline.

4) Timing constraints - Any limitations on program timing need to be defined for planning purposes.

5) Links between visits - visits that must consecutively may constrain the planning process.

6) Visit duration - estimates of the visit duration may serve as a constraint (e.g., for long visits and lengthy target observations).

Once the TAC approves the science programs for a JWST observing cycle, these constraints on programs defined at the Phase I deadline will help the STScI schedulers construct the intial LRP. After the Phase II program definition deadline, scheduling constraints for aperture position angle, links between visits, and visit duration will be refined for many programs, allowing for a more robust LRP to be generated. As is presently done for HST, JWST science programs will execute based on a detailed short term plan that is constructed using the initial LRP as a guideline.

References

Fullerton et al. 2007, JWST Technical Report, JWST-STScI-01257
Sparks et al. 2009, JWST Technical Report, JWST-STScI-001667

Table 1: JWST Instrument Modes and Observing Templates

<table>
<thead>
<tr>
<th>Instrument Name</th>
<th>Instrument Modes</th>
<th>Wavelength Range</th>
<th>Observing Template</th>
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</thead>
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<td>NIRCam</td>
<td>Coronography</td>
<td>~1.0 to 5.0μm modules for Imaging</td>
<td>Imaging Template</td>
</tr>
<tr>
<td></td>
<td>Grism Spectroscopy</td>
<td>simultaneous long and short wavelength images</td>
<td>Coronography Template</td>
</tr>
<tr>
<td></td>
<td>Tunable Filter</td>
<td>Imaging Template</td>
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<td>FGS TFI</td>
<td>Coronography</td>
<td><del>0.7 to 5.0μm Spectral Imaging at R</del>100</td>
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<td>Non-Redundant Mask</td>
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<tr>
<td></td>
<td>IFU Spectroscopy</td>
<td>Spectroscopy at R~100, 1000 or 3000</td>
<td>IFU Template</td>
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BetaDrizzle: A Redesign of the MultiDrizzle Package

A. S. Fruchter, W. Hack, N. Dencheva, M. Droettboom, P. Greenfield

Space Telescope Science Institute, Baltimore, MD 21218

Abstract. We describe substantial changes to MultiDrizzle intended to allow users to easily and accurately align and combine HST images taken at multiple epochs, and even with different instruments. In the first part of this program, the correction of the time-dependent distortion of ACS was introduced into MultiDrizzle to provide distortion corrections good to a few hundredths of a pixel level. Now, we are undertaking a more profound change in the underlying method that MultiDrizzle (and HST) use to represent astrometry and geometric distortion. As part of this change we have temporarily renamed MultiDrizzle as BetaDrizzle to avoid confusion between the programs. BetaDrizzle uses new simple extensions of the fits format that allow us to fully represent the ACS and WFC3 distortions in the header of the image, meaning that a calibrated image needs no other files to describe its astrometry. Users will also be able to easily pass astrometric header information to each other or back to the archive for use by others. The presence of the full astrometric solution in the header means that users can fit the astrometry of one image to another image, or to an external catalog, without drizzling the image. Precise coordinate information can be extracted directly from the .flt image itself.

1. Introduction

When we started this program we had one guiding principle: Astrometry should be like any other calibration; once an image is calibrated you should not need to carry around separate calibration files. All of the necessary information should be in the image and its headers. This lead us to several more specific goals:

- The best available geometric distortion information should be incorporated into HST image headers.
- A user should be able to easily align an image and update its WCS to match another image or catalog
- Users should be able to easily exchange astrometric solutions with each other and/or the archive

To accomplish these goals we have extended the FITS standard and have created several new pieces of software to extend the capabilities and accuracy of MultiDrizzle (Koekemoer et al. 2002). The FITS extensions allow users to incorporate (compressed) versions of the astrometric information in their headers and to keep several different WCS solutions in an image at the same time. The upgraded MultiDrizzle software set, which we call BetaDrizzle for now to avoid user confusion, provides a means for users to control this WCS information, to update it as needed, and to transfer astrometric solutions to each other or the archive.
Figure 1: A comparison of distortion error with and without time-dependent distortion (TDD) correction. The stellar positions from two ACS/WFC images of 47 Tuc are compared. Both images were taken in 2005, but at a relative rotation of about 90 degrees. The blue points represent offsets between the same stars in the two different images after stellar positions have been fit for a possible rotation, shift and scale change, but without the application of the time-dependent distortion correction. The red points show the result of the same fit after time-dependent distortion correction is applied. While small systematic errors remain (up to perhaps 0.05 pixels), the peak-to-peak distortion error has been reduced nearly an order of magnitude. The TDD correction is added to the image by modifying the World Coordinate System matrix to reflect the TDD.
2. A (Relatively) Simple Image Convention

One of the greatest difficulties faced by users of Drizzle (Fruchter and Hook 1999) and Multidrizzle is aligning large sets of images either of the same field taken at different epochs and thus different orientations and guide stars, or of adjacent fields to create a mosaic. The standard way of handling this has been through pixel shifts. However, for larger fields – particularly mosaics – it is preferable to think in sky coordinates. Therefore the first part of our program has been to redesign the HST image headers so that all distortion information is included in the headers. To do this we are using an extended version of the Simple Image Polynomial convention (Shupe et al. 2005) which has already been used by Spitzer:

\[
\begin{pmatrix}
  x \\
  y
\end{pmatrix} = \begin{pmatrix}
  CD_{1,1} & CD_{1,2} \\
  CD_{2,1} & CD_{2,2}
\end{pmatrix} \begin{pmatrix}
  u + f(u,v) + LT_x(u,v) \\
  v + g(u,v) + LT_y(u,v)
\end{pmatrix}
\]

Here \((u,v)\) represent the input pixel positions, and \((x,y)\) are output sky coordinates. \(f\) and \(g\) are polynomials with no term smaller than second order – thus the linear terms of the distortion are entirely contained in the CD matrix. \(LT_x\) and \(LT_y\) are look-up tables which follow the convention of Paper IV proposed as a FITS standard (Calabretta et al. 2004). The distortion of ACS, for instance, can be described to a few tenths of a pixel with only a quadratic polynomial. However, to get distortion errors below 0.1 pixel with ACS one needs two sets of corrections beyond a low order polynomial. One needs higher order corrections, which can be added by a lookup table (Anderson 2002), and one needs to adjust the skew of the CD matrix with time (Anderson 2007). The distortion correction look up tables have been in use with the ACS for some time – but as full size images rather than an interpolated table. Here we return them to the much smaller look-up tables, which enables us to put the information in the header. Note that in the SIP convention all of the linear terms of distortion are included in the World Coordinate System (WCS) matrix. Our software now also removes the time dependent distortion of the ACS by modifying the skew of the WCS when the SIP coefficients are created for the first time by the STScI pipeline (see Figure 1).

In addition to these primary corrections, very small corrections to \((u,v)\) will be allowed prior to the SIP implementation above. For instance, both the ACS and WFPC2 have very small periodic changes in pixel column width (\(\ll 0.1\) pixel). Detector distortions such as these are best thought of as separate from and subject to the optical distortions, which are represented fully in the extended SIP format.

For the ACS this means the user will be trading more than 128 MB of external reference files (the D GEO – or differential geometry – files) for about an extra 50kB of header data. WFPC3 at present needs no calibration beyond the SIP polynomial; however, we expect that higher order corrections will be added in the future. In the case of WFPC2 data, which is now statically archived, the user will still need to retrieve the D GEO file, which contains a correction for a small lithographic errors in the detector. BetaDrizzle will incorporate this information into the header of the WFPC2 image the first time it is run. Figure 2 shows what a present image header looks like when it comes out of the archive.

3. Off With Their Headerlets

Incorporating the full astrometric solution in the header allows a new approach to aligning images. One does not need to Drizzle images before aligning them. One can find the pixel location of sources in the flt images and then accurately convert these pixel locations to sky coordinates. The image sky coordinates can then be compared either to a catalog of coordinates derived from an astrometric survey (such as 2MASS) or with a catalog from another image. Because we have already removed any distortion in the image, we now only need to solve for a shift, rotation and possible scale change to bring the images into
Figure 2: A listing of an ACS image showing the new extended FITS format we describe in this paper. Extensions 0 through 6 should already be familiar to ACS users. Extension 7 is the file for the detector distortions, and extensions 8 through 11 are the replacement for the present DGeo file. Extension 12 is a binary file which keeps track of the astrometric solutions and their distortion files. This binary table allows us to handle multiple astrometric solutions (and their potentially different correction tables) in a single header. The image shown here has only one astrometric solution – the original solution provided by the archive.

alignment. The solution derived can then be directly applied to the image header. Removing the requirement of drizzling the image not only speeds up the process of aligning images it can improve alignment accuracy, as drizzling a single image can introduce small astrometric errors when the original images are undersampled. A task presently called "TweakReg" has been written which can in a single call find the sources in an HST image, match these to sources in a catalog (obtained either from another image or from the literature) and update the image header of the image so that the sky coordinates of the objects in the image match those in the catalog.

Our approach also enables another important leap. Once one has obtained an astrometric solution for an image, the header now contains all the information that is needed to describe that solution. If one were to package the astrometric information into a small FITS file, one could imagine passing that solution directly to another user. Our software will allow just this. We call the small astrometric FITS files headerlets. Users will be able to exchange these headerlets among themselves or back and forth with the archive. Thus, for instance, if a Multi-Cycle Treasury program derives good astrometric solutions for all of its images, it will be able supply headerlets for all of those solutions to the archive and other users will be able to download them.

There is, however, no reason why an image should be limited to only one astrometric solution in its header – a single image could be fit to multiple (and slightly discrepant) astrometric catalogs. We will allow multiple astrometric solutions in the header and we will provide a simple means for the user to choose which of these solutions to make the default. If all of these solutions use the same SIP coefficients (as will generally be the case), then the presently available DS9 software can already correctly switch between the multiple solutions.
Figure 3: One of the advantages of the new Teal GUI. In this case the user has decided not to do separate Drizzles of the images. In the first image she has turned off the separate drizzles (and the parameters associated with that choice are grayed out). In the second image she has turned off the median step. While this sort of parameter handling is now common in, for instance, the Mac OS X operating system, it is something IRAF still cannot do.

4. A New Engine and Chassis

BetaDrizzle has been built upon a philosophy of modularity. BetaDrizzle has two primary Python scripts that call modular components. The modular components are themselves written in Python or C, if that language would be more efficient (Drizzle itself has been converted from Fortran to C). The first primary script handles the coordinate transformations using the full distortion model, while the second is responsible for resampling (drizzling) the image. The new modular BetaDrizzle will enable the user to more easily customize the program to the specific needs of their data, while making the standard use of the software even simpler.

BetaDrizzle also employs a new graphical user interface, called Teal, to handle parameter input. Teal behaves in many ways like the "epar" mechanism of IRAF, but has new features that will allow for a better and simpler user experience. For instance, as shown in Figure 3, Teal can grey out parameters depending upon choices made by the user. Teal also makes it easy for the user, or the Institute, to create "configuration" files – that is sets of parameters – that can be named and saved, and written to handle specific instruments and situations. This will enable the Institute to provide better tailored parameter files and for the user to adapt those files for his or her use.

5. Conclusions

In this short paper we have provided a basic introduction to program to substantially advance the ability of users to combine HST images. The header philosophy and programs we have developed will allow users to determine offsets between images directly from their flt files without the need for drizzling, and to update the image headers to reflect the new astrometry. The full astrometric solution, including all distortion corrections, will be in
the image header, and users will be able to pass astrometric solutions back and forth to each other or the archive through the use of headerlets. The BetaDrizzle program itself is modular and will use easily modifiable configuration files making it far simpler for users to adapt it to their particular needs. And BetaDrizzle uses a new GUI, Teal, that is a substantial advance beyond the present IRAF epar interface. We believe that all of these changes should make accurately combining HST images far simpler and less error prone.

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Part 3. Contributed Posters
Cosmic Origins Spectrograph Commissioning: Overview of COS Servicing Mission Observatory Verification

Charles (Tony) Keyes

Space Telescope Science Institute, Baltimore, MD 21218

Abstract.

The Cosmic Origins Spectrograph (COS) was installed on the Hubble Space Telescope (HST) during Servicing Mission 4 (SM4) in May 2009. COS is designed and optimized to obtain spectra of faint objects at moderate spectral resolution ($R \approx 16,000$) with two channels: FUV, covering wavelengths from 1150 to 1450 Å and NUV, covering 1700 – 3200 Å. Two low resolution gratings ($R \approx 1500$) cover the 1900 – 2050 Å (FUV) and 1650 – 3200 Å (NUV) wavelength regions. An imaging capability is also available on the NUV channel. As part of the HST Servicing Mission Observatory Verification (SMOV) program, an extensive period of checkout, fine-tuning, and preliminary characterization began after the installation of COS. The COS SMOV program was a cooperative effort between the Space Telescope Science Institute and the Instrument Definition Team based at the University of Colorado. Nearly 2800 COS exposures in 34 separate observing programs were obtained during the course of SMOV. Early activities included an initial instrument functional checkout, high-voltage turn-on and initial characterization of the detectors, NUV and FUV channel focus and alignment, wavelength calibration, and target acquisition verification and assessment. Once this initial period was completed, science-related calibrations and verifications were performed in order to prepare the instrument for normal operations. These activities included flux calibration, detector flat field characterization, spectroscopic performance verification, high S/N operation, and thermal and structural stability measurements. We present the design, execution and results of the SMOV program, including the interrelationships and dependencies between the various tasks, and how the pre-launch plan was adjusted in real-time due to changing conditions.
On-Orbit Performance of the COS Detectors

David Sahnow
Department of Physics & Astronomy, The Johns Hopkins University, Baltimore, MD 21218

Jason McPhate
Space Sciences Laboratory, University of California, Berkeley, CA 94720

Thomas B. Ake, Derck Massa
Space Telescope Science Institute, Baltimore, MD 21218

Steven V. Penton, Eric Burgh
Center for Astrophysics and Space Astronomy, University of Colorado, Boulder, 80309

Wei Zheng
Department of Physics & Astronomy, The Johns Hopkins University, Baltimore, MD 21218

Abstract. The Cosmic Origins Spectrograph includes two microchannel plate detectors. The FUV channel uses a two-segment cross delay line (XDL) device, while the the NUV channel uses a MAMA, similar to those used by other HST instruments. Since SM4, both detectors have been characterized in detail, using data collected as part of specific calibration programs and in the course of normal science observations. We will describe the on-orbit performance of both detectors since the installation of COS, and describe ongoing detector characterization work.

1. Overview

The Cosmic Origins Spectrograph uses a separate detector for each of its two channels. The performance of both detectors has been monitored extensively since the installation of COS during SM4 in May 2009.

For the FUV channel, an open-faced microchannel plate (MCP) detector with a cross delay line anode (XDL) is used. The layout is similar to the detectors used for the Far Ultraviolet Spectroscopic Explorer (FUSE), containing two independently-controlled 85 × 10 mm segments placed end-to-end with a small gap between them. The MCPs are curved to match the focal surface of the spectrograph. A CsI photocathode on the front of the MCPs provides sensitivity in the FUV from below 900 Å to ∼1750 Å.

Photons reaching either MCP generate a cloud of ∼10^7 electrons which is collected by the anode. The position of the charge cloud on each segment is digitized to 14 bits in the spectral direction (X) and 10 bits in the cross-dispersion direction (Y), giving a digitized region of 16384 × 1024 pixels, each 6 × 24 µm; in addition, a 5 bit pulse height (total charge collected by the anode) is calculated. The active area of the MCP covers approximately 14,000 × 400 pixels, and a typical science spectrum covers the full active area in X, and < 50 pixels in Y. Details of the FUV detector system are discussed by McPhate et al. (2000) and...
On-Orbit Performance of the COS Detectors

Figure 1: FUV Segment A count rate as a function of time during an orbit that skims the SAA (top), and during one that is farther from the SAA (bottom).

Vallerga et al. (2002); some of the early on-orbit performance characteristics are described by McPhate et al. (2010).

For the NUV channel, a Multi-Anode Microchannel Plate Array (MAMA), originally built as a backup for STIS, is used. The MAMA is a windowed detector, with 1024 × 1024 pixels, each 25 μm square. It has sensitivity from ∼1600 Å to > 3000 Å. NUV spectra are imaged on to the detector in three stripes, each covering the full pixel range in X, and <25 pixels in Y.

Most COS data are saved in TIME-TAG mode, where the time of arrival of each photon event is recorded with a resolution of 32 msec, along with the position (and in the case of FUV, the pulse height). For count rates greater than 20,000 to 30,000 counts per second, data is typically saved in ACCUM mode, which discards the timing information (and pulse height for the FUV), saving only a two-dimensional image of the counts collected during the exposure. For the FUV channel, this image includes only a subset of the active area, while for the NUV it includes the full detector.

Special detector calibration observations have been made since installation of COS, and in the case of dark rate monitoring, continue weekly. Other regular monitoring of the detectors is made using data from normal science observations.

2. Dark Rates

FUV detector dark rates measured on the ground were very low, with rates of approximately 0.4 counts cm$^{-2}$ s$^{-1}$. Typical dark rates on-orbit away from the South Atlantic Anomaly (SAA) during the first year on orbit for both segments are ∼1.8×10$^{-6}$ counts pixel$^{-1}$ s$^{-1}$ (∼1.25 counts cm$^{-2}$ s$^{-1}$), equivalent to ∼2.6×10$^{-4}$ counts s$^{-1}$ per X resel in a spectrum with the default extraction height. These rates have remained stable with time since SM4.
When *HST* passes though the SAA, observations stop and the detector high voltage is lowered in order to prevent damage to the detector. Elevated dark rates of up to 30 times higher than normal (Figure 1) have been observed during exposures made when skimming the edge of the SAA. In order to minimize the observing time with higher background, the SAA model was shifted 6 degrees to the west in May 2010.

The spatial distribution of background counts on Segment A is quite uniform, independent of pulse height thresholding or proximity to the SAA (Figure 2). For segment B, however, there are a number of bright spots in the region where the spectra fall when all pulse heights are included. These features disappear when the appropriate pulse height thresholding (used by default in the calcos pipeline for TIME-TAG data) is applied, as shown in Figure 3. Their relative intensity varies depending on proximity to the SAA. There was very little change in other detector features (hot spots, dead spots, etc.) between the pre- and post-launch data.

The NUV dark rate is fairly uniform across the detector. The initial on-orbit rate was $\sim 6 \times 10^{-5}$ counts pixel$^{-1}$ s$^{-1}$, which was below the expected on-orbit value. Since SM4 the rate has steadily increased, and by mid-2010 had reached $\sim 3.5 \times 10^{-4}$ counts pixel$^{-1}$ s$^{-1}$. Additional information on the trend of the NUV dark rate as a function of time is given in Zheng et al. (2010).

### 3. Flat Fields

FUV flat fields taken before launch were not able to improve the signal-to-noise ratio of spectra, so calcos currently does not apply a flat field for this channel. On orbit spectra of white dwarfs have been combined in an attempt to create a useful flat field for flight data. Although analysis has shown that two-dimensional flats constructed from these data do not
Figure 3: Dark counts for FUV Segment B with no pulse height thresholding (top), and with the default thresholding used by calcos (bottom). Using the appropriate thresholding minimizes the effects of the extra features near the middle of the segment.

give a significant improvement in the signal-to-noise ratio, one-dimensional flats (i.e., those summed over the cross-dispersion direction in the region of the spectra) show promise at doing so, and are currently being evaluated. Details of the analysis performed so far are described by Ake et al. (2010a).

Before launch, NUV flat field exposures were taken using both the internal flat field lamp and an external deuterium lamp. These data were used to create a high signal-to-noise flat field. On orbit, a series of exposures of the flat field lamp were collected during SMOV. In order to preserve the life of calibration lamps, a lower signal-to-noise ratio flight flat, with only enough counts to allow a comparison between the ground and flight flats, was obtained. A comparison shows that the two are identical, so they were combined into a single 'superflat' for use in the calcos pipeline (Figure 4; Ake et al., 2010b).

4. Sensitivity

Initial measurements of the COS effective area were consistent with the component throughputs determined on the ground, including a substantial sensitivity below the nominal MgF$_2$ cutoff at $\sim$1150 Å (McCandliss et al, 2010). Since then, the FUV channel has shown a wavelength-dependent sensitivity decline, at a rate of up to 12% per year, with the largest declines at longer wavelengths. This appears to be a detector effect, possibly related to the photocathode, and its cause is still under investigation.

Sensitivity monitoring measurements of both channels have been made monthly through much of Cycle 17. The details of the time dependent sensitivity monitoring, and the results, are discussed in these proceedings (Osten et al. 2010a), and in an STScI Instrument Science Report (Osten et al. 2010b).
Figure 4: NUV flat field used in the calcos pipeline, created by combining on-orbit and ground data. This flat only covers the region of the detector where the science spectrum is imaged, and includes a vignetting correction on the left side.
Figure 5: On-orbit cumulative image map for FUV Segment B. The bright stripe centered near Y=550 is where the science spectra fall during normal observations. The wavelength calibration lamp spectra are seen near Y=630, and Lyman-α airglow lines through the unused aperture at a variety of central wavelengths and FP-POSs are visible near the top and bottom of the segment. The most exposed areas are at the positions of the airglow lines in the science spectra.

5. FUV Thermal Stability

The FUV detector does not have physical pixels, but instead uses the digitization of timing signals to determine the location of each event. Since these calculations are sensitive to thermal variations of the detector hardware, ‘stim pulses’ are injected into the data stream in order to permit calcos to make a geometric correction for any thermal changes. On-orbit monitoring of the stim pulses has shown that normal orbital temperature variations are properly corrected by the calcos pipeline. However, there have been several instances where the detectors have been powered off for several days. The stim pulses in the first COS exposures collected after these events sometimes appear at incorrect locations, leading to an error in the thermal correction. Allowing the detector to reach a stable operating temperature before collecting science data, eliminates this problem.

6. FUV Gain Sag

The FUV detector has begun to show the effects of gain sag, which is a localized decrease in the number of electrons output by the MCPs due to their previous illumination history. The total number of counts that have fallen on the detector is spatially non-uniform (Figure
Figure 6: Modal gain as a function of total counts collected for Segment B. As the cumulative counts increase, the pulse height decreases nonlinearly. The Lyman-\(\alpha\) airglow regions at the far right show the most degradation.

5), and, as a result, the effect of gain sag is also very non-uniform. Regions of the detector that have seen the highest number of counts now show a substantial drop in the peak of their pulse height distribution.

Figure 6 shows the typical pulse height for all of the pixels on Segment B as a function of the total number of counts that have fallen on that pixel on-orbit. The pixels that have seen the most counts, which are those where geocoronal Lyman-\(\alpha\) falls, are clearly showing a decrease in the pulse height.

Work is ongoing to determine the implication of the gain sag on future observations. The FUV channel was designed to allow five lifetime positions on the detector, so the spectrum can be moved to more pristine areas. In addition, the high voltage on a segment can be adjusted in order to increase the MCP gain.

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Cosmic Origins Spectrograph: On-Orbit Performance of Target Acquisitions

Steven V. Penton

Center for Astrophysics and Space Astronomy - Astrophysics Research Lab,
University of Colorado, Boulder, CO, 80309

Abstract. COS is a slit-less spectrograph with a very small aperture (R=1.25"). To achieve the desired wavelength accuracies, HST+COS must center the target to within 0.1" of the center of the aperture for the FUV channel, and 0.04" for NUV. During SMOV and early Cycle 17 we fine-tuned the COS target acquisition (TA) procedures to exceed this accuracy for all three COS TA modes; NUV imaging, NUV spectroscopic, and FUV spectroscopic. In Cycle 17, we also adjusted the COS-to-FGS offsets in the SIAF file. This allows us to recommend skipping the time consuming ACQ/SEARCH in cases where the target coordinates are well known. Here we will compare the on-orbit performance of all COS TA modes in terms of centering accuracy, efficiency, and required signal-to-noise (S/N).

1. Introduction

There are four COS Target Acquisition (TA) modes:

- **ACQ/SEARCH**: performs a search in a spiral pattern by executing individual exposures at each point in a square grid. This mode can use either dispersed-light or imaging exposures.

- **ACQ/IMAGE**: obtains an NUV image of the target field, moves the telescope to center the object, and obtains a second NUV image as a confirmation of the centering.

- **ACQ/PEAKXD**: determines the cross-dispersion (XD) centroid of a spectrum in the direction perpendicular to dispersion and moves the telescope to center the XD direction.

- **ACQ/PEAKD**: centers the target spectrum in the along-dispersion (AD) direction by executing individual exposures at each point in a linear pattern along the AD axis.

Coordinate accuracy and target brightness will determine your choice of TA strategy and optional parameters. **ACQ/IMAGE** is often the fastest and most accurate centering mode, but, depending on target brightness, can take longer than the spectroscopic TA sequence of **ACQ/PEAKXD+PEAKD**. TAs on targets with positional uncertainties larger than 0.4" need to begin with an **ACQ/SEARCH** to ensure the target is in the COS aperture.

This document summarizes two more detailed COS TA documents: COS TIR 2010-03(v1) “COS Target Acquisition Guidelines, Recommendations, and Interpretation”, and COS TIR 2010-14(v1) “On-Orbit Target Acquisitions with HST+COS”. Interested observers should use these documents, and the COS Instrument Handbook, as their primary sources of information on COS TA.

2. Centering Requirements

Wavelengths assigned to COS data are required to have an absolute uncertainty of less than ±15 km/ms in the medium resolution modes, ±150 km/s in G140L mode and ±175 km/s in G230L mode. In the XD direction, the requirement is to be centered to within ±0.3", however, our goal is ±0.1" for FUV flat-fielding purposes. Since the AD requirement is in
units of km/s, it is detector and wavelength dependent as defined and shown in Equation 1.

$$\Delta AD(\text{arcsec}) = \frac{\Delta D_{\text{velocity requirement}} \times \lambda}{c \times \text{dispersion} (\AA/p) \times \text{platescale} (p/\text{arcsec})}$$  \hspace{1cm} (1)$$

Assuming that the wavelength error budget is split evenly between the COS TA and wavelength scale accuracy, the allowable TA centering errors, in arcseconds ("), are given in Table 1. The strictest requirement for each grating is highlighted in bold.

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3. HST Blind Pointing

By reverse engineering the actions performed by COS TAs during SMOV, we determined a bias in the blind pointing of HST+COS (left panel of Figure 1). The mean blind pointing offset in [AD,XD] coordinates was [-0.22±0.47, -0.17±0.49]", indicating a offset in the FGS-to-COS alignment. The outer black dashed circles in the Figure 1 represents the R=1.25" aperture, the red dotted circle is the centering limit to transmit maximum flux, the blue dashed circle is the FUV centering goal (0.106"), and the small black solid circle is the NUV centering goal (0.41"), In March 1, 2010 an adjustment was made to the Science Instrument Aperture File (SIAF) to correct for this blind-pointing bias. The right panel of Figure 1 shows the blind pointing estimate for observations taken after March 1, 2010. As expected, the HST blind pointing accuracy is not a function of TA mode employed (the blind pointing accuracies of each mode is given in Figure 1.), dominant FGS (red squares), or whether the target was a guest observer (GO), or calibration (CAL) program.

4. ACQ/SEARCH

As envisioned prior to launch, ACQ/SEARCHs should be able to center a target to within 0.1-0.2" in both AD and XD given that the target was within the box on the sky contained by the outer dwell points. This expectation was based on simulations based upon the predicted, and not the observed PSF. The observed on-orbit PSF is noticeably asymmetric and contains a much larger percentage of the light in the extended wings (see Figure 2). The PSF asymmetries and extended wings, along with the extended transmission function of the aperture tend to feed incorrect information into the ACQ/SEARCH centering algorithm.

Based upon the centerings required by TA procedures following ACQ/SEARCHs, the initial pointing accuracies of ACQ/SEARCH can be deduced. As detailed in COS TIR-2010-03(v1), imaging ACQ/SEARCHs show average [AD,XD] centering errors of magnitude [0.29±0.24]", while NUV and FUV spectroscopic ACQ/SEARCHs show average centering errors of [0.10±0.07]" and [0.19±0.15]", respectively. On-orbit, a single ACQ/SEARCH should, therefore, not be expected to center the target to better than 0.3" in either AD and XD, and depending on where the target falls in the pattern, it could easily be 0.4" off-center. This is why we recommend that all COS TAs follow up their initial ACQ/SEARCH with either an ACQ/IMAGE, ACQ/PEAKXD+PEAKD, or a second 2 × 2 ACQ/SEARCH. These pointing statistics
Figure 1: HST Blind Pointing: GO observations are highlighted by green circles, calibration exposures by red circles. If the dominant FGS was #3 a red box is added. The left panel is before the March 1, 2010 SIAF adjustment, the right panel is after the COS-to-FGS correction.

include only those ACQ/SEARCHs with S/N > 25 in the brightest dwell point. This is the minimum recommended S/N for ACQ/SEARCH, and the suggested S/N is 40. Higher S/Ns do not improve the final target centering as ACQ/SEARCH is always followed by additional TA procedures.

Because of the improvement in the HST 'blind pointing' due to the FGS-to-COS adjustment on March 1, 2010, targets with very well known coordinates are no longer required to precede ACQ/IMAGEs or spectroscopic ACQ/PEAKXD+ACQ/PEAKD sequences with ACQ/SEARCHs. Table 2 shows our updated ACQ/SEARCH usage guidelines. SCAN-SIZE equal to zero implies that the observer is bypassing ACQ/SEARCH altogether and proceeding directly to an ACQ/IMAGE or ACQ/PEAKXD+ ACQ/PEAKD.

<table>
<thead>
<tr>
<th>SCAN-SIZE</th>
<th>Target Coordinate Uncertainty (&quot;)</th>
<th>STEP-SIZE</th>
<th>CENTER method</th>
</tr>
</thead>
<tbody>
<tr>
<td>0</td>
<td>(\sigma \leq 0.4)</td>
<td>1.767&quot;^{a}</td>
<td>FW</td>
</tr>
<tr>
<td>2</td>
<td>0.4 (\leq \sigma \leq 0.7)</td>
<td>1.767&quot;^{b}</td>
<td>FW</td>
</tr>
<tr>
<td>3</td>
<td>0.7 (\leq \sigma \leq 1.0)</td>
<td>1.767&quot;^{b}</td>
<td>FWF</td>
</tr>
<tr>
<td>4</td>
<td>1.0 (\leq \sigma \leq 1.3)</td>
<td>1.767&quot;^{b}</td>
<td>FWF</td>
</tr>
<tr>
<td>5</td>
<td>1.3 (\leq \sigma \leq 1.6)</td>
<td>1.767&quot;^{b}</td>
<td>FWF</td>
</tr>
</tbody>
</table>

Table 2:

*a*Default STEP-SIZE value and the largest value which samples the search area without gaps.

*b*If the target coordinate uncertainty is on the low edge of the given range, you can reduce the STEP-SIZE slightly (e.g., 1.5") to improve centering accuracy by sacrificing the total area of the sky covered by the search pattern.
Figure 2: The left figure shows the on-orbit PSF at the COS aperture, the pre-launch estimate is on the right. The red circles are the size of the aperture. The pre-launch PSF has a central peak and a symmetric outer ring of radius 0.75″. The on-orbit PSF is not symmetric and has considerable power (> 5%) in extended wings and rings.

5. ACQ/IMAGE

First, a wavelength calibration lamp (WCA) exposure determines the location of the apertures on the NUV detector, then an image of the sky is taken. After measuring the positions of the target and the WCA, the slew required to center the target is determined and performed. Depending on target flux, this can be performed with (PSA or BOA) + (MIRRORA or MIRRORB). PSA/BOA images land on about the same location on the detector, but MIRRORA and MIRRORB (a reflection off the order sorter in front of MIRRORA) images land at different locations. Figure 3 below shows the on-orbit WCA, initial target locations, and the TA subarrays used in determining spot locations. The dashed boxes are proposed subarrays for an upcoming adjustment. This adjustment is required due to the rising NUV background rate. In this figure, only ACQ/IMAGEs after the FGS-to-COS adjustment of March 1, 2010 are shown.

Comparing the final target locations in the confirmation images to those commanded, we can measure the on-orbit centering accuracy of ACQ/IMAGE. The final [AD, XD] centering accuracy of all S/N > 25 ACQ/IMAGE to date is an impressive [-0.007 ± 0.016″, -0.009 ± 0.012″]. This is the minimum recommended S/N for ACQ/IMAGE, and the suggested S/N is 40 for PSA TAs and S/N of 60 for BOA TAs. Higher S/Ns will improve the final target centerings.

6. ACQ/PEAKXD

ACQ/PEAKXD sequences flash the wavelength calibration lamp (WCA), followed by a brief target spectrum. The WCA position determines the desired XD target location. The slew required to put the spectrum at this XD location is calculated and performed. For NUV TAs, one specifies which stripe to use (STRIPE=A, B, or C). For FUV TAs, one must specify which segment(s) to use (SEG=A, B, or BOTH). To limit detector and Geocoronal backgrounds from contributing to the XD location determinations, subarrays are used; one for the NUV and one or two for each of the FUV segments (as shown in Figure 4).

FUV ACQ/PEAKXDs are complicated by the:

1. distortions present in the raw coordinate frame (no thermal or geometric correction)

2. wavelength dependent XD profiles (different targets are centered differently)
Figure 3: Detector coordinate locations of NUV Imaging TA subarrays. Solid lines are current subarrays, dashed boxes are proposed Cycle 19 replacements. The two distributions on the left are the WCA locations for MIRRORA and MIRRORB images. The two distributions on the right are the initial target locations for PSA/BOA + MIRRORA and MIRRORB images.

3. mapping of segment-B spectra onto the segment-A coordinate system (each segment has different raw digital element XD sizes and zero point offsets).

These complications can cause objects with different spectral energy distributions (SEDs) to be centered differently. This offset can be as much as 0.3″, and is the most problematic for G160M TAs.

NUV ACQ/PEAKXDs are complicated by:

1. an additional error is present if STRIPE is not equal to B (±0.06″)
2. the stripes are tilted on the detector (different targets are centered differently (±0.05″)
3. each grating has a unique plate scale, but the average is currently used (±0.05″).

These NUV effects are smaller than the those seen in FUV TAs. The minimum recommended S/N for ACQ/PEAKXD is 40 for both BOA and PSA TAs. Higher S/Ns will improve the final target XD centerings.

7. ACQ/PEAKD

ACQ/PEAKD is a 1D along-dispersion (AD) version of ACQ/SEARCH. The transmission of the COS apertures is essentially flat within the central 0.5″ then tails off in a non-linear, but approximately symmetric profile.

The upper left panel of the left figure of Figure 5 shows the normalized ACQ/PEAKD results. The upper right panel of this figure shows the normalized counts after alignment based upon the chosen centering method. The lower left panel highlights the NUV results, while the lower right panel of the left figure gives the FUV results. The centering results for all appropriate observations were merged to a composite profile representing the average on-orbit AD profile. The average correcting maneuver was 0.018±0.127″. This profile was used to create simulations for all ranges of the parameters SCAN-SIZE and STEP-SIZE, with a detector background appropriate for a 15 second observation. As shown in the right
On-Orbit Performance of COS TAs

Figure 4: FUV TA subarrays superimposed on FUV raw detector coordinate images. Segment-A is on top, segment-B in on the bottom. Note the significant XD geometric distortion “kink” near X/10=400–500.

Figure of Figure 5, there are several ACQ/PEAKD centering options that are exceeding the FUV (0.106") and NUV (0.041") centering requirements.

8. Detector Backgrounds/SAA

The NUV detector background has been rising since launch, and is now 300 counts/sec for the entire detector. NUV ACQ/SEARCH and WCA ACQ/IMAGE subarrays are being optimized to minimize impact on COS TA. The FUV detector background has been steady since launch at 2-3 counts/sec/segment. No FUV subarrays currently need to be optimized for COS TA due to background effects. Both detectors show large background rate excursions near the SAA. The COS SAA defining contours were modified in May 2010 to prevent COS TAs from occurring near the SAA.

9. COS Focus

By monitoring the FWHM of the confirmation ACQ/IMAGEs we are tracking the imaging performance of COS, and hence, the stability of the HST+COS focus. The FWHM of the confirmation images from all COS ACQ/IMAGEs are shown in Figure 6. Non-point sources show larger FWHMs, so the focus is monitored by the narrowest measurements. No significant change in imaging performance has been detected for either the PSA or BOA in either the along-dispersion (AD) or cross-dispersion (XD) directions.

10. Conclusions and Recommendations

By evaluating the on-orbit actions performed by the COS TA flight software (FSW), we have been able to demonstrate the excellent performance of all COS TA modes during SMOV and Cycle 17. In this article and in COS ISR 2010-14(v1), we have updated Cycle 18 COS TA parameter choices and strategy recommendations. During SMOV and Cycle 17 many COS FSW TA parameter values were adjusted to our best determinations of the values for early operations. As outlined in COS TIR 2010-03(v1), there are still some outstanding TA issues and further TA optimizations are planned for Cycle 18 and beyond.
Figure 5: ACQ/PEAKD observations and simulation results. The upper left panel of the left figure shows the normalized ACQ/PEAKD results. The upper right panel of this figure shows the normalized counts after alignment based upon the chosen centering method. The lower left panel highlights the NUV results, while the lower right panel of the left figure gives the FUV results. The centering results for all appropriate observations were merged to a composite profile representing the average on-orbit AD profile. This profile was used to create simulations for all ranges of the parameters SCAN-SIZE and STEP-SIZE, with a detector background appropriate for a 15 second observation. The right figure shows that there are several ACQ/PEAKD centering options that are exceeding the FUV (0.106") and NUV (0.041") centering requirements. ACQ/PEAKDs with CENTER=RTB (return-to-brightest) never achieve these requirements.

Figure 6: COS focus monitoring. Neither the BOA ACQ/IMAGE images (top, circles) or the PSA images (bottom, triangles) show any change in AD (blue) or XD (red) FWHM on-orbit.
References


Wavelength Calibration of the Cosmic Origins Spectrograph

Cristina Oliveira\(^1\), Stéphane Béland\(^2\), C. Tony Keyes\(^1\), Alessandra Aloisi\(^1\), Sami Niemi\(^1\), Steve Osterman\(^2\), and Charles Proffitt\(^1,3\)

Abstract. Cosmic Origins Spectrograph (COS) calibration programs 11487 and 11474 were carried out during Servicing Mission Observatory Verification (SMOV) to characterize the wavelength scales of the FUV and NUV detectors, respectively. Science data were obtained with all the FUV (G130M, G160M, and G140L) and NUV (G185M, G225M, G285M, and G230L) gratings through the Primary Science Aperture (PSA), simultaneously with PtNe lamp line spectra, obtained through the Wavelength Calibration Aperture (WCA). These data were partially calibrated with the COS pipeline and then, in conjunction with STIS data of the same target, used to derive offsets between the PSA and WCA wavelength scales. These offsets are then used to place the dispersion relations derived from thermal vacuum 2003 (TV03) in the on-orbit frame of reference, so that correct wavelengths can be assigned to the on-orbit COS data.

1. Introduction

The COS FUV channel contains one low resolution (G140L) and two medium resolution (G130M, G160M) gratings, providing wavelength coverage from the Lyman limit to \(\sim 2250\) Å, while the NUV channel contains one low resolution (G230L) and three medium resolution (G185M, G225M, and G285M) gratings, providing coverage from 1670 to 3560 Å.

Spectra of external targets are obtained through the Primary Science Aperture (PSA) while wavelength calibration spectra are obtained through the Wavelength Calibration Aperture (WCA). There are two PtNe line lamps on-board, and wavecal data can be obtained concurrently with science data (TAGFLASH mode) or either before or after a science exposure is obtained. Wavecal data is used by the COS pipeline (CalCOS) to correct for any offsets between the wavecals and the lamp template reference files due to non-repeatability of the grating mechanism positions. These offsets are then applied to the science data as well, before wavelengths are assigned (COS Data Handbook, Shaw et al. 2009).

During thermal vacuum 2003 (TV03) science (PSA) and wavecal (WCA) data were obtained at all central wavelengths of the FUV and NUV gratings, using an internal and an external PtNe line lamp. These data were used to derive dispersion relations for all the FUV and NUV modes (for FP-POS = 3 only), by cross-correlating the observed spectra, in pixel space, with a PtNe wavelength line list. For the FUV a linear dispersion relation was used for the medium resolution gratings (G130M and G160M) and a second order polynomial was used with the low dispersion grating (G140L). A second order dispersion relation was used for all the NUV gratings. Because of the different optical paths between the PSA and WCA, the dispersion relations derived for each aperture are different.

\(^1\)Space Telescope Science Institute
\(^2\)University of Colorado
\(^3\)Computer Sciences Corporation
The dispersion solutions for the PSA are not adequately represented by a simple offset of the WCA dispersion solutions. A separate dispersion solution is needed per segment, for each central wavelength of each grating, for each aperture. Because the WCA and PSA have different dispersion solutions we cannot use WCA data obtained on-orbit to determine directly the on-orbit dispersion solutions for the PSA. Instead we use the PSA dispersion solutions derived from TV03 data to wavelength calibrate PSA data obtained on-orbit. However, the TV03 dispersion solutions cannot be directly applied to on-orbit data without placing the TV03 dispersion solutions on the on-orbit frame of reference, which implies deriving the on-orbit offsets between the PSA and WCA spectra, as described below. Note that all the FUV data needs to be thermally and geometrically corrected, assuring that the spectra are all in the same reference frame, before any other analysis.

Programs 11487 (FUV) and 11474 (NUV) were executed during SMOV to obtain spectra of external targets, that would allow us to determine the on-orbit offsets between the PSA and WCA data. These offsets are then used to place the dispersion relations derived in TV03 in the on-orbit frame of reference, so that correct wavelengths can be assigned to the on-orbit COS data, using the TV03-derived dispersion coefficients. As a first approximation we determine an average value, constant in wavelength for each setting, for the offsets between the PSA and WCA apertures. Future refinements to the FUV wavelength calibration will take into account the second order effect resulting from the fact that the PSA and WCA dispersion solutions are different, and so the offsets between the PSA and WCA are wavelength dependent.

Programs 11488 and 11475, companion programs to programs 11487 and 11474, respectively, were executed during SMOV with the purpose of obtaining wavelength calibration spectra through the WCA.

2. Observations

All the visits in program 11487 (COS FUV Internal/External Wavelength Scales) used a NUV target acquisition sequence consisting of ACQ/SEARCH, followed by ACQ/SEARCH, ACQ/PEAKXD, and ACQ/PEAKD to achieve maximum centering accuracy. All the exposures were obtained in TIME-TAG mode and used special lamp flash parameters (available in ENG mode only) to ensure that multiple lamp flashes were present in each exposure (particularly for the longer exposures) and that any mechanism drift, if at all present, could be easily corrected. Visits 50 and 73 observed NGC 330–B37, a B super-giant in the SMC, which has been observed with STIS/E140M. The analysis presented here use data from these two visits.

Program 11488 (COS Internal FUV Wavelength Verification) obtained internal wavelength calibration spectra using the default PtNe lamp (Lamp 1) with each FUV grating, at each central wavelength setting, and for each FP-POS position. Each exposure had a duration of 120 sec. In order to allow the mechanism drift to settle after a grating motion, a 1800 sec exposure (containing multiple flashes of the lamp) was taken at the beginning of each visit.

All the visits in program 11474 (COS NUV Internal/External Wavelength Scales) used the same target acquisition strategy as described above for the FUV to achieve maximum centering accuracy. All the exposures (except for G285M/2979 and G285M/3035 in Visit 01) were obtained in TIME-TAG mode and used the special lamp flash parameters described above for the FUV.

Visit 01 observed HD 187691 (F8V), a CORAVEL radial velocity standard (Udry et al. 1999), with $v_{\text{rad}} = 0.0 \pm 0.3$ km/s, which has also been observed by STIS (E230M, 0.2x0.06, exposure o52601020, covering [2274,3118] Å). Visit 02 observed another radial
velocity standard, HD 6655 (F8V), which has $v_{\text{rad}} = 19.5 \pm 0.3 \text{ km/s}$. The analysis presented here uses data obtained in the two visits described above.

3. Analysis

To assess the impact of mechanism drift in the wavecal exposures obtained in programs 11488 and 11475 (FUV and NUV programs used to update the lamp template reference file), each exposure was divided into 30 sec time slices, for each segment or stripe, and each time slice was cross-correlated with the first time slice. The maximum drift for any wavecal exposure was 1.3 and 0.9 pixels (FUV and NUV, respectively), but most were within 0.5 and 0.3 pixels (FUV and NUV, respectively), much smaller than the resolution element of 6 pixels for the FUV and 3 pixels for the NUV, and so no drift correction was applied to the wavecal exposures.

After being corrected for thermal and geometric distortions, the wavecal data from program 11488 (FUV) were extracted into 1d spectra (counts vs pixel) and then the 1d spectra were used to produce a new FUV lamp template reference file, containing one entry per FP-POS for each grating/cenwave/segment combination. In addition, each of the 1d spectra corresponding to FP-POS = 3 was cross-correlated with the 1d spectra for the same grating/cenwave/segment mode and FP-POS = 3 in the TV03-based lamp template reference file, in order to determine their separation, $\text{WCA}_{\text{TV03}} - \text{WCA}_{\text{SMOV}}$.

Similarly to the FUV, the wavecal data from program 11475 (NUV) were used to produce a new NUV lamp template reference file, containing one entry per FP-POS for each grating/cenwave/stripe combination and in conjunction with TV03 data, $\text{WCA}_{\text{TV03}} - \text{WCA}_{\text{SMOV}}$ was determined.

This value is needed so that the wavelength calibration is performed properly, as all the wavecal data obtained on-orbit is compared with a lamp template reference file that was also updated from on-orbit data. In addition, since the dispersion solutions from TV03, which are applied to the on-orbit data, and the lamp template reference file for TV03 are only for the FP-POS = 3 position, the separation between the $\text{WCA}_{\text{TV03}} - \text{WCA}_{\text{SMOV}}$ for other FP-POS is not necessary (see below).

3.1. NGC 330–B37

The rawtag files for each of the science exposures obtained in Visits 50 and 73 of program 11487 were partially calibrated with CalCOS by setting all the calibration switches to $\text{OMIT}$ except for $\text{TEMPCORR}$, $\text{GEOCORR}$, $\text{IGEOCORR}$, $\text{DOPPCORR}$, and $\text{WAVECORR}$, which were set to $\text{PERFORM}$. The new lamp template reference file, described above, was used in this calibration, so that the exposures are placed in the on-orbit frame of reference. To produce 1d science and wavecal spectra, the $x_{\text{full}}$ and $y_{\text{full}}$ columns of the resulting corrtag files, which are corrected for temperature and geometric distortions, doppler smearing, and mechanism drift for the science spectra, and for temperature and geometric distortion, and mechanism drift for the wavecal spectra, were extracted using $\text{Cedar}$, each segment independently, using the default extraction parameters used by CalCOS. The resulting 1d files contain then counts vs pixel.

The on-orbit separation, in pixels, between each segment of the PSA and WCA data, was determined by first cross-correlating the 1d WCA spectra with a PtNe line list, so that wavelengths can be assigned to each of the PtNe emission lines seen in pixel space. Each emission line, $i$, has then a pixel coordinate, $\text{WCA}_{\text{SMOV},i}$.

The heliocentric correction is only applied by the pipeline when wavelengths are being assigned. Since our analysis of the science data was done in pixel space, this correction was not applied. However, since for all the exposures analyzed here $|v_{\text{Helio}}| \leq 2 \text{ km/s}$ this effect can be neglected as it is less than 0.5 pixels for both the M and L gratings.
The next step is to identify the pixel positions in the 1d PSA spectra, corresponding to the wavelengths of the PtNe emission lines, identified in the previous step, PSA_{SMOV,i}. In the case of data obtained in Visits 50 and 73 for NGC 330–B37 this was accomplished by using the STIS/E140M spectrum of this target. For each feature corresponding to the wavelength being considered, the STIS spectrum was used as a guide, so that the pixel position of the same feature could be looked up in pixel space, in the 1d PSA COS spectrum. For each grating/cenwave/segment, this procedure was applied to as many features as possible so that one can determine PSA_{SMOV,i} - WCA_{SMOV,i} across each segment of the FUV detector. The on-orbit separation between the PSA and WCA data is then calculated, for each grating/cenwave/segment, by calculating the mean of \( (PSA_{SMOV,i} - WCA_{SMOV,i}) = PSA_{SMOV} - WCA_{SMOV} \).

3.2. HD 187691 and HD 6655

The rawtag files for each of the science exposures obtained in Visit 01 and Visit 02 of program 11474 were partially calibrated with CalCOS by setting all the calibration switches to OMIT except for DOPPCORR and WAVECORR, which were set to PERFORM. The new lamp template reference file, described above, was used in this calibration, so that the exposures are placed in the on-orbit frame of reference. To produce 1d science and wavecal spectra, the xfull and yfull columns of the resulting corrtag files, which are corrected for doppler smearing and mechanism drift for the science spectra, and for mechanism drift for the wavecal spectra, were extracted using Cedar, each stripe independently, using the default extraction parameters used by CalCOS. The resulting 1d files contain then counts vs pixel.

The heliocentric correction is only applied by the pipeline when wavelengths are being assigned. Since our analysis of the science data was done in pixel space, this correction was not applied. However, since for all the exposures analyzed here \( |v_{\text{Helio}}| \leq 2 \text{ km/s} \) this effect is less than 0.5 pixels for both the M and L gratings.

The on-orbit separation, in pixels, between each stripe of the PSA and WCA data, was determined by first cross-correlating the 1d WCA spectra with a PtNe line list, so that wavelengths can be assigned to each of the PtNe emission lines seen in pixel space. Each emission line, \( i \), has then a pixel coordinate, WCA_{SMOV,i}.

The next step is to identify the pixel positions in the 1d PSA spectra, corresponding to the wavelengths of the PtNe emission lines, identified in the previous step, PSA_{SMOV,i}. In the case of data obtained in Vis 01 for HD 187691 this was accomplished by using the STIS/E230M spectrum of this target. For each feature corresponding to the wavelength being considered, the STIS spectrum was used as a guide, so that the pixel position of the same feature could be looked up in pixel space, in the 1d PSA COS spectrum. For each grating/cenwave/stripes this procedure was applied to as many features as possible so that one can determine PSA_{SMOV,i} - WCA_{SMOV,i} across the detector. The on-orbit separation between the PSA and WCA data is then calculated, for each grating/cenwave/stripes, by calculating the mean of \( (PSA_{SMOV,i} - WCA_{SMOV,i}) = PSA_{SMOV} - WCA_{SMOV} \).

HD 6655, observed in Visit 02 of program 11474, has not been observed by STIS, and so this procedure cannot be applied as described above. However, since HD 6655 and HD 187691 have the same spectral type, and both have accurately known radial velocities, one can simulate the STIS spectrum of HD 6655 by using the STIS spectrum of HD 187691 and taking both of the radial velocities into account. The simulated HD 655 STIS spectrum can then be used to derive PSA_{SMOV,i}, as described above.

4. Wavelength Calibration in CalCOS

Wavelength calibration by the COS pipeline is described in detail in the COS Data Handbook (Shaw et al. 2009), here we only present a summary.
The COS pipeline determines the dispersion-direction shift that needs to apply to the science data (PSA) to correct shifts due to the non-repeatability of the OSM mechanism by comparing the wavecal data (WCA) obtained concurrently (if in tagflash mode) with the data contained in the lamp template reference file. This reference file contains one entry per FP-POS for each of the available modes, and so the wavecal data is compared with the lamp data at the same FP-POS. The lamp template reference file contains also a column that gives the separation in pixels between the lamp template at FP-POS = 3 and the lamp templates at the other FP-POS. The shift determined by comparing the wavecal data to the lamp plus the shift read from the column of the lamp template reference file are applied to the data, producing the $x_{\text{full}}$ pixel coordinate.

The COS pipeline computes then the wavelength from the $x_{\text{full}}$ pixel coordinate, $x_{\text{full}}$, as $x_{\text{prime}} = x_{\text{full}} + d_{\text{PSA}}$, where $d_{\text{PSA}} = d_{\text{TV03}} - d$. $d_{\text{TV03}}$ is the separation between the PSA and WCA data in TV03, PSA_{TV03} - WCA_{TV03}, and $d$ is the separation between the PSA and WCA data on-orbit (SMOV), taking into account that the WCA data in SMOV is offset from the WCA in TV03: $d = PSA_{\text{SMOV}} - WCA_{\text{SMOV}} - (WCA_{\text{TV03}} - WCA_{\text{SMOV}})$. The wavelengths are then computed from $\lambda = a_0 + a_1 \times x_{\text{prime}} + a_2 \times x_{\text{prime}}^2$, where $a_0$, $a_1$, and $a_2$ are the dispersion coefficients derived from TV03 data. Each segment/stripe of each grating/cenwave mode has its own dispersion solution.

5. Results

5.1. FUV Results

The offsets derived in this analysis, $d_{\text{PSA}}$, were used to update the FUV wavelength dispersion reference file which was then used to process COS data obtained in three SMOV programs (independent from the programs used for wavelength calibration), so that the accuracy of the wavelength scales can be determined.

All the COS data were processed through CalCOS (version 2.11f) and all the calibration switches were set to PERFORM, except for TDSCCORR, STATFLAG, and BRSTCORR, which were set to OMIT. STIS data were retrieved from the MAST archive and used to evaluate the COS wavelength scales. To facilitate the comparison between the COS and STIS wavelength scales, the STIS/E140H data used for comparison with the COS M resolution data and the STIS/E140M data used for comparison with the COS/G140L data, were convolved with the COS LSF appropriate to the COS grating being compared and appropriate to the wavelength in question (see Ghavamian et al. 2009).

Figure 1 compares wavelength calibrated COS G130M/1291 data for SK 155 (program 11489, FP-POS = 3; COS External Spectroscopic Performance - Part I) and STIS/E140H data for the same target (red). STIS data convolved with the LSFs appropriate to the wavelengths and settings displayed (see Ghavamian et al. 2009) are also shown (green). For segment A the COS and STIS wavelengths are in good agreement; for segment B the COS wavelengths are under-predicted by $\sim 4$ pixels at 1190 Å while at 1260 Å they are under-predicted by $\sim 2$ pixels.

Figure 2 compares wavelength calibrated COS G160M/1623 data for SK 155 (program 11489, all FP-POS combined) with that of STIS E140H for the same target (red). STIS data convolved with the LSF appropriate to the wavelengths and settings displayed are also shown (green). For segment A the COS wavelengths are over-predicted by $\sim 3$ pixels, for segment B the COS wavelengths are in good agreement with those of STIS at $\sim 1526$ Å, but they are under-predicted by $\sim 3$ pixels at $\sim 1609$ Å.

Figure 3 compares the wavelength scales of G140L/1230 and G140L/1105, using COS data of NGC 330–B37 (program 11487, COS FUV Internal/External Wavelength Scales, FP-POS=3), with that of STIS E140M (red), for two wavelength regions of segment A. STIS data convolved with the LSFs appropriate to the wavelengths and settings displayed
Figure 1: Comparison between wavelength calibrated COS G130M/1291 data for SK 155 (black) and STIS/E140H data for the same target (red). STIS data convolved with the LSF appropriate to the wavelengths and settings displayed are also shown (green). For segment A the COS and STIS wavelengths are in good agreement; for segment B the COS wavelengths are under-predicted by $\sim$4 pixels at 1190 Å while at 1260 Å they are under-predicted by $\sim$2 pixels. COS data is from program 11489, for FP-POS = 3.

Figure 2: Comparison between wavelength calibrated COS G160M/1623 data for Sk 155 (black) and STIS/E140H data for the same target (red). STIS data convolved with the LSF appropriate to the wavelengths and settings displayed are also shown (green). For segment A the COS wavelengths are over-predicted by $\sim$3 pixels; for segment B the 1526 Å wavelength region the COS wavelengths are in good agreement with those from STIS, while at 1609 Å they are under-predicted by $\sim$3 pix. COS data is from program 11489, FP-POS = 1,2,3,4 (combined).
Figure 3: Comparison between wavelength calibrated COS G140L/1230 (left panel) and G140L/1105 (right panel) data for NGC 330-B37 (black, program 11487, FP=3) and STIS E140M data for the same target (red). STIS data convolved with the LSF appropriate to the wavelengths and settings displayed are also shown. For G140L/1230 the COS wavelengths are over-predicted by $\sim 3$ pix for 1335 Å and under-predicted by $\sim 1$ pix for 1609 Å. For G140L/1105 the COS wavelengths are under predicted by $\sim 2$ pix for 1335 Å and they agree well with those of STIS at 1609 Å.

Evaluating the COS FUV wavelength scales in the figures discussed above is made complicated by two issues that affect the COS FUV data. One of this issues is due to the wings of the LSF, which fill the cores of the saturated lines, affecting also the depths of weaker non-saturated lines (for more on the LSF see Ghavamian et al. 2009). We tried to minimize this effect by convolving the STIS E140H (SK 155) and E140M (NGC 330–B37) data with the COS LSFs. The other issue is related to the fact that the shadows from the gain efficiency (GE) grid wires produce features that are similar to absorption lines and that change the profiles of real absorption lines (at the time this data was processed grid wires were not flagged by the COS pipeline, but they are currently).

Even though the figures discussed above seem to indicate that the accuracy of the wavelength scales is within the specifications defined in Figure 4, the limited number of datasets on which the updated wavelength scales have been tested makes any conclusions preliminary. More work is required to further refine our evaluation of the FUV wavelength scales in the future.

5.2. NUV Results

The offsets derived in this analysis, $d_{PSA}$, were used to update the wavelength dispersion reference file, which was then used to process COS data obtained in two SMOV programs (independent of the SMOV wavelength calibration programs), so that the accuracy of the wavelength scales can be determined.

Visit 01 of COS program 11477 (COS NUV External Spectroscopic Performance - Part 2) observed the symbiotic star AG Draconis, a spatially unresolved point source with an emission line spectrum, on Sep 7 2009. Exposure labm01eq was obtained with G225M/2390
Figure 4: COS specifications for FUV wavelength accuracies. The internal error includes the accuracy of the wavelength scale, the dispersion relation, aperture offsets, distortions and drifts. The error goal includes contributions from external sources, which are dominated by target mis-centering in the aperture. Error goal given per exposure.

<table>
<thead>
<tr>
<th>Grating</th>
<th>Error Goal ((1\sigma)) (km s(^{-1}))</th>
<th>Error Goal ((1\sigma)) (pixels)</th>
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<td>G130M</td>
<td>15</td>
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<td>150</td>
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at FP-POS=3 (TAGFLASH, 110 sec); exposure labm01dqq was obtained with G285M/2739 also at FP-POS=3 (TAGFLASH, 70 sec). AG Draconis has been previously observed by STIS with the E230M grating (exposure o6ky01030, obtained with the 0.2x0.2 aperture and 2707 cenwave; 374 sec) and so the STIS observations can be used to compare the STIS wavelength scales to those of COS.

The two COS exposures of AG Draconis were calibrated with CalCOS (version 2.11f) by setting all the calibration switches to \textit{PERFORM} except for \textit{STATFLAG} and \textit{TDS_CORR} which were set to \textit{OMIT}. The resulting spectra, where then compared with the STIS spectra of AG Draconis, which were retrieved from the MAST archive. No convolution of the STIS data, with the COS LSF, was performed. Data with intrinsically broad lines was used to evaluate the G225M and G285M wavelength scales; in the case of G230L, the STIS data used for the comparison of the wavelength scales is of poorer resolution.

Figure 5 shows one wavelength region for each of the three COS stripes of the G225M/2390 AG Draconis observation. Overplotted in red in the STIS/E230M spectrum of this target. There is good agreement between the COS and STIS wavelengths for stripes A and B, for stripe C the COS wavelengths are under-predicted by \(\sim 2\) pixels.

Figure 6 shows one wavelength region for each of the three COS stripes of the G285M/2739 AG Draconis observation (in black), again with the STIS/E230M spectrum of this target overplotted in red. For both stripes A and C the COS wavelengths are over-predicted by \(\sim 1\) pixel, while for stripe B the COS wavelengths are under-predicted by \(\sim 1\) pixel.

To evaluate the G230L wavelength scales we used COS data of AzV18, observed in program 11472 (COS NUV Dispersed-light Acquisition Algorithm Verification), with the G230L/2950/FP-POS=3 setup (exposure laa004p1q, observed in TIME-TAG mode for 60 sec on Aug 11, 2009). This target has also been observed by STIS, with the G230LB grating, through the 52x2 aperture (exposure o6df06010). Even though the G230LB grating is of lower resolution than the G230L grating, comparing the two spectra is still useful to evaluate the G230L wavelength scales.

Figure 7 displays the Mg\(\text{II}\) \(\lambda\lambda 2796, 2803\), and Mg\(\text{I}\) \(\lambda 2852\) regions in the COS observation of AzV18 (G230L/2950/strip B, in black), with the STIS/G230LB spectrum overplotted in red. For this stripe the COS wavelengths are under-predicted by \(\sim 1–2\) pix-
Figure 5: Comparison between STIS/E230M (red) and COS/G225M/2390 (black) wavelength scales using AG DRAC data. COS data is from program 11477, for FP-POS = 3.

Figure 6: Comparison between STIS/E230M (red) and COS/G285M/2739 (black) wavelength scales using AG DRAC data. COS data is from program 11477, for FP-POS = 3.
Figure 7: Comparison between STIS/G230LB (red) and COS/G230L/2950 (black) wavelength scales using AzV 18 data. COS data is from program 11472, for FP-POS = 3.

d. Even though the figures presented here seem to indicate that the accuracy of the wavelength scales is within the specifications defined in Figure 8 (except for the G185M wavelength scales which have not been updated yet), the limited number of datasets on which the updated wavelength scales have been tested makes any conclusions preliminary. More work is required to further refine our evaluation of the NUV wavelength scales, in the future.

6. Future Work

These updated wavelength scales are currently in use by the CalCOS pipeline. We have only compared the COS wavelength scales to those of STIS for a limited number of datasets both in the FUV and NUV. However, all the cases presented here seem to indicate that the accuracy of the wavelength scales is within the specifications defined in Figures 4 and 8. There are a few cases where localized distortions of up to 10 pixels have been seen in the FUV. The nature and frequency of these localized distortions is under investigation. Two Cycle 17 calibration programs (11997 for FUV and 11900 for NUV) are used to monitor the offsets between the PSA and WCA apertures. Preliminary analysis of the data obtained in these programs indicates that the offsets between the PSA and WCA apertures are stable. We are in the process of analyzing data obtained in other SMOV, GTO, and GO programs that will allow us to further refine our evaluation of the wavelength accuracies.

References

Shaw, B. et al., "COS Data Handbook", version 1.0, (Baltimore:STScI)
Figure 8: COS specifications for NUV wavelength accuracies. The internal error includes the accuracy of the wavelength scale, the dispersion relation, aperture offsets, distortions and drifts. The error goal includes contributions from external sources, which are dominated by target mis-centering in the aperture. Error goal given per exposure.

<table>
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<td>G230L</td>
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Abstract. During the SM4 Servicing Mission Observatory Verification (SMOV) we discovered that the on-orbit shape of the COS LSF with the HST optical telescope assembly (OTA) exhibits broad wings. The wings are caused by mid-frequency wave-front errors (MFWFEs) that are produced by the zonal (polishing) errors on the HST primary and secondary mirrors; these errors could not be simulated during ground testing. The MFWFEs are particularly noticeable in the FUV. The on-orbit LSF has up to 40% of its total power distributed in non-Gaussian wings (as compared to 26% for a Gaussian). The power in these wings is largest at the shortest wavelengths covered by the COS FUV medium-resolution gratings (∼1150 Å). The effect decreases with increasing wavelength but has a non-negligible effect on encircled energies even at the longest wavelengths. We have calculated optical models incorporating the MFWFE effects into the LSF for the whole spectral range covered by the FUV and NUV medium-resolution gratings. We show that for the FUV, the convolution of these model LSFs with high-resolution STIS echelle spectra yields an excellent match to the on-orbit COS spectra of the same targets. The model LSFs are available online and can be used by COS observers to assess the impact of the MFWFE broadening on their COS spectra. In ground tests it was shown that COS can spatially resolve two equally bright objects separated by 1" in the cross-dispersion direction in the FUV. Using the FUV spectra of white dwarfs acquired during the Cycle 17 COS Spectroscopic Sensitivity Monitoring program, we show the on-orbit spatial resolution (as defined by the full-width half maximum of the spectrum along cross-dispersion) meets this specification, though in a wavelength-dependent manner. The wavelength dependence is primarily due the astigmatism introduced by the FUV gratings in cross-dispersion. The spatial resolution also depends on the central wavelength (CENWAVE) setting used, with spatial resolution monotonically improving with increasing CENWAVE settings.

1. Introduction

The Cosmic Origins Spectrograph (COS) was installed during the most recent servicing mission of the Hubble Space Telescope (SM4) and is the most sensitive ultraviolet spectrograph flown to date. With its medium-resolution gratings (G130M and G160M, covering 1150 Å - 1800 Å) this channel was designed to reach a spectroscopic resolving power of at least 20,000 (15 km s$^{-1}$) across 80% of its passband (STE-63, 2004). The COS near ultraviolet channel was designed to cover the 1750 Å - 3200 Å spectral range with a sensitivity 2-3 times larger than STIS. It utilizes medium-resolution gratings (G185M, G225M and G285M) with similar spectroscopic resolving power requirements to the FUV.
Thermal vacuum measurements showed that for the G130M and G160M FUV gratings, the line spread function (LSF) of an unresolved line was well approximated by a Gaussian profile with a FWHM of approximately 6.5 pixels, corresponding to $\Delta\lambda = 0.065$ Å at 1300 Å and $\Delta\lambda = 0.079$ Å at 1600 Å. The LSFs of the NUV gratings, on the other hand, are expected to have broad wings due to the response of the MAMA detector in the NUV (similar to what is seen for the STIS MAMA LSFs).

The FUV spatial resolution was measured during the 2003 thermal vacuum testing of COS. Spectra of a PtNe lamp were obtained through a pinhole array, with 3 of the pinholes seen in a single exposure. The instrument specification for spatial resolution required that 90% of the energy of a single point source centered in the Primary Science Aperture (PSA) fall within a range of 300 µm or less in the cross-dispersion direction (STE-63; 2004), corresponding to 1.1″. In tests with the G160M grating, the PtNe spectra from each pinhole were spatially almost completely unblended, with 99% of the counts falling within 12 pixels (300 µm). The ground test indicated that the spatial resolution was considerably better than 1″ in G160M.

2. On-Orbit Results

Analysis of stellar spectra obtained during SMOV indicates that mid-frequency wavefront errors (MFWFEs) created by zonal (polishing) errors on the HST OTA produce non-Gaussian wings in the on-orbit COS LSF, both broadening the core of the profile and lowering its amplitude. These features were mapped via a phase-retrieval analysis of WFPC2 imagery by Krist & Burrows (1995; see Figure 1). However, the MFWFEs are not corrected by the optics of COS or any other HST instrument. Therefore, the beam entering COS on-orbit is slightly different from the beam fed into the instrument during thermal vacuum testing.
COS Spectral and Spatial Resolution

Figure 2: Left: Calculated LSFs for the COS FUV medium-resolution gratings, as observed through the PSA aperture. Models including the effects of the mid-frequency wavefront errors are shown in solid. The dotted line shows a Gaussian LSF with a 6.5 pixel FWHM (the nominal prediction from ground testing, which did not include the wavefront errors). Right: same models, but for the NUV channel. Note the presence of non-Gaussian wings even in the absence of the MFWFEs, induced by the response of the MAMA NUV detector. The MFWFE contribution to the LSF wings is smaller in the NUV than the FUV, particularly for $\lambda > 2500 \, \text{Å}$.

3. LSF Models: FUV and NUV

We computed model LSFs for the COS gratings from the expected aberration content of the COS + HST OTA system, OTA pupil geometry, OTA MFWFEs as determined by Krist & Burrows (1995), and estimates of the point response function of the detectors. The LSFs for the FUV channel were produced for each grating by first matching monochromatic images generated by a Code V optical model of the COS + HST OTA system to emission line images from thermal vacuum testing. In particular, a mean detector-induced blur kernel was estimated by matching a suite of images over the spectrum. The model was then used to generate Zernike aberration coefficients, which, together with the OTA pupil function and the MFWFEs, were employed to compute the expected PSF at a number of wavelengths for each grating. The PSFs were then convolved with the estimated detector blur kernel and integrated in the cross-dispersion direction to form the LSFs. Figure 2 shows that these FUV LSFs are characterized by prominent wings, broader cores and lower central peaks than the nearly Gaussian LSFs computed without the OTA MFWFEs.

We also produced model LSFs for the COS NUV medium-resolution gratings that incorporate MFWFEs. The relative contributions of the wings from the MAMA detector and the wings produced by the MFWFEs are illustrated in the right panel of Figure 2, where we compare model NUV LSFs with and without the MFWFEs included.

We have produced optical models of the COS LSF, taking into account the MFWFEs (as well as the appropriate detector responses). We have made these model LSFs available to the astronomical community for use in analyzing COS science spectra and for the plan-
Figure 3: Right: Closeup views of prominent absorption features in the COS G130M (Segment B) spectrum of Sk 155 (bottom panels). For comparison, the same features in the STIS E140H spectrum (R \sim 114,000) are shown (top panels). Left: The STIS E140H spectrum convolved with an R = 20,000 Gaussian is overplotted in blue, while the convolution of the STIS data with the MFWFE LSF model appropriate to the wavelength range is shown in red. The Gaussian model produces a noticeably poorer match to the COS spectrum.

4. Comparison to On-Orbit COS Data

In Figure 3 we compare the COS spectrum from SMOV (black solid line) of the O9 Ib star Sk 155 in the LMC with a STIS E140H high resolution spectrum (R \sim 114,000) that has been convolved with an R = 20,000 Gaussian LSF (blue dashed line). Also shown in Figure 3 is the STIS spectrum convolved with an MFWFE LSF model appropriate for the wavelengths displayed (solid red line). This figure clearly shows that the rounded, filled-in absorption cores are not properly reproduced by an R = 20,000 Gaussian LSF, which systematically underpredicts the flux at line centers and produces more boxy shapes for the broad, saturated absorption lines than is observed. Similar trends are observed in data from the G160M grating.

5. Impact on COS Science

Science observations intending to use the full resolution of the FUV G130M and G160M and the NUV G185M medium-resolution gratings will be most seriously affected. The impact of the on-orbit LSF on the COS science observations are as follows:

• At a given signal-to-noise, weak, narrow features \((b \leq 35 \text{ km s}^{-1})\) will be more difficult to detect.

• Closely spaced spectral features may blend and become more difficult to isolate kinematically.

• Studies requiring measurement of accurate line profile shapes (as well as saturated lines) will also require full consideration of the LSF.

• Spectral purity will be reduced, resulting in decreased contrast between line cores and wings.

6. FUV Spatial Resolution

In Figure 4 we show the fitted widths of the cross-dispersion (spatial) profiles for COS FUV spectra. The spectra are of white dwarfs observed during the Cycle 17 FUV sensitivity monitoring program (G130M/G140L: WD0947+857; G160M: WD1057+719). Although the spatial profiles are double-peaked at the shortest wavelengths (G130M, Segment B), a single Gaussian FWHM is still useful as an approximate measure of the spatial resolution. The main results from Figure 4 are:

1. For a given FUV grating, the spatial resolution depends strongly on central wavelength setting (CENWAVE). The lowest spatial resolution (i.e., the broadest spatial profiles) occurs at the shortest CENWAVE settings, with longer CENWAVE settings giving progressively better spatial resolution. The resolution meets or exceeds specification for all gratings and central wavelengths save for the G130M/1291 combination.

2. At a given CENWAVE, the spatial resolution tends to be highest at the extreme wavelength ends of the medium resolution spectra, while for G140L the spatial resolution is highest near the center of the spectrum. The resolution ranges from 0.5” to 1.8” for G130M, 0.3” to 0.8” for G160M and 0.4” to 1.7” for G140L. While the FUV holographic gratings are shaped to correct for the spherical aberration of HST and are grooved to both correct the astigmatism along dispersion and provide nearly constant spectral resolution with wavelength, the astigmatism in the spatial (cross-dispersion) dimension is not corrected. This feature, along with the modified Rowland circle layout of the spectrograph, results in the width variations seen in Figure 4.

The small fluctuations seen in the fitted FWHM of the spatial profiles (Figure 4) are dominated not by the statistical noise, which is much smaller than the observed fluctuations, but rather the residual structure on the FUV detectors (e.g., grid wires, dead spots, small residual distortions left after geometric corrections by CalCOS, etc). These structures remain in the spectra because CalCOS does not currently apply a flat-fielding correction to the FUV data. Some of the fluctuations are also due to the presence of locations of spectral absorption lines.

Additional systematic trends can be seen in each of the profile plots in Figure 4. For Segment B spectra from both G130M and G160M, the spatial profiles are narrowest at the shortest wavelengths and become progressively broader at longer wavelengths. The opposite trend is seen for the Segment A G130M spectra, while all the profiles of the Segment A G160M spectra reach a common minimum near 1700 Å, before broadening again at the longest wavelengths. The profiles in the G140L Segment A spectra (we do not consider the G140L Segment B data here) reach a minimum width between 1400 Å and 1550 Å, and progressively broaden out to shorter and longer wavelengths.

The systematic variations in shapes and widths of the spatial profiles are the result of the modified Rowland Circle layout of the COS FUV channel. The aspheric FUV gratings
Figure 4: The spatial resolution of the COS FUV channel, as determined by Gaussian fits to the cross-dispersion profiles of white dwarf spectra. Results are shown for Segments A and B for G130M and G160M, and Segment A for G140L. The color of each curve corresponds to a different central wavelength setting. To obtain the plots, spectra were summed at 1 Å intervals along dispersion (∼ 100 pixels), then fit with a Gaussian along cross-dispersion (XD). The wiggles in the curves are primarily the result of flat field structure in the data (grid wires, dead spots, etc) which are not yet removed by CalCOS.
used in COS are shaped to correct for the spherical aberration of HST (in both dispersion and cross-dispersion). On the other hand, the holographic rulings on the gratings simultaneously disperse the light from the HST OTA and correct for the astigmatism inherent in the Rowland circle layout of the FUV channel (the MFWFEs are of course not corrected). However, the gratings only correct the astigmatism along dispersion. The remaining astigmatism in cross-dispersion is not corrected, resulting in the wide variations in profile FWHM (and hence spatial resolution) seen in Figure 4. The FWHM variations are similar to those seen in FUSE spectra, where a Rowland circle layout is also used.

The spatial and spectral focus vary independently in the COS FUV channel: for a given FUV grating, COS is designed to produce spectra having nearly constant spectral resolution with wavelength (corresponding to constant focus along dispersion). This is accomplished by placing the COS FUV detector along the Rowland circle and approximately matching the curvature of the detector to match that of the Rowland circle. This leaves the spatial (cross-dispersion) focus to vary along the detector. The surface of constant spatial focus intersects the detector at two points, one on Segment A and one on Segment B. By design, these focus points correspond to wavelengths of 1150 Å and 1450 Å for G130M and 1350 Å and 1750 Å for G160M, which is approximately what is observed in Figure 4. Elsewhere, the spatial profile is not at optimum focus. This is adequate for the majority of cases where the science targets are isolated point sources, and the flux is merely collapsed along cross-dispersion to produce the final x1d spectra. However, in a moderately crowded field (for example, several point sources located ~ 0.5″ apart from one another in cross-dispersion) the two sources may overlap spatially at one CENWAVE setting, but not another. Therefore, the extent of the spatial profile may require consideration for some observations. The overlap between spectra is even more important if one or more of the sources is extended.

References


COS Data Processing Improvements Based on HST SMOV Results

Thomas B. Ake\textsuperscript{1}, P. Hodge, A. Aloisi, R. I. Diaz, P. Ghavamian, C. Keyes, D. Massa, S. Niemi, C. Oliveira, R. A. Osten, C. R. Proffitt\textsuperscript{1}, B. York

Space Telescope Science Institute

D. Sahnow

Johns Hopkins University

S. Béland, S. Penton, E. Burgh, K. France

University of Colorado

J. McPhate

University of California, Berkeley

Abstract. After the Cosmic Origins Spectrograph (COS) was installed onboard the Hubble Space Telescope (HST) in May 2009, it underwent an extensive calibration and characterization check-out during the Servicing Mission Observatory Verification (SMOV) period. The results from this program were used to update reference files and make changes to CALCOS, the COS data processing software. Improvements to the standard data products are discussed. For the FUV channel, we have begun development of a flat-field correction. As an intermediate step, grid wire shadows are now ignored when combining FP-POS exposures. Pulse-height filtering has been activated to reduce background features. For the NUV channel, vignetting corrections are incorporated in the flat field file. For both channels, improvements have been made in the wavelength scales, flux calibration, and data quality flagging. Additional data are included in the FITS products to allow users to perform customized processing.

1. Introduction

This paper describes several areas in which CALCOS and the COS calibrated data products have been improved since SMOV due to our better understanding of the instrument, changes to the reference files, and updates to the CALCOS software. Some improvements were planned before launch and were merely awaiting the arrival of SMOV data. Others were in response to investigations of the on-orbit behavior of the instrument. We highlight several advancements that involved either changes to the CALCOS software or changes to the way the software operates in standard pipeline processing. The paper is organized as follows. In Section 2, we explain changes to the handling of the grid wire shadows in merged X1DSUM spectra for the FUV detector, and, in Section 3, the filtering of TIME-TAG events by pulse height amplitude. In Section 4 we discuss improvements to the wavelength calibration. Section 5 presents new capabilities for customized processing and off-line tools for working with COS data, and Section 6, other improvements and plans for future work. Topics that

\textsuperscript{1}Computer Sciences Corporation
are described in more detail elsewhere at this workshop are: FUV flat fields (Ake et al. 2010), spectral resolution (Ghavamian et al. 2010), wavelength calibration (Oliveira et al. 2010), changes in flux calibration (Osten et al. 2010), far ultraviolet performance (Osterman et al. 2010), target acquisition (Penton et al. 2010), detector performance (Sahnow et al. 2010), and dark count rate (Zheng et al. 2010).

A description of the CALCOS processing steps and reference files can be found in Chapter 3 of the COS Data Handbook (http://www.stsci.edu/hst/cos/documents/datahandbook). The latest information on CALCOS changes and dates of installations can be found at http://www.stsci.edu/hst/cos/pipeline/CALCOSReleaseNotes.

2. FUV QE Grid Wires

The FUV detector employs a wire grid above the microchannel plate (MCP) stacks to improve quantum efficiency. The wires cast shadows on the MCPs that appear as regularly spaced depressions in extracted spectra (X1D files), about 20% deep every 835 (segment A) or 841 (segment B) pixels. The affected locations can be identified in the CALCOS data products by a data quality flag, DQ=4.

Prior to launch, it was expected that the grid wire shadows could be removed with a flat field, but a methodology for flat fielding the FUV detector is still being developed. In the meantime, X1DSUM processing has been modified so that the grid wire regions are not included in the average (Figure 1). This is accomplished by including DQ=4 in the default value for the SDQFLAGS keyword, the parameter that controls which data will be ignored when FP-POS exposures are merged into the X1DSUM spectrum (Section 6 and Table 3). The thermal and geometric distortion corrections must be accurate to correct grid shadows with a flat field, but excluding the flagged regions is effective regardless of the shadow shape. The regions can also be made slightly wider than the actual shadows to accommodate misregistration. If data were included regardless of the DQ flags, the grid wire shadows would be reduced in depth, but would appear in more places. By excluding DQ=4 when forming the average, the signal-to-noise ratio is lower than in neighboring regions, but the grid wire shadows are effectively removed. For observations with two FP-POS positions, the contributions from only one exposure are present in the X1DSUM in the shadows; for four FP-POS steps, typically three exposures contribute to those locations. If no FP-POS stepping was performed, the X1DSUM spectrum would be a “sum” of only one exposure, and the spectrum will have gaps where the shadows lie. The data in the shadow regions are still available in the individual X1D spectra.

3. FUV Pulse Height Filtering

The FUV detector in TIME-TAG mode transmits a pulse height value between 0 and 31 for each event. The pulse height amplitudes (PHA) from photons have a different distribution from background events. Noise events typically have very large or very small PHAs. The PHACORR step in CALCOS flags TIME-TAG events that have PHAs outside specified limits. Such events are excluded when constructing spectra, thereby reducing the background. Prior to launch, CALCOS was set to accept all PHA values, so no filtering would occur.

During SMOV, FUV segment B was found to exhibit artifacts that resemble emission lines. These features have PHA=0, which is outside the normal range for true photon events. To eliminate these features and reduce detector noise, CALCOS now uses PHA values 4–30 for TIME-TAG exposures. See Figure 2.

When an observation is taken in ACCUM mode, the individual events and information about them (such as PHA) are not preserved, so ACCUM observations cannot be filtered by CALCOS. Thus the features in segment B may be present in the extracted spectra. Since
Figure 1: FUV G130M spectra of WD0320–539, offset by an arbitrary amount, taken at two FP-POS settings (X1D spectra) and coadded as an X1DSUM spectrum product. Only two lines, O I $\lambda$1302 and Si II $\lambda$1304, are real. The top two spectra are the individual exposures. Violet bars indicate grid wire regions. The blue spectrum is the average of the two (old X1DSUM). The shadows are reduced by a factor of 2, but occur in twice as many places. The red spectrum is the current CALCOS processing, where the grid wire regions in the individual spectra are ignored in the coaddition.

ACCUM mode exposures are taken only for high count rate objects, the artifacts will likely be overwhelmed by the target spectrum, but may affect a low flux area such as the bottom of an absorption line. The regions where these features are found are listed in the FUV bad pixel table and are flagged by CALCOS with DQ=4096.

Since PHA filtering can eliminate valid photons with pulse heights at the extremes of their distribution, the filtering process changes the flux calibration by a small amount, as can be seen in Fig. 2. For the time being, CALCOS is using the same sensitivity curves for TIME-TAG and ACCUM data, i.e. with no PHA filtering. CALCOS will be modified in the future to select flux calibrations based on the range of pulse heights that were used in the PHACORR step.

4. Wavelength Assignment

CALCOS uses PtNe lamp data, which are taken during TIME-TAG exposures or, for ACCUMs, just before or after the exposure, to convert X pixel values to wavelengths through WAVECORR. The lamp data are extracted as spectra and are cross-correlated with reference file spectral templates to compute offsets to the grating dispersion relations. The offsets are then used as a correction in assigning wavelengths to the pixels.

Prior to obtaining on-orbit data during SMOV, the reference files for the lamp templates and dispersion relations were based on the default FP-POS=3 setting. An exposure for any FP-POS setting was cross-correlated with the one template for that central wavelength. During SMOV, sufficient data were obtained at all four FP-POS settings to create template lamp spectra separately for all FP-POS settings for every central wavelength. This then allows better registration of an exposure’s wavecals with the wavelength reference frame.
COS Data Processing Improvements Based on HST SMOV Results

Figure 2: FUV G130M spectrum of Lin 358 showing pseudo-emission features in segment B. Upper panel: COUNTS image showing areas (blue) in the FUV bad pixel table where the artifacts occur. Lower panel: Extracted spectra with (red) and without (black) event pulse height filtering, illustrating how the filtering removes the contaminating features and lowers the extracted counts somewhat. The effects of the artifacts range from distorting some line profiles to inducing false emission lines. Note that CALCOS cannot eliminate these features in ACCUM exposures.

In addition, for the NUV detector prior to SMOV, the PtNe data from the three stripes were added together (pixel by pixel, i.e. without regard to wavelength) to form a single collapsed spectrum prior to cross-correlation. This was done because it was originally thought that the shifts of all the stripes would be nearly the same, and combining the spectra would help to fill in regions where one stripe had few lines with good signal. But in fact the shifts can differ by several pixels from one stripe to another, and these differences lead to incorrect wavelengths as well as misalignments between FP-POS spectra in X1DSUM files. The current version of CALCOS uses a separate lamp template for each FP-POS position (with known offsets from the FP-POS=3 dispersion relation), and for NUV, each stripe is processed separately.

The wavelength calibration has been modified to make it more robust as well as more accurate. The shift between the wavecal spectrum and template lamp spectrum is now found by minimizing $\chi^2$, rather than by maximizing the cross correlation. When a bright emission line is partially truncated by an edge of the NUV detector, the distorted line shape can bias the shift that is determined. CALCOS checks for such truncated lines and sets the wavecal data to zero in their vicinity. The improvement in wavelength accuracy for COS spectra is illustrated in Figure 3. See COS ISRs 2010–05 and –06 for details (Oliveira, 2010a and 2010b).

5. Customized Processing

CALCOS is not only used in the STScI pipeline, but runs in the STSDAS environment. This allows users to customize their own data processing. Also it is expected that GOs will develop their own tools to work either in concert with or in addition to CALCOS processing. Some updates have been made to CALCOS since SMOV to facilitate user processing.
The corrected time-tag event list (CORRTAG file) is the primary intermediate product most useful for customized processing. Two modifications to CALCOS have been made for this purpose so far. CALCOS can now accept CORRTAG files as input, not just raw files. This lets a user do some of the processing by setting some but not all of the calibration switch keywords to PERFORM and running CALCOS on the raw file. Then calibration switches can be reset in the CORRTAG header, the CORRTAG data can be modified (e.g. by applying one’s own flat field), and then CALCOS can be rerun with the modified CORRTAG file as input. The CORRTAG file itself has also been modified; a new wavelength column has been added to the table. After wavelength calibration is computed as part of preparing for spectral extraction, CALCOS assigns a wavelength to each event in the file. This will allow users to combine and extract spectra directly without converting the data into images. Table 1 lists the columns in the updated version of the CORRTAG file.

There are two new tasks (in addition to CALCOS) in the hstcos package in STSDAS in PyRAF. The xidcorr task extracts a 1-D spectrum (or spectra, if NUV), given a CORRTAG file as input. This task creates FLT and COUNTS files from the CORRTAG file and then extracts the spectrum from the FLT and COUNTS files, using the same code as in CALCOS. The splittag task takes a CORRTAG file and splits it by time into multiple files in the same format, i.e. CORRTAG. Then running CALCOS on the output files from splittag will generate a spectrum for each of those time intervals.

6. Other Improvements

The COS reference files are updated periodically as calibration studies are completed. Table 2 lists the reference files used by CALCOS, sorted by their installation dates, since SMOV and through July 2010. Many files created pre-SMOV and presently in use were not expected.
### Table 1: Corrected Event List (CORRTAG) Columns

<table>
<thead>
<tr>
<th>Parameter</th>
<th>Type</th>
<th>Description</th>
</tr>
</thead>
<tbody>
<tr>
<td>TIME</td>
<td>FLOAT</td>
<td>Time of event from beginning of exposure (seconds)</td>
</tr>
<tr>
<td>RAWX</td>
<td>INT</td>
<td>Uncorrected position along dispersion (pixel)</td>
</tr>
<tr>
<td>RAWY</td>
<td>INT</td>
<td>Uncorrected position cross dispersion (pixel)</td>
</tr>
<tr>
<td>XCORR†</td>
<td>FLOAT</td>
<td>Thermally and geometrically corrected RAWX (FUV)</td>
</tr>
<tr>
<td>YCORR†</td>
<td>FLOAT</td>
<td>Thermally and geometrically corrected RAWY (FUV)</td>
</tr>
<tr>
<td>XDOPP</td>
<td>FLOAT</td>
<td>XCORR corrected for HST Doppler shift at this time</td>
</tr>
<tr>
<td>XFULL</td>
<td>FLOAT</td>
<td>XDOPP shifted to wavecal frame</td>
</tr>
<tr>
<td>YFULL</td>
<td>FLOAT</td>
<td>YCORR shifted to wavecal frame</td>
</tr>
<tr>
<td>WAVELENGTH‡</td>
<td>FLOAT</td>
<td>XFULL wavelength assignment (Å)</td>
</tr>
<tr>
<td>EPSILON</td>
<td>FLOAT</td>
<td>Deadtime, flat field correction factor</td>
</tr>
<tr>
<td>DQ</td>
<td>INT</td>
<td>Data quality flag</td>
</tr>
<tr>
<td>PHA</td>
<td>BYTE</td>
<td>Pulse height amplitude (0–31)</td>
</tr>
</tbody>
</table>

†For NUV, XCORR, YCORR are the same as RAWX, RAWY since no correction is needed
‡Newly added column

to change. Some processes (bad time and burst corrections) have not been needed. Time dependent sensitivity corrections began in July 2010.

Other enhancements have been incorporated into CALCOS since SMOV.

- Corrections for NUV vignetting have been integrated into the flat field file (Ake, Burgh & Penton 2010); in the future vignetting may be handled with its own reference file.

- Data quality flags have been redefined to mark additional detector characteristics (Table 3). One particular change is that the DQ=4 flag, which had only been applied to FUV grid wire locations, is now used for the NUV to flag regions corrected for vignetting. While CALCOS should exclude wire shadows when creating an FUV XIDSUM spectrum since no flat field is available, for the NUV channel the vignette regions can be used in the sum since they are corrected by its flat field. The difference in processing is controlled by having the detector shadow bit in SDQFLAGS set to 1 for the FUV and 0 for NUV.

- The data quality extension of the FLT and COUNTS files are more accurately shifted by the wavecal offset and Doppler shift, which is important when combining FP-POS stepped spectra.

- The usefulness of the G140L grating below 1150 Å has been demonstrated during SMOV, requiring wavelength calibration changes since the PtNe lamps do not extend into that region.

- Several keywords have been added to preserve such information as the number of events flagged as bad or exposure time lost due to FUV bursts, and keywords have been added for spectral-location information.

Further improvements are planned for the future.

- The acceptable range of pulse heights will need to be changed to track the changes in FUV gain due to exposure to bright sources. Currently a single pulse height range is applied for an entire FUV detector segment; however, the gain varies depending on position on the detector, so the PHA filtering should also vary with position. Filtering
<table>
<thead>
<tr>
<th>File Type</th>
<th>FUV File Name</th>
<th>NUV File Name</th>
<th>Description</th>
</tr>
</thead>
<tbody>
<tr>
<td>Pre-SMOV</td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>FLATFILE</td>
<td>n9u20182l_flat.fits</td>
<td></td>
<td>Flat field image (unused)</td>
</tr>
<tr>
<td>GEOFILE</td>
<td>s7g1700cl_geo.fits</td>
<td></td>
<td>FUV geometric correction</td>
</tr>
<tr>
<td>BADTTAB</td>
<td>s7o1739kl_badt.fits</td>
<td></td>
<td>FUV bad times (unused)</td>
</tr>
<tr>
<td>BPIXTAB</td>
<td>s7g1700pl_bpix.fits</td>
<td>s7g1700ql_bpix.fits</td>
<td>Bad pixel locations</td>
</tr>
<tr>
<td>BRFTTAB</td>
<td>s7g1700cl_brf.fits</td>
<td></td>
<td>FUV reference frame</td>
</tr>
<tr>
<td>BRSTTAB</td>
<td>s7g1700ql_burst.fits</td>
<td></td>
<td>FUV burst rejection (unused)</td>
</tr>
<tr>
<td>DEADTAB</td>
<td>s7g1700ql_dead.fits</td>
<td>s7g1700ql_dead.fits</td>
<td>Deadtime correction</td>
</tr>
<tr>
<td>TDSTAB</td>
<td>t2314312l_tds.fits</td>
<td>t2314314l_tds.fits</td>
<td>Time dependent sensitivity (unused)</td>
</tr>
<tr>
<td>WCPTAB</td>
<td></td>
<td>t2314313l_wcp.fits</td>
<td>Wavelength calibration parameters</td>
</tr>
</tbody>
</table>

**Sep 11, 2009**
FLATFILE | t9b18111l_flat.fits | Flat field image
XTRACTAB | t9b1608jl_1dx.fits | Spectrum extraction regions
WCPTAB   | t9b18112l_wcp.fits | Wavelength calibration parameters

**Sep 14, 2009**
DISPTAB  | t9e1307kl_disp.fits | t9e1307ll Disp. fits | Wavelength dispersion relations

**Sep 17, 2009**
FLUXTAB  | taj18183l.phot.fits | Sensitivity curves

**Oct 19, 2009**
FLUXTAB  | taj18183l.phot.fits | Sensitivity curves

**Jan 29, 2010**
DISPTAB  | u1t1616ml_disp.fits | u1t1616pl_disp.fits | Wavelength dispersion relations
LAMPTAB  | u1t1616nl_lamp.fits | u1t1616ol_lamp.fits | PtNe lamp templates
WCPTAB   | u1t1616ql_wcp.fits | u1t1616ol_wcp.fits | Wavelength calibration parameters
PHATAB   | u1t1616llpha.fits  | Pulse height screening

**Mar 10, 2010**
BPIXTAB | u381724gl_bpix.fits | Bad pixel locations

**Apr 14, 2010**
XTRACTAB | u4d1930sl_1dx.fits | u4d1930tl_1dx.fits | Spectrum extraction regions

**Apr 29, 2010**
FLUXTAB  | u4t18348l_phot.fits | Sensitivity curves

**Jun 28, 2010**
LAMPTAB  | u6s1320vl_lamp.fits | PtNe lamp templates
DISPTAB  | u6s1320ql_disp.fits | Wavelength dispersion relations
DISPTAB  | u6s1320plDisp.fits | Wavelength dispersion relations

**Jul 13, 2010**
TDSTAB   | u7d20377l_tds.fits | u7d20378l_tds.fits | Time dependent sensitivity
Table 3: Original and New COS Data Quality (DQ) Flag Assignments

<table>
<thead>
<tr>
<th>Bit</th>
<th>Decimal</th>
<th>Old Description</th>
<th>New Description</th>
<th>SDQFLAGS†</th>
<th>FUV</th>
<th>NUV</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>1</td>
<td>Reed-Solomon error</td>
<td>Reed-Solomon error</td>
<td></td>
<td>0</td>
<td>0</td>
</tr>
<tr>
<td>2</td>
<td>2</td>
<td>Brush mark</td>
<td>Detector flaw</td>
<td></td>
<td>0</td>
<td>0</td>
</tr>
<tr>
<td>3</td>
<td>4</td>
<td>Grid wire</td>
<td>Detector shadow</td>
<td></td>
<td>1</td>
<td>0</td>
</tr>
<tr>
<td>4</td>
<td>8</td>
<td>Spectrum near detector edge</td>
<td>Spectrum near detector edge</td>
<td></td>
<td>1</td>
<td>1</td>
</tr>
<tr>
<td>5</td>
<td>16</td>
<td>Dead spot</td>
<td>Dead spot</td>
<td></td>
<td>1</td>
<td>1</td>
</tr>
<tr>
<td>6</td>
<td>32</td>
<td>Hot spot</td>
<td>Hot spot</td>
<td></td>
<td>1</td>
<td>1</td>
</tr>
<tr>
<td>7</td>
<td>64</td>
<td>Burst</td>
<td>Burst</td>
<td></td>
<td>0</td>
<td>0</td>
</tr>
<tr>
<td>8</td>
<td>128</td>
<td>Out of bounds</td>
<td>Pixel out of bounds</td>
<td></td>
<td>1</td>
<td>1</td>
</tr>
<tr>
<td>9</td>
<td>256</td>
<td>Fill data</td>
<td>Fill data</td>
<td></td>
<td>0</td>
<td>0</td>
</tr>
<tr>
<td>10</td>
<td>512</td>
<td>PHA too low</td>
<td>PHA out of bounds</td>
<td></td>
<td>0</td>
<td>0</td>
</tr>
<tr>
<td>11</td>
<td>1024</td>
<td>PHA too high</td>
<td>Unused</td>
<td></td>
<td>0</td>
<td>0</td>
</tr>
<tr>
<td>12</td>
<td>2048</td>
<td>Bad time interval</td>
<td>Bad time interval</td>
<td></td>
<td>0</td>
<td>0</td>
</tr>
<tr>
<td>13</td>
<td>4096</td>
<td>Wavelength below 900 Å</td>
<td>Background feature</td>
<td></td>
<td>0</td>
<td>0</td>
</tr>
<tr>
<td>14</td>
<td>8192</td>
<td>Wrinkle (segment A)</td>
<td>Low gain area</td>
<td></td>
<td>0</td>
<td>0</td>
</tr>
<tr>
<td>15</td>
<td>16384</td>
<td>S distortion (segment A)</td>
<td>Unused</td>
<td></td>
<td>0</td>
<td>0</td>
</tr>
</tbody>
</table>

†An SDQFLAGS bit=1 excludes data with the corresponding DQ flag from the X1DSUM spectrum

by pulse height affects the flat field and flux calibration; CALCOS will need to check (via a header keyword) that the reference files are consistent with the PHA filtering currently being applied to the data.

• The FUV flat field calibration will likely be modified. Currently CALCOS divides by a 2-D image, as is done for most instruments. There is evidence that a 1-D flat for the relevant spectral extraction region works better, however, and if further investigation shows this to be the case, CALCOS will be modified to use this format for the flat field.

• CALCOS may be modified to check the high voltage, and calibration would be limited if the high voltage is out of bounds.

• The algorithm for detecting FUV bursts is based on what was done for FUSE (Far Ultraviolet Spectroscopic Explorer). If bursts are actually present in COS data, the burst-detection algorithm should be improved and customized for COS.

7. Summary

CALCOS and the COS data products continue to improve with time. Prior to SMOV, a prioritized list was created for CALCOS enhancements and bug fixes, and plans were in place to update reference files derived from on-orbit observations. A great deal was learned about the instrument during SMOV, and valuable calibration data were obtained during this period. Several new reference files were prepared based on SMOV data, including lamp template spectra, dispersion relations, spectrum extraction regions, and time-dependent flux calibration. Besides the planned changes to the software and data products, some unexpected instrument characteristics, such as the low-gain artifacts in the FUV detector and vignetting of the NUV channel, were corrected. Other updates, like FUV flat fielding and calibrations below 1150 Å, are still on the horizon.
References

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Osterman, S. et al. 2010, this volume
Penton, S. et al. 2010, this volume
Sahnow, D. et al. 2010, this volume
Zheng, W. et al. 2010, this volume
Trend of Dark Rates of the COS and STIS NUV MAMA Detectors

W. Zheng\textsuperscript{1,2}, C. R. Proffitt\textsuperscript{2,3}, D. Sahnow\textsuperscript{1,2}, T. B. Ake\textsuperscript{2,3}, C. Keyes\textsuperscript{2}, P. Goudfrooij\textsuperscript{2}, P. Hodge\textsuperscript{2}, C. Oliveira\textsuperscript{2}, A. Bostroem\textsuperscript{2}, C. Long\textsuperscript{2}, and A. Aloisi\textsuperscript{2}

Abstract. The dark rate of the STIS NUV MAMA detector was about an order of magnitude higher after SM4 repair than anticipated, with an initial rate of $0.01$ count sec\textsuperscript{-1}pixel\textsuperscript{-1}. Measurements over the past year show a dual-component exponential decline with e-folding timescales of $\approx 30$ and 300 days. The most recent measurements show a rate of $2.8 \times 10^{-3}$ count sec\textsuperscript{-1}pixel\textsuperscript{-1}. The dark rate of the COS NUV detector started at a very low value of $6 \times 10^{-5}$ count sec\textsuperscript{-1}pixel\textsuperscript{-1}, and has displayed a steady increase, approaching the ground-tested level of $3.7 \times 10^{-4}$ count sec\textsuperscript{-1}pixel\textsuperscript{-1}. Still, the rate of COS NUV detector is considerably lower than that of its STIS counterpart. The rates for both detectors are sensitive to detector and tube temperatures, and the rate fluctuations can be fit with an empirical model.

1. Introduction

During the Hubble servicing mission in 2009, the new COS instrument was installed and the STIS revived. As part of the Cycle 17 calibration programs, we have collected dark images regularly and have monitored the dark rates of the FUV and NUV detectors. In general, the dark rates of the FUV detectors, both COS and STIS, are steady and consistent with pre-SM4 estimates. The NUV detectors, however, display significant changes. We model the measurements of dark rates with existing parameters that were derived for STIS (Proffitt, et al. 2010) over its many year operation. Note that the COS NUV detector is a spare device for its STIS counterpart, considered as nearly identical. However, their dark rates follow opposite trends.

2. Observations

The STIS dark images were obtained in ACCUM mode. Since all the data points are used, the dark rates include contributions from cosmic rays and other effects. The COS dark images, on the other hand, were obtained in TIME-TAG mode, thus offering time resolution. During the SAA (South Atlantic Anomaly) passages, the dark rate becomes considerably higher than average. We used the time variation of dark rates to identify and exclude data during SAA passages. When the average dark rate rises to a value more than twice the normal dark rate or declines from such a high level, a flag was set for SAA. Further iterations over rates at different intervals yielded an improved identification of the SAA passage, and the data during that period were excluded from our study.

\textsuperscript{1}The Johns Hopkins University
\textsuperscript{2}Space Telescope Science Institute
\textsuperscript{3}Computer Sciences Corporation
Figure 1: STIS NUV dark rate vs. modified Julian date. Two moderate surges, at the beginning of STIS operation and at modified Julian date (MJD) ≈ 51500, took place when the detector was turned on after a long period when the detectors were off. The surge during SM4 (MJD ≈ 55000) is unprecedented in terms of its large scale and slow recovery.

3. STIS NUV

The NUV MAMA dark rate is believed to be dominated by a phosphorescent glow from the detector window. Impurities in the window contain meta-stable states that are populated by cosmic-ray impacts. At lower temperatures, the reduced de-excitation rates result in a build-up of populated meta-stable states, leading to a large, but temporary, increase in the dark rate once the detector is warmed again. Previous experience with STIS had suggested that this increase in dark rate would dissipate within a few weeks. However, after its post-SM4 recovery, the NUV MAMA showed a dark rate about a factor of 10 higher than the typical pre-failure rate of about $1 \times 10^{-3}$ count sec$^{-1}$pixel$^{-1}$ (Figure 1), and it has since come down very slowly. Even now, more than a year after the STIS repair, the mean dark rate is approximately 2.5 times its typical pre-failure values.

The temperature of the STIS MAMA detector never reaches equilibrium, as the MAMA high-voltage power supplies are cycled frequently. As a result, the dark rate fluctuates considerably. Our earlier model of the NUV MAMA window glow as a single population of meta-stable states with a single band gap and only one time-scale for de-excitation does not work for the post-SM4 data. We model the dark rate after SM4 with the following parameters: an exponential function of instrument temperature and a dual-component exponential decay over time. The best fit is (in units of count sec$^{-1}$pixel$^{-1}$):

$$\text{Dark rate} = 2.9 \times 10^{24} \exp\left(-18646/(T_c + 273.16)\right) \ast \exp((t - t_0)/29) + 4.9 \times 10^{15} \exp\left(-12515/(T_c + 273.16)\right) \ast \exp((t - t_0)/398),$$

where $T_c$ is OM2TUBET (temperature of the STIS MAMA tube) in Celsius, $t$ is modified Julian date, and $t_0 = 55050$. The fitting result is shown in Figure 2. The rapid term resembles some previous surges in dark rate prior to 2004, but there is another population
Figure 2: STIS NUV dark rate since SM4, plotted in filled circles. The upper curve is a model fit of temperature parameter OM2TUBET and dual decay components. The lower curve is an extrapolation of the model fit from data prior to SM4.

that evolves very slowly over time. Fitting with other temperature parameters yields similar results, as these telemetry data follow similar trends.

4. COS NUV

The dark rate of the COS NUV detector (Dixon et al. 2010) started from a very low level, at $\approx 6 \times 10^{-6}$ count sec$^{-1}$pixel$^{-1}$, which is considerably lower than the pre-launch prediction. Since SM4, the rate has increased approximately linearly over time, as shown in Figure 3. The fluctuations appear to be correlated with the detector temperature. We use a single-temperature-component model to fit the data after MJD = 55020 (when the detector conditions became stable), and the results suggest that

Dark rate = $1.1 \times 10^9 \exp(-8791/(T_c + 273.16)) \ast ((t - t_0)/163 + 1.5)$,

where $T_c$ is LNTUBET, temperature of the COS MAMA tube, and $t_0 = 55100$.

5. Conclusion

Both the NUV MAMA detectors of COS and STIS show considerable but slow changes in dark rate. The current measurements are useful in updating the COS and STIS ETC (Exposure Time Calculator) parameters. The COS NUV MAMA detector still exhibits a considerably lower dark rate than the STIS detector. More data are needed in the next year to confirm at which level their values will stabilize, and to understand the cause for such changes.
Figure 3: COS NUV dark rate vs. MJD, plotted in filled circles. The curve is a model fit with tube temperature parameter LNTUBET.

References

Sahnow, D., et al. 2010, this proceedings
Monitoring of the Wavelength Calibration Lamps for the Hubble Space Telescope

Ilaria Pascucci\textsuperscript{1,2}, Charles R. Proffitt\textsuperscript{1,3}, Parviz Ghavamian\textsuperscript{1}, David Sahnow\textsuperscript{1,2}, Cristina Oliveira\textsuperscript{1}, Alessandra Aloisi\textsuperscript{1}, Charles Keyes\textsuperscript{1}, Steven V. Penton\textsuperscript{4}

\textit{Space Telescope Science Institute, Baltimore, MD, USA 21218}
\textit{Johns Hopkins University, Department of Physics & Astronomy, Baltimore, MD, USA 21218}
\textit{Science Programs, Computer Sciences Corporation, Baltimore, MD, USA 21218}
\textit{University of Colorado, Boulder, CO, USA}

Abstract. The Space Telescope Imaging Spectrograph (STIS) and the Cosmic Origins Spectrograph (COS) are the two optical-UV spectrographs on board the Hubble Space Telescope. To determine the wavelength scale for individual science observations, internal arc lamp spectra accompany most observations of external targets. Here we present a detailed analysis of the changes in the COS and STIS internal lamp fluxes and spectra over time, and also compare our results to pre-launch ground testing, and to laboratory accelerated aging testing of similar lamps. We find that the STIS LINE lamp has faded by a factor of $\sim 15$ in the very short FUV wavelengths (1150-1200\AA) over the 13-year period on which STIS was in space, a much steeper fading than predicted from accelerated aging tests in the laboratory. We also find that all STIS lamps have faded during the period in which the spectrograph was not operational (2004-2009) thus pointing to on-orbit conditions as an additional and important cause of lamp fading. We report that the COS P1 lamp output appears to decline with usage with a similar slope as the LINE and HITM1 lamps on STIS. Finally, we recommend switching from the LINE to the HITM2 lamp for a more efficient wavelength calibration of the STIS settings covering the very short FUV wavelengths.

1. INTRODUCTION

Ultraviolet (UV) and visible spectroscopy are prime tools to unravel the properties of stars, planets, galaxies, and of the interstellar and intergalactic matter. The Hubble Space Telescope (HST) has two instruments performing sensitive UV-optical spectroscopy.

The Space Telescope Imaging Spectrograph (STIS, Kimble et al. 1998) is a versatile imaging UV-visible spectrograph. STIS operated on-orbit from Feb 1997 until a malfunction in August 2004, but resumed operations in May 2009 after a successful repair during Hubble Servicing Mission 4 (SM4). During SM4, a new UV instrument, the Cosmic Origins Spectrograph (COS, Green et al. 2003) was also installed. For wavelength calibration, STIS contains three hollow cathode Pt/Cr-Ne arc lamps, while COS contains two similar Pt/Ne lamps. These lamps provide a continuous distribution of suitable emission lines for wavelength calibration from UV to optical wavelengths, as needed for STIS (with Pt emission lines dominating the spectra out to 3200Å, Ne lines beyond 5000Å, and Cr lines in between). To allow the wavelength scale for individual science observations to be corrected for non-repeatability and drifts in the STIS and COS optical alignment, most observations
of external targets are accompanied by one or more wavelength calibration spectra. This provides a large data set for monitoring on-orbit lamp performance over time.

There are two main mechanisms that can cause a reduction in the spectral output of these arc lamps (lamp fading) with time. The first one is deposit of metals sputtered from the interior of the cathode into the mica disks and other lamp surfaces. In the process metals trap neon atoms, gradually decreasing the gas pressure in the lamp. This, in turn, requires an increasing turn-on voltage to operate the lamps, an effect observed in aging tests of two Pt/Cr-Ne lamps similar to those installed on STIS (Kerber et al. 2004) and of a Pt/Ne lamp (in particular an air-aged lamp) similar to those in COS (Nave et al. 2008). Unfortunately, both STIS and COS were not designed to report lamp turn-on voltage as part of the telemetry. Therefore, we cannot use this diagnosis to test the lamp aging. The second mechanism of lamp fading has to do with the aging of the barium getter: a lower efficiency in removing impurities such as H, N, O, and C results in decreasing turn-on operating voltage which suppresses the lamp line spectrum. The most obvious sign of the getter aging would be the appearance of strong H lines in the discharge.

An additional factor that could mimic lamp aging has been recently proposed for COS lamps (and might also apply to the STIS lamps). The beam from the lamps grows larger with time, resulting in less and less light making it through the aperture.

In the following, we use total count rates and spectra to measure how the spectral output of the STIS and COS lamps has changed with time and as a function of wavelength. With STIS we have a time baseline of 13 years (Sect. 2.), while with COS we have about 6 months (Sect. 3.). We compare our results with pre-launch ground testing, and with laboratory testing of similar lamps and discuss what mechanisms could be responsible for the lamps fading. We also provide recommendations for the exposure time/lamp combinations to perform adequate wavelength registration with STIS.

2. THE STIS WAVELENGTH CALIBRATION LAMPS

On STIS there are three Pt/Cr-Ne lamps for wavelength calibration. These lamps are designed as the LINE, HITM1, and HITM2 lamps. The LINE lamp is located on the Insertion Mechanism platform. When the lamp is used the Calibration Insert Mechanism (CIM) is inserted into the light path and all external light is blocked. The HITM1 and HITM2 lamps are located on the Hole In The Mirror platform and can be used both for wavelength calibration and target acquisition (though the HITM2 lamp is considered the spare lamp and has not been used for target acquisition). When the HITM lamps are used, light from the external sky still falls on the detector unless the STIS external shutter is closed. The different locations and optics of the lamps have been chosen to allow for a range of fluxes on the detector and thus optimize lamp exposure times throughout the large wavelength range covered by the STIS spectrograph with its various grating apertures. The LINE lamp, being the brightest of the 3 Pt/Cr-Ne lamps on the detector, is being used for the least sensitive modes on STIS, especially the FUV echelle modes. The optics of the HITM1 and HITM2 lamps are such that their lower flux on the detector is more suited for the wavelength calibration of more sensitive modes on STIS, such as the optical low-resolution modes.

The flux of one lamp relative to another is also varying as a function of wavelength. This is due to intrinsic lamp-to-lamp variations in the FUV, which are a strong function of the transmissivity of the MgF2 lamp window, and to the optical path of the lamp. In particular, the HITM lamps illuminate a UV beamsplitter whose transmissivity varies from the UV to the visible. Hence, the HITM2 lamp which illuminates the beamsplitter in reflection is much brighter than the HITM1 lamp in the visible region of the spectrum due to the increased transmissivity of the reflected path at visible wavelengths. Pre-launch
ratios of the lamp brightness at a 10 mA operating current are reported in the STIS System Handbook and User’s Guide and compared to current flux ratios in Sect. 2.2.

Between installation in February 1997 and the end of 2009 the LINE lamp has been used for a total of 145.8 h, the HITM1 lamp for 97.2 h, and the HITM2 lamp for 3.3 h. The expected lifetime of Pt/Cr-Ne lamps is 15,000 mA h or 1,500 h at 10 mA, which is the typical lamp operating current on STIS. Therefore, even the LINE lamp has been used just 1/10 of its expected lifetime. Aging tests simulating the working conditions of the STIS lamps do not report major changes in appearance of the emerging spectra and in the intensity of the lines over the lamp lifetimes (Kerber et al. 2004). In contrast, we show here that the lamps output has changed with time and that the change is wavelength-dependent (Sect. 2.1.). These changes are much larger than the sensitivity degradation of the STIS detectors which is estimated to be <2% per year throughout the full UV-optical wavelength range (Sect. 2.2.). In addition to lamp outputs, we present current lamp flux ratios and total count rates for individual lamps (Sect. 2.2.). These changes in lamp flux with time and usage could be due to the aging of the lamps and/or to changes in the optics.

2.1. Lamps Fading as a Function of Wavelength

Analysis of spectra from the STIS Pt/Cr/Ne lamps obtained before and after SM4 hinted at a fading of the lamps especially at the very shortest FUV wavelengths. Following up on this, we have extracted (using the calstis pipeline) and analyzed all lamp spectra obtained since 1997 in the G140L/STIS grating. The main results from this analysis are shown in Figs. 1, 2, and 3.

Figs. 1 compares total counts per second in the G140L grating over the full wavelength range covered by the grating (1150-1730 Å) for each of the three lamps. We have scaled the counts to the same slit width of 52×0.1″ for all three lamps and restricted the analysis to exposures obtained with the 10 mA current setting. The decline in the LINE flux can be fitted by a second order polynomial and appears to be steeper than that of the HITM1 flux. The LINE flux is still a factor of ~4 higher than the HITM1 flux. Fig. 2 shows that the fading is wavelength-dependent and that the LINE lamp flux has decreased exponentially with time at the very shortest FUV wavelengths (1150-1200 Å). The LINE flux at these wavelengths is currently less than the HITM1 flux, see also Sect. 2.2. and Fig. 4. We also show that the exponential decay in the LINE lamp flux is steeper than the exponential decay suffered by the HITM1 lamp flux at the same short FUV wavelengths (see Fig. 3). Fluxes from the two lamps appear to decay similarly up to an usage of ~500 mA -h, at which point the LINE lamp flux starts to decrease faster. The same figure and Fig. 2 show that their Post-SM4 fluxes are well below the fluxes just before the STIS malfunction in August 2004. The fading of the LINE lamp is again more dramatic post-SM4, a factor of 4 versus a factor of 1.8 for the HITM1 lamp.

2.2. Current Lamp Flux Ratios

We have recently used 4 internal HST orbits to acquire lamp exposures with each of the LINE, HITM1, and HITM2 lamps in the low-resolution STIS gratings G140L, G230LB, G430L, and G750L. In this way, we are able to get the full spectrum of the lamps brightness from the FUV to the optical wavelengths. We set the lamp current to 10 mA and used the same narrow (52×0.1″) slit aperture when possible. Exposure times varied between 5 and 400 seconds depending on lamp and grating (for more details see the calibration program PID 12079).

Because the STIS low-resolution gratings cover a large wavelength region (between several hundreds to thousands Å), we analyzed both total counts within each grating as well as total counts within three narrower wavelength regions. We have chosen these regions to correspond to the following STIS medium-resolution gratings to sample the short, middle, and long wavelength portions of each low-resolution grating: G140M-1173, G140M-
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1470, G140M-1741, G230MB-1713, G230MB-2146, G230MB-3115, G430M-3165, G430M-4451, G430M-5471, G750M-5734, G750M-8311, and G750M-10363. Flux ratios for the LINE over the HITM1 lamps are plotted in Fig. 4 and summarized in Table 1 together with line flux ratios from the other lamps.

We find that the LINE flux remains higher by about an order of magnitude at all wavelengths except at the shortest FUV wavelengths. At these short wavelengths, corresponding to the G140M-1173 setting, the HITM1 flux is currently 1.5 times higher than the LINE flux while the HITM2 flux is 3 times higher. The HITM2 flux is about 1.5 times higher than the HITM1 flux at all wavelengths.

We have also extracted and analyzed in the same way deep lamp exposures obtained in 1997, right after the installation of STIS (calibration program PID 7722). Fig. 4 compares the current LINE/HITM1 ratios with the pre-launch and the 1997 ratios. These ratios are very similar over the full wavelength range covered by STIS except at the shortest FUV wavelengths. The LINE/HITM2 ratios show a similar behaviour (see Table 1).

To visualize the decrease in the LINE flux in absolute terms we have plotted in Fig. 5 the 1150-1200˚A portion of the G140L spectra obtained in our 2010 lamp monitoring program and those obtained in 1997. The LINE flux between 1997 and 2010 has decreased by a factor of ~15 in absolute terms in the short FUV wavelengths, while the HITM1 and HITM2 have decreased by a factor ~3. We note that the STIS sensitivity degradation has been measured to be typically less than 2% per year throughout the full UV-optical wavelength range1, implying a flux loss of only ~25% between 1997 and 2010 as due to sensitivity degradation. The much steeper fading of the lamps reported here has to do with either lamp aging and/or changes in the optics.

2.3. Impact on Wavelength Calibration Exposures

Based on the analysis above, we expect that the STIS grating/wavelength settings most affected by the LINE lamp fading are those covering only the shortest FUV wavelengths. Note that the LINE/HITM1 flux ratio is very close to the pre-launch and 1997 ratios already for the middle wavelength region (1443-1497˚A) of the G140L grating, so any lamp fading should affect more strongly the lamp output at wavelengths ≤1400˚A. Among all available STIS settings the G140M-1173˚A is definitively the one that is most impacted given that it covers only from 1145˚A out to 1201˚A. To show the impact of the LINE fading on the detectability of Pt, Cr, and Ne lines, we present in Fig. 6 two typical lamp exposures obtained for the wavelength calibration of G140M-1173˚A spectra (the 1997 lamp spectrum is from PID 7667 while the 2009 spectrum is from PID 11860). Lamp exposure times were set in 1997 to be 46 seconds and are still kept to this value. Fig. 6 clearly shows that the number of detected lines in the 2009 spectrum has dramatically decreased and only a few bright lines, all at wavelengths ≥1170˚A, are still detected with current lamp exposure times. Our 2010 lamp monitoring program (PID 12079, see Sect. 2.2.) demonstrates that the HITM2 lamp is more than 3 times brighter than the LINE lamp at these short wavelengths, while the HITM1 lamp is just 1.4 times brighter (see Table 1). Thus, the most efficient way to recover lines at the shortest FUV wavelengths seems to be a switch from the LINE to the HITM2 lamp for the wavelength calibration of the STIS settings covering wavelengths below 1400˚A.

3. THE COS WAVELENGTH CALIBRATION LAMPS

COS has two Pt/Ne hollow-cathode wavelength calibration lamps suitable for determining the wavelength scale of all available spectroscopic modes, except for those covering only

1http://www.stsci.edu/hst/stis/performance/sensitivity/
wavelengths below $\sim$1180Å$^2$. Both lamps have been tested during SMOV and found to perform as expected (PID 11467, PI D. Sahnow). Only one of them (P1) is currently used for wavelength calibration and is expected to be used until it fails, at which point operations will be switched to the other lamp. The second lamp (P2) is used for target acquisition.

Light from the Pt-Ne lamp reaches the spectrograph through the WCA (wavelength calibration aperture), while science exposures use either the PSA (Primary Science Aperture) or the BOA (Bright Object Aperture). Because it was found in thermal vacuum testing 2003 that the COS optical selection mechanisms cause internal and external spectra to drift in the spectral direction, most COS lamp exposures are acquired contemporaneously to science observations to guarantee an accurate wavelength registration. Lamp exposures for TIME-TAG/PSA data are taken in the so-called TAG-FLASH mode with flashing of the P1 lamp ranging between 5 to 30 seconds in duration. As mentioned above the P2 lamp is used during target acquisition to provide a reference point between the location of the WCA and the detector pixel coordinates$^3$. Hence, the lamps on COS will be used much more extensively than any other lamp on previous spectrographs, see also Penton et al. (2010) for estimates of the lamps’ usage.

3.1. Preliminary Trends on Lamp Fading

We have extracted and analyzed all TIME-TAG exposures obtained in the G140L-1105 setting since the installation of COS, a total of 80 exposures obtained with the lamp set to medium current (10 mA). This setting covers the 1120-2246Å FUV region and thus overlaps with the 1150-2000Å wavelength region where STIS lamps have shown the steepest fading. For each of the corrected TIME-TAG images produced by the calcos pipeline we performed a sum in dispersion direction and divided the total counts by the lamp exposure time to obtain count rates. The signal-to-noise of the COS lamp spectra is not as high as that in the STIS lamp spectra due to the much shorter exposure times used in the COS TIME-TAG mode. This results in more scattered count rates when summing the counts just in the 1150-2000Å wavelength region. For this reason, we have decided to report count rates over the full wavelength range covered by the G140L-1105 setting. The lamp usage is a running cumulative total of the lamp flash times (for P1), converted into hours, and then multiplied by 10mA.

Another issue that needs to be taken into account for COS is the sensitivity degradation of the FUV detector. Monitoring has shown that the FUV sensitivity is degrading much faster than the STIS FUV sensitivity and that the degradation is wavelength-dependent (see the STScI Analysis Newsletter, May 2010$^4$). In the G140L-1105 setting the sensitivity degradation is estimated to be $\sim$-9% per year (ISR by Osten et al. in prep.), in contrast to $\sim$2% sensitivity loss per year for STIS. Therefore, in comparing the COS with the STIS lamp fading we have removed the trend in the COS detector sensitivity loss. Fig. 7 shows exactly this comparison with the output of the STIS and COS lamps versus usage in the G140L setting. We note that the larger scatter in the COS count rates could be due to orbits partially impacted by the elevated background during passage through the South Atlantic Anomaly. All three lamps, two on STIS and one on COS, show the same decline in output with usage in the first $\sim$150 mA-h. As noted in Sect. 2.1., both STIS lamps faded at a similar rate up to an usage of $\sim$500 mA-h after which the LINE lamp started to fade faster.

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$^2$The G140L-B settings are calibrated using the G140L-A portion of the spectra

$^3$http://www.stsc.edu/hst/cos/documents/handbooks/current/cos_cover.html

$^4$http://www.stsci.edu/hst/cos/documents/newsletters/stan0510.html
Accelerated aging tests have been performed in the laboratory on three Pt-Ne lamps similar to those installed in COS. All lamps lasted between 800 and 1000 h and showed that the radiance of the lamps changed by less than 20% during the first 800 h of use (Nave et al. 2008). Fig. 7 uses spectra taken on-board and shows a similar change but just in the first 25 h of use for all three lamps. We measure a reduction of $\sim 80\%$ for the STIS/LINE lamp output in less than 150 h of use in these very short FUV wavelengths (Fig. 3).

These results demonstrate that monitoring of the lamps output will be important to ensure a proper wavelength calibration with the STIS and COS spectrographs.

4. CONCLUSIONS

We have measured the output of the wavelength calibration lamps on the STIS and COS HST spectrographs versus time. Our main findings can be summarized as follows:

- The STIS LINE lamp, which was the brightest at the beginning of operation in 1997, has faded dramatically at the shortest FUV wavelengths (1150-1200Å) and is now fainter than the other two STIS lamps
- The STIS LINE lamp output remains higher than the HITM1 and HITM2 lamp outputs by factors of $\sim 5-10$ at all wavelengths $\geq 1300Å$
- The fading of the STIS LINE lamp between 1997 and 2010 in absolute terms is factor of $\sim 15$ at the very short FUV wavelengths, much greater than what expected from accelerated aging tests in the laboratory
- We have evidence that all STIS lamps fade even when they are not used indicating that on-orbit factors are impacting the lifetime of the lamps
- The 6-month monitoring of the COS P1 lamp shows that the lamp output fades similarly to the STIS lamps as a function of lamp usage

References

Green, J. C., Wilkinson, E., Morse, J. A. 2003, SPIE, 5164, 17
Kerber, F. et al. 2004, SPIE, 5488, 679
Penton, S. V. et al. 2010, SPIE, 7011, 70113
Monitoring of the Wavelength Calibration Lamps for the Hubble Space Telescope

Figure 1: Total counts versus time for the three wavelength calibration lamps on STIS. The LINE lamp flux (black) has decreased faster than the HITM1 flux (red) but remains brighter by a factor of about 4 when total counts are computed on the full wavelength range covered by the G140L setting (1150-1730 Angstrom). The LINE lamp flux data (from 1997 until now) are best fitted with a second order polynomial (blue dashed line). The HITM2 flux (green) was measured twice: one, right after the installation of STIS in 1997, the second time in 2010 with our lamp monitoring program (PID 12079). Note that numbers have not been scaled to take into account the detector sensitivity loss since this is small for STIS (<2% per year, Sect. 2.2.).

Figure 2: Like Fig. 1 but for the short wavelength portion of the G140L grating (1150-1200Å). The LINE lamp flux (black) has decreased exponentially with time (blue dashed curve, fit to all LINE lamp data from 1997 until now) and it is currently lower than HITM1 lamp flux (red) in this wavelength region. Our 2010 lamp monitoring program (PID 12079) shows that the HITM2 lamp has currently the largest flux among the three wavelength lamps (see Table 1). Note that the HITM2 lamp has also faded after SM4 even if it was not used during STIS operations.
Figure 3: Total counts normalized to the maximum versus usage of the LINE (black) and HITM1 (red) lamps in milliAmper-hour. Note the faster exponential decay of the LINE output ($\propto 10^{-0.7 \times \text{use}}$) in comparison to the HITM1 output ($\propto 10^{-0.3 \times \text{use}}$) in the 1150-1200 Å wavelength range. Note also that the LINE lamp decline became steeper for usage $\geq 200$ mA-h.

Figure 4: Ratio of the LINE and HITM1 lamps in four low-resolution gratings (G140L, G230LB, G430L, and G750L) which cover the full wavelength range offered with STIS. Circles are flux ratios from our 2010 Lamp Calibration Program (PID 1209). Different colors correspond to ratios sampling different wavelength regions within the same grating (see main text). 'X' symbols show lamp flux ratios from the 1997 Lamp Calibration Program (PID 7722) for the whole wavelength region (black) and for the 1145-1201 Å region for the G140L (red). Flux ratios for the other lamps are summarized in Table 1. Note that the LINE/HITM1 flux ratio in the G140L short wavelength band (1145-1201 Å region) has decreased below 1.
Figure 5: Plots comparing the 1997 (left panel) and 2010 (right panel) lamp fluxes for the very short wavelengths of the G140L setting. These figures demonstrate how the LINE lamp has faded at these wavelengths in comparison to the other two wavelength calibration lamps. The broader line profiles in the 2010 HITM1 and HITM2 spectra compared to the 1997 spectra are due to the larger slit width employed for the wavelength calibration (0.1″ in 2010 versus the 0.05″ slit in 1997.)

Figure 6: LINE lamp spectra obtained as part of the standard wavelength calibration for G140M-1173Å observations (PID 7667 for the 1997 spectrum and PID 11860 for the 2009 spectrum). The exposure time for standard calibration spectra is set to 46 seconds. Note that only a few bright lines are still detected in the 2009 lamp spectra and no lines are detected shortward of ∼1170Å.
Figure 7: Like Fig. 3 but including the output of the COS wavelength calibration lamp (blue circles). Since the STIS/HITM2 lamp was only used twice its output is not reported here. The COS lamp outputs have been corrected for the FUV sensitivity degradation which is estimated to be $\sim 9\%$ per year for the G140L-1105 setting, see text. The COS lamp outputs also show more scatter than the STIS lamps. This could be due to differences in the background for orbits close to the South Atlantic Anomaly (the STIS MAMAs cannot be used during any part of an orbit that is even partially impacted by the Anomaly).

Table 1: 2010 STIS lamp flux ratios measured in the full wavelength region covered by each grating and in three smaller regions covering the short, middle, and long wavelength portions of each setting. Lamp spectra were obtained as part of our Cycle 17 calibration program (PID 12079).

<table>
<thead>
<tr>
<th>Lamps</th>
<th>G140L</th>
<th>G230LB</th>
<th>G430L</th>
<th>G750L</th>
</tr>
</thead>
<tbody>
<tr>
<td>LINE/HITM1</td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>Full $\lambda$ range</td>
<td>4</td>
<td>5</td>
<td>7</td>
<td>22</td>
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<tr>
<td>Short $\lambda$ range</td>
<td>0.7</td>
<td>5</td>
<td>5</td>
<td>25</td>
</tr>
<tr>
<td>Middle $\lambda$ range</td>
<td>4</td>
<td>4</td>
<td>7</td>
<td>11</td>
</tr>
<tr>
<td>Long $\lambda$ range</td>
<td>4</td>
<td>5</td>
<td>10</td>
<td>10</td>
</tr>
<tr>
<td>LINE/HITM2</td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>Full $\lambda$ range</td>
<td>3</td>
<td>6</td>
<td>5</td>
<td>8</td>
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<tr>
<td>Short $\lambda$ range</td>
<td>0.3</td>
<td>2</td>
<td>5</td>
<td>12</td>
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<tr>
<td>Middle $\lambda$ range</td>
<td>3</td>
<td>5</td>
<td>4</td>
<td>3</td>
</tr>
<tr>
<td>Long $\lambda$ range</td>
<td>5</td>
<td>7</td>
<td>4</td>
<td>1</td>
</tr>
<tr>
<td>HITM1/HITM2</td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>Full $\lambda$ range</td>
<td>0.6</td>
<td>1</td>
<td>0.6</td>
<td>0.4</td>
</tr>
<tr>
<td>Short $\lambda$ range</td>
<td>0.4</td>
<td>0.4</td>
<td>1</td>
<td>0.5</td>
</tr>
<tr>
<td>Middle $\lambda$ range</td>
<td>0.7</td>
<td>1</td>
<td>0.6</td>
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<tr>
<td>Long $\lambda$ range</td>
<td>1</td>
<td>1</td>
<td>0.4</td>
<td>0.1</td>
</tr>
</tbody>
</table>
Updated Status and Performance of the STIS CCD

M. A. Wolfe, A. Aloisi, R. C. Bohlin, P. Goudfrooij, R. A. Osten

*Space Telescope Science Institute, Baltimore, MD 21218*

C. R. Proffitt

*Computer Science Corporation*

W. V. Dixon

*Johns Hopkins University, Baltimore, MD 21218*

D. Lennon

*Space Telescope Science Institute, Baltimore, MD 21218*

T. R. Gull, B. E. Woodgate

*NASA Goddard Space Flight Center, Greenbelt, MD*

D. J. Lindler

*Sigma Space Corporation*

**Abstract.** A description is provided of the overall performance of the STIS CCD after HST Servicing Mission 4 (SM4) and during Cycle 17 calibrations. Most aspects of CCD performance are found to be consistent with extrapolations of the trends seen prior to the failure of STIS in August 2004. The gain values for 1 and 4 have not measurably changed from pre-SM4 measures. The read noise through Amp D, as determined from unbinned bias images taken during Cycle 17, is slightly higher than pre-SM4 values. As is expected due to the on-orbit radiation environment the dark current continues to increase. The spurious charge has increased as well and has a slope that results in a smaller value at the top of the chip (near the readout amplifier) when compared to the center of the chip. The charge transfer inefficiency (CTI) also continues to increase although the increasing dark current will noticeably minimize actual CTI losses for typical exposure times exceeding several hundred seconds. The post-SM4 CTI measured values agree well with extrapolation of the trend seen through Cycle 17 calibrations. Changes in the sensitivity of STIS CCD modes have been modest.

1. **Dark Normalization**

After the failure of Side-1 electronics and the transition to Side-2, the CCD detector temperature could not be held at a constant value because Side-2 lacks a working thermistor on

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1Space Telescope Science Institute, Baltimore, MD 21218

2European Space Agency
the CCD chip. A thermistor located on the CCD housing acts as a proxy to monitor, not control, the CCD thermal variations. Figure 1 depicts the fluctuating housing temperature over the one year since SM4. A scaled measure of the net solar illumination (solid line in Figure 1) appears to account for most of the housing temperature variations and hence can be used as a predictor for the CCD dark count. The measured dark current correlates well with the housing temperature and the fluctuations in the dark current can be scaled using equation 1:

\[
\text{normalized dark image} = (\text{dark image}) \times (1 + \text{slope} \times (T_O - T))
\]

where normalized dark image is the result, dark image is the dark image being normalized, slope is a constant scaling factor, $T_O$ is the reference temperature, and $T$ is the housing temperature of the exposure. Table 1 shows the old and new slopes and reference temperatures used to normalize the darks.

![Housing Temperature Trend](image)

Figure 1: Note that there is less variation in the measured housing temperature well after the SMOV4 testing period. During a portion of the SMOV4 period, a SIC&DH safing led to suspension of STIS and resultant instrument cooling. The solid curve is an appropriately scaled integrated solar flux (average solar radiation times sun illumination time) impinging on HST. After end of SMOV4, variations of the housing temperature appear to track the integrated solar flux variations.

<table>
<thead>
<tr>
<th>Dates</th>
<th>Slope</th>
<th>Reference Temperature ($T_O$)</th>
<th>Dark Exp Time</th>
</tr>
</thead>
<tbody>
<tr>
<td>07/2001 - 08/2004</td>
<td>0.070/°C</td>
<td>18.0 °C</td>
<td>1100 s</td>
</tr>
<tr>
<td>06/2009 - Present</td>
<td>0.056/°C</td>
<td>22.0 °C</td>
<td>1100 s</td>
</tr>
</tbody>
</table>
2. Sensitivities

STIS sensitivities for the CCD low resolution grating spectral regions are presented in Table 2 for 2010 and 2004 relative to the initial 1997 measures. Sensitivities for the less fully sampled medium resolution gratings follow the same trends.

<table>
<thead>
<tr>
<th></th>
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<th></th>
</tr>
</thead>
<tbody>
<tr>
<td>G230LB</td>
<td>0.949</td>
<td>0.920</td>
</tr>
<tr>
<td>G430L</td>
<td>0.981</td>
<td>0.964</td>
</tr>
<tr>
<td>G750L</td>
<td>0.988</td>
<td>0.978</td>
</tr>
</tbody>
</table>

3. Charge Transfer Inefficiency

Note that the value quoted for the CTE is in terms of charge transfer inefficiency (CTI). The CTI slope quoted in Table 3 is for the linear equation (Goudfrooij et al. 2009) that evaluates the CTI as time progresses. This time-dependent equation is:

\[ \text{CTI}(t) = \left( \text{CTI}_0 \right) \times \left( 1 + \alpha (t - t_0) \right) \]  

where \( \text{CTI}(t) \) is the value of the CTI at a specific time \( t \), \( \text{CTI}_0 \) is the CTI extrapolated to zero background, \( \alpha \) is the slope of the time dependence, \( t \) is the time in years, and \( t_0 \) is 2000.6. Note the value for \( \alpha \) can be found in Table 3. The value for \( \alpha \) is derived from the Internal Sparse Field Test (Goudfrooij et al. 2006).

Another CTI test is the Extended Pixel Edge Response (EPER). This test, while not being an absolute measure of CTI, produces a relative measure of CTI thus allowing a robust measure of time trends (Janesick 2001). Figure 2 plots the CTI as a function of time for both the parallel (top plot) and serial (bottom plot) directions. The parallel CTI increased noticeably more than predicted from a linear extrapolation from pre-SM4 measures while the serial CTI increased marginally. Note that the serial CTI is about an order of magnitude less than the parallel CTI measure. However, the achievable operating temperature for the STIS CCD is \( \sim -83 \, ^\circ C \), well within the temperature range where there is a strong dependence upon temperature. This is demonstrated in Figure 3 which shows the ratio of the derived EPER CTI value, relative to the expected value based upon the linear fit shown in Figure 2, plotted against housing temperature. The parallel CTI appears to show a 20 to 40% increase beginning abruptly at temperatures above 20 \( ^\circ C \). The serial CTI temperature dependence is noticeable, but dominated by measurement noise and/or additional contributions from other sources.

4. CCD Parameters

Values for various parameters that characterize the performance of the STIS CCD are given in Table 3. The slopes in both the dark current and spurious charge are thought to be due to the accumulation of radiation damage (see Goudfrooij et al. 2009 and the references therein). The slope in the dark current along CCD columns results in a dark current near the detector readout amplifier that is about 37.9% lower than the mean over the whole detector. The spurious charge is approximately 36.3% lower at the top of the chip (row 900) than it is in the center of the chip (row 512) for gain = 1. The same behavior for the
Figure 2: The time dependence of the EPER CTI in the parallel (top plot) and serial (bottom plot) directions. Note that the parallel CTI is much greater than the CTI in the serial direction.
Figure 3: This plot shows the relationship between the ratio of the derived values (from the stis_eper.cl script) to linear fit values (shown in Figure 2) against the CCD housing temperature. This figure shows that for data taken post SM4 that the ratios are much greater than 1, suggesting a relationship between the CCD housing temperature and parallel CTI (top plot). The bottom plot shows that the dependence of CTI on housing temperature is not as strong for the serial direction.
gain = 4 spurious charge is also present but the percentage difference between the top of the chip and the center is 6.36%.

### Table 3: Updated CCD Parameter Values

<table>
<thead>
<tr>
<th>Parameter</th>
<th>Pre-SM4 Value</th>
<th>Current Value</th>
</tr>
</thead>
<tbody>
<tr>
<td>Dark Current (Top of Chip)</td>
<td>0.0043 e⁻/pixel/s</td>
<td>0.0102 e⁻/pixel/s</td>
</tr>
<tr>
<td>Dark Current (Center of Chip)</td>
<td>0.0070 e⁻/pixel/s</td>
<td>0.0165 e⁻/pixel/s</td>
</tr>
<tr>
<td>Number of Hot Pixels</td>
<td>19816 (0.1 e⁻/s level)</td>
<td>44394 (0.1 e⁻/s level)</td>
</tr>
<tr>
<td>Spurious Charge (Top of Chip, Gain = 1)</td>
<td>0.55e⁻/pixel</td>
<td>0.80e⁻/pixel</td>
</tr>
<tr>
<td>Spurious Charge (Center of Chip, Gain = 1)</td>
<td>0.75e⁻/pixel</td>
<td>1.26e⁻/pixel</td>
</tr>
<tr>
<td>Spurious Charge (Top of Chip, Gain = 4)</td>
<td>5.15e⁻/pixel</td>
<td>6.33e⁻/pixel</td>
</tr>
<tr>
<td>Spurious Charge (Center of Chip, Gain = 4)</td>
<td>5.35e⁻/pixel</td>
<td>6.76e⁻/pixel</td>
</tr>
<tr>
<td>Read Noise, Gain = 1</td>
<td>5.40 e⁻</td>
<td>5.54 e⁻</td>
</tr>
<tr>
<td>Read Noise, Gain = 4</td>
<td>7.70 e⁻</td>
<td>7.94 e⁻</td>
</tr>
<tr>
<td>Fractional increase in rate of CTI (α)</td>
<td>0.218/year</td>
<td>0.216/year</td>
</tr>
<tr>
<td>EPER CTI at ≈ 12,000 e⁻ signal level</td>
<td>4.9×10⁻⁵</td>
<td>9.2×10⁻⁵</td>
</tr>
<tr>
<td>Gain = 1</td>
<td>1.0 e⁻/DN</td>
<td>1.0 e⁻/DN</td>
</tr>
<tr>
<td>Gain = 4</td>
<td>4.032 e⁻/DN</td>
<td>4.016 e⁻/DN</td>
</tr>
</tbody>
</table>

### 5. Observational Recommendations

Note that the top of the CCD has reduced values of dark current and spurious charge. Furthermore, the effects of charge transfer inefficiency are mitigated at the top of the chip due to fewer charge transfers. This significantly reduces the strength of cosmic ray and CTI tails. It is, therefore, recommended that observations (if possible) be done at the top of the chip at row 900. Observations at row 900 can accomplished by using the long slit E1 pseudo-apertures of 52X2E1, 52X1E1, 52X0.2E1, 52X0.1E1, and 52X0.05E1. Additionally, because of the total number of hot pixels, which is 5.7% of all the pixels, appropriate dithering of CCD images is becoming increasingly important.

### References

Removing the Pattern Noise from all STIS Side-2 CCD data

Rolf A. Jansen, Rogier Windhorst, Hwihyun Kim
Arizona State University, School of Earth & Space Exploration, Tempe, AZ 85287

Nimish Hathi\textsuperscript{1}, Paul Goudfrooij\textsuperscript{2}, & Nicholas Collins\textsuperscript{3}
\textsuperscript{1}University of California, Riverside, CA 92521; \textsuperscript{2}Space Telescope Science Institute, Baltimore, MD 21218; \textsuperscript{3}Wyle Information Systems, McLean, VA 22102, and NASA/Goddard Space Flight Center, Greenbelt, MD 20771

Abstract. When HST/STIS resumed operations in July 2001 using its redundant “Side-2” electronics, the read-noise of the CCD detector appeared to have increased by $\sim 1 \text{ e}^-$ due to a superimposed and highly variable “herring-bone” pattern noise. For the majority of programs aiming to detect signals near the STIS design limits, the impact of this noise is far more significant than implied by a mere $1 \text{ e}^-$ increase in amplitude of the read-noise, as it is of a systematic nature and can result in $\sim 8 \text{ e}^-$ relative deviations (peak-to-valley). We discuss the nature of the pattern noise, and summarize a method to robustly detect and remove it from raw STIS CCD frames. We report on a Cycle 16/17 Archival Calibration Legacy program to (semi-)automatically remove the herring-bone pattern noise from all raw, unbinned Side-2 STIS/CCD frames taken between 2001 July and 2004 August — representing a gain in effective sensitivity of a factor $\sim 3$ at low S/N. We also present some trends in the characteristics of the noise pattern.

1. Nature of the Pattern Noise

The superimposed noise signal, due to analog-digital cross-talk or a grounding issue in the STIS Side-2 circuitry, is not a spatial signal, but a high frequency signal in time. That signal manifests itself as a spatial “herring-bone” pattern (Fig. 1) that can drift erratically — even during the relatively short time it takes to read the CCD. The pattern tends to be locally semi-coherent, however, and is best described as a modulated $\sim 14$–$18 \text{ kHz}$ wave. The amplitude of that high-frequency wave is modulated by the superposition of three $\sim 1 \text{ kHz}$ sinusoidal waves with phases that are shifted $120^\circ$ from one another, and which have amplitudes of $3$–$5 \text{ e}^-$ (see Fig. 2a and b).

Since a $14$–$18 \text{ kHz}$ frequency corresponds to a spatial period of $2.5$–$3.2$ pixels, the values of adjacent pixels along a row tend to be affected by offsets of opposite signs (Fig. 2a), resulting in relative deviations of up to $\sim 8 \text{ e}^-$ (peak-to-valley). Adjacent pixels along columns experience offsets that are shifted in phase by amounts that vary from region to region in a single frame, and also from frame to frame. The resulting impact on Side-2 CCD data is therefore far more serious than implied by a mere $1 \text{ e}^-$ increase in the amplitude of the read-noise, and is partly systematic in nature.

2. Removing the pattern noise

Brown (2001) introduced a method to filter out the pattern noise by noting that the sequen-
Figure 1: A section of a raw, unbinned STIS/CCD bias frame, taken in July 2001. This section features the highly variable “herring-bone” noise pattern, several (vertical) columns and individual pixels with elevated bias level, as well as three regions affected by cosmic ray hits.

Figure 2: (a) The noise pattern is not a spatial signal, but results from a high-frequency signal in time. The difference of two adjacent pixels can be affected by up to $\sim 8 e^-$ (peak-to-valley). Apart from the $\sim 16$ kHz (2.8 pixel) pattern in this example, three sinusoidal waves—with a frequency of $\sim 1$ kHz and phases that differ by $120^\circ$—define an envelope on the amplitude of the high-frequency primary pattern. (b) The pattern can be semi-coherent over tens to hundreds of pixels. The $\sim 1$ kHz signal is likely associated with an onboard power supply, oscillator, or clock.
tial charge shifts during read-out of the CCD allow one to convert a 2-D image into a timed signal. That time-series may be Fourier transformed to the frequency domain, where one can search for the frequencies responsible for the noise pattern, and then suppress them in various ways. This works well in images or portions of images where few bright and/or sharp (spatially very concentrated) features are present, but requires manual definition of the frequency limits of the filter. If the filter is chosen too wide, or if many genuine high-frequency non-periodic signals (e.g., stars, spectral lines, cosmic ray events) are present, ringing may occur (see, e.g., figures 1b and 6b of Brown 2001).

Jansen et al. 2003 noted that the problem of automatically and robustly finding the frequencies that correspond to the pattern is greatly reduced if the genuine background and science signals are modeled and subtracted first. The resulting residuals image, ideally, only contains photon noise, read-noise, and the herring-bone pattern. In practice, since the model won’t be (and does not need to be) perfect, there are systematic residuals of genuine features in the data as well. But the contrast of the herring-bone pattern has become much higher than in the original image. This means that, in the frequency domain, one can blindly run a peak finding routine with much relaxed constraints on the frequency interval (or alternatively on much poorer data — e.g., very long spectroscopic exposures that are riddled with cosmic ray hits) and still correctly find, fit, and filter out the pattern frequencies. Also, since most of the power from genuine signal has been removed prior to constructing the power spectrum, the problem of ringing is effectively avoided.

We further improved the method by replacing the power at frequencies associated with the noise pattern with white noise at a level and amplitude that matches the “background” power in two intervals that bracket the affected frequencies. In the original method, such frequencies were suppressed using multiplicative filters or windowing functions, or were set to zero. Replacement with white noise is less likely to introduce artefacts due to the absence of power at frequencies that should have some, or which may result when many adjacent frequencies have identical or zero power. The resulting modified power spectrum is inverse Fourier transformed, converted to a 2-D image, and added to the previously fitted “data model” to produce a CCD frame from which the pattern noise is completely removed.

Figure 3 [next page]: Overview of the autofilet procedure. (a) Section of the raw STIS/CCD bias frame of 1. (b) A data “model” constructed for this section, containing most of the signal (as fitted to the image lines and columns) and also all pixels deviating from that fit by more than 3σ, or by more than 0.5σ when adjacent to a pixel that deviates by more than 3σ. The difference of the original image section and this model, i.e., the residuals image, is converted to a time-series and Fourier transformed to frequency space. (c) Portion of the power spectrum centered on the frequencies responsible for the herring-bone pattern. After finding this peak, an estimate of its width (resulting from the erratic drift in frequency of the pattern during the time it takes to read the CCD) is obtained by fitting a Gaussian. All power within ±3σ of the peak’s central frequency is then replaced by white noise that matches the noise in the two bracketing regions located 4–7σ away. The resulting modified power spectrum is inverse Fourier transformed and converted back into a 2-D image, to which the model of panel b is added. (e) The resulting pattern-subtracted, cleaned frame. Note, that there is no “ringing” around bright regions affected by cosmic ray hits. The difference between panels a and e, i.e., an image of the detected noise pattern, is shown in (d). (f) Comparison of the distribution of pixel values in the raw and pattern-subtracted bias frames. Whereas the noise in the raw bias frame is distinctly non-gaussian near the mean pixel value and has a σ ∼ 5.5e−, after removal of the inferred herring-bone pattern the remaining noise closely resembles white noise with a significantly smaller standard deviation σ ∼ 4.0e−. Autofilet therefore successfully reproduces the nominal “Side-1” CCD read-noise observed prior to July 2001.
Figure 3:
Removing the Pattern Noise from all STIS Side-2 CCD data

Figure 4: Comparison of a weekly “superbias” reference frame retrieved from the HST Archive and one constructed from pattern-subtracted biases. While the “herring-bone” patterns vary from one frame to the next, they are not sufficiently random to cancel out completely when averaging multiple frames. In the left panel, significant residuals from the pattern noise are seen even when more than 100 individual frames are averaged. The frame constructed from our pattern-subtracted biases (right) is free of such residuals. Indeed, in the bottom panel, the pixel histogram of the STScI/OPUS bias reference frame shows a broader distribution of pixel values, while our frame approximates the theoretically expected gaussian distribution. The observed tail toward higher pixel values results from hot and warm pixels, mostly located along discrete detector columns.
Our optimized Fourier filtering method, briefly outlined above and summarized in Fig. 3, was implemented in IDL procedure autofilet.pro. Several auxiliary shell-scripts provide input and allow batch processing of multiple CCD frames, while a compiled program, fits2mef, generates multi-extension FITS datasets that are compatible again with calstis. A comparison of the pixel histograms of original and cleaned bias frames (Fig. 3f) demonstrates that the noise in the pattern-subtracted frames approximates the theoretically expected distribution very closely and matches the nominal "Side-1" CCD read-noise that was observed prior to July 2001.

3. Archival Calibration Legacy program AR 11258

As part of AR 11258, all raw, unbinned, full-frame Side-2 STIS/CCD data sets taken between 2001 July and 2004 August (each containing one or more individual frames) were retrieved from the HST Archive and processed at ASU using autofilet to remove the herring-bone pattern noise. The 75345 cleaned frames were quality verified and merged back into 47192 multi-extension FITS files and delivered to STScI. The removal of the pattern noise represents a gain in effective sensitivity of up to a factor $\sim$3 at low S/N, if one uses superbias (Fig. 4) and superdark frames generated from pattern-cleaned frames in calstis. The cleaned datasets are available from the Hubble Legacy Archive:

$$http://archive.stsci.edu/pub/hlsp/stis-herringbone/$$

Autofilet, all auxiliary software, and instructions for its use are available for download from the lead author’s web page:

$$http://www.public.asu.edu/~rjansen/stis2/stis2.tar.gz$$

For each successfully cleaned frame, we logged the detected peak frequency, frequency drift width, and peak power, as well as selected information from the FITS headers of each dataset and frame for the purpose of a trending analysis. Three examples are shown in Fig. 5. A more detailed description of Autofilet, program AR 11258, and our trending analysis is forthcoming (Jansen et al. 2010).

Acknowledgements

This work was funded by grants HST-AR-11258 and HST-GO-9066 from STScI, which is operated by AURA under NASA contract NAS5-26555. We thank Bruce Woodgate for getting us started. We would not have had the same success without the prior work by Thomas M. Brown.

References

Removing the Pattern Noise from all STIS Side-2 CCD data

Figure 5: Noise pattern trends. (a) Detected peak power in the frequencies associated with the “herring-bone” pattern noise. The DARKs show that pattern detection contrast depends on the spatial density of genuine (or cosmic ray induced) strongly peaked signals. (b) We find little change with time in the amplitude of the pattern noise. (c) The average frequency associated with the pattern noise has decreased by $\sim 6\%$ from 2001 July till 2004 July, a trend that continues also after the successful repair of STIS during SM4. At any given epoch there is a wide range of $\sim 1$–3 kHz in pattern-frequency measured in individual CCD frames, but frames taken in close succession tend to show similar pattern-frequencies. Some of the larger excursions in frequency may be associated with monthly anneals.
Post - SM4 Flux Calibration of the STIS Echelle Modes

Azalee Bostroem, A. Aloisi, R. C. Bohlin, C. R. Proffitt, R. A. Osten, D. Lennon

Space Telescope Science Institute, Baltimore, MD

Abstract. Like all STIS spectroscopic modes, STIS echelle modes show a wavelength dependent decline in detector sensitivity with time. The echelle sensitivity is further affected by a time-dependent shift in the blaze function. To better correct the effects of the echelle sensitivity loss and the blaze function changes, we derive new baselines for echelle sensitivities from post-HST Servicing Mission 4 observations of the standard star G191-B2B. We present how these baseline sensitivities compare to pre-failure trends.

1. Introduction

The Space Telescope Imaging Spectrograph includes 2 medium and 2 high resolution ultraviolet (UV) echelle gratings, E140M, E230M, E140H, and E230H, with resolving powers of \( R \sim 45,800, \) 30,000, 114,000, and 114,000, respectively. These gratings provide the highest resolution spectra available on the Hubble Space Telescope (HST). There are multiple central wavelength settings or modes, available for each grating. A few central wavelengths in each grating are referred to as prime modes and the rest are denoted secondary modes. See the Chapter 13 in the STIS Instrument Handbook for a complete list of prime and secondary central wavelengths.

The absolute flux calibration of the echelle modes has proved particularly difficult. The initial calibration of the 12 prime modes was performed in 1998 (Bohlin, 1998). In 2006, prior to the reprocessing of STIS data, all echelle modes were recalibrated (Aloisi et al., 2007). Between 1998 and 2006, the characterization of many STIS features improved significantly. The 2006 sensitivity calibration includes these improvements as well as a correction for shifts in the blaze function (BF). It is also the first direct calibration of the secondary echelle modes and the first flux calibration of all echelle modes since the switch to Side 2 electronics in May 2001.

The STIS detector sensitivity has been shown to decline smoothly with wavelength over time. Additionally, large temperature changes (such as the switch from Side 1 to Side 2 electronics) can cause elements within the instrument to shift, creating unpredictable shifts in the echelle BF. For this reason, when STIS was restarted in May 2009, after HST Servicing Mission 4 (SM4), a program was undertaken to characterize the sensitivity of the echelle modes for post-SM4 data.

2. Observations

All observations were taken as part of the Cycle 17 calibration program 11866. The HST primary standard star G191-B2B, a DA white dwarf, was observed in all modes using the \( 0.2'' \times 0.2'' \) aperture. Unlike previous calibrations, all data was taken in the zero monthly offset position (monthly offsetting was disabled for echelle modes in August 2002), creating a more uniform data set. All observations were taken between November 25, 2009 and January 6, 2010. Exposure times were set to obtain a peak signal-to-noise (S/N) ratio of...
30 in each central wavelength for the E140M, E140H, and E230M grating and a S/N ratio of 20 for the E230H grating. For the three most commonly used Cycle 17 FUV modes (E140H/1234, E140M/1425, E140H/1343), exposure times were chosen to obtain a S/N ratio of 100.

3. Analysis

This analysis closely follows that of Bohlin (1998) and Aloisi et al (2007). First the sensitivity curves for each order are found. The sensitivity curves are then combined with the blaze function to create the photometric throughput (PHT) tables, which are used by CALSTIS in the absolute flux calibration of STIS data.

3.1. Finding the Sensitivity Curves for Each Order

The sensitivity curve of each order in a central wavelength is determined by performing three spline fits: one across each order (in the dispersion direction), one to the average sensitivity of each spectral order, and one across the spectral orders for each node defined in the first fit.

The observed count rate is divided by the model flux spectrum, which has been interpolated to the observed wavelength scale. This is done for each order, m, of an extracted spectrum (x1d file), creating a sensitivity curve for each order. A 7 (9 for E230M) node spline fit is made to the median smoothed sensitivity spectrum, placing the nodes, N(i,m), (where i is the node number within an order) at constant x pixel locations (where the x axis is defined horizontally in the same direction as the increasing wavelengths within a given order). The median smoothing smoothes many narrow absorption lines which are in the observed spectrum, but not in the model spectrum. In some cases, this smoothing is insufficient and the line must be masked or the entire order excluded from the final fit. Additionally, pixels with bad data quality flags are excluded from the fit.

The sensitivity values of these nodes are averaged to create an average sensitivity for each order. A spline fit is then performed to the average values and the fit used as the average sensitivity for each order, A(m).

Next, the nodes are normalized by the average sensitivity of a given order, N_{norm}(i, m) = N(i, m)/A(m). All normalized nodes at a single x pixel location are fit, smoothing vertically along the detector. The normalization brings all node values at a given x pixel location to approximately the same sensitivity, making them easier to fit. A spline fit is used again, with closely spaced nodes for the end orders and a few widely spaced nodes for the middle orders. This fit primarily smooths out any lingering abnormal features (such as absorption lines or bad data). The values of this spline fit at the original nodes, N(i, m), for each order are multiplied by the average to obtain the original sensitivity and are then written to a file with the corresponding wavelengths.

Finally, if more than one observation exists for a given grating and central wavelength, then the two files are combined and weighted by exposure time.

3.2. Blaze Function

A blaze function (BF) is an optical throughput curve that describes the change in sensitivity of a grating across a detector. Two of the variables that affect the BF are the incident light angle on the grating and the angle of the grating with respect to the detector. In general, the BF is optimized so that its peak falls on the center of the detector. However, the change in position of a spectrum on a detector is indicative of a change in grating position or incident light angle, and therefore means a shift in the BF. In a static system, the BF shift is simply a function of detector location: observed wavelength (OBSW) and cross-dispersion location (OBSY). However, STIS is more complicated. The mode select mechanism (MSM) is a
nutating wheel, allowing a grating to move in all three dimensions. Hence, the BF does not shift by the same amount as the spectrum. Additionally, there appears to be a drift of the BF with time (and thus a time dependence) as well as an offset which is thought to be the result of the flexing of the instrument with large thermal changes. The BF shift equation, as defined in Aloisi et al. (2007), is:

$$B_{\text{SHIFT}} = B_{\text{SHIFT VS } X} \Delta x + B_{\text{SHIFT VS } Y} \Delta y + B_{\text{SHIFT VS } T} \Delta T + B_{\text{SHIFT OFFSET}}$$

(1)

where

$$\Delta x = (\text{REFWAV} - \text{OBSW})/\text{DISP}$$

$$\Delta y = \text{OBSY} - \text{REFY}$$

$$\Delta t = \text{OBSDATE} - \text{REFMJD}.$$ 

REFWAV and REFY are the wavelength and cross-dispersion position, respectively, of the reference spectral order at pixel $x = 512$. OBSDATE is the observation date in Modified Julian Days (MJD) and REFMJD is the MJD of the observation used to infer the sensitivity of the echelle mode. $B_{\text{SHIFT VS } X}$ and $B_{\text{SHIFT VS } Y}$ are functions of detector location and vary only with grating. $B_{\text{SHIFT VS } T}$ and $B_{\text{SHIFT OFFSET}}$ vary with grating, order, and Side (Side1, Side2, or Side2 repaired electronics). The temporal components of the BF shift ($B_{\text{SHIFT T}}$ and $B_{\text{SHIFT OFFSET}}$) reflect changes with time since the sensitivity curves were created. As we created new sensitivity files, these were 0.

3.3. Creating the PHT Table

The sensitivity curves must be corrected to meet the PHT table sensitivity definitions. Specifically, the sensitivity curves are derived from count rates observed through a $0.2'' \times 0.2''$ aperture and extracted using a 7 pixel high extraction region. The net counts are corrected to an infinite extraction region and aperture. Additionally, the sensitivity of the detector has decreased over time at a rate that is wavelength dependent. The sensitivity curves are measurements of the degraded sensitivity as of Cycle 17. The PHT table assumes sensitivity measurements from the beginning of STIS operations in 1997. The time-dependent changes in the sensitivity are therefore removed from the sensitivity curves. The sensitivity is also converted to throughput. The correct throughput values and the BF shift coefficients are then written to the PHT table. One table is made for each detector.

4. Results

A PHT table is tested by running the STScI calibration pipeline program, CalSTIS, on data from the SMOV program 11403 and the Cycle 17 program 11860. These programs periodically observe the HST secondary standard star BD+28D4211 to monitor changes in sensitivity of the STIS detectors. The programs were chosen as they observe a source with a reference spectrum and have observations which are taken at regular intervals. CALSTIS is run twice on observations from August and September 2009 and April and June 2010, once with the current PHT table (currently used by OPUS for On the Fly Reprocessing) and once with the new PHT table. The flux from both these calibrations is plotted together with the reference spectrum. While the overall sensitivity level is expected to be the same (since it is corrected with the time-dependent sensitivity files) the shape and location of the blaze function is expected to have changed, causing the data calibrated with the current PHT table to show arcs in each order. The data calibrated with the new PHT table should follow the reference spectrum closely. This result can be seen in Figure 1, where the green line is the data calibrated with the new PHT table, the red line is the data calibrated with
the current PHT table, and the black line is the reference spectrum. The inset plot shows that the new PHT table spectrum matches the reference spectrum much better than the current PHT table spectrum.

Figure 1: A plot of the flux calibration of the HST standard star BD +28D4211A. This data was taken as part of the sensitivity monitoring program in June 2010 using the E140M grating. In the plot, a clear comparison can be drawn between the reference spectrum (black line) and a spectrum calibrated with the new PHT table (green line) and the current PHT table (red line). Inset is the same plot for the wavelength range 1550-1610. In this plot, the BF is clearly being calibrated incorrectly with the current PHT table as the BF arc of each order is apparent in the red line but not in the green line.

4.1. E140H Anomaly

Sensitivity monitoring data from the E140H grating showed a much lower sensitivity than anticipated in August 2009. Since the anomaly was first detected, the sensitivity has steady increased to close to expected levels. A comparison of April and June data show that the sensitivity is continuing to increase. As the echelle flux calibration data was taken between November 2009 and January 2010, it is possible that it has been impacted by the anomaly. The anomaly would affect the flux calibration by artificially increasing the flux above the reference spectrum. Figure 2, shows the behavior of the anomaly over time. While the spectrum calibrated with the new PHT table is systematically higher than the spectrum...
calibrated with the current PHT table, there is no evidence that the new PHT table is over correcting the flux above the reference spectrum. Additional monitoring data obtained in September 2010 will be used to confirm the validity of the new PHT table prior its implementation as a reference file.

Figure 2: Sensitivity monitoring over time of the HST standard star BD+28D4211. The black line is the reference spectrum, the green line is the data calibrated with the new PHT table, and the red line is data calibrated with the current PHT table. The date each observation was taken is given at the top of each plot. From these plots, the E140H anomaly is clearly visible, as is its evolution over time towards the expected sensitivity.

5. Conclusion

The absolute flux calibration of the STIS echelle modes post-HST SM4 was undertaken to accurately correct the blaze function and the decreased sensitivity of the detectors over time. Observations of the white dwarf G191-B2B were taken and used to create sensitivity curves. These curves were corrected and combined with the blaze function files to create new PHT tables. The new PHT tables were used to calibrate observations of the star BD+28D4211 and results were compared to results from the current PHT tables. It was found that the new PHT tables provide a better absolute flux calibration than the current PHT tables.
through a better characterization of the blaze function shift and sensitivity curves. These files will be implemented in On the Fly Reprocessing of post HST-SM4 data.

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The Space Telescope Imaging Spectrograph Flat Fielding

Sami-Matias Niemi
Space Telescope Science Institute, Baltimore, MD 21218

Charles R. Proffitt and Thomas B. Ake
Space Telescope Science Institute / CSC, Baltimore, MD, 21218

Daniel Lennon
Space Telescope Science Institute / ESA, Baltimore, MD, 21218

Abstract. The Space Telescope Imaging Spectrograph (STIS) has two types of reference files that are related to flat-fielding. A low-order flat field (the so-called L-flat) consists of a map of the large scale (> 64 pixels) variations in the sensitivity across a detector and contains the wavelength-dependent, low spatial frequency information about the uniformity of the detector. A pixel-to-pixel flat field (the so-called P-flat) contains the high-frequency pixel-to-pixel variations and the small scale blemishes like dead and bright pixels, thinning differences, and dust motes. The main features of the P-flats are the high frequency odd-even pattern seen in the high-resolution multi-anode microchannel array exposures and the small scale blemishes from dead and bright pixels, dust motes, etc. seen in the charge-coupled device (CCD) images. The L-flats are not expected to change significantly as a function of time. We therefore concentrate on P-flats and describe the generation of new post Servicing Mission 4 (SM4) STIS P-flats. The programs for taking data and software used to generate P-flats are described briefly. Application and the effectiveness of the new P-flats are also discussed. New post-SM4 FUV P-flats do not show any structural changes, hence no new flats have been made available, while the analysis of post-SM4 NUV P-flats is underway. New post-SM4 P-flats for the STIS CCD were delivered in December 2009 for use in pipeline processing.

1. Introduction

The Space Telescope Imaging Spectrograph (STIS) has three different detectors: two Multi-Anode Microchannel Arrays (MAMA) and a charge-coupled device (CCD). The MAMAs are used for far ultraviolet (FUV) and near ultraviolet (NUV) observations, while the CCD is reserved for optical wavelengths. The STIS internal calibration sub-system includes a Krypton (Kr), a Deuterium (D$_2$) and a Tungsten (W) lamp, which provide rather uniform illumination for the FUV, NUV and CCD detectors, respectively. In general, STIS flat fields rely on the flux of these internal calibration lamps.

Since the installation of STIS, several methods to map the pixel-to-pixel variations in the detector response have been explored. Complications in generating P-flats arise because the spectrum of each calibration lamp is not a completely uniform continuum.

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1 University of Turku, Department of Physics and Astronomy, Tuorla Observatory, Väisäläntie 20, Piiikkiö, Finland
without emission features, and because the STIS slit aperture bars occult some regions of the detector. Moreover, normalization of the lamp and slit profiles, are best done in geometrically rectified space. Hence, several different procedures must be applied when the lamp spectra are normalized to show the true underlying pixel-to-pixel variation of the detector.

During each cycle, flat fielding related calibration programs for all three detectors are executed. Data from these programs are used to monitor possible changes in the pixel-to-pixel response that may arise as a function of time. The data are also used to create new P-flats. However, a new P-flat is delivered to the pipeline only if it can be shown that the new P-flat will improve the calibration. The current super flats in the pipeline are of high signal-to-noise ratio ($SNR$) and should not be replaced by a lower $SNR$ flat, unless improvement in the calibrated output files can be demonstrated.

This proceeding paper is organized as follows. We start in Section 2. by briefly describing the three programs used to obtain flat fielding data. In Section 3. we describe the techniques applied to make a pixel-to-pixel flat field, while Section 4. is reserved for the discussion of the Cycle-17 findings. We conclude in Section 5. by briefly discussing future STIS flat field programs.

2. Data

2.1. FUV: Program 11861

Data for the Cycle-17 FUV P-flat come from program 11861, P.I. Lennon. The program obtains a set of FUV-MAMA flat field observations with the Kr lamp with sufficient counts to construct P-flats for all FUV modes. Approximately 11 visits are required to construct a P-flat with $SNR \sim 100$ per low resolution pixel. The Cycle-17 calibration program called for obtaining flats with G140M at five SLIT-STEP positions to illuminate regions of the detector normally shadowed by the slit fiducial bars. A central wavelength of 1470 Å and the $52 \times 0.1$ slit were used, resulting in approximately 284,000 counts per second in the initial exposure of the sequence. This declined to $\sim 264,000$ counts per second by the final exposure of the program. The data, when combined, can be used to achieve a signal-to-noise of $\sim 100$ or slightly below per low resolution pixel.

2.2. NUV: Program 11862

Data for the Cycle-17 NUV P-flat come from program 11862, P.I. Lennon. This program obtains a set of NUV-MAMA flat field observations with the D$_2$ lamp with sufficient counts to construct pixel-to-pixel flat fields (P-flats) for all modes. Approximately 11 visits are required to construct a P-flat with $SNR \sim 100$ per low resolution pixel. The Cycle-17 calibration program calls for obtaining flats with G230M at five SLIT-STEP positions to illuminate regions of the detector normally shadowed by the slit fiducial bars. A central wavelength of 2338 Å and the $52 \times 0.5$ slit are being used, resulting in approximately 275,000 counts per second when the decline in lamp flux from one lamp exposure to the next is being ignored. The data, when combined, can be used to achieve a signal-to-noise of $\sim 100$ or slightly below per low resolution pixel.

2.3. CCD: Program 11852

Data for the Cycle-17 CCD P-flat come from program 11852, P.I. Proffitt. This program uses the internal Tungsten lamp and it has been split into three different types of visits. The first type of visits execute every second month for a single grating and central wavelength (G430M and 5216 Å, respectively) but with two gain settings (1 and 4). Note that these visits execute with the clear aperture (50CCD) in the light path. The second type of visits uses the same grating and central wavelength as above, but this time the $52 \times 2$ slit
is positioned in the light beam. Due to the fiducial bars five different slit positions are used to provide illumination over the whole detector. The slit observations are required to provide illumination for the edges normally shadowed when the clear aperture is being used. Additionally, program 11852 executes also a third type of visits to monitor changes in the dust motes. These visits are done with G430L at five slit positions and are appropriate for monitoring because the illumination angle and magnification are different for L and M gratings.

3. Making of P-flats

3.1. MAMA

Experience with pre-flight and on-orbit monitoring flats (see e.g. Shaw, Kaiser & Ferguson 1998) show that the flat field characteristics are in large measure color- and mode-independent, so that high-quality P-flats constructed with the G140M settings should suffice for all FUV-MAMA spectroscopic and imaging programs. The same is true for NUV-MAMA when using the G230M setting (Shaw, Kaiser & Ferguson 1998). This simplifies both taking data and creating a new P-flat as data obtained with a single grating is enough, albeit several slit steps are needed to provide illumination over the whole detector.

MAMA P-flats can be created using several different techniques and over the years since the installation of STIS, procedures for creating new P-flats have evolved. Lately, the process of creating a new P-flat has followed the procedures outlined in Brown & Davies (2002) and use an iterative technique. We have adopted this method, and below we briefly outline the key points of their method.

The first step towards creating a new MAMA P-flat is to combine all suitable images by weighting each pixel with the number of input files after invalid data have been masked. The detector regions that are occulted by the occulting bars are masked out and excluded when files are being combined. The region occulted for each slit step position is $\sim 60$ pixels high in the $y$ direction. However, it is possible to move the occulting bar so that the whole detector can be illuminated when using large enough slit steps. Even so, there will still be regions in the corners and at the edges that are vignetted and cannot be fully corrected, especially for the NUV-MAMA, where the two top corners are highly occulted. The FUV detector has in addition a repeller wire in front of it that casts a curving horizontal shadow, but this shadow is correctable.

After suitable illumination patterns have been combined to make a single frame, the combined image must be normalized. The combined image is dominated by lamp and slit features in the horizontal and vertical directions, respectively, and the odd-even pattern of the detector. The normalization can be done by dividing the combined data by the lamp and slit profiles. However, to best characterize the lamp and slit features the image should be geometrically corrected. This is best done using CalSTIS, which corrects for the geometric distortion and rotation. In a rectified frame, the lamp and slit profiles can be obtained by collapsing the image in $y$ and $x$ directions, respectively.

Due to the vignetting present at the edges of the MAMAs, the lamp and slit profiles start to roll over artificially. Hence, the lamp and slit profiles must be extrapolated where they start to roll over. The point of extrapolation is the only place in the procedure where the user must make a choice where the artificial roll over starts and the extrapolation occurs. After the extrapolated profiles are available they can be smoothed by a Gaussian with FWHM = 1.5 pixels. After the smoothed profiles are available the two dimensional profiles are created and the geometric distortion correction is reverted. Finally, the combined frame can be divided by the two dimensional lamp and slit profiles.

The previously described normalization process is repeated once to obtain an accurate normalization. However, some low-level residuals remain after the lamp and slit profiles have been divided out. To remove all features larger than 64 pixels (taken into account
by the L-flat in case they are real) the normalized image is convolved with a Gaussian of FWHM = 64 pixels. The convolution is done iteratively, with masking of bad pixels, to prevent the bad pixels from throwing off the end result. Finally, after all of this processing is done, the P-flat is written to the first extension of a FITS file while the statistical errors and data quality flags are recorded in the second and third extensions, respectively.

3.2. CCD

Except for fringing effects at the longer wavelengths, the pixel-to-pixel flat field for the spectral CCD modes is independent of wavelength but does evolve with time (Bohlin, 1999a). Below \( \sim 6200 \, \text{Å} \) where there is no fringing, the STIS CCD has \( \sim 0.8\% \) root mean square (RMS, for definition see Eq. 1) intrinsic structure, which is independent of wavelength. In an extracted spectrum, the intrinsic structure is reduced to about 0.3\%. Thus, a high precision flat field correction becomes important for observational data with around 16,000 counts per pixel in the CCD image or around 100,000 counts per pixel in the extracted spectrum.

Bohlin (1999b) showed that for the STIS CCD medium resolution grating (M-mode) data can be used to generate a low dispersion grating (L-mode) P-flat; however, the dust motes of the M-mode data appear shallower and their centroids are shifted relative to the L-mode data due to the different optical magnifications. The L-mode dust motes have not been updated since 1999 as they have been shown to be relatively stable. However, since SM4 the dust motes have changed; in some cases by more than \( \pm 2\% \). Thus, a new L-mode P-flat may be generated in the future.

Due to the reason mentioned above, generating new CCD P-flats follows a slightly more complicated scheme than in case of the MAMAs discussed in the previous Section. Additionally, the process of generating new CCD P-flats is not as robust as in case of MAMAs, but the users have several choices when removing the lamp and slit profiles. However, unlike in the case of the MAMAs, the procedures to produce CCD P-flats have stayed almost the same since the installation of STIS, originally outlined in Bohlin (1999a) and revised by Niemi (2010). Niemi (2010) describes new software that was developed to create STIS CCD P-flats and to streamline the process, to improve the speed, and to provide a user friendly command line interface. Below we briefly summarize the methods used when generating a new STIS CCD P-flat.

Before anything else is done to the tungsten lamp data, the standard bias and dark removal should be performed. This, together with cosmic ray rejection, can be done using CalSTIS when the CRCORR keyword has been set to perform. After each individual exposure has been cosmic ray rejected with the standard CalSTIS technique, similar illumination patterns (i.e. the same grating and aperture or slit wheel position) can be co-added. Note, however, that one should use a fairly high (e.g. 20) sigma limit for the cosmic ray rejection when co-adding, so as not to reject any real detector features that may be relatively sharp and confused with a cosmic ray if a low sigma limit were applied. The co-adding of the same illumination pattern files can be done using CalSTIS and will produce several files depending on how many slit step positions and apertures were used when obtaining the data.

After the same illumination pattern files have been co-added, all invalid data in each file have to be masked. For the clear aperture (50CCD), the edges of the image should also be masked, while for the 52 \( \times \) 2 aperture the fiducial bar positions must be masked. Again, as in case of the MAMAs, the fiducial bar positions change together with the slit wheel position and therefore depend on the actual slit wheel encoder position. Because the position of the fiducial bars imaged onto the detector can be calculated from the slit wheel encoder position, pre-calculated slit wheel positions can be used to illuminate the whole detector. After all regions that hold no valid data, such as fiducial bar positions, have been masked, the dust motes and blemishes smaller than \( \sim 50 \) pixels should be masked. The
dust mote locations are not expected to change, thus the old information has been found to be valid year after year. Saturated pixels with $DN > 32000$ or $> 30000$ for gain 1 and 4, respectively, should also be masked.

After a mask for each co-added and cosmic ray rejected file has been generated, the next step is to normalize the lamp spectrum present in each file. Before the data can be normalized, possible sharp features should be removed. To achieve this each column of the data can be median filtered with a width of, e.g., 13 pixels. This step is not absolutely necessary and can be skipped if unwanted results are noted. Independently of whether the median filtering is applied or not, the low frequency structure and the spectrum of the lamp must be removed. To achieve this, a spline function can be fitted to all unmasked points, so that the overall low frequency form can be removed (if this is real, it will be corrected by the L-flat). Note that the number of nodes for a spline is in principle a free parameter and can be varied for the best possible outcome. Details of the removal of the lamp profile depend on whether a slit was in the light path or not, hence the following procedure is different for 50CCD and the slit data.

To remove the lamp profile in the case of the clear aperture (50CCD) data, one can start by fitting a spline to the unmasked and possibly median filtered data in the column direction. A number of spline nodes ranging from 20 to 25 has been found to give the most robust results. Note that this technique allows the number of nodes to be varied from one column to another, however, this was not used to minimize the difference in the treatment of each column. After each column of the 50CCD data has been fit and normalized to unity, each row can be fit with splines and normalized again to unity to remove possible discontinuities that might have been left after the column fits. Note that for STIS CCD data the slit profile (in row direction) is removed simultaneously when the lamp profile is being normalized, hence this step is not necessary but it is recommended. 13 nodes have been found to give the most robust results in the row direction and this value was used for the new CCD P-flats.

For the $52 \times 2$ observations, removal of the lamp profile, i.e. the column fitting, is done in a fashion similar to that for the clear aperture, however, the fiducial bars cause at least one longer gap in the data in the column direction. Because the fiducial bars render the valid data region shorter, the best number of nodes for this case is in general 13. Note that the row fitting is not appropriate for the $52 \times 2$ Tungsten spectra, which have a steep increase in signal with wavelength. Instead, we apply the spline fitting only in column direction for the $52 \times 2$ data. The error incurred from skipping the row fitting is generally less than 0.2% (Bohlin 1999a), because the slit profile gets normalized simultaneously when the lamp profile is divided out. Note that due to the masking of dust motes, detector blemishes, fiducial bars, etc. the spline fit nodes are in general unevenly spaced, and will therefore introduce an uneven weighting to the valid data when the splines are being fitted.

When robust spline fits for every column (and row if applicable), have been generated, the original, unmasked image can be divided by the smooth fits. This will remove unwanted low and mid frequency features, as well as the spectra in the lamp data, leaving only the pixel-to-pixel variations (see Fig. 4) and normalizing the image to unity. The masked points with no valid information are set to unity. Furthermore, if more than 55% of a column or a row contains invalid data, the P-flat pixel values of the entire column or row are set to unity.

After the individual P-flat images have been created and normalized, a data quality array can be populated to identify and ignore saturated pixels and invalid regions when co-adding individual P-flats to generate the final master P-flat. A statistical RMS uncertainty image is also created, and can be used to weigh pixels when separate P-flats are co-added to create a single super P-flat for a given mode. The data quality (DQ) array and the RMS uncertainty (err) extensions are populated appropriately (DQ = 512, err = 0) for invalid data, so that when different illumination pattern P-flat files are being combined for the final P-flat, these rows, columns, and pixels can be treated appropriately.
In case new L-mode P-flats are generated from the medium dispersion data the dust motes have to be added “artificially”. The information describing the dust motes is stored in a reference file that can be multiplied with the normalized M-mode P-flat file. The dust motes of the medium dispersion gratings are not appropriate for L-modes, because the dust motes are illuminated from a different direction, which causes them to shift slightly. Due to the different optical magnification in the low and medium dispersion gratings, the dust motes are also slightly smaller for L-modes.

4. Results

4.1. FUV MAMA

Figure 1 shows the current STIS FUV-MAMA P-flat reference file that is being used for the pipeline calibration. Note the strong hexagon and Moiré patterns present in the STIS FUV flat. To assess how the old reference file compares to the new FUV-MAMA P-flat, which was generated in July 2010, we studied how the structures between different flats have changed. The time evolution in the P-flats was studied by comparing the new P-flat to the reference file in the pipeline. To further characterize potential structure that may have grown over a longer time period we also compared the new P-flat to old P-flats of previous cycles. To look for time evolution in the P-flats, we divided the P-flat of one cycle by the P-flat of another, and calculated the RMS deviation

\[
\text{RMS} = \sqrt{\frac{1}{(N-1)} \sum_{i} (x_i - \bar{x})^2}
\]

in a region near the center.

Brown & Davies (2002) used the same procedure to compare P-flats of cycles seven to 10. They found that for the FUV-MAMA, the RMS deviation in the region near the center range from 3 to 4.4% in high resolution mode. They did not take out the Poisson contribution, which in most cases is \(\sim 1\%\) or less. However, they conclude that the RMS is from 10 to 90% higher than one would expect if the only limitation on the P-flats was the Poisson noise. They also conclude that the P-flats are likely to be somewhat sensitive to the mode and central wavelength used, and that the time evolution they note is real.

Our analysis, using the P-flat generated from the Cycle-17 data, shows that the RMS in the region near the center varies from 1 to 3% (after the Poisson noise contribution has been removed) for the high resolution mode. This is in agreement with the findings of Brown & Davies (2002) and is therefore expected. The variation in the low resolution P-flat between the last cycle and the new P-flat and the current reference file and the new P-flat is about 0.8 and 0.9%, respectively. As the calibrated STIS data are delivered, in most cases, in low resolution format, we can conclude that the time evolution since STIS broke has been as expected and the RMS change is less than one per cent.

Figure 2 shows a ratio between the P-flat generated from the Cycle-17 data and a P-flat generated from program 8428. The image is in low resolution format and shows that the RMS deviations are about 1.5%. Note however that the Poisson contribution has not been removed and is fairly high, as neither of the P-flats have \(SNR \geq 100\). No clear structure was noticed in any of the differential P-flats. We therefore conclude that the post-SM4 P-flat does not show any new structure and the time evolution compared to the old P-flats and to the reference file in the pipeline is as expected. As the \(SNR\) of the Cycle-17 P-flat is not enough to support science that requires a very high \(SNR\) we, at this point, choose not to replace the high \(SNR\) P-flat reference file used in the pipeline calibration.
Figure 1: Example of a STIS FUV-MAMA P-flat. Note the strong hexagon and Moiré patterns visible.

Figure 2: Differential STIS FUV-MAMA P-flat. The image shows a ratio between the new P-flat and an old flat produced from program 8428 data.
4.2. NUV MAMA

We do not yet have an estimate for the time evolution in the NUV flat, as we are still in the process of obtaining Cycle-17 data. However, a new P-flat will be generated from Cycle-17 data as soon as the data become available and it will be compared to the reference file currently in the pipeline, as well as, to the old reference files. Since the FUV-MAMA did not show any structural changes, we do not expect major time evolution in the NUV-MAMA flat field structure either.

4.3. CCD

New STIS CCD P-flats (see Fig. 4) were produced in November 2009 after which we thoroughly tested them. For testing the new P-flats we used post-SM4 data (GO proposal 11567, P.I. Proffitt) that were taken with the G230MB grating through a long slit 52 × 0.2. This data set is well suited for testing as these data are of high signal-to-noise (SNR ∼ 300) and moreover these observations were dithered along the slit. The observations used are of variable stars of beta Cep type (HD024760 and HD109668). There does not appear to be any detectable photometric or velocity variation between the individual exposures. The 1-dimensional extracted spectral stripes, which are all of the same target, had been recorded on different places on the detector and should therefore tell us readily if the P-flat correction is appropriate by minimizing the scatter between the spectra at different detector positions. We also compared the extracted 1-dimensional spectra (1dx) that were calibrated with the new P-flat to a situation where no flat fielding was performed to assess whether the P-flat can improve the signal-to-noise ratio of the data or not, and to ensure that the P-flat does not generate any spurious features.
Figure 4: Example of a STIS CCD P-flat. The dust motes and small scale blemishes have been marked with red circles.

Figure 5 shows a ratio between the new and the old P-flat based on data taken in 2003. No systematic differences are noted in the ratio image and the differences are in general less than ±1%. We also made a ratio image of the new and the old error arrays to assess the signal-to-noise ratio of the new P-flat. The new P-flat was found to have slightly (in general within few %) larger errors across the detector. This difference is due to the number of files, i.e. the number of electrons recorded, used to generate the new P-flat, which is slightly lower than for the old P-flat. Even so, the signal-to-noise ratio of the new P-flat is high enough so that it can be used for high signal-to-noise data without degrading the data quality.

As no systematic and unexplainable differences between the old and the new P-flat were noted, we tested the new P-flat on high signal-to-noise post-SM4 data. Because no accurate stellar models are available for HD024760 or HD109668, we used the spectrum observed at the last dither position as a template for the testing. All values quoted are always with respect to the template spectrum so only relative differences should be compared.

The new P-flats can reduce the residual scatter and the RMS deviation up to ∼22% in comparison to the old reference files when high signal-to-noise ratio spectra recorded at different dither positions are compared. No spurious features were detected when the new P-flats were used. The new P-flat was found to have only a modest impact on the signal-to-noise ratio of the extracted 1-dimensional spectra; adopting the new P-flat improved the signal-to-noise ratio of the extracted spectra < 1% compared to the old P-flat or if no flat fielding was applied. The improvements on high signal-to-noise data range from modest to moderate, numerical values are listed in Table 1. The RMS deviation in extracted 1-dimensional spectra (x1d) is from ∼14 to up to ∼24% smaller in case of the spectra that were calibrated using the new P-flat instead of the old reference file. The RMS deviation was found to be from ∼5 to ∼19% smaller when extracted 1-d spectra were calibrated with the new P-flat in comparison to the situation where no flat fielding was performed. Thus, it was found that the old P-flat, in case of post-SM4 data, actually increased the scatter.
relative to not applying any P-flat. In general, we find that the new P-flat reference file can reduce the scatter between the spectral stripes of the same target that were recorded on different detector positions by a moderate amount.

Table 1: The improvements on high signal-to-noise data with the new P-flat.

<table>
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<th>Root Name</th>
<th>RMS Deviation</th>
<th>Smaller (%)</th>
<th>SNR</th>
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<td>14.3</td>
<td></td>
<td>318.9</td>
</tr>
</tbody>
</table>

Figure 6 shows an example of a high signal-to-noise spectra that have been calibrated using the new and the old P-flat reference files while all other calibration flags have been kept the same. Two different objects are present in the figure, thus, shapes of the line profiles and flux levels are not identical. The first spectrum was used as a template and the remaining spectra were linearly interpolated to the same wavelength scale as the template. The wavelength region from \(\sim 2028\) to \(\sim 2032\) Å is typical and shows how the new P-flat reference file (top panel) reduces the scatter of the different extraction positions shown with different colors.

We can also estimate the signal-to-noise ratio of the extracted 1-dimensional spectra using the residuals. The signal-to-noise ratio of the spectra that had been calibrated using the new P-flat reference file was found to be marginally higher than in case of the old
reference file. However, for the high signal-to-noise data used for testing, the difference was found to be $\leq 1\%$ in all cases. Even so, one should keep in mind that a robust estimation of the signal-to-noise ratio is more complicated than implemented in our simple recipe. Thus, the actual impact may be larger than presented here. Finally, after thorough testing, the new STIS CCD P-flat field reference files were delivered to CDBS in December 2009 and adopted for the pipeline calibration.

5. Future Activities

In future Cycles the COS/STIS Team will most likely continue to monitor the flat field structure and time evolution in all three STIS detectors. For the CCD it is likely that the Team will continue to produce new P-flats each Cycle. For the MAMAs the Team will likely want to continue to monitor the time evolution. However, due to the relatively small time evolution, at least in the FUV-MAMA P-flats, it may not be necessary to produce new P-flats every Cycle. The primary reason not to make a new P-flat every Cycle is that both the Kr and D$_2$ lamps fade significantly with use. It would be unfortunate to end up in a situation during the expected lifetime of STIS that the lamp does not produce enough photons to make new P-flats if it were noted that the changes in the flat field have become significant. Hence, monitoring but not producing high $SNR$ P-flats may be advisable in case of the MAMAs.

References

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Calibration of B,V,R,I Photometry for the Wide Field Channel of the HST Advanced Camera for Surveys

Richard Shaw

NOAO, Tucson, AZ

Abstract.

We present new observations of two Galactic globular clusters, PAL4 and PAL14, using the Wide-Field Channel of the Advanced Camera for Surveys (ACS) on board the Hubble Space Telescope, and reanalyze archival data from a third, NGC2419. We matched our photometry of hundreds of stars in these fields from the ACS images to existing, ground-based photometry of faint sequences which were calibrated on the standard B,V,R,I system of Landolt. These stars are significantly fainter than those generally used for HST calibration purposes, and therefore are much better matched to supporting precision photometry of ACS science targets. We were able to derive substantially more accurate photometric transformation coefficients for several commonly used ACS filters compared to those published by Sirianni, et al. (2005), owing to the use of a factor of several more calibration stars which span a greater range of color. We find that the inferred transformations from each cluster individually do not vary significantly from the average, except for a small offset of the photometric zeropoint in the F850LP filter. Our results show that the published prescriptions for the time-dependent correction of CCD charge-transfer efficiency appear to work very well over the 3.5 yr interval that spans our observations of PAL4 and PAL14 and the archived images of NGC2419.
Post-SM4 ACS/WFC Bias Striping: Characterization And Mitigation

N. A. Grogin, P. L. Lim, and A. Maybhate

Space Telescope Science Institute, Baltimore, MD 21218

R. N. Hook

Space Telescope European Coordinating Facility, D-85748 Garching, Germany

M. Loose

Markury Scientific, Thousand Oaks, CA 91360

1. Introduction

The Advanced Camera for Surveys (ACS) CCD Electronics Box Replacement includes a SIDECAR Application-Specific Integrated Circuit (ASIC) that exhibits a low frequency noise (1mHz to 1Hz) on the bias and reference voltages it generates for the Wide Field Channel CCDs (Loose 2011).

This noise contribution, first seen in thermal-vac testing, does not matter for bias voltages going to the CCD since it is cancelled out by correlated double sampling (CDS). However, there is one reference voltage from the ASIC that is used to offset the signal, applied after the CDS stage. Here, the noise does not cancel out, and manifests as a slow moving variation of the baseline.

In practice, we observe a “striping” noise in all post-SM4 Wide Field Channel (WFC) images, virtually uniform across both amplifier readouts (the entire 4096 columns) of each WFC CCD. An example of this phenomenon is shown in Figure 1, where we have taken a post-SM4 bias frame and subtracted the (~stripe-free) calibration superbias to remove fixed bias structure and thereby highlight the striping noise.

2. Characterization

Because of the uniformity of the striping across WFC rows (standard deviation between amplifiers on the same CCD is at the 0.3e− level), it is straightforward to characterize and remove this low-level 1/f noise from individual WFC bias frames, as seen in Figure 2. The amplitude distribution of the striping noise, shown in Figure 3, is well approximated by a Gaussian with $\sigma_G = 0.75e^−$: less than 20% of the WFC readnoise. The distribution has an enhanced negative tail that results in an overall standard deviation of $\sigma \approx 0.9e^−$. Figure 4 illustrates the nearly 1/f power spectrum (crosses), as averaged from the striping patterns extracted from an ensemble of 16 post-SM4 bias frames. The corresponding WFC readnoise power spectrum is represented in this figure (blue dots) by an ensemble of 16 simulated white-noise-only images with $\sigma = 4e^−$. We superpose an arbitrarily normalized 1/f profile (dotted red line) for comparison.

Averaging $N$ bias images diminishes the 1/f noise nearly as $N^{-1/2}$ (Fig. 5, dashed blue line). A non-linear power-law fit to the noise dampening (Fig. 5, dotted red line) yields a
best-fit relation of:

\[ \frac{d(\log_{10} \sigma)}{d(\log_{10} N)} = -1/\left[2.1562 - 0.1408 \log_{10} N + 0.1700 (\log_{10} N)^2\right]. \]

Because of this dampening, the overall noise in the post-SM4 WFC superbias calibration images approaches pre-2007 levels without the need for frame-by-frame stripe removal.

We have investigated the long-term stability of the bias striping characteristics by analysing subsets of biases segregated according to the \( \sim \) monthly ACS anneal cycle. During the first year of post-SM4 WFC operation, the histogram of bias striping amplitude appears to be highly stable. Figure 6 plots the Gaussian width of the striping noise from month to month over the span July 2009 – June 2010. There is very little change, and at the \( \pm 0.02e^{-} \) level. Such variation is negligible when plotted relative to the \( 4e^{-} \) readnoise.

The bias striping is not well estimated by the limited WFC overscan regions alone, complicating stripe removal from individual non-bias frames such as science exposures and the calibration darks. A WFC superdark calibration image (average of \( \approx 24 \) individual dark exposures) is effectively stripe-free. However ACS/WFC science programs rarely comprise so many frames (at a fixed orientation on the sky) to average away the striping noise. Image structure along the Y-axis in a science (or dark) exposure runs the risk of being mistakenly removed along with the bias stripes.

3. Mitigation Methodology

We designed and tested three different codes for automated bias-stripe removal from science images, each performing row-by-row correction with different algorithms for modeling the stripes in the presence of increasingly complex astronomical scenes.

The first method was a minimally altered translation of the code already made available by STScI to correct a qualitatively similar bias anomaly appearing in later WFPC2 WF4 images (Biretta & Gonzaga 2005). This “widestreak” algorithm, described in detail by Maybhate et al. (2008), iteratively created a global mask of presumed non-sky pixels, then proceeded with a row-by-row determination of the non-masked pixel mean, and finally corrected each row by the difference between the row mean and the global mean of unmasked pixels.

The second method treated each row of the CCD entirely independently, and determined the sky-stripe level of a given row by fitting a Gaussian to the \( \sigma \)-clipped histogram of pixel intensities along the row. Each row was then corrected so that all had the same Gaussian-fit centroid.

The third method also used a row-by-row sky fitting algorithm, but in this case the sky-stripe level was determined in a manner similar to SExtractor’s (Bertin & Arnouts 1996): iterative \( \sigma \)-clipping to reject non-sky pixels, then a hybrid mean & median estimator applied to the remaining pixels to compensate for slight distortion of the pixel histogram due to low-level flux from astronomical sources and cosmic-ray hits. In detail, the algorithm performed \( \leq 15 \) iterations of \( 2\sigma \) clipping, and then determined the sky-stripe level as \( 2.5 \times \text{Median} - 1.5 \times \text{Mean} \).

Our test suite comprised four pre-failure WFC programs spanning a broad range in sky level and scene complexity. These include the large galaxy NGC 1300 (Fig. 7), the crowded stellar field 47 Tucanae (Fig. 8), and long-exposure “blank field” images with wide-band filters (Fig. 9) and narrow-band filters (Fig. 10). We seeded each of the pre-failure images (from 8 to 40 in the four programs) with 30 independent realizations of the striping noise as measured from post-SM4 bias frames. To better assess the stripe recovery as a function of background noise level, we repeated the simulations with twice the signal level in the pre-SM4 frames — treating these GAIN=1e^{-}/DN images as 2e^{-}/DN iamges. Figure 11 shows the total range of image complexity probed by our simulations, as summarized by \( \sigma \)-clipped image mean and standard deviation. Because the post-SM4 ACS/WFC readnoise
Post-SM4 ACS/WFC Bias

(dotted line in Fig. 11) is lower than the pre-failure readnoise, our simulations were unable to probe stripe removal effectiveness all the way down to current readnoise levels, even with short exposures in narrow-band filters.

4. Mitigation Testing

We rated performance of striping mitigation by comparing the standard deviations of the residual (improperly-corrected) striping for each algorithm. Figure 12 shows the results for each of the four astronomical scenes, plotted against image complexity as summarized by the scene standard deviation. In these plots, method #1 results are shown in red, method #2 in green, and method #3 in blue. In some cases, an algorithm identically recovered 30 different stripe seedings of a given image to the same fidelity, resulting in pile-up of symbols on Figure 12.

Overall the top-performing code, and the one we have chosen for public release to the ACS/WFC user community, was method #3. Not surprisingly, the best results were achieved for the narrow- and wide-band “blank field” images, where the injected striping noise was reduced to $0.3e^-$ and $0.4e^-$, respectively. The CCD-spanning “large galaxy” case had uneven results, worsening the scatter to $2e^-$ in half the trials. All the codes failed badly for the crowded stellar field 47 Tuc, often worsening the scatter to $\gg 10e^-$. Due to these instances of poor performance, de-striping is not yet incorporated into the automated ACS calibration pipeline (“calacs”). In September 2010, STScI is making available the stripe-removal code as a stand-alone Python script.

5. Conclusions

Although the eye easily picks out the row-correlated striping in post-SM4 ACS/WFC images, we emphasize that the standard deviation of this striping noise is $< 25\%$ of the WFC readnoise. Our post-SM4 observations of WFC photometric calibration fields show no measurable impact upon aperture photometry of compact sources with local background estimation.

In the case of science programs where low-level, row-correlated noise may be a severe systematic contamination, such as weak-lensing or faint-isophotal analyses, we recommend the use of the STScI-tested de-striping algorithm. This algorithm may also be generically useful for cosmetic improvement of image mosaics, as can be seen for the large galaxy NGC 6217 observed in the H$\alpha$ filter (Fig. 13) as the post-SM4 first-light target.

References

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Figure 1: A single post-SM4 ACS/WFC bias frame, which has been superbias-subtracted to remove fixed bias structure. The linear stretch shown at bottom is in units of DN, for gain setting 2.0 e⁻/DN. The row-correlated 1/f “striping noise” is easily detected by the eye, though the standard deviation of this noise is less than 25% of the pixel-to-pixel read noise.
Figure 2: Identical post-SM4 bias frame as shown in Figure 1, but now with the de-striping algorithm applied. Given the near-perfect stripe removal, only pixel-to-pixel read noise remains, along with a few bad columns and sparse, CTE-trailed cosmic ray hits that occurred during read-out.
Figure 3: (Left) Stripe intensities from 319 ACS/WFC bias frames (4096 stripes in each) in the first year after SM4. The best-fit Gaussian to the peak (red dashed line) has a width of $\sigma_G = 0.74e^-$. The full distribution, with negative skewness, has $\sigma \approx 0.9e^-$. 

Figure 4: (Right) Power spectrum of the Y-axis bias striping, averaged from 16 bias frames. Simulated white-noise spectrum matched to the $4e^-$ WFC readnoise is shown as blue dots.

Figure 5: (Left) Data points illustrate the $\approx N^{-1/2}$ reduction in striping noise measured from $N$ averaged post-SM4 bias frames. The STScI calibration superbiases and superdarks currently mitigate striping noise via large-$N$ averaging, rather than frame-by-frame stripe removal.

Figure 6: (Right) Long-term stability in the Gaussian width of the striping amplitude, $\sigma_G$, as measured from bias frames in successive $\sim$monthly ACS anneal cycles following SM4.
Figure 7: Representative ACS/WFC large galaxy field (NGC 1300) exposures, from pre-failure Program 10342, used in the testing of optimal de-striping algorithms. The upper two exposures are 380 sec with F435W (left) and 340 sec with F814W (right). The lower two exposures are 380 sec with F658N (left) and 340 sec with F555W (right).
Figure 8: Representative ACS/WFC crowded stellar field (47 Tuc) exposures, from pre-failure Program 10775, used in the testing of optimal de-striping algorithms. The upper two exposures are 3 sec (left) and 50 sec (right) with F606W. The lower two exposures are 3 sec (left) and 50 sec (right) with F814W.
Figure 9: Representative ACS/WFC wide-band blank field exposures, from pre-failure Program 10809, used in the testing of optimal de-striping algorithms. The upper two exposures are 513 sec (left) and 514 sec (right) with F606W. The lower two exposures are 513 sec (left) and 514 sec (right) with F814W.
Figure 10: Representative ACS/WFC narrow-band blank field exposures, from pre-failure Program 10843, used in the testing of optimal de-striping algorithms. The upper two exposures are 357 sec with F660N. The lower two exposures are 1437 sec with F658N.
Figure 11: Image complexity among the astronomical scenes chosen to test stripe-removal algorithms. Simulations were run twice, at (actual) GAIN=1e−4/DN and at 2e−4/DN.

Figure 12: Fidelity of recovered bias striping for the three estimation methods (labeled red; green; blue) described in the text, measured as standard deviation of residual striping for high-noise (squares) and low-noise (crosses) realizations of: a large galaxy (upper left); a crowded stellar field (upper right); a wide-band blank field exposure (lower left); and a narrow-band blank field exposure (lower right). The standard deviation of the simulated striping, $\sigma = 0.9e^-$, is shown as the horizontal dotted line in each plot.
Figure 13: The ACS/WFC post-SM4 first-light mosaic, galaxy NGC 6217 in F658N (Hα), with extreme stretch to contrast corrected (right) and uncorrected (left) bias striping.
ACS/WFC Crosstalk

A. A. Suchkov\textsuperscript{1}, N. A. Grogin, \textsuperscript{2} M. Sirianni\textsuperscript{3}, E. Cheng\textsuperscript{4}, A. Waczyński\textsuperscript{5}, \& M. Loose\textsuperscript{6}

Abstract. The ACS/WFC detector consists of two CCDs, each of which is read out through two amplifiers. While reading each quadrant of the detector, the electronic crosstalk between the amplifiers induces faint, typically negative, mirror-symmetric ghost images on the other three quadrants. The effect is strongest for high-signal offending (source) pixels. Analysis of the pre-SM4 crosstalk showed that its impact on ACS/WFC science was insignificant and can be ignored in most science applications. In this report, we analyze crosstalk after SM4. Crosstalk due to low-signal offenders is much weaker than before SM4 and does not produce ghosts similar to those seen in pre-SM4 images. For high-signal offending pixels, we find substantial differences between the gain=1 $e^-$/DN and gain=2 $e^-$/DN cases. For the default gain setting of 2, the crosstalk is similar to what it was before the SM4, up to 5–8 $e^-$ per pixel on the same CCD. For gain=1, the crosstalk is 100 $e^-$ per pixel for saturated offending pixels on the same CCD, which is more than an order of magnitude above the pre-SM4 level. The crosstalk from saturated pixels is 20–30 $e^-$ per pixel on the other CCD, which is also much higher than it was before SM4. For gain=1.4, the crosstalk is very similar to that at gain=2.

1. Introduction

The ACS/WFC detector has four amplifiers, through which the four quadrants of the detector are read separately and simultaneously. As the quadrants are read out, electronic crosstalk between the amplifiers can be induced. As a result, an imaged source in one quadrant may appear as a faint, mirror-symmetric ghost image in the other quadrants. The ghost image is often negative, so bright features on the “offending” quadrant show up as dark depressions on the “victim” quadrants. Giavalisco (2004a, 2004b) analyzed the pre–SM4 ACS/WFC crosstalk using image frames of external targets. He found that crosstalk ghosts from extended offending sources are observed as faint, negative images, $\sim 2$ $e^-$ relative to the background, characterized by approximately constant surface brightness (here and below the ghost and offending signal values are given as electrons per pixels). He also noticed that sources in some quadrants apparently are more effective in generating ghosts than sources located in other quadrants, and some quadrants are more susceptible to crosstalk than others are.

\textsuperscript{1}Johns Hopkins University, Baltimore, MD 21218
\textsuperscript{2}Space Telescope Science Institute, Baltimore, MD 21218
\textsuperscript{3}European Space Agency/ESTEC, N2200 AG Noordwijk, Netherland
\textsuperscript{4}Conceptual Analytics LLC, Maryland
\textsuperscript{5}NASA, GSFC, Maryland
\textsuperscript{6}Markury Scientific Inc., California
Crosstalk turned out to be stronger for high-signal offending pixels, and gain=1 images were found to be more affected by crosstalk than images taken at gain=2. In general Giavalisco (2004a, 2004b) concluded that crosstalk ghosts have very little impact on ACS/WFC photometry and can be ignored in most cases, especially if the observations are performed using gain=2. Therefore, the gain=2 setting was recommended as a preferable observation mode as it mitigates the crosstalk effects on ACS/WFC data (Giavalisco 2004b). In this paper we analyzed ACS/WFC crosstalk after Servicing Mission 4 in May 2009, when the ACS/WFC electronics was replaced (see also Suchkov et al. 2010).

2. Crosstalk in new electronics ground tests

The performance requirements for the new ACS-R electronics identified crosstalk as one of the limiting criteria for the CCD Electronics Box (CEB). Crosstalk was carefully tested at multiple stages of the ACS-R program. Special test equipment was developed to facilitate crosstalk testing at early stages of electronics development by simulating the CCD video signal with arbitrary image patterns. In complex multichannel instruments like ACS, the primary sources of crosstalk are signal couplings through circuit board layout, wiring, emission, common supply lines and grounding. All these aspects were carefully considered during the design phase. It was a challenging task because crosstalk parameters are difficult to model. Depending on the coupling mechanism, crosstalk can manifest itself as a ghost image of either positive or negative magnitude, and may not be proportional to the source signal. It can also produce additional noise when digital lines couple to the analog sections of signal processing chain. Coupling from digital lines can also produce crosstalk that is not proportional to signal magnitude but shows up in a particular spatial location or data value. Crosstalk had been measured in engineering models of the ACS-R CEB and verified to comply with requirements. During the instrument integration and testing, it was observed that crosstalk can be strongly amplified when the signal significantly exceeds the normal operating range. Special circuitry was developed and implemented to limit signal levels at sensitive interfaces to minimize this crosstalk source. The final test results showed that the crosstalk stays within the requirements specification, below the 0.1% level for ghost-to-signal ratio. Note that the ACS-R program, unlike normal flight projects, was unable to verify full instrument performance during ground testing because the detector assembly was aboard HST and unavailable. Instead, a substitute detector (the ACS5 flight-spare unit) was used for final performance verification. It is therefore possible that the observed on-orbit crosstalk may differ from ground test results due to differences in wiring harnesses, pre-amplifier boards, and detector characteristics.

3. Quantifying crosstalk in ACS/WFC images

We characterized crosstalk using ACS/WFC dark frames and external images obtained during the Servicing Mission Observatory Verification (SMOV) that took place after SM4. Post-SM4 science observations with ACS/WFC are performed with the default gain=2, which was thoroughly tested and characterized during SMOV. We had a substantial number of gain=2, 1000 sec dark frames and a few dark frames for gain=1.4. Only a limited number of gain=1 images of external targets were obtained in the SMOV crosstalk assessment program (ID 11371). Therefore, in this paper the crosstalk characterization for gain=1 is more limited than for gain=2. In dark frames, the offending source pixels are produced by cosmic rays and hot pixels. In external data, crosstalk is generated by the observed target, cosmic rays, and hot pixels. A ghost overlapping a target image on a victim quadrant is hard to isolate, because it is difficult to distinguish brightness variations due to the ghost and the target itself. A more reliable and much easier way to isolate and quantify crosstalk is to measure ghost images against the sky or dark background in external or dark
frames, respectively. Consequently, we used the following analysis strategy: In each detector quadrant, we first isolate the background, which we define as an area with pixels below a pre-set value. This value is selected using the quadrant’s pixel distribution. Frames suitable for crosstalk characterization have pixels that follow a quasi-gaussian distribution peaked near the mean of the distribution. From the pixel histogram, we identify the lower limit of pixel values associated with imaged objects and then compute the mean and standard deviation of all pixels below that limit. The pixels within $\pm 2\sigma$ of the mean are then used to compute the mean background within each quadrant against which crosstalk ghosts are measured. Next, we identify in all quadrants the offending pixels within a specified signal range and map them onto the backgrounds of the victim quadrants. Finally, we compute the mean pixel values in the mapped area of each victim quadrant, and calculate the crosstalk ghost as the difference between the victim quadrant’s mean background and the mean of the mapped area.

4. Results

Unlike the distinct crosstalk ghosts seen in pre-SM4 images, no ghosts from low-signal sources were seen in post-SM4 gain=2 and gain=1 images. We inferred that the new ACS/WFC electronics are less susceptible to negative crosstalk at low signal levels. We detected crosstalk from high-signal source pixels in post-SM4 images obtained with both gain=1 and gain=2. We found remarkable differences between the two gains as we analyzed ghosts due to saturated stellar images. For gain=1, the ghost are $\sim 100 \, \text{e}^{-}$ in the quadrant on the same CCD, whereas 30–40 $\, \text{e}^{-}$ ghosts appear in the quadrants on the other CCD. These values represent a dramatic increase from the pre-SM4 level.

At this point we can only speculate why the gain=1 crosstalk is highly elevated. It may be caused by saturation of the pre-amplifier internal to the SIDECAR ASIC. At gain=1, the pre-amplifier outputs are highly saturated for large input signals, which substantially increases the power consumption and enhances crosstalk. However, we found that unsaturated pixels above $\sim 25000 \, \text{e}^{-}$ also produce ghosts, albeit at a lower level; we estimated that the ghost effect for high-signal sources below saturation is about $\sim 10 \, \text{e}^{-}$. It is possible that improvements in the gain=1 crosstalk can be implemented after suitably designed on-orbit testing.

Because gain=2 is the more common mode of operation, it has been more extensively tested on the ground and on-orbit. Unlike the gain=1 case, there are no significant changes in the gain=2 high-signal crosstalk after SM4. It is observed only among quadrants on the same CCD, as noted in pre-SM4 data, and its depth of 5–8 $\, \text{e}^{-}$ is similar to that before SM4. The crosstalk signal dependence is essentially linear, with post-SM4 slopes in the range 5e-5 to 7.5e-5. Crosstalk is about two times higher when the offender is the CCD’s left quadrant. It is seen only between the quadrants on the same CCD and becomes noticeable only for signals above 20000 $\, \text{e}^{-}$, with a maximum of 5–8 $\, \text{e}^{-}$ for signals near the pixels full-well depth. This is seen in Table 1, which lists ghost values (and their errors) as obtained from 18 dark frames.

The ghost estimates become slightly better after the frames are corrected for a quasi-regular bias spatial noise detected in post-SM4 bias frames. The noise represents a low-level signal ($\sigma \leq 1 \, \text{e}^{-}$) slowly varying across the CCDs rows but remaining homogeneous across the columns, and is observed on bias frames as stripes running parallel to the CCD’s rows. Because of the uniformity of the striping across the CCD rows, it is straightforward to remove bias stripes (de-stripe) from WFC bias or dark frames by performing row-by-row correction with a simple algorithm for minimizing the mis-estimation due to bias structure. Table 1 shows the results for both original and de-striped frames.
Table 1: Post-SM4 crosstalk at gain=2 for offending signal > 65000 e−.

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<thead>
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<th>average error, e⁻</th>
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</table>

5. Conclusions

- Negative crosstalk due to low-signal sources is much weaker than it was before SM4 and can be safely ignored, both for gain=1 and gain=2.

- For the current default gain=2, the post-SM4 crosstalk induced by high-signal pixels remains at pre-SM4 levels (ghost-to-signal ratio ∼ 0.01% for high-signal offending pixels) and thus can be ignored in most scientific applications of ACS/WFC.

- The gain=1.4 crosstalk is very similar to that of gain=2.

- The gain=1 crosstalk due to high-signal sources increased dramatically after SM4 and is now ~10 times stronger than the gain=2 crosstalk.

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References

The NIRSpec Calibration Concept

Guido De Marchi

European Space Agency

Abstract.
NIRSpec is the main near-infrared spectrograph on board the James Webb Space Telescope, a segmented off-axis telescope in deep space. Besides the traditional ‘fixed slits’ for long-slit spectroscopy, NIRSpec offers both integral field unit and multi-object spectroscopy capabilities. The highest level of multiplexing is provided by the Micro-Shutter Array, a fixed grid of micro-shutters allowing up to 100 objects to be observed simultaneously over a large field of view (10' square) and over nearly a factor of 10 in wavelength, since NIRSpec’s detector response ranges from 0.6 to 5 micron. The combination of these factors results in a number of unique challenges for an efficient calibration of the instrument. In this paper we present a high-level description of the calibration and outline some of the challenges that it entails.
NIRSpec - the JWST Multi-Object Spectrograph

Guido De Marchi
European Space Agency

Abstract.
The Near-Infrared Spectrograph (NIRSpec), to be launched in 2014 on the James Webb Space Telescope, will be the first slit-based multi-object spectrograph used in space, and is designed to simultaneously provide spectra of faint objects in the range 0.6 - 5.0 micron at resolutions of $R=100$, $R=1000$ and $R=2700$. The all-reflective wide-field optics, a novel MEMS-based programmable micro-shutter array slit selection device, and large format low-noise HgCdTe detectors combine to allow simultaneous observations of $\sim 100$ objects within a 3.4x3.4 arcmin$^2$ field. A 3x3 arcsec$^2$ integral field unit and five fixed slits are also available. The NIRSpec sensitivity is expected to allow detection of a continuum flux of 20 nJy (AB$\sim$28) in $R=100$ mode, and a line flux of $6 \times 10^{-19}$ erg s$^{-1}$cm$^{-2}$ in $R=1000$ mode, both at S/N $\sim 3$ in 104 seconds. NIRSpec is being built for the European Space Agency by EADS Astrium as part of the European contribution to the JWST mission, with the micro-shutter and detector arrays being provided by NASA/GSFC.
Dithering with JWST’s NIRCam

Jay Anderson

Space Telescope Science Institute, Baltimore, MD 21218

Abstract.

HST allows users great flexibility in designing their own dither patterns, but JWST will follow a different model. Users will have to select their dither pattern from a set of pre-designed templates. With few exceptions, all observations will be required to dither, and no two exposures will be taken through the same filter at the same pointing. This report summarizes the dithering strategies that have been developed for NIRCam, JWST’s workhorse near-IR camera.

1. Introduction

Users of HST have great flexibility in designing their dither patterns. They can decide to make use of the canned patterns provided by the Institute, or they can design their own dither patterns. They even have the flexibility of not dithering at all, which can make cosmic-ray rejection easier but does nothing to mitigate defects.

While many HST programs have benefitted from this dithering flexibility, many other programs have suffered from insufficient attention to dithering. The archive in particular has suffered both from programs that chose not to dither and from programs that were not dithered in a way that the association-oriented pipeline can deal well with.

Except in rare circumstances, such as exo-planet-transit studies, users of JWST will be forced to dither. This means that no two images will be taken at the same pointing with the same filter. This compulsory dithering will allow automatic mitigation of bad pixels and will allow improved sampling and better field coverage (by imaging over the gaps). This must-dither requirement will help ensure a high-quality archive and a better pipeline product.

The responsibility for the users, then, is to determine which of the canned patterns meets their needs. This will involve various compromises between sub-pixel sampling, defect avoidance, L-flat mitigation, gap coverage, etc. The responsibility for the Institute is to ensure that the patterns provided are sufficient to accomplish a wide variety of scientific goals and also to ensure that the users have the information needed to choose the right pattern for their science.

2. NIRCam

NIRCam is the workhorse near-IR imager for JWST. It has been designed by Marcia Rieke at the University of Arizona and is being built by Lockheed-Martin in Palo Alto. It will have a di-chroic to allow simultaneous observations in two channels. The short-wave channel (SWC) will sense from 0.7 to 2.4 $\mu$m with 0.034" pixels. The long-wave channel (LWC) will sense from 2.4 to 5 $\mu$m with 0.065" pixels. The SWC will be Nyquist-sampled at 2 $\mu$m and undersampled blueward of this. It will be severely undersampled at 0.7 $\mu$m. NIRCam will also have a set of defocusing lenses, which will allow it to sense the telescope’s wavefront error so that JWST’s mirrors can be kept in near-perfect alignment.
Figure 1: The NIRCam SWC footprint.

NIRCam’s field will span $5.17' \times 2.25'$ with 10 detectors: 8 in the SWC and two in the LWC. The field will be divided between two identical modules (A and B) with a $40''$ gap between them. The four SW detectors in each module will have small $5''$ gaps between them. Figure 1 shows the NIRCam SWC footprint.

3. Choices in dithering

The choice of dithering is only one of many decisions that must be made. The first concerns the target and the desired S/N. For an input S/N, the ETC will report the necessary total exposure time. This total exposure time will often have to be broken up into several independent exposures, either because it exceeds the maximum $4000 \text{ s}$ for a given exposure or because dithering will enhance the science.

Once users determine how many different exposures to take, they must then choose a dither pattern. The dithers must be divided among primary dithers and secondary dithers. **Primary** dithers involve moving the target an appreciable distance across the detector. They are used to mitigate field lost to the gaps, L-flat errors, detector defects, etc. **Secondary** dithers involve moving the target a very small amount. These small dithers are able to carefully place the target at a variety different locations relative to the pixel grid, so that a reconstructed stack can sample the scene better than a single exposure. The secondary dithers are small enough that the pixel-phase coverage will remain coherent across the entire $5' \times 2'$ field.

The dither pattern will be specified by the number of primary dithers ($N_P$) and the number of secondary dithers ($N_S$). A full set of secondary dithers will be executed about each of the primary offsets. As such, the total number of exposures $N_T$ will simply be

$$N_T = N_P \times N_S.$$ 

If the number of possible primary and secondary dithers is quantized too coarsely, then the total number of exposures will be quantized even more coarsely, which could have a severe impact on our ability to plan an optimal set of observations. We therefore must design primary and secondary patterns with as much granulation as possible. This will be one of our aims in constructing useful patterns.
4. Primary Dithers

There will be three general strategies for Primary dithers, depending on the size of the field of interest. We can do (1) full-field primary dithers, (2) single-module primary dithers, or (3) single-chip primary dithers.

4.1. Full-Field Primary Dithers

The main priority of the full-field primary dithers will be to image over the gaps, both the 40′′ gap between modules and the 5′′ gap between SWC chips. (See Figure 1.). Since the NIRCam footprint has both vertical and horizontal gaps, it will take at least three dither positions to image over them and ensure that a given point in the field will be covered by at least one image.

In many deep-imaging programs, such as the GOODS, COSMOS, or the UDF, we will want to achieve much more than single-exposure coverage, so we have designed primary patterns with \( N_P \) between 3 and 45 dithers that will aim to cover the 5′×2′ NIRCam footprint as evenly as possible. The coverage for the first three patterns is shown in color in Figure 2. The 3-point pattern gets 71% of the field in 2/3 exposures, 18% in 3/3 and 11% in 1/3. The 9-point pattern gets 14% of the field in 5/9, 52% in 6/9 and 34% in 7/9.

In Figure 3, we show how nicely the patterns inter-lock when the pointings are tiled together for mosaics. To provide just one example, the 45-point pattern covers 85% of the field with between 30 and 32 out of the 45 exposures. This is extremely even coverage, and should make for a very uniform data set. The SWC detectors cover 69% of the footprint, so the best we can hope for is even coverage in 31 exposures. Dither patterns have been devised with \( N_P \)'s of 3, 6, 9, 15, 21, 27, 35, and 45.

4.2. Single-Module Primary Dithers

A second set of primary dithers has been designed to image objects that fit within the roughly 2′×2′ field of a single module. The dithers will move the target around the center of the FOV, ensuring that a given point in the field will not fall in the 5′′ gap in too many exposures. One can choose any number of \( N_P \) for this, from 1 to 16.
Figure 3: The mosaic coverage achieved for the eight available full-field dithers.
4.3. Single-Chip Primary Dithers

The third set of primary dithers is designed to image small sources — sources smaller than an individual chip. This pattern will ignore the gaps between the chips and will simply move the source around the detector. Sources that are small can afford to be moved around more than larger sources. The user can dial in anywhere from $N_p = 1$ to 25. Figure 4 shows a schematic of the pattern.

5. Secondary Dithers

Most of the filters in NIRCam’s SWC will be somewhat undersampled by the $0.034''$-spaced pixels. The LWC filters will also be undersampled blueward of 4.0 microns, but not as severely. Moving the target around the detector with large dithers will randomize the pixel-phase at which it is observed. This will improve the sampling somewhat, but only in a random-walk kind of way. It would be better to take some observations at deliberate pixel-phase offsets with respect to each other. Good pixel-phase coverage will improve significantly our ability to reconstruct the scene. That is the purpose of the secondary dithers.

The distortion in NIRCam is specified to be less than 2%, which means that the plate scale at the edge of the field will be within 2% of that at the center of the field. This means that if we have a dither of 10.5 pixels at the center of the field, the dither at the edge will be at worst 10.7 pixels. So, if we keep our secondary dithers less than 10 pixels, then the pattern should not differ by more than 0.2 pixel in pixel-phase spacings from the center of the field to the edge. On the other hand, if we dither by much more than 10 pixels, then the pixel-phase pattern achieved at the edge of the field will be qualitatively different from that at the center.

An additional goal of the secondary dithers is to mitigate the impact of single-pixel artifacts. So, in addition to the fractional-pixel dithers, we will also dither by several integer pixels to ensure that a single bad pixel will not impact a small target in more than one exposure.

If $N_S$ is less than about 10, we can fit all of the secondary dithers within a $10 \times 10$-pixel box and both maintain pixel-phase coherence across the detector and mitigate bad pixels. If $N_S$ is greater than this, then we must either give up on the bad-pixel mitigation or the pixel-phase coherence. We deal with these two cases below.

5.1. Patterns for $N_S$ from 1 to 9

We have designed secondary-dither patterns for $N_S$ of 1 to 9 that will sample the pixel-phase space as evenly as possible. The pixel-phase diagrams are shown in Figure 5. The goal of the even sampling is to ensure that no matter where a particular object falls with respect
to the pixels in the first dither, the other dithers will provide complementary high-frequency information.

Undersampling will be more severe in the SWC than in the LWC, but it will still be important to improve the LWC sampling. It is not trivial to optimize them simultaneously, since the SWC and LWC have pixels that are almost exactly a factor of two different in size. Nevertheless, we have optimized the patterns for the SWC, but have simultaneously ensured that the resulting patterns for the LWC are acceptable.

5.2. Patterns for $N_S$ from 10 to 64

When more than 9 secondary dithers are needed, we cannot confine them to a 10×10-pixel box without the source landing on the same pixels more than once. So, they must be spaced out over a larger region. This compromises the pixel-phase coherence from the center to the edge of the field, but provides continued improvement in mitigating bad pixels.

This problem can be solved in a similar way to what was done for the UDF: the dithers will be bunched into groups of 4. We take each group of four with the small-number pattern for $N_S=4$ above, which doubles the native sampling of the pixels. The dithers within each group span less than 10 pixels, so the phase remains coherent among them. The various groups get spread out over many tens of pixels, so the phase between the groups can drift across the detector. However, since the phase within each group is coherent, we are guaranteed to get excellent ×2-super-sampled coverage at worst.

For targets that are smaller than the entire NIRCam footprint, the dithers among the groups will be able to maintain more coherence, so we have offset the dithers to that we will get the evenest possible pixel-phase coverage at the target location. The accuracy of achieving dithers may be limited to 7 mas (about 0.20 of a NIRCam pixel), but there is no harm in spreading things out as well as possible.

The top panels of Figure 6 show the patterns at the center of the field for the SWC and for the LWC. The bottom panels show the total dithers (fractional pixels plus integer offset). Note that the integer shifts have been chosen to minimize the travel time through the pattern.

Users can choose any number of secondary dithers, from 1 to 64 and by executing the first $N_S$ dithers, we will get the best possible pattern.

6. Choosing the Best Pattern

The NIRCam team at STScI and at the University of Arizona will be putting together a NIRCam Dithering Handbook that will describe these patterns in detail and will provide many examples of what strategy works best for specific science objectives. The Handbook will also include a discussion of how many secondary dithers are needed to fully sample the scene. This number will in general be different for different filters: a 2-point pattern may be sufficient to sample F150W images, but an 8-point pattern may be needed to recover all the structure in F070W images.

Finally, the Handbook will contain a discussion of how various combinations of $N_P$ and $N_S$ will sample the scene. While the primary dithers will have incoherent pixel phases among them, this random coverage will provide some mitigation of undersampling, so that if one is taking a lot of primary dithers, it may not be necessary to take as many secondary dithers. Primary dithers do a better job mitigating the impact of chip gaps and large artifacts, so users will want to take as many primary dithers as possible.
Figure 5: The pixel-phase coverage for the secondary dither patterns with $N_S$ from 1 to 9. The dotted line shows the outline of a pixel and the filled circles are the pixels from the $N_S$ dithers that would fall within the bounds of that pixel. The open circles show where relative to this pixel the neighboring pixels would sample.
Figure 6: Top: the pixel-phase coverage for the 64-point general secondary dither pattern for the SWC (left) and LWC (right). Bottom: the actual dithers in the pattern, including the integer-pixel shifts.
The JWST Tunable Filter Imager and its Calibration

A. R. Martel, A. Fullerton
STScI/HIA

S. Holfeltz
STScI

R. Doyon, M. Beaulieu, D. Lafrenière
Université de Montréal

J. Hutchings
HIA

N. Rowlands
COM DEV

Abstract. The JWST Tunable Filter Imager (TFI) is a sensitive camera that shares the optical bench of the Fine Guidance Sensor (FGS). It provides JWST with the capability of imaging a 2.2' × 2.2' field-of-view through narrow-band filters over a wavelength range of 1.5−5.0 μm, with a non-functional region between 2.5 and 3.2 μm. The narrow-band filters are characterized by spectroscopic resolving powers of ∼75 to ∼120 for central wavelengths selected by the user. The TFI is also capable of high-contrast imaging with a choice of four coronagraphs and an aperture mask. Here, we provide a brief summary of the technical specifications and scientific capabilities of the TFI and present an overview of the upcoming ground calibration and on-orbit commissioning activities.

1. Description and Capabilities

The primary components of the integrated TFI are identified in Fig. 1 (left panel) and the optical layout is shown schematically in Fig. 1 (right panel). The overall characteristics of the instrument are summarized in Table 1. The heart of the TFI is a Fabry-Pérot etalon (Fig. 2, left panel). It consists of a pair of flat, semi-transparent surfaces that are held parallel to each other under active control. After multiple reflections within the etalon cavity, only those wavelengths that interfere constructively are transmitted. By changing the spacing (“gap”) of the etalon, the wavelength of constructive interference at some reference position (e.g., the optical axis of the instrument) can be tuned to any desired value within the functional wavelength range. Since the response of an etalon is periodic, a blocking filter (in the Filter Wheel: Fig. 3, right panel) is moved into the light path to reject light from adjacent orders of interference.

The operational wavelength range of the TFI is ∼1.5−5.0 μm but there is a non-functional region over ∼2.5−3.2 μm. This region is unusable for scientific observations. The TFI will be routinely operated in the third order on the blue side of this region and in
Table 1: TFI Characteristics

<table>
<thead>
<tr>
<th>Imaging</th>
<th>Wavelengths:</th>
</tr>
</thead>
<tbody>
<tr>
<td>1.5 – 2.5 µm</td>
<td>1. Full FOV: 2.2′ × 2.2′</td>
</tr>
<tr>
<td>3.2 – 5 µm</td>
<td>2. Bright-object subarrays: 16&lt;sup&gt;2&lt;/sup&gt; to 512&lt;sup&gt;2&lt;/sup&gt; pixels</td>
</tr>
<tr>
<td></td>
<td>3. High angular-resolution imaging: 7-hole aperture mask with non-redundant spacing that provides 75 mas resolution at 4.6 µm</td>
</tr>
</tbody>
</table>

<table>
<thead>
<tr>
<th>Coronagraphy</th>
<th>Occulters:</th>
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</thead>
<tbody>
<tr>
<td></td>
<td>0.58″, 0.75″, 1.5″, 2.0″</td>
</tr>
<tr>
<td></td>
<td>Lyot masks:</td>
</tr>
<tr>
<td></td>
<td>MASK21, MASK66, MASK71</td>
</tr>
</tbody>
</table>

<table>
<thead>
<tr>
<th>Detector (HAWAII-2RG)</th>
<th>Dark rate:</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>&lt; 0.0137 e⁻/sec</td>
</tr>
</tbody>
</table>

| Gain: |
| 1.5 e⁻/DN |

<table>
<thead>
<tr>
<th>Filter Wheel</th>
<th>Blocking filters:</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>1-F158M, 2-F175M, 3-F200M, 4-F229M, 5-F246M, 6-F349M, 7-F425M, 8-F470M, 9-CLEAR</td>
</tr>
</tbody>
</table>

| Pupil Wheel | 1-MASK21, 2-CLEAR, 3-FLAT, 4-LINE, 5-MASK71, 6-MASK66, 7-FND, 8-OPAQUE, 9-MASKNR |

Figure 1: Left: Light from the optical telescope element (OTE) is incident on the pick-off mirror (POM) and directed through the collimator, pupil wheel, filter wheel, etalon, camera, and onto the detector. Right: Optical layout of the TFI.
the first order on the red side. Other orders can be accessed for engineering purposes. The resolving power of the TFI is $\sim 100$ over its full wavelength range.

The Pick-Off Mirror (POM) assembly includes a focus mechanism which can be translated with a stepper motor. The four coronagraphic occulters are engraved on the POM and are paired to three Lyot apodization masks located in the Pupil Wheel. Internal calibration units consist of miniature incandescent lamps mounted behind the Dual Wheel assembly. They produce internal flat fields through the FLAT element of the Pupil Wheel and monochromatic spots centered at five different reference wavelengths through the LINE element. The intensity level of the lamps is adjustable.

Astronomical sources can be imaged in narrow, user-selected bandpasses over the entire field-of-view (FOV) or in a suite of subarrays (bright objects, coronagraphic occulters, and aperture masking with the non-redundant mask (NRM)). The TFI science team will exploit all the capabilities of the TFI in their guaranteed observing time, which includes “First Light” programs such as the detection of high-redshift ($z > 12$) Ly$\alpha$ emitters and programs to characterize the assembly of galaxies and the orbits and atmospheres of exoplanets.

2. Ground Calibration

The primary goal of the ground-based calibration effort is to perform a comprehensive characterization of the TFI flight model which will serve as a baseline for on-orbit calibrations and data processing. This pre-launch verification of the performance of the instrument and its detector will provide a good understanding of the fully integrated instrument in a simulated on-orbit environment and will allow the acquisition of calibration files that would be difficult or impossible to obtain in orbit. It also permits the early detection, diagnosis, and resolution of problems, both with the hardware and the flight software.

The ground-based calibration of the TFI will make use of the large vacuum and cryogenic chamber at the David Florida Laboratory (DFL) in Kanata, Canada. In this chamber, the instrument will be subjected to the full range of temperatures that it will experience from launch to normal operation ($35 - 40$ K) at the L2 Lagrangian point. The external stimulus is provided by the Optical Ground Support Equipment (OGSE) built by ABB Bomem (Fig. 3). This apparatus can illuminate the detector with nine broad-band and/or monochromatic point sources over the detector’s field-of-view.

A wide array of calibration items will be addressed in the final cryogenic calibration campaign of the TFI in the spring of 2011 when the instrument will be in its final flight
Figure 3: The OGSE consists of three arrays of telescopes that provide point-source illumination for the fields of view of the two Guider channels (top section; 16 telescopes) and the TFI (lower section; 9 telescopes). The assembly shown here resides inside the cryogenic vacuum chamber. Fiber optic cables couple each of the telescope to the light source, which is located outside the chamber.

configuration. The tests have been devised by the COMDEV, STScI, and U de M teams and will include: measurements of the detector dark current, noise, non-linearity and cross-talk, the transmission of the FLAT, LINE, and OPAQUE elements, the focus, and end-to-end system throughput; characterization of the detector persistence, the blocking filter profiles, the neutral density filter and the NRM; the acquisition of internal flat fields; the wavelength calibration of the etalon and measurement of its detuning and parallelism. A library of calibration reference files will be assembled from these activities to process the on-orbit science images through the JWST pipeline.

3. On-orbit Commissioning

The on-orbit commissioning of the TFI is scheduled over a six-month period immediately following launch. The commissioning will encompass a series of engineering and science activities to verify the integrity and performance of the instrument while it is cooling down and once it has reached its operating temperature. An aggressive plan is being devised to characterize fully the instrument in orbit. The plan may be augmented with programs that directly address issues or problems that were uncovered during the ground calibration at DFL. The activities will include: a basic checkout of the instrument electronics, calibration lamps, detector noise, dark current, and electronic gain; calibration of the etalon wavelength, control electronics, and detuning; adjustment of the focus; updates to the plate scale, geometric distortion, coordinate system, photometric zero-points, and location of the coronagraphic occulters; characterization of stray light contamination and PSF wings; verification of the coronagraph target acquisition and performance of the Lyot masks and the NRM. The Early-Release Observations will be taken as soon as feasible during commissioning, followed by the science programs of the TFI science team and of the guest observers in Cycle 1.
JWST Fine Guidance Sensor Calibration

P. Chayer, S. Holfeltz, E. Nelan

Space Telescope Science Institute, 3700 San Martin Drive, Baltimore, MD, 21218, USA

J. Hutchings

Herzberg Institute of Astrophysics, 5071 West Saanich Rd., Victoria, BC, V9E 2E7, Canada

R. Doyon

Département de physique and Observatoire du Mont Mégantic, Université de Montréal, C.P. 6128, Succ. Centre-Ville, Montréal, QC, H3C 3J7, Canada

N. Rowlands

COM DEV Ltd, 303 Terry Fox Drive Suite 100, Ottawa, ON, K2K 3J1, Canada

Abstract. The primary function of the Fine Guidance Sensor (FGS) is to provide continuous pointing information to the Observatory. The FGS will image two separate regions of the JWST field of view onto two independent 2k x 2k infrared focal planes arrays, which will cover a wavelength range of 0.6 to 5 \( \mu \)m. Once in fine guiding, the FGS will provide continuous pointing information to a precision of \( \leq 5 \) milli-arc seconds at an update rate of 16 Hz for \( J \approx 18.5 \) magnitude stars. In order to fulfill its functions, i.e., Identification, Acquisition, Track, and Fine Guide, the FGS will require calibration. In this paper, we present a brief overview of the FGS design, its operations, and its calibration.

1. Introduction

The Fine Guidance Sensor (FGS) is an near-infrared (NIR) instrument that will provide the data for fine pointing and attitude stabilization to the James Webb Space Telescope (JWST) Attitude Control System (ACS). By measuring and providing accurate positions of a guide star to the ACS, the FGS will be used to stabilize the line of sight in order to achieve the pointing stability that is required by the science instruments on board JWST. The Canadian Space Agency (CSA) is contributing the FGS to the JWST Observatory in collaboration with COM DEV Space Systems of Ottawa Canada which is CSA’s prime contractor. CSA is also contributing a NIR narrow band imager called the Tunable Filter Instrument (TFI) that will be part of the JWST science instrument collection (see Martel et al., these proceedings). We describe in this paper the design of the FGS and its operations, and present some aspects of its calibration.

2. FGS Optical Design

Figure 1 illustrates the optical layout of the FGS with respect to the Optical Telescope Element (OTE). The FGS is located behind the primary mirror where it is mounted on the Integrated Science Instrument Module (ISIM) structure. The FGS pick-off mirror, which is
placed just past the OTE focus, picks up the light that comes from two adjacent fields of view (FOV) and feeds it to the relay Three Mirror Assembly (TMA). The light is collimated by the TMA and is then focused by a fold mirror onto two NIR detectors. Both FOVs cover $2.3' \times 2.3'$ each.

Figure 2 shows a drawing of the FGS optical assembly and illustrates the optical sub-system components that are mounted on the FGS aluminum bench. The bench is attached to the ISIM structure via three bipod kinematic mounts. The TFI optical assembly is underneath the FGS optical assembly and shares the same optical bench (see Figure 1 of Martel et al., these proceedings). In Figure 2, the baffles are not drawn in order to show clearly the pick-off mirror, the Fine Focus Mechanism (FFM) with its fold mirror, and the detector assembly. The FGS focus is adjusted after the TMA by moving a fold mirror mounted on the FFM. A calibration lamp is located behind a small aperture in the center of the tertiary mirror of the TMA and provides a light source that can illuminate both detectors.

The FGS Focal Plane Arrays (FPA) consist of two $2048 \times 2048$ pixels $5 \, \mu m$ cutoff HAWAII-2RG HgCdTe detectors from Teledyne Imaging Systems (TIS) (see, e.g., Loose et al. 2007). Four rows and columns surrounding the detectors are not sensitive to light and are used as reference pixels. This leaves a total of $2040 \times 2040$ sensitive and active pixels that can detect light spanning a wavelength range of $0.6 \, \mu m \leq \lambda \leq 5 \, \mu m$ (Rowlands et al. 2008). TSI also provides two application specific integrated circuits (ASIC) that control the FPAs. Each ASIC is paired with an FPA. The ASICs, dubbed SIDECAR, for System for Image Digitization, Enhancement, Control and Retrieval, are placed on the FGS bench next to the detector assembly.
3. FGS Operations

3.1. FGS Functions

The FGS Flight Software (FSW) uses the FGS for taking full frame and sub-window images in order to provide data to the ACS, which will position science targets within the FOV of a science instrument and stabilize the line-of-sight during an observation. During normal science observations, a sequence of functions is executed by the FGS FSW under the control of the Activity Description (AD) scripts that are running on the ISIM Command & Data Handling Unit. For instance, here are the four functions that the FGS will execute after the Observatory has slewed to and settled to a desired attitude:

**IDENTIFICATION:** During this process, the FGS FSW takes full frame images, extracts the positions of stars observed in the FOV, compares them to the positions of known stars that were uploaded from the ground, identifies the guide star, measures its position and reports it to the AD.

**ACQUISITION:** First, a 128 × 128 sub-window (~8.7" × 8.7") is placed on the guide star and a series of images are taken. The images are passed to the FGS FSW that measures the average position of the guide star. Then a 32 × 32 sub-window (~2.2" × 2.2") is centered on this new guide star position and a second series of images are taken in order to measure the position of the guide star more accurately.

**TRACK:** The AD commands the FGS FSW to center a 32 × 32 sub-window on the guide star. This time the sub-window is continuously read out and the FGS FSW computes the centroid of the guide star. These centroids are passed to the ACS at a rate of 16 Hz for line of sight stabilization of the spacecraft. The guide star location is continuously reported such that the sub-window location is updated in case the guide star is still moving across the FOV. Once the jitter and drift have been reduced, an 8 × 8 sub-window (0.54" × 0.54")
is centered on the guide star and the FGS FSW computes more accurate centroids that are passed to the ACS at a rate of 16 Hz.

**FINE GUIDING:** The AD commands the FGS FSW to transition directly from Track to Fine Guiding by using a fixed window size of $8 \times 8$ pixels centered on the guide star. The FGS FSW computes the guide star's centroids and sends them to the ACS every 64 msec. Once in Fine Guiding, the image motion that is expected over a typical duration of science observation should be $< 7$ mas.

### 3.2. Guide Star Selection

The ground system will select guide stars and reference objects for JWST observations. Guide stars are used for fine guiding the telescope and reference objects are used in on-orbit pattern matching to identify the guide star. Stars with $12.1 \leq J \leq 18.6$ will be used as guide stars for fine guiding; reference objects may be even fainter. Up to three guide star candidates will be selected for each visit, dramatically reducing the risk of guide star acquisition failure. Guide stars and reference objects will be selected from the Guide Star Catalog II (GSC-II) and the Two Micron All Sky Survey (2MASS) Point Source Catalog (see, Lasker et al. 2008 and Skrutskie et al. 2006).

Although the GSC-II is based on a photographic survey, it has been selected as one of the best sources of guide stars for the JWST mission, because it is the deepest and most complete all-sky survey available. GSC-II contains approximately a billion objects ($\sim 0.3$ billion stars) that were observed in three photographic pass bands: $B_J$ ($\lambda_{\text{eff}} = 0.47 \, \mu m$), $R_F$ ($\lambda_{\text{eff}} = 0.64 \, \mu m$), and $I_N$ ($\lambda_{\text{eff}} = 0.85 \, \mu m$). As for the 2MASS point source catalog, it contains 0.47 billion objects over the entire sky in the $J$ (1.25 $\mu m$), $H$ (1.65 $\mu m$), and $K_S$ (2.16 $\mu m$) band passes and is complete down to $J \approx 16$. The resolution ($\geq 2$ arcseconds) of the 2MASS data is not sufficient to ensure that all objects are truly point sources. Sources at the bright end of the catalog ($J < 12.1$) are too bright for guiding and the faint limit of the catalog ($J \approx 17$) is brighter than the expected faint limit for JWST FGS guide stars ($J \approx 18.5$).

The on-board guide star identification and acquisition process uses a small star table provided by the ground system containing the positions and predicted detector count rates of the guide star and reference objects expected to be in the FGS FOV. As the FGS is a broadband camera with a wavelength range of $\sim 0.6 \, \mu m$ to $5.0 \, \mu m$, the magnitudes of the GSC-II objects, $B_J$, $R_F$, and $I_N$, must be transformed into $J$, $H$, $K_S$ if 2MASS magnitudes are not available, and then extrapolated from $K_S$ to 5 $\mu m$. These optical and NIR magnitudes are then convolved with the telescope aperture, the FGS throughput, and detector response to derive the predicted electron count rate.

Transformations of the GSC-II optical magnitudes into the near infrared $J$, $H$, and $K_S$ have been derived from color-color diagrams for stars common to both GSC-II and 2MASS (Chayer et al. 2006). At high galactic latitudes ($|b| > 45$ degrees), where only 48% of GSC-II stars have 2MASS matches within 1.75 arcseconds, the effects and variations of reddening are negligible, and consequently, the magnitude transformations are fairly accurate (Holfeltz et al. 2009). At low galactic latitudes ($|b| < 30$ degrees), where highly variable reddening hinders accurate magnitude transformations and the GSC-II catalog is subject to confusion, 87% of GSC-II stars have 2MASS matches within 1.75 arcseconds, eliminating the need to predict their NIR magnitudes.

### 4. FGS Calibration

The calibration of the FGS affects many more aspects of the observatory than the accuracy and precision of pointing. The FGS must be calibrated in order to plan and schedule visits, select guide stars, optimize operational parameters, and execute spacecraft maneuvers, in addition to guiding JWST. The characterization and calibration of the FGS started during
instrument assembly, integration and ground testing. These measurements and analytic calculations will serve as the initial calibration database. These initial calibrations will be verified or updated during on orbit commissioning. Some calibrations may only be possible on orbit; initial data for these will likely be obtained during the commissioning phase. During routine on orbit operations, after commissioning is complete, the calibration plan is likely to consist mostly of periodic calibration monitoring and updating of the calibration database and operational instrument settings. Below is a partial list of the planned FGS calibrations. For more information, the reader is referred to the FGS Calibration Plan prepared by Nelan 2007 for CSA.

**PSF CHARACTERIZATION:** The effective point spread function (PSF) and its variation across the FGS field of view will be characterized; (the FGS PSF is under-sampled and cannot be directly measured). This calibration is needed for the geometric distortion correction and to verify the centroid non-linearity correction applied to the measured guide star position. Calibration data will be obtained by dithered observations of an astrometric field in the Large Magellanic Cloud (LMC). Dithering compensates for the under-sampled FGS pixels. This process will be iterative, since the commanded sub-pixel dithers cannot be actualized before the effective PSF in the guiding FOV is calibrated.

**PHOTOMETRIC CALIBRATION:** The magnitude transformations predicting FGS count rates from GSC-II and 2MASS magnitudes will be calibrated. Relatively accurate predicted FGS count rates are needed to select and identify guide stars. Predicted count rates are used in the guide star selection process to determine if a candidate is bright enough to serve as a guide star. Predicted count rates are used on orbit to pattern-match the predicted and observed scenes, and identify the guide star. To obtain the necessary calibration data, we will bootstrap from NIRCAM observations of the astrometric LMC field.

**GEOMETRIC DISTORTION:** The geometric distortion will be determined across the FGS FOVs to $\leq 5$ mas. Knowledge of the geometric distortion is needed to support fine pointing and accurate offset maneuvers. Calibration data to determine the geometric distortion across the FGS FOV will be obtained by dithered observations of the astrometric LMC field at various orientations.

**SPACECRAFT TO FGS ALIGNMENT:** The alignment of the star tracker coordinate frame relative to the FGS coordinate frame will be measured. Knowledge of the relative alignment is needed to place guide stars in the FGS FOV for identification & acquisition and set the orientation of ISIM focal plane. First, the star trackers must be aligned relative to each other by comparing catalog and observed positions of stars in the Bright Star Catalog (BSC). Next, the alignment of the star trackers with respect to the FGS is determined by simultaneously observing several BSC stars in the star trackers and LMC stars in the FGS.

**ISIM TO FGS ALIGNMENT:** The relative alignment of each ISIM instrument with respect to the FGS will be measured. Knowledge of the relative alignments is needed to fine point the instruments and support accurate offsets. The initial ISIM-to-FGS alignment will be measured on the ground during ISIM assembly. Accurate on-orbit calibration data will be obtained by simultaneously observing the astrometric LMC field with the FGS and each science instrument at selected positions and rolls.

**FLAT FIELD:** The FGS flat field will be determined. Knowledge of the pixel-to-pixel non-uniformity is needed for accurate measurement of guide star positions in the FGS field of view. Limited flat field data will be collected during ground tests. On-orbit flat field calibration data will be collected from observations of star fields and the internal flat lamp.

References

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Nelan, E. 2007, FGS Calibration Plan, JWST-STScI-001050
Detecting Cosmic Rays in Infrared Data

Rachel E. Anderson and Karl Gordon

Space Telescope Science Institute, Baltimore, MD 21218

Abstract. Cosmic rays are a known problem in infrared astronomy, causing both loss of data and data accuracy. The problem becomes even more extreme when considering data from a high radiation environment such as in orbit around Earth, or outside the Earth’s magnetic field altogether, unprotected, as the James Webb Space Telescope will be. To find the best method to correct for this disturbance we develop three cosmic ray detection methods: a 2-point difference method, a deviation from the fit method, and a y-intercept method. These methods will be applied to simulated non-destructive read ramps with varying combinations of slope, number of frames, number of cosmic rays, and cosmic ray frame number and strength.

1. Introduction

Advances in technology that allow us to see deeper, build more complex systems, and even send telescopes further into space, have challenged us to continue to improve our cosmic ray (CR) detection methods. Infrared detectors, such as those that will be on the James Webb Space Telescope (JWST) require a lower operating temperature, therefore JWST will be placed at the L2 point outside of the earth’s magnetic field. Here the observatory will be largely unprotected making CRs a serious problem. Furthermore, newer infrared telescopes will have a lower read-noise, and thus will detect fainter CRs than before (and most CRs are faint!). Finally, long observing time will be necessary to complete many of the scientific goals of the JWST, causing CRs to be an even larger problem. A study by Offenberg et al. (1999) calculated that in every 1000 seconds, up to 10% of the field of view of the JWST will be affected by CRs. However, according to a more recent study by Rauscher et al. (2000), this rate is predicted to be 20%. Clearly, a reliable method to detect and correct for CRs is needed.

Correcting for CRs is not a new problem, however with the recent development of non-destructive reads there are more solutions available. Instead of utilizing charge-coupled devices (CCD’s), infrared arrays are now made with an infrared material plus silicon Complementary Metal Oxide Semiconductor (CMOS) multiplexer. These do not use charge transfer as a CCD does, and instead collect charge at each pixel and have an amplifier and readout multiplexer at each pixel. This technology allows us to continue to non-destructively read the array as we integrate charge on a pixel, and choose when we want to reset. By integrating the charge on each pixel, we can calculate the slope of this ramp in counts per time to get the flux of the sky (Rieke 2007).

To best learn how to detect and correct for CRs in non-destructive read ramps, we simulate mid-infrared data to test various techniques for finding CRs. This will be described in Section 2. There are three proposed CR detection techniques that we explore here; these methods will be described in Section 3. The first is a two-point difference method (Subsection 3.1.), the second is a deviation from the fit method (Subsection 3.2.), and the
third is the y-intercept comparison (Subsection 3.3.). Concluding remarks can be found in Section 4..

2. Simulating Non-Destructive Read Ramps

In order to test how well a CR detection method works, we have to test it on a known CR. For that reason we have simulated non-destructive read ramps with photon-noise and read-noise added in, to which we can either add CRs with known frame numbers and known amplitudes or leave the ramps clean (CR free). We build these ramps with the slope, zero point, gain, dark current, integration time and frame time as inputs. At time, \( t = 0 \), the counts are equal to the y-intercept. This is not considered part of the ramp \( (t = 0 \) would be when the reset occurred, but we start reading at time equal to one frame time). For the first frame, we add the expected number of photons to the y-intercept (counts at \( t = 0 \)), which would be the slope times the frame-time. To this we add the correlated noise. This is a random number pulled from a Gaussian with the photon-noise as the standard deviation of that Gaussian. We calculate the photon-noise by taking the square-root of the electrons added. We then move onto the next frame, and calculate the counts there by adding the expected photons to the previous count, and then adding the correlated noise as before. When all frames are populated, we add the uncorrelated noise, the read-noise, to each frame. If we decide to add one or more CRs to the ramp, we simply add the electrons when we add the expected counts. The uncertainties for these ramps are calculated by adding the photon-noise (Poisson noise is used) and the read-noise in quadrature.

The inputs used for these ramps are given in Table 1.

| Slope (DN/s) | 10.0 |
| Zero Point (DN) | 3000.0 |
| Number of Frames | 10 |
| Frame Time (s) | 27.7 |
| Gain (e-/DN) | 7.0 |
| Dark Current (e-/s) | 0.02 |
| Read-noise (e-/sample) | \( 16.0/\sqrt{8} \) |

Table 1: Parameters used to create non-destructive read ramps.

3. Cosmic Ray Detection Methods

All three CR detection methods have a unique algorithm, however, all follow the same procedure to calculate a slope once the CRs are found. This is outlined here:

- Detect CRs.
- Calculate the slope for the resulting ramp segments before and after the CR events (we will refer to these as ‘semi-ramps’ from here on, following Robberto (2008)).
- Calculate the final slope of the entire ramp. If there is one CR or more, do this by taking the weighted average of the slopes of the semi-ramps (see Regan 2007).

Anytime a slope is calculated for the above process, or for the following methods, we do so using linear regression for data with correlated and random uncertainties (in consideration of the photon-noise and the read-noise respectively) following Gordon et al. (2005). Furthermore, the largest outlier is always found first.
Detecting Cosmic Rays in Infrared Data

3.1. Two-Point Difference Method

For the two-point difference method we calculate the two-point differences between the counts in each set of adjacent frames. The largest outlier is flagged as a CR given that it fulfills the rejection criteria:

\[
\frac{|d_i - \mu_d|}{\sigma_d} > r_t
\]

where \(d_i\) is the difference between the science data \(i\) and \(i-1\), \(\mu_d\) is the median, \(\sigma_d\) is the uncertainty in the \(d_i\)'s, and \(r_t\) is the rejection threshold. The median is used because it is more robust than the mean when there are outliers in the data (Press et al. 1986, Offenberg et al. 1999). Furthermore, we use the absolute difference so that we remove outliers in both directions and do not bias the data (Offenberg et al. 1999, Offenberg et al. 2001). However, if larger rejection thresholds are used, and therefore there are no longer as many outliers (i.e. only picking up the strongest CRs), then one-sided clipping would work just as well (Windhorst et al. 1994). Once one CR is identified, the faulty \(d_i\) (that which includes the CR) is removed, and the process is repeated on the remainder until no more are found. This method is depicted in Figure 1.

When calculating the uncertainty in the \(d_i\)'s, \(\sigma_d\), the most obvious solution would be to use the standard deviation. However we have found that this does not work well for ramps with low number of frames (\(\sim 5\)). An improvement would be to use the photon-noise and
read-noise added in quadrature. We can calculate the photon-noise as Poisson noise, but since we are dealing with the 2-point difference we use the charge accumulated since the last frame rather than the total charge in a frame (also helps to avoid biases towards the end of the ramp). This charge can be estimated by a linear fit to the data multiplied by the frame time, but in the case of a CR this would not quite be accurate. Rather we use the median of the two point differences, $\mu_d$ to avoid contamination by a CR. Therefore:

$$\sigma_d = \sqrt{n^2_p + 2 \times n^2_r}$$  \hspace{1cm} (2)

where the ‘2’ is because of the read-noise from each of the two frames, $n_r$ is the read-noise and $n_p$ is the photon-noise defined as $n_p = \sqrt{\mu_d}$.

### 3.2. Deviation from the Fit Method

To apply the deviation from the fit method, we fit a line to all of the counts in a ramp and for each frame, calculate the difference to the fit as a ratio to the uncertainty in the counts:

$$diff_n = \frac{x_i - f_i}{\sigma_i}$$  \hspace{1cm} (3)

where $x_i$ is the counts in frame $i$, and $f_i$ is the fit at that sample, $\sigma_i$ is the calculated noise for $x_i$.

We then take the first difference of these ratios, and look for the largest. If it satisfies the criteria:

$$|diff_{i+1} - diff_i| > r_t$$  \hspace{1cm} (4)

then frame $i$ is flagged as a CR. The ramp is then split into the semi-ramps, and this method is applied again to the resulting semi-ramps.

If a sample contains a CR hit, it’s high value will bias the fit upwards, so that the one sample by itself doesn’t necessarily stick out as high as it should above the fit. This also means that the sample preceding the hit (or maybe even more than one sample) will usually lie below the fit. So by looking at the change in distance from the fit from one sample to the next, you’re more sensitive to picking up real hits. This is shown in Figure 2. Notice in the figure that the background is not at zero as it was for the 2-point difference method, but $\sim 10$. The reason is that for the 2-point difference method we subtract the median (which would exclude the CR) while here we subtract the fit which includes the CR in its calculation.

The noise in a frame, $\sigma_i$, is calculated from the photon-noise and read-noise added in quadrature. Instead of using the Poisson noise for the photon noise (as for the 2-point difference method, we found that this resulted in more false CR detections at the end of the ramp where there are higher counts, thus greater Poisson noise), we give an equal weighting to each sample by calculating the photon noise as the square root of the average number of counts added to each frame. Therefore $n_p = \sqrt{m \times t_f}$ where $m$ is the slope calculated for the fit, and $t_f$ is the frame time.

### 3.3. Y-Intercept Method

For the y-intercept method, we step through each frame and assume that there is a CR there, and fit a line to the semi-ramps before and after. Every time the x-axis is shifted so that the y-intercept is located at the frame number of the assumed CR. Then, we take the ratio of the absolute differences in the two y-intercepts ($b_1$ and $b_2$) to the expected uncertainty. After doing this for every frame, we look at the frame with the largest ratio. If this ratio is larger than a given $r_t$, then we flag that frame as a CR. The process is repeated.
Figure 2: The deviation from the fit method is depicted here. The black line and points are the data, and the blue dashed line is the linear fit made up of points, $f_i$, corresponding to the data points, $x_i$. In the lower panel the green dotted line is the ratio of the difference between the fit and the data to the uncertainty, $diff_i$. The red line is the 2-point differences of the $diff_i$’s. This is what will be compared to the rejection threshold, $r_t$. 
on the resulting semi-ramps. This equation is shown in Equation 5, and the method is shown in Figure 3.

\[ \frac{b_2 - b_1}{\sigma_y} > r_t \] (5)

The expected uncertainty, \( \sigma_y \), is calculated by adding the uncertainty in the y-intercept calculation from both the read-noises and from the photon noise in quadrature. The uncertainty in the y-intercept calculation from the read-noise comes from the random errors in the linear fit algorithm described by Gordon et al. (2005). The photon noise that goes into this measurement is calculated by taking the weighted average of the two slopes (from the two semi-ramps), and multiplying that by the frame time to get the average number of counts added to each frame, and taking the Poisson noise of that. Therefore,

\[ \sigma_y = \sqrt{n_p^2 + n_{r_1}^2 + n_{r_2}^2} \] (6)

where \( n_r \) is the read-noise, and \( n_p = \sqrt{\bar{m} \times t_f} \).

4. Conclusion

Each of the three methods described above find the CR, evidence being the strong peak in their detection ratios at the location of the CR (Figures 1, 2, 3). However, how well each method finds a CR depends on the rejection threshold used. Without considering the results, there are a few things we can acknowledge. The 2-point difference method takes the least amount of computations, while the y-intercept method can claim the most. Furthermore, once a CR is found, the 2-point difference method is the least affected. With the 2-point difference method, instead of splitting the ramp into two semi-ramps, we can simply remove the 2-point difference that includes the CR. With the other two methods the ramp is split into two semi-ramps and our calculations of the slope and y-intercept are not as robust.

These three CR detection methods have been tested on simulated non-destructive read ramps, adjusted, corrected, and refined, in order to present the optimum version here. However, we ask ourselves which of the three is the best? The goal of our future work is to answer this question.

To compare these methods, we look at simulated ramps with 5, 10, 20, 30, and 40 frames, with CRs of 20 different strengths from 0 to 140 DN (beyond that all CRs were found with all methods), and with slopes of 0.1, 1.0, 5.0, 10.0, 30.0, 40.0, 50.0, 60.0, and 70.0 DN/s, with 1, 2, or 3 CRs, and with the CR located in the beginning, middle, and end of the ramp. We then simulate 10,000 non-destructive read ramps for each combination, and apply each method to each ramp to compare the results. For each of the methods described above, the rejection threshold, \( r_t \), is chosen such that the rate of false detections is the same for all methods. In this way we can best compare how will they find the CRs. These results will be published in a future paper.

References


Figure 3: The y-intercept method is shown here. In (a) we show the method as it would look when assuming the CR is in frame 2. The black line and points are the data, and the blue dashed lines are the liner fits to the semi-ramp before and after the assumed CR. The two y-intercepts, $b_1$ and $b_2$ are highlighted with red dots and labeled. In (b) we are assuming the CR is in frame 4 (which indeed it is). In plot (c) we show the ratio of the absolute difference in the y-intercepts to the uncertainty, which must be above the rejection threshold, $r_t$, to be counted as a CR. These are the results after stepping through the entire ramp assuming a CR was in each frame.


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WFC3/UVIS Detectors: On-orbit Performance

S. M. Baggett, T. Borders, S. Deustua, R. L. Gilliland, J. S. Kalirai

*Space Telescope Science Institute*

R. A. Kimble

*Goddard Space Flight Center*

V. Kozhurina-Platais, J. MacKenty, A. Rajan

*Space Telescope Science Institute*

Abstract.

WFC3 is a fourth-generation instrument installed on HST in May 2009. Designed as a panchromatic camera, it features both a UVIS and an IR channel, selectable via a channel select mechanism (CSM), allowing sequential imaging. Both detectors are performing well on-orbit and are now in standard science operations. This poster presents some of the UVIS detector highlights as well as remaining issues being addressed.

1. Introduction

The WFC3 UVIS detectors consist of two 2Kx4K e2v thinned, back-illuminated devices, separated by about 35 pix or 1.4”. The field of view is 160”x160”, with a spectral range of 200-1000 nm. The CCDs reside in a hermetic package which includes a vent tube to provide for outgassing and a thick housing wall to provide radiation shielding. Four 2-stage thermo-electric coolers (TECs) cool the inner radiation shield and window while a 4-stage TEC cools the CCDs to -83C. Figure 1 shows a photo of the UVIS detector package before integration into the instrument.

Figure 1: UVIS detector package.
The chip design includes a charge injection feature to provide a mechanism for mitigating the charge transfer efficiency degradation which will occur on-orbit. Standard operations for WFC3/UVIS is to read out the full-frame via four amps with serial physical and virtual overscan regions and a parallel overscan area; binning or subarray readouts are available options. Figure 2 illustrates the chip layout.

![Figure 2: WFC3/UVIS CCD readout configuration.](image)

The devices exhibit low readnoise (about 3e-) and dark current (about 3e-/hr in 2010); the gain is about 1.6 e-/DN. The 16-bit A-to-D converter provides an ADC maximum of 65535 with a full well is 63,000-72,000 e-. The chips are photometrically linear to 1% up to 100x saturation (chip 2), somewhat less in chip 1 (about 15% low at 100x full-well) but it appears to be correctable to 2%. The cosmetics of these devices is excellent: less than 0.05% of the science pixels are dead; pixels deviating more than 10% from the average QE comprise less than 0.2% of the chips.

2. Quantum Efficiency

The QE as measured in ground tests was excellent and particularly high in the UV (solid lines, Figure 3). On-orbit data have shown that the total integrated throughput is even higher than expected, by 5-20% (UVIS) and by about 20% (IR); the dashed lines in Figure 3 are the resulting QE if the entire gain is due to QE though it is likely other instrument components are contributing as well (Kalira et al., 2009).

An example flatfield is shown in Figure 4. The minor vignetting at the lower left is an artifact of the ground test optical stimulus. The large wedge-shaped feature, with vertex originating in the lower right quadrant and flaring out to the upper left into the other quadrants is an expected result of reflections between the windows and the detectors, which are positioned at a 20 deg angle to each other. On-orbit low-frequency corrections to the flats (about 1-5%) are in progress; an initial alpha release of new flatfields are available on the main STScI WFC3 web page.

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1alpha releases of flatfields and other calibration files are available from [http://www.stsci.edu/wfc3](http://www.stsci.edu/wfc3)
3. Radiation Damage and Mitigation

The low-earth orbit radiation environment is slowly degrading the performance of the WFC3 CCDs as energetic particles damage the silicon lattice of the detectors and form charge traps. As a result of this damage, hot pixels increase, the charge transfer efficiency (CTE) declines, and the median dark current grows over time (at the rate of about 0.5 e-/hr/pix/yr).
Figure 5 illustrates the gradual on-orbit increase in the number of hot pixels. Each vertical line represents an anneal procedure, where the chips are warmed to 20°C; the heating process repairs a large fraction of the hot pixels, giving the plot its characteristic saw-tooth shape. About 1000 new hot pixels are generated every day; the anneal repairs between 80% and 90% of those but a small portion remain hot and accumulate over time. Frequent superdarks are generated for use in the calibration pipeline and can provide some correction; any residual hot pixel effects can be mitigated by dithering observations (Dahlen et al. 2010).

Figure 6 illustrates the charge transfer inefficiency as measured via the EPER (extended pixel edge response) method. Approximately once a month, a set of flatfields are taken with the internal lamp using a special readout format which provides a significant number of extra overscan rows. Any traps in the CCD will cause a fraction of the flatfield charge to be captured and trail out later into the overscan during readout, appearing as deferred charge. An evaluation of the charge in this tail provides a means of tracking the relative CTE degradation, typically plotted as (1-CTE) versus signal level in the flat. Each line in the Figure represents the results from a monthly EPER measurement; the slope has remained stable but as expected, the intercept (CTI) is slowly increasing. The EPER analysis is primarily useful for tracking the relative changes in CTE; for calibrating CTE effects on an absolute scale and determining the impact to science targets, external observations are necessary; these are being acquired as part of the standard calibration plans.

To mitigate CTE, the WFC3 chips have a charge injection (CI) capability. Effectively an "electronic preflash", CI fills the radiation-induced traps; based on ground tests, the additional noise due to CI is only 15 e- rms for 10,000 injected electrons, considerably lower than the Poissonian noise from an optical pre- or postflash image. CI is expected to be made available as an option to observers once the CTE degradation becomes more significant, probably in 2011 or 2012.

4. Low Level Image Effects

Several low level effects are present in the WFC3 UVIS images: potential slight decrease in QE (hysteresis), crosstalk, and fringing.
The hysteresis was first detected during ground testing, when occasional images showed a "bowtie" pattern (at left in Figure 7) or sometimes spots where targets had been located in previous images. The levels of the features at the operating temperature of -83°C were low, usually 0.1% to 0.2%. Tests performed on WFC3-like devices in the Goddard Space Flight Center Detector Characterization Lab (DCL) were able to reproduce a similar effect, a small reduction in QE, by cooling the CCDs without illumination. Furthermore, the lab tests revealed that 1) a uniform QE deficit could be present after cooling (i.e., no characteristic bowtie shape thus making it more difficult to detect), 2) the QE of the chips could be completely restored by "flashing" the devices with a saturated flatfield, and 3) the restoration, or pinning, of the QE remained effective for days.

On-orbit, the internal lamps are used to regularly supply the pinning exposure. Monitoring of the QE levels has shown evidence of slightly lower QE immediately after the detectors are cooled following the monthly anneal procedure. However, the pinning exposure performed as part of the anneal, before any science data are acquired, successfully quenches the effect: frequent monitoring between anneal procedures shows no evidence of hysteresis (Figure 7; see also Baggett and Borders, 2009).

Common in devices where more than one chip section is read out simultaneously, crosstalk (CT) is also seen in the WFC3 images. In the case of the UVIS channel, a target in one quadrant generates a negative mirror image in the other quadrant within the
same chip (i.e., the CT does not cross chips). The CT is linear and low level, about 7x 10\(^{-5}\) to about 1x 10\(^{-4}\) and can be corrected in post-readout image data processing. Such a correction may be added to the calibration pipeline.

Finally, we note that as back-thinned CCDs, the chips also exhibit fringing caused by multiple reflections between the layers of the CCD detector. The amplitude of the fringes is a strong function of the silicon layer thickness and the spectral energy distribution of the light source. Apparent mostly at wavelengths longward of 700nm, the amplitude with monochromatic input increases gradually with wavelength, and can reach levels higher than 50% at the longest wavelengths (Wong, 2010 as well as Wong, this conference). The fringing amplitudes will of course be a strong function of the target spectrum. For example, continuum sources in broad filters will effectively smooth out fringing effects but that same filter can show strong fringes when illuminated by sources with strong spectral lines or SEDs much narrower than the filter bandpass. For sources with SEDs similar to the calibration lamp, the fringes can be corrected by the flat-fielding process; the fringe pattern has been shown to be very stable, as long as the wavelength of light on a particular part of the CCD stays constant, so fringing can be corrected if an appropriate flat field is available. Tools for correcting fringing are under development.

The on-orbit geometric distortion calibration is obtained from observations of the Large Magellanic Cloud (LMC). Measured star positions are compared to those in the LMC standard astrometric catalog (derived from ACS/WFC observations) to obtain distortion coefficients. The distortion induced by the WFC3 optics is significant, elongating pixels by about 7\% and varying the projected pixel area by about 7\% over the FOV, affecting positions and photometry (see Kozhurina-Platais et al., 2009, as well as Kozhurina-Platais et al., this conference).

Shown in Figure 8 is the residual distortion map for both chips: the accuracy of the calibration is such that the boundaries of the mask used for photolithography during the chip manufacture are evident. The size and direction of the vectors indicate the magnitude and direction of the residual; the longest vector is 0.08 pix (UVIS1) and 0.05 pix (UVIS2); the scale is magnified 2000x. Discontinuities between the vectors coincide with the lithographic edges (red lines) as measured from flatfield images. The pattern is effectively removed by the pipeline flatfields.

5. Conclusions

The WFC3/UVIS detectors have shown excellent performance on-orbit, with low readnoise and dark current, superb QE particularly in the UV, great cosmetics, and extraordinary linearity even well beyond saturation. The astrometric solutions are of sufficient accuracy that the lithography pattern used during the chip manufacturing process is clearly evident in the solution residuals. Radiation damage is causing increased number of hot pixels though these can be controlled to some extent by regular anneals. The on-orbit radiation damage is also slowly increasing the dark current and degrading the CTE; charge injection will be a future option for mitigating the latter. Finally, some low-level effects remain, such as crosstalk or fringing, though these are generally calibratable; hysteresis is being well-controlled via periodic pinning exposures.

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Figure 8: Geometric distortion residuals.


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For further information:
http://www.stsci.edu/hst/wfc3
http://wfc3.gsfc.nasa.gov
Anomalies and Artifacts of the WFC3 UVIS and IR Detectors: An Overview

M. J. Dulude, A. Rajan, A. Viana, S. Baggett, and L. Petro

Abstract. Since first light, a number of anomalies or artifacts have been found in the WFC3 UVIS and IR detectors. These include trails from passing satellites, “ghost” images from reflected light, scattered light, and a handful of detector idiosyncrasies. In this paper, we present a “rogues gallery” detailing the various types of anomalies found so far, their likely causes and possible remedies.

1. Introduction

The HST WFC3 IR/UVIS detector is a dual-channel instrument with a HgCdTe IR side and a CCD UVIS side. The details of the instrument design can be found in Section 5 of the WFC3 Instrument Handbook. The WFC3 instrument was designed to take advantage of lessons learned from previous HST instruments in order to optimize detector performance. However, there are still unavoidable abnormalities in the detector response particularly under extreme conditions. Some of these anomalies are well understood while others are still being investigated. Further refinements will be published in future Instrument Science Reports.

2. UVIS Stray Light

The UVIS stray light anomaly is characterized by diffuse a horizontal and/or vertical strip of abnormally bright pixels. The horizontal feature is brighter and more common that it’s vertical counterpart, and sometimes contains diffuse ‘blobby’ structures, as seen in Figure 1. The anomaly is believed to be caused by light from a moderately bright source slightly outside the detector field of view. This is supported by the fact that the location of the anomaly is highly dependent on telescope pointing. This anomaly is very rare. It has only been seen in less than 1% of all full-frame UVIS external images.

3. UVIS Ghosts

“Ghost” image artifacts are produced by light reflecting off various elements along the light path of the WFC3/UVIS detector. They were predicted and characterized prior to launch in a number of works (See Stiavelli, Sullivan and Fleming (2001), Brown & Lupie (2004a), Brown & Lupie (2004b), Bond & Brown (2005), Brown (2005), and Brown (2007)). These anomalies can be classified into three groups: CCD ghosts, window ghosts, and filter ghosts.

CCD ghosts are caused by light reflecting off the surface of the UVIS CCD and the detector and Dewar windows. CCD ghosts manifest themselves as large, diffuse donuts or figure-8 shaped features. These features are widely separated from the source star. They occur along a diagonal line up and to the left of the source star (see Figure 2), and hence occur when bright sources are placed in the lower right quadrant of the detector. Generally, CCD ghosts contain ~2-3% of the source signal.
Window ghosts are caused by light reflecting between window surfaces. They are characterized by a series of small diffuse donuts in the immediate vicinity of the source star, as seen in Figure 3. Overall, window ghosts contain ~0.1-0.4% of the source signal.

Filter ghosts are caused by light reflecting off the surfaces of layers in the filters. The location, morphology and impact of filter ghosts varies according to the filter, as different filters were manufactured with different internal structures (see Figure 4). Generally, filter ghosts manifest themselves as either a series of compact points in the immediate vicinity of the source, or as a series of donuts (which can be indistinguishable from window ghosts) in

Figure 1: Brighter, blobby horizontal and fainter vertical stray light artifacts in a full-frame UVIS image.

Figure 2: Full-frame external UVIS image with two CCD ghost artifacts (circled in white) from the bright star near the lower right corner of the image.
the immediate vicinity of the source. Although the filter ghost brightness varies according to the filter, and there are a handful of notable exemptions (see Brown (2007) for more details), most filter ghosts contain roughly 0.1% - 0.3% of the source signal.

The effects of ghosts can be mitigated using several different techniques. In general, dithering and/or rolling *HST* is the simplest solution. Additionally, CCD ghosts can be avoided by keeping bright sources out of lower right quadrant. For more complex situations, deconvolution algorithms can assist in ghost mitigation as well. For more details, see Bond & Brown (2005).

Figure 3: Window ghosts (circled in white) can be seen emanating from the eleven o’clock position of the two brightest stars

Figure 4: F656N image with the donut-shaped filter ghosts circled in white.
4. IR Banding

Banded images exhibit a rectangular region containing pixels with brightness levels that are significantly different (typically ± 3-5 DN) from values in the rest of the image. This region is vertically centered and extends all the way across the image horizontally into the reference pixels. The banded region is “bookended” on top and bottom by single row of pixels with “discontinuous” brightness levels (see example images in Figure 5, accompanying brightness profiles in Figure 6). Finally, although the vertical width of the band does vary from image to image, it only does so by very specific quantized steps – All observed banded regions have a vertical width of either 512, 256, 128 or 64 pixels.

Figure 5: Examples of banded images. Left: 64-pixel-wide band in a SPARS50 full-frame external science image. Right: 128-pixel-wide band in a SPARS10 256x256 subarray dark calibration image.

Figure 6: 3-σ clipped robust mean brightness profile along y-axis of the full-frame external science image (left panel) and 256x256 dark calibration image (right panel) in figure 5. Note the central banded region and the discontinuous rows that bound it.

One of the most puzzling properties of this anomaly is “induced” banding. Under the right conditions, it seems that banding can be induced in almost any IR image by the exposure (or exposures) that immediately preceded it in time. Assuming the first image is
smaller than the following image and that the time interval between the two images is less than an hour, there is a strong possibility of banding in the following image.

Although images with induced banding have little to nothing in common, the images that immediately precede them do. In nearly every documented case of induced banding, an IR subarray image whose size exactly matches the vertical width of the induced band was taken in the previous hour. For example, a 128x128 subarray dark was taken several minutes before the image in the right panel of Figure 5 – which has a 128-pixel wide band, and several 64x64 subarray images were taken just minutes prior to the banded full-frame science image illustrated in the left panel of Figure 5 – which has a 64-pixel wide band.

Calibration is another open issue. It is not fully understood how banding affects external science images, and if the effects can be reduced or eliminated by calibration. Further complicating the issue is the fact that many (but not all) subarray dark calibration files exhibit strong banding. (Table 1 summarizes our preliminary banding survey of subarray dark calibration frames) As the effects of banding on calibrated images and dark calibra-

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Table 1: Preliminary results of the banding survey of WFC3/IR subarray dark calibration files

In summing, banding still not well understood and very much an open issue. The behavior of the banding anomaly is complex enough that additional analysis is required to fully determine the root cause.

5. IR Blobs

According to Pirzkal, et al. (2010), blobs are small circular blemishes with typical radii of 10-20 pixels that will, in general, reduce the flux from a star by 5 to 10%, and in some cases as much 15-20% (see Figure 7, left panel for an example image). Blobs are thought to be caused by material deposited on the Channel Select Mechanism (rather than the IR detector itself).

The locations of the blobs is completely static and their locations are well determined (see Figure 7, right panel), and the WFC3 calibration code (CALWF3) has already incorporated a mask to automatically flag affected pixels. Thus, the effects of blobs can be mitigated by simply avoiding them or dithering around them.
6. IR Snowballs

Snowballs can be described as ‘fuzzy’ blobs of bright pixels with saturated cores whose occurrence (in terms of both when and where) is totally random. Each snowball affects between 15 and 35 pixels, saturating between 1 and 13 central pixels. Overall, they seem to occur at a rate of between 0.4 and 0.8 per hour per full-frame image. The cause of snowballs is not fully understood (see Hilbert (2009) and McCullough (2009) for more details and further discussion).

Due to the transient nature of snowballs, they are largely removed by up-the-ramp signal fitting. The resulting calibrated image usually contains a small patch of pixels with non-physical (negative) values at the site of the snowball. An example of a calibrated image with a snowball can be found in Figure 8. In general, the best mitigation strategy is to simply take more than one exposure, and possibly dithering as the chances of a snowball striking the same pixel in two exposures is slim.

Figure 8: Section of a calibrated flt.fits image with the remains of a snowball circled in white.
7. IR Scattered Earth Light

Scattered Earth light is an anomaly most often seen in IR grism observations, but can occur in IR direct (non-spectral) images as well. As illustrated in Figure 9, this anomaly is characterized by a diffuse region of bright background of variable width that extends from the side of the image.

![Figure 9: Example IR grism image with scattered earth light.](image)

This anomaly occurs when the telescope is pointing near the bright Earth limb. Continuous Viewing Zone (CVZ) observations are therefore the most susceptible, because the telescope can be pointing near the bright limb for extended periods. Grism observations most often suffer from scattered Earth light because grisms have very large overall throughputs when compared to the IR filters. It should also be noted that wide-band filters are susceptible as well, as the throughputs for these filters are also quite large. Thus, non-CVZ and medium- and narrow-band filter images are least affected.

8. UVIS and IR Satellite Trails

Satellite passes occur when Earth-orbiting objects pass through WFC3’s field of view during an exposure. This is an unavoidable event that typically occurs once per few tens to hundred images. Observations with longer exposure times are naturally more susceptible as they present a larger interval for a pass to occur.

Figure 10 shows affected UVIS and IR images. The path is always straight and randomly oriented. The width of the trail varies from image to image, but typically is approximately 10-15 pixels. Affected pixels are almost always saturated. This is of special concern to IR observers. The effects of bright and saturated satellite passes will linger beyond the initial image due to persistence effects.

The best way to protect against, or minimize the effects of satellite passes is to take more than one exposure, dither and/or use MultiDrizzle to produce a final product.
Figure 10: Satellite passes as seen in UVIS (left panel) and IR (right panel)

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WFC3/UVIS Linearity: Low to Near Saturation

A. Rajan, S. Deustua and R. Gilliland

Abstract. We present measurement results of the linearity of the WFC3 UVIS detectors by sampling over the response curve from very low counts up to near saturation. We use data from the Cycle 17 calibration proposal 11925. The in-orbit program obtained exposures of NGC 1850 to verify the linearity response and to characterize the linearity at very low counts between a few hundred and a few thousand electrons. Our results indicate that the CCDs are linear to <1% between a few thousand and ~65,000 electrons.

1. Target and Observations

We observed the star cluster NGC 1850 in the Large Magellanic Cloud using the Wide Field Camera 3 UVIS detector in October 2009. The observations were all carried out in F467M filter. For the purpose of this study we observed the cluster in 2 sets with two exposures 10s which were cosmic ray cleaned and combined. And single images of 20s, 50s, 100s, 300s and 500s. Each set of 7 images were observed twice, see Fig. 1. In the first set the cluster is centered on UVIS1 (left) and in the second set cluster is centered on UVIS2 (right).

Figure 1: The two pointings used in our program are shown here. The image on the left centers the cluster on UVIS and the image on the right is centered on UVIS2.
2. Results: Linearity Near Saturation

To test the linearity of the detector we decided to use a combination of the longest and the shortest exposures in our dataset. We have also carried out these tests with the shorter exposures and find that the results do not vary.

Cosmic rays can skew the results for the shorter exposures and thus before we did our tests we first cosmic ray cleaned all the images. To do this we made use of MULTIDRIZZLE (Fruchter & Hook 2002), and one of the bi-products of the code is a CR cleaned image in the FLT frame. We then generated a catalog of stars (using SExtractor) using the 500s exposures as reference frames. These catalogs were used to carry out photometry (using DAOPHOT) on all the images.

![Graph showing linearity near saturation](image)

Figure 2: Brightest pixels in each star in the short exposure (10s) plotted against the long exposure (500s).
Figure 3: The upper panel plots the photometry in a 3 pix aperture in the short exp vs the long exp. The bottom panel plots the ratio of long and short exposures vs. short. At low levels (<2000e^-) there is a clear deviation from the expected value of 50 (green line above).

In Fig. 2 we plot the brightest pixel of each star in the long exposure (500s) against the short exposure (10s) for all the stars in the catalog. Further the stars are differentiated based on their position on the detector i.e: UVIS1 or UVIS2. The black points shown in the figure are the observations. The green line is the expected value i.e: Short Exp*50, which
is what we expect if the detector were perfectly linear since the long exposure is a factor of 50 greater than the short exposure. To ensure that we are working within the linear regime we enforced a physical cutoff wherein only stars with counts $<55000e^-$ were included in our calculations. Note that this value is lesser than originally estimated saturation value for the UVIS detector, please refer to the WFC3 ISR 2010-10.

3. Results: Linearity at Low Counts

The main goal of this program was to test the linearity of the detector in the low count regime. During thermal vacuum testing we were able to test the medium to high count regime satisfactorily and the results are summarized in the WFC3 ISR Deustua et al 2010 (in prep). Fig. 3 plots the aperture photometry carried out on the stars, the results in the upper panel appear similar to the one seen in Fig. 2, but when we plot the difference of the observed and the expected (lower panel in Fig. 3) we see that the detector deviates from the linear (given by the green line) at levels below 2000 $e^-$ (NOTE: this is the sum in an aperture of 3 pix).

3.1. CTE Losses

One possible solution to the deviation from linearity at low counts is charge transfer efficiency (CTE) problem, which all the previous HST detectors have suffered from. CTE basically means that electrons in stars are inefficiently transported from the region they are measured to the readout electronics. The effect is much more problematic for faint stars than bright stars and while WFC3 was expected to suffer from CTE, we were not expecting to see it so early.

![CTE losses](image)

Figure 4: CTE measurement for stars having sum greater than 2000$e^-$. The data however clearly shows how going to progressively fainter stars affects the linearity of the detector.

We plot in Figures 4, 5, 6 the ratio of the exposures used in our test normalized to one versus the Y-position of the star on the detector. What these plots indicate is that the further you are from the detector readout (which occur at 0 and 4192 in the plots) the worse the linearity appears to get. This is the classical signature of CTE damage. Stars
with aperture sums greater than 2000e$^-$ appear not to be affected, while stars under 500e$^-$ can lose up to 6% of their counts. Having seen that the detector appears to show on-orbit CTE damage, the WFC3 team is undertaking a careful study during Cycle 18 to mitigate this effect.

4. Summary

We have met the initial goal of this program which was to ensure that the detector is linear up to saturation, which occurs at $\sim$65000e$^-$. At count levels near saturation the detector is
linear to better than 1%. In addition we have studied the low count regime using the faint stars in the cluster and found first on-orbit evidence for CTE damage to the WFC3 UVIS detector. There will be follow-up studies in Cycle 18 to measure the CTE and to continue monitoring the linearity of the detectors using different calibrators.

References

Gilliland, R. L. 2010, WFC3 ISR 2010-10, "WFC3 UVIS Full Well Depths, and Linearity Near and Beyond Saturation"
WFC3 UVIS Full Well Depths, and Linearity Near and Beyond Saturation

Ronald Gilliland

*Space Telescope Science Institute, Baltimore, MD 21218*

**Abstract.**
Maps of full well depth ranging from 63-72K electrons are presented. Both ACS and STIS showed excellent linearity beyond saturation when the gain used sampled full well – one simply needs to sum over all the pixels bled into to recover accurate photometry. I show that WFC3/UVIS photometry requires a simple calibration to obtain good results beyond saturation. With this calibration it is possible to recover photometry for saturated point sources to near 1% even for stars that are 7 magnitudes beyond saturation.

This work is presented in a contemporaneous WFC3 ISR, these proceedings are used only to provide a pointer to Gilliland, Rajan and Deustua (2010) which has the same title as this paper.

**References**

Measurement of the Count-Rate Non-linearity in WFC3-IR Detectors

Robert J. Hill
NASA Goddard Space Flight Center, Greenbelt, MD

Abstract.
The Detector Characterization Laboratory at NASA/GSFC has investigated the count-rate non-linearity (reciprocity failure) characteristics of 1.7 micron cut-off HgCdTe devices that are very similar to the WFC3 IR detector. The reciprocity failure follows a power law behavior over the range of fluxes tested (0.1-10^4 photons/second). The slope of the power law varies among detectors, ranging from 0.3-1% dex at 1.1 micron, which is much smaller than the 6% dex effect observed with the HST NICMOS 2.5 micron cut-off detectors. Reciprocity failure is nevertheless an important effect in the calibration of WFC3 data, as well as in other applications in which there is a large difference in flux between the photometric standards and the scientific sources of interest. Furthermore, the variation among detectors demonstrates that a measurement made with the WFC3 IR detector is necessary in order to provide a proper calibration of the WFC3 IR data.
WFC3 Low-Frequency Flat Field Corrections

Jennifer Mack

Space Telescope Science Institute, Baltimore, MD 21218

Abstract.

Multiple dithered observations of the globular cluster Omega Centauri (NGC 5139) have been used to measure inflight corrections to the WFC3 UVIS and IR ground flat fields for a subset of key filters. To obtain an adequate characterization of the flat field over the detector field of view (FOV), 9 pointings were obtained for each filter using a 3x3 box dither pattern with steps of approximately 25% of the FOV. By measuring relative changes in the brightness of a star over different portions of the detector, low-frequency spatial variations in the detector response (L-flats) have been used to correct the flat fields obtained during ground testing. The broad wavelength range covered by these observations allow an interpolation of the L-flat correction for the remaining wide, medium and narrow-band filters, assuming a simple linear dependence with pivot wavelength. Initial results indicate that the required L-flat corrections are ±1.5% (standard deviation) in the IR and ±1.0% in the UVIS, and that the photometric response for a given star after applying the L-flat correction is now stable to better than 1% for any position in the field of view. Followup observations of the same field at multiple orientations will be used to verify the accuracy of the L-flat solutions and to quantify any temporal changes in the detector response while in orbit.
WFC3: On-Orbit Measurement of the Point Spread Function on the WFC3 Detectors

Linda Dressel and George Hartig

Abstract. Knowledge of the point spread function (PSF) on a detector is important for scientific applications such as photometric measurements and analysis of blended sources. It is also important for ensuring that good optical alignment of the instrument is achieved and maintained. The quality of the PSF on HST instruments varies over the detector and varies with time. The focus changes throughout an orbit as the optical bench "breathes", and the average focus changes slowly over time. Locally, the detected PSF is affected by pixelation and charge diffusion or interpixel sensitivity variation. The PSF is also dependent on wavelength. All of these factors both complicate PSF analysis and make it more challenging to provide useful PSFs to observers. Here we review the on-going measurement and analysis of the PSF on the WFC3 detectors.

1. Introduction

Calibration programs to measure and monitor the PSF as a function of time and of position on the WFC3 detectors began during the SMOV period and have continued into cycle 17. In Section 2, we examine stellar cluster exposures from the CAL/OTA program 11877 and show that PSF variations over the course of an HST orbit (due to "breathing" of the Optical Telescope Assembly) are greater than the gradual mean variation that occurs over the course of months. In Section 3, we use similar stellar cluster exposures from the CAL/WFC3 program 11918, which all happen to have approximately the same breathing, to determine the values of a variety of PSF metrics, and to display two of those metrics as a function of position on the WFC3/UVIS detector. In Section 4, we show deep PSFs on the WFC3/UVIS and WFC3/IR detectors made by combining short to long exposures of the isolated star GD 153, made in the CAL/WFC3 programs 11438 and 11439. A brief summary of future work is given in Section 5.

2. Dependence of PSFs on Breathing

Cycle 17 CAL/OTA program 11877 periodically observes the open cluster NGC 188 with WFC3/UVIS filter F410M. We have measured the PSFs in the full frame exposures (2x15 sec) taken in the first 10 visits. An example of the ~30 stars analyzed in one visit is shown in Figure 1.

The mean FWHM and the error in the mean for the PSFs measured in each visit are shown as a function of time in Figure 2. The scatter is due to the change in focus that results from the "breathing" of the Optical Telescope Assembly, due to heating loads and altitude histories. The breathing must be taken into account when analyzing long-term changes in focus and instrument stability.

Predictions of the breathing model are shown for each of the 10 visits in Figure 3, expressed as changes in the spacing between the HST primary and secondary mirrors in microns over the course of two orbits. An extrapolation of the long-term variation is included.
The squares indicate the time at which the WFC3/UVIS exposures were made. A range of nearly 6 microns has been sampled by the exposures. Larger excursions occur occasionally.

The mean FWHM and the error in the mean for the PSFs measured in each visit have been replotted as a function of the modelled breathing in Figure 4. There is clearly a dependence of the size of the PSF on the focus over the range typically encountered during an orbit. Short exposures will generally sample different degrees of defocus, while long exposures will blend the effects of the changing focus.

### 3. Monitoring of Image Quality

Monitoring of image quality by observing open cluster NGC 188 with two filters per detector began with SMOV WFC3 programs 11436, 11437 (1 visit each) and continues with cycle 17 CAL/WFC3 programs 11918, 11920 (3 visits each). The values of the PSF metrics for a visit depend on the OTA breathing sampled by the exposures, as shown in Figures 3 and 4. Analysis of image stability must therefore take breathing into account.

A summary of PSF measurements is shown Table 1 for exposures made with filter F621M on 1 Aug 2009 (program 11436, visit 01) and 21 Nov 2009 (program 11918, visit 01). Each visit included four F621M exposures (2x30 sec) taken with steps ~25 arcsec between exposures, producing ~40 measurable stellar images per exposure. The mean, minimum value, maximum value, and standard deviation are shown for each PSF metric and for the modeled breathing of the exposures.

<table>
<thead>
<tr>
<th></th>
<th>EE.15</th>
<th>EE.20</th>
<th>EE.25</th>
<th>sharpness</th>
<th>peak</th>
<th>FWHM (pix)</th>
<th>ellip</th>
<th>breathing (microns)</th>
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<tr>
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<td>mean</td>
<td>0.587</td>
<td>0.709</td>
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<td>0.139</td>
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<td>0.110</td>
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<tr>
<td></td>
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<td>0.614</td>
<td>0.731</td>
<td>0.809</td>
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<td>0.182</td>
<td>1.925</td>
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<tr>
<td></td>
<td>std</td>
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<td>0.011</td>
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<td>0.066</td>
<td>0.027</td>
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<td>11918</td>
<td>mean</td>
<td>0.601</td>
<td>0.720</td>
<td>0.799</td>
<td>0.060</td>
<td>0.149</td>
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<td>min</td>
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<td>0.053</td>
<td>0.117</td>
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<td>0.016</td>
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</table>

The first three columns of metrics show the encircled energy, EE.NN, which is the fraction of flux within a diameter of 0.NN arcsec. The next column gives the sharpness, which is the sum of the pixel values in the unity-normalized PSF times itself. The next column, "peak", is the fraction of flux in the peak pixel. The full width half maximum (FWHM) and ellipticity (ellip) are given in the following columns. The final column (breathing) gives the estimated relative spacing of the primary and secondary mirrors from the breathing model. The encircled energies and "peak" have been adjusted by 5% to include the flux in the wings measured in the UVIS PSF Wings program 11438. (See the following Section).

The PSFs in the later visit are more compact because they were taken when the breathing produced better focus. Fortuitously, all of the exposures in that visit are at nearly the same focus, so they can be combined to investigate PSF characteristics as a function of location on the detector, as shown in Figure 5.
4. Characterizing PSF Wings

Figures 6 and 7 show images of deep PSFs made from exposures of the isolated star GD153 in the SMOV PSF wings programs (11438, 11439). The images are \( \sim 20 \) arcsec on a side, uncorrected for geometric distortion, displayed with a log stretch of 6 dex. Field objects are also visible. Faint ghosts due to window reflections can be seen in the UVIS images.

The UVIS detector is tilted along a diagonal with respect to the principal ray, resulting in a rhomboidal projection on the sky and slightly non-perpendicular diffraction spikes in the image. The IR detector is tilted about the X axis, resulting in a nearly rectangular projection on the sky and diffraction spikes that appear as an elongated X in the image.

5. Future Work

Analysis of the SMOV programs on image quality and PSF wings has been presented in WFC3 ISRs 2009-37 and 2009-38. Cycle 17 CAL/WFC3 data is being analyzed as described here, along with other data as needed, to characterize the changes in PSFs on the timescale of orbits (due to breathing), to monitor the evolution of typically observed PSFs on the timescale of months (due to gradual changes in focus), and to characterize the dependence of PSF properties on location on the detector. The results will be used to inform users of PSF properties and to assess the possible need for optical re-alignment of the WFC3 optics.

References

Hartig, G.F. 2009, STScI ISR WFC3 2009-38
Hartig, G.F. 2009, STScI ISR WFC3 2009-37
Figure 1: In this typical WFC3/UVIS F410M exposure of the open cluster NGC 188 from program 11877, circles indicate the locations of stars used to determine PSF metrics.
Figure 2: Mean FWHM (pixels) and error in the mean for stars in WFC3/UVIS F410M exposures from program 11877 as a function of date of observation.
Figure 3: Modeled breathing (changes in the spacing between the HST primary and secondary mirrors in microns) as a function of time over the course of two orbits for each of the first ten visits in program 11877. Squares indicate the times of the WFC3/UVIS F410M 30 sec exposures analyzed here. Time intervals of 1 to 5 weeks between the visits have been removed to make the structure of the curves visible.
Figure 4: Mean FWHM (pixels) and error in the mean for stars in WFC3/UVIS F410M exposures from program 11877 as a function of modelled breathing (microns of mirror spacing) at the time of the exposure.
Figure 5: PSF metrics for filter F621M as a function of position on the WFC3/UVIS detector. Color represents Encircled Energy within 0.25 arcsec diameter (violet: low to red: high). Symbol size (larger) represents FWHM (larger).
Figure 6: Deep WFC3/UVIS PSFs near amp B for filter F275W (left) and near amp A for F625W (right), ∼20 arcsec on a side, uncorrected for geometric distortion, displayed with a log stretch of 6 dex. Ghosts due to window reflections are apparent, with orientation dependent on location on the detector.

Figure 7: Deep WFC3/IR PSFs for filters F098M (left) and F160W (right), ∼20 arcsec on a side, uncorrected for geometric distortion, displayed with a log stretch of 6 dex. Field objects are visible in these deep exposures of an “isolated” star, GD153.
An example reduction of WFC3/IR slitless data

M. Kümmel, H. Kuntschner, J. R. Walsh

Space Telescope - European Coordinating Facility, Karl-Schwarzschild-Str. 2, D-85748 Garching b. München, Germany

H. Bushouse
Space Telescope Science Institute, 3700 San Martin Drive, Baltimore, MD 21218, USA

A. Straughn
Astrophysics Science Division, Goddard Space Flight Center, Code 665, Greenbelt, MD 20771, USA

Abstract. The G102 and G141 grisms in the WFC3 IR channel together cover the wavelength range 0.8 – 1.6 μm with resolutions of $R \sim 210$ and $R \sim 130$, respectively.

In this contribution we present an example reduction of G102 and G141 slitless data taken during the WFC3 Early Release Science programme. All core tasks are done by the spectroscopic extraction software aXe, which was specifically designed to handle data from the HST slitless modes. The reduction scheme includes a sky background subtraction with a master sky image to achieve a homogeneous, flat background for the spectral extraction and the most recent on-orbit calibrations. Moreover we present a new method for co-adding data from individual slitless images that, similar to MultiDrizzle in direct imaging, allows the rejection of deviating pixels from e.g. cosmic ray hits. As a result of applying this processing we show the spectra of some emission line galaxies down to $\text{mag}_{AB}(F140W) = 24.5$, demonstrating the remarkable efficiency and capability of the WFC3 NIR grisms.

1. The data

The data that is the basis for this work has been taken as part of the Wide Field Camera 3 (WFC3) Early Release Science (ERS, PI: R. O’Connell, PID: 11359) programme in October 2009. The ERS grism field, which is located in the GOODS South field (see Straughn et al. 2010), was observed in two orbits per grism (G102 and G141) together with the corresponding direct imaging taken with the filters F098M and F140W. For every grism image, an associated direct image was taken within the same pointing to allow the projection of the position and the sizes of the dispersed objects to the grism image using the WCS header information (Kümmel et al. 2009b) with high accuracy. Table 1 lists all data taken in the ERS grism field.

2. Combining the direct images

To generate a list of objects whose spectra shall be extracted from the grism images, we combine the direct images with the MultiDrizzle software (Koekemoer et al. 2005, Koekemoer et al. 2010). The WFC3/IR grisms are very sensitive to the detection of faint emission lines of compact objects. In extreme cases when the spectrum is dominated by the emission
line flux (and thus very low continuum flux), these objects can appear very faint in direct imaging. To detect these objects, we first multidrizzle the F098M and the F140W images independently and then co-add these images to form a deep detection image. The STSDAS task tweakshifts (in the dither-package) was used to improve the registration of the direct images. The residual shifts were on the order of $\sim 0.5$ pixels.

Figure 1 shows the F098M/F140W combined detection image. While the overall image quality is very good, some persistence effects (Long et al. 2010) from the slitless images taken immediately before each direct image can clearly be identified (on the right of the bright star in the lower left corner of the image; around the middle of the left edge of the image, resulting from a bright star outside of the direct image FOV).

<table>
<thead>
<tr>
<th>G102</th>
<th>expt. [s]</th>
<th>F098M expt. [s]</th>
<th>G141 expt. [s]</th>
<th>F140W expt. [s]</th>
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</tbody>
</table>
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Figure 2: The G141 image $ib6o23rsq$ (left), the G141 master sky image (middle) and the background subtracted image $ib6o23rsq$ (right). The cut levels are $[0.6, 1.15]$, $[0.9, 1.1]$ and $[-0.05e/s, 0.5e/s]$, respectively.

3. Source detection and object lists

The object lists for the direct images were compiled using the SExtractor software package (Bertin & Arnouts 1996). For the spectral extraction, the “multi”-colour information in the F098M and F140W data can improve the estimate of the mutual overlap of spectra (see Kümmel et al. 2009b). In order to retrieve this multi-colour information, we have used SExtractor in the so called “dual-image mode” (see the SExtractor manual), using the combined image for object detection and the F098M and F140W multidrizzled images for retrieving photometric information in matched apertures.

The resulting photometric catalogs were combined and set to the correct format using custom made Python scripts. Spurious detections in the image corners or at the traces coming from persistence effects (Long et al. 2010) were removed. The crosses in Fig. 1 mark the positions of all $\sim 550$ objects in the final list. The spectra of these objects were extracted from the G102 and G141 data.

4. Spectral extraction

4.1. The aXe spectral extraction software

Slitless spectroscopic data has some specific problems (Walsh et al. 2010) that are not addressed in the standard spectral extraction tasks offered in public data reduction packages (e.g., IRAF, IDL or MIDAS). The aXe software (Kümmel et al. 2009b) was designed to extract spectra in a consistent manner from all the slitless spectroscopy modes provided by WFC3 and the Advanced Camera for Surveys (ACS). aXe is distributed as a PyRAF/IRAF package with several tasks which can be successively used to produce extracted slitless spectra. aXe is distributed as part of the STSDAS software package. Within STSDAS, aXe is located under the sub-packages ‘analysis.slitless.axe’, and the version 2.1 of aXe described here is part of STSDAS 3.12. aXe has successfully been used in several large science programs such as GRAPES (Pirzkal et al. 2004), PEARs (Pirzkal et al. 2009) and within the Hubble Legacy Archive programme (Kümmel et al. 2009a).

4.2. Configuration and calibration files

While specifically developed for HST slitless data, by design aXe was intended to be instrument-independent. This aim was achieved by placing all the instrument specific parameters and sensitivities into external configuration and calibration files. The G102 and G141 configuration and calibration files used to reduce the ERS grism data have been de-
WFC3/IR grism example reduction

# iraf.iolprep(mdrizzle_image='F140W_drz.fits', input_cat='F098140_clean.cat')
#
# iraf.axeprep(inlist="G141.lis", configs="WFC3.IR.G141.BCK.conf", backgr="YES", 
# backims='G141_Apr09.fits', mfwhm=3.0, norm='NO')
#
# iraf.axecore(inlist="G141.lis", configs="WFC3.IR.G141.V1.0.conf", back="NO", 
# extrfwhm=4.0, drzsfwhm=3.0, backfwhm=0.0, slitless_geom="NO", 
# orient="NO", exclude="NO", cont_model="gauss", model_scale=3.0, 
# inter_type="linear", interp=1, np=15, niter_med=3, niter_fit=3, 
# kappa=3.0, smooth_length=10, smooth_fwhm=1.0, spectr="YES", 
# adj_sens="NO", weights="NO", sampling="drizzle")
#
# iraf.drzprep(inlist="G141.lis", configs="WFC3.IR.G141.V1.0.conf", 
# opt_extr="YES", back="NO")
#
# iraf.axedrizzle('G141.lis','WFC3.IR.G141.V1.0.conf', 4.0, 3.0, back='NO', 
# driz_separate='YES', clean='YES', combine_type='median', 
# combine_maskpt=0.7, combine_nsigmas="4.0 3.0", combine_nlow=0, 
# combine_nhigh=0, combine_lthresh="INDEF", combine_hthresh="INDEF", 
# combine_grow=1.0, blots_interp='poly5', blots_sinscl=1.0, 
# driz_cr_snr="5.0 4.0", driz_cr_grow=1, driz_cr_scale="1.2 0.7", 
# makespc='YES', adj_sens='NO', opt_extr='YES')

Figure 3: A python script used to run the entire aXe reduction of the ERS G141 data

rived using on-orbit HST data and are described in Kuntschner et al. (2009a), Kuntschner et al. (2009b) and Kuntschner et al. (2010). The calibration of the ACS grism is given in Pasquali et al. 2006.

For the spectral extraction of many objects over a large FOV, such as in the ERS grism field, the accurate subtraction of the sky background is critically important. Using \( \sim 100 \) publicly available, deep slitless science images for each grism, we have compiled master sky images for G102 and G141. For each grism the science images were scaled to the same average sky level and then combined to the master sky image. Object signatures were masked and removed using kappa-sigma rejection techniques (Kümmler et al. 2010). Figure 2 shows an original G141 image, the G141 master sky image and the background subtracted G141 image from the ERS data on the left, center, and right panels, respectively. The numerous dead pixels appear as white (\textit{value} = 0.0) spots. The background subtracted G141 image from the ERS data on the right panel does not show any gradients or inhomogeneities, thus illustrating the high quality of the master sky image.

4.3. Using aXe

aXe extracts the slitless spectra semi-automatically. After an interactive preparation of the detection image and object catalog, the extraction runs automatically, and the user can concentrate on optimizing the parameters of the various aXe tasks in repeated runs on the data. In addition to co-adding the direct images and preparing the object list (Sect. 2. and 3.), the setup includes organizing the input such as preparing and listing for each grism image the associated direct image. The actual aXe reduction, e.g. for the ERS G141 data, is then accomplished by running the sequence of tasks, as listed in Figure 3. The purpose of the individual tasks is:

- \texttt{iolprep}: The coordinates of all objects in the object list are projected into the coordinate system of every single direct image. Partial object lists covering the FOV of each direct image are generated.
Figure 4: The reduction and source spectrum extraction of WFC3/G102 data with aXe (observations of the planetary nebula Vy2-2). The left panels show an example of a G102 dispersed image (top) together with its associated direct image in the F098M filter (bottom). The right panels show the combined, aXedrizzled stamp image from the entire dataset (top) and the extracted one-dimensional spectrum (bottom) with the identification of prominent emission lines.

- **axeprep**: The spectral reduction is prepared. For WFC3 grism data, this task performs the sky background subtraction.

- **axecore**: Performs the spectral extraction on the single grism images. That is marking the spectral beams for all objects, estimating the contamination from neighboring objects, collecting all spectral beams in a special transport format and flat-fielding all beams.

- **drzprep**: Prepares the subsequent aXedrizzle step by re-assembling the spectral beams and computes the coefficients to co-add the individual 2D grism stamp images.

- **axedrizzle**: Co-adds for each object the individual spectral beams to a deep 2D grism stamp image with constant dispersion and extracts the spectra from these stamp images. There is an option to detect and flag deviant pixel values as MultiDrizzle does in direct imaging before extracting the final spectrum from the co-added grism stamps.

Figure 4 illustrates the aXe extraction on the basis of the G102 calibration data taken for the planetary nebulae Vy2-2. The position on the F098M direct image (lower left panel) is transported to the G102 grism image (upper left panel). There the spectral beam is marked and the pixel values are flat fielded. aXedrizzle co-adds the stamp images from the individual grism images to a deep, 2D grism stamp image (upper right panel) and extracts the final, deep 1D spectrum from it (lower right panel). Some prominent emission lines are identified in the spectrum.
Figure 5: The comparison of a 2D drizzled G141 image produced with the “basic” aXedrizzle (left column) and the new aXedrizzle that can reject deviant pixels (right column). Unnoticed hot or cosmic ray affected pixels are detected and masked out in the new aXedrizzle.

4.4. aXedrizzle with pixel rejection

In WFC3 IR data, cosmic ray hits are detected and rejected during the “up-the-ramp” fitting of the individual detector reads. However this procedure can miss flagging some of the CR affected pixels. In direct imaging, MultiDrizzle (Koekemoer et al. 2006, 2010) can detect deviant pixel values when combining dithered data.

Similar to MultiDrizzle we have extended aXe to combine spectral images, in the axedrizzle task, to be able to detect deviating pixel values. This is accomplished by individually drizzling the 2D spectra, median-combining them, blotting them back and then identifying cosmic rays by comparing the original 2D spectra with the blotted versions.

Figure 5 shows a comparison of a deep, aXedrizzle combined 2D grism stamp image reduced with the “basic” aXedrizzle and the new aXedrizzle in the left and right columns, respectively. There is a cluster of pixels with quite high values in the data, which are just co-added in the “basic” aXedrizzle procedure and result in an artificial emission feature at ~15800Å. In the new aXedrizzle on the right side, these pixels are identified as deviant and rejected, and the artificial emission feature no longer exists in the 2D grism stamp and the extracted spectrum.

As in MultiDrizzle the pixel rejection in aXedrizzle improves the results only in case of properly registered and dithered data, e.g. if there are astrometric shifts in the grism data that are not recorded in the WCS, aXedrizzle would reject many good pixels as cosmic rays and thus result in vastly different and wrong spectra. A careful comparison of the results obtained with and without rejecting pixels in aXedrizzle should be an important and integral part of the reduction process.

4.5. Viewing the results

The results of the aXe extraction outlined in the previous sections are ~550 spectra covering a large range of signal-to-noise ratios, plus auxiliary data such as the 2D grism stamp images produced by aXedrizzle. The visual checking of such a large number of spectra is very tedious.
The first column with an index number is followed by the object ID number, the object brightness and a column with the positional information. The further columns show the direct image stamp, the 2D grism stamp image and the extracted spectrum in [e/s] and physical units.

For this purpose we have developed the tool aXe2web\(^1\) (Walsh & Kümmel 2004), which produces web pages that show for each source in the extraction catalogue a postage stamp image from the direct image, the 2D grism stamp image and a 1D spectrum in units of detected counts and flux calibrated. An example of such a web page is given in Figure 6. These pages are a useful tool to provide a quick overview of the extracted spectra, to assess the quality of the extraction process and to find interesting objects (e.g., high redshift galaxies, emission line objects, etc.) which can be highlighted for further study or interactive spectrum extraction.

In order to facilitate the navigation within a dataset, an overview and an index page accompany the object pages which carry all the information. The overview page contains, for each object, the basic information sequence number, reference number, X, Y, RA, Dec and magnitude. The index page includes a table with the ordered reference number of all objects. Direct links from both the overview page and the index page lead to the corresponding locations of the objects in the object pages.

\(^1\)see: http://www.stecf.org/software/slitless\_software/axe/axe2web.php
Figure 7: Selected examples of spectra extracted from the ERS grism data set. The G102 (red) and G141 (blue) spectra are plotted together without any scaling. The smooth transition illustrates the high accuracy of the flux calibration. The redshifts were derived from the identifications of at least two emission lines as marked in the panels. Each panel also shows a typical 1-σ error bar for G102 and G141.

5. Results

Figure 7 shows selected spectra from the ERS grism data set. The object identifications are derived from the object numbers in the SExtractor catalog. All four objects are emission line galaxies with at least two line identifications in the spectral range covered by the WFC3/IR grisms. Three sources had already been identified as emission line galaxies in the PEARS survey (Pirzkal et al. 2004), and one object (ID258) was identified as emission line object in the WFC3 data. An in-depth analysis of all emission line objects in the ERS grism data is given in Straughn et al. (2010). Although no scaling has been applied to the data, the transition from the G102 (red) to the G141 (blue) is remarkably smooth, which illustrates the high accuracy of the flux calibration and the quality of the spectral extraction.

Even at a rather shallow depth of 2 HST orbits per grism, the ERS grism observations yield spectra of hundreds of faint objects and allow the detection of emission line galaxies down to $m_{AB}(F098M) = 24.5$. These results demonstrate the remarkable efficiency and capability of the WFC3/IR grisms for measuring galaxy properties to faint magnitudes.

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WFC3/IR Filter Wedge

Tiffany Borders\(^1\), Elena Sabbi\(^1\), and John MacKenty\(^1\)

Abstract. In order to verify if any of the WFC3/IR imaging filters presents variations in its thickness (filter wedge) we observed the young and bright star cluster NGC 1850 through all the WFC3/IR filters consecutively within 1 orbit. We found that, with the exception of the F098M and F126N filters, the positions of stars observed though different filters differ on average by less than 0.14 ± 0.06 pixels.

1. Introduction

A key feature of the Wide Field Camera 3 (WFC3) is its panchromatic wavelength coverage. The optical design features two independent channels, one operating in the wavelength range 200 to 1000 nm (the UVIS channel), and the other sensitive to the 800 to 1700 wavelength range (the IR channel). WFC3/UVIS is equipped with 62 wide-, medium-, and narrow-band filters in the UVIS channel, plus one low-dispersion grism, while WFC3/IR has 15 wide-, medium-, and narrow-band filters, 2 grisms, and an opaque element.

Thickness variations in a filter can modify the intensity distribution of the light beam, an effect called filter wedge. Filter wedge introduces a slight shift in positions when comparing observations of astronomical sources observed without moving the telescope. If the two faces of a filter are not coplanar (or wedged) the displacement of the light beam will cause a rigid shift of the entire field of view.

During Cycle 17 proposals 11913 and 11923 (P.I. MacKenty) were executed to measure the relative displacement of stars in the cluster NGC 1850 caused by filter wedge in the IR and UVIS respectively. This report summarizes the results obtained for the IR channel using the data acquired in proposal 11913. Results for the UVIS tests will be presented in a forthcoming ISR (Borders et al. 2010).

2. Observations

The IR channel is equipped with five broad-band filters (F105W, F110W, F125W, F1450W, and F160W), four medium-band filters (F098M, F127M, F139M, and F153M0), and six narrow-band filters (F126N, F128N, F130N, F132N, F164N, and F167N), plus two low-dispersion grisms (G102 and G140). Filters and grisms are housed in an 18 slot wheel, called the Filter Select Mechanism (FSM), with the 18th slot reserved to the opaque element (or BLANK). For IR observations, the selected filter is rotated into the light beam. The FSM can rotate in both directions in the IR channel and always takes the shortest path to a new filter position. Figure 1, from the Wide Field Camera 3 Instrument Handbook for Cycle 18, shows how the filters are housed in the FSM. The first household on the right side of the BLANK element housing the F160W filter is defined as position 1 of the FSM while filter F164N is in position 17.

We observed the young and bright star cluster NGC 1850 in the Large Magellanic Cloud using all the IR filters in order to verify if any of the WFC3/IR filters are affected

\(^1\)Space Telescope Science Institute, 3700 San Martin Dr, Baltimore, MD 21218

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by wedge. NGC 1850 has been chosen because it has a low amount of extinction and hosts several bright red and blue stars, which makes it a good target for both the IR and UVIS channels.

For each filter in the IR channel we acquired one short exposure in the MULTIAC-CUM mode. To minimize shifts between consecutive observations (e.g. due to guide star (re)acquisition) all the images were acquired within the same orbit. Finally to minimize the read-out overheads all data were acquired using a default 512 x 512 sub-array. The data were acquired starting with the F105W filter (corresponding to position 5 in Figure 1) and then rotating the FSM clockwise toward the F125W filter. We also acquired an extra F105W image at the end of the orbit to verify the telescope had not moved during the observations (i.e. because it lost the Guide Star). A log of the observations are reported in Table 1 including the image name, filter, read mode, number of samplings, and total exposure time in seconds.

3. Analysis

All raw images were processed through the IRAF \textit{calwf3} (version 1.7) for calibration. The calibration steps included the overscan correction, dark file subtraction, and flat-field correction (DQICORR, BLEVCORR, DARKCORR, and FLATCORR turned on). The reference files employed were t291659mi\_bpx.fits, t2c16200i\_ccd.fits, q911321mi\_osc.fits, t5s1754di\_drk.fits and sca2025ti\_pfl.fits, for the bad pixel table, ccd table, overscan table, superdark file, and the pixel to pixel flat field file, respectively. For the IR channel, the pipeline applies an “up-the-ramp” fit through the non destructive reads to remove the cosmic rays and recover the information associated with saturated stars, and finally it converts the images from counts to count rates (refer to Appendix E of the Wide Filed Camera 3 Instrument Handbook for a detailed description of the calibration steps performed by \textit{calwf3}).

We measured the position of the stars on each image using DAOPHOT in IRAF and the sources were identified using DAOFIND in IRAF. Aperture photometry was performed using the IRAF task PHOT. To refine the positions of the detected sources for each image
Table 1: Log of observations for proposal 11913 including file name, filter, read mode, number of samples, and the total exposure time in seconds. Data were taken on February 6th, 2010.

<table>
<thead>
<tr>
<th>Image Name</th>
<th>Filter</th>
<th>STEP</th>
<th>NSAMP</th>
<th>Total Exposure Time (sec)</th>
</tr>
</thead>
<tbody>
<tr>
<td>ibcg01fqiq</td>
<td>F105W</td>
<td>RAPID</td>
<td>6</td>
<td>17.549</td>
</tr>
<tr>
<td>ibcg01fhiq</td>
<td>F098M</td>
<td>RAPID</td>
<td>6</td>
<td>17.549</td>
</tr>
<tr>
<td>ibcg01fwq</td>
<td>F140W</td>
<td>RAPID</td>
<td>6</td>
<td>17.549</td>
</tr>
<tr>
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<td>F153M</td>
<td>RAPID</td>
<td>8</td>
<td>23.458</td>
</tr>
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<td>23.458</td>
</tr>
<tr>
<td>ibcg01fwhq</td>
<td>F127M</td>
<td>RAPID</td>
<td>6</td>
<td>17.549</td>
</tr>
<tr>
<td>ibcg01flq</td>
<td>F128N</td>
<td>RAPID</td>
<td>10</td>
<td>29.323</td>
</tr>
<tr>
<td>ibcg01foq</td>
<td>F129N</td>
<td>RAPID</td>
<td>10</td>
<td>29.323</td>
</tr>
<tr>
<td>ibcg01fpoq</td>
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<td>RAPID</td>
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</tr>
<tr>
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<tr>
<td>ibcg01fpoq</td>
<td>F170W</td>
<td>RAPID</td>
<td>6</td>
<td>17.549</td>
</tr>
</tbody>
</table>

We made an empirical PSF, by selecting ∼10 well isolated stars, and then used ALLSTAR to perform PSF-fitting on each image. This method also allowed us to reject non-stellar objects such as cosmic rays (CRs) and hot pixels that were not removed by calwf3, as well as galaxies and blended stars. In order to reject non-stellar objects we used two of the statistical parameters provided by ALLSTAR to evaluate how close a detected source is to the empirical PSF. In particular, we used the $\chi^2$ of the residual and kept only those sources with $\chi^2 < 1.5$. We also used the sharpness parameter to discard those objects that were either too sharp (sharpness $<-0.2$) or too broad (sharpness $>0.2$) to be a star.

We matched together all the catalogs and, for each star, we measured the shift between the X and Y positions on the original frame and the reference image (ibcg01fqiq.fits) in the F105W filter (see Table 1). A 3$\sigma$ clipping rejection was used to reject possible false matches, due to the difference in the crowding conditions of the cluster on the various images. The difference along the X and Y directions are shown in Figures 2, 3, 4, and 5. With the exception of filters F098M and F126N, where the shifts with respect of the reference image is more than 0.3 pixels, the average shift is 0.14 ± 0.06 pixels. The average shifts are listed in Table 2 for each filter.

To verify if there is any dependence with wavelength we plotted the shifts along the Y axes (upper panel), the X axes (middle panel), and the total shifts (lower panel) as a function of wavelength shown in Figure 6. No obvious relation between the total shift and the wavelength are found. We did, however, find a moderate dependences with wavelength along both the X and Y axes, consistent with the tilted angle of ∼8.6° between the RCP (which applies the spherical-aberration correction) and the field center chief ray.

4. Summary and Conclusions

During Cycle 17 we observed the young star cluster NGC 1850 (calibration proposal 11913) to verify and quantify if any of the WFC3/IR imaging filters are affected by wedge. By observing the cluster in all the WFC3/IR filters within 1 orbit we were able to directly measure the displacement of the sources over the detector in each filter.
Figure 2: Difference between the position of the stars in the images acquired with the F098M, F140W, F153M and the F139M filters, and the position on the reference image (F105W). The red lines highlight the 0,0 (no shift) position.

Figure 3: Same as Figure 2 but for filters F127M, F128N, F130N, and F132N.
Figure 4: Difference between the position of the stars in the images acquired with the F126N, F167N, F164N and F160W, and the position on the reference image (F105W). The red lines highlight the 0,0 (no shift) position.

Figure 5: Same as Figure 4 but for filters F125W, F110W, and F105W.
Figure 6: Average shifts of the stars measured in each image with respect to the reference image (F105W) as a function of wavelength along the Y axes (Top Panel), X axes (Middle Panel), and the total shift (Bottom Panel).

With the exception of the F098M and F126N filters, which show a displacement relative to the reference image greater than 0.3 pixels, the average shift for all the other filters is 0.14 ± 0.06 pixels. We also find that the displacement of the detected sources increases along the X axes and decrease along the Y axes as a function of wavelength. This behavior is consistent with the tilted angle of about 8.6° between the RCP (which applies the spherical-aberration correction) and the field center chief ray.

References

WFC3: Internal Monitoring of UVIS Charge Transfer Efficiency

V. Kozhurina-Platais\textsuperscript{1}, S. Baggett\textsuperscript{1}, B. Hilbert\textsuperscript{1}

Space Telescope Science Institute, Baltimore, MD, 21218

Abstract. We present the results of one year of Charge Transfer Efficiency (CTE) monitoring of the WFC3 UVIS detector, based on data acquired during the monthly internal Extended Pixel Edge Response (EPER) observations. We present an algorithm for CTE assessment and fit a power-law to CTE measures versus signal level. We find that at each signal level, CTE declines linearly over time and CTE losses are worst at the lowest signal levels. We compare WFC3/UVIS in-flight results to similar ACS/WFC in-flight results 400 days after installation.

1. Introduction

It is well-known that there is a gradual loss of Charge Transfer Efficiency (CTE) for HST CCDs in the environment of space: radiation damage to the crystalline structure of the detector creates defects which trap and release charge. A degradation in CTE was observed in all HST CCDs: WFPC2 (Whitmore \textit{et al.} 1999, Dolphin 2000), STIS (Gilliland \textit{et al.} 1999; Goudfrooij & Kimble 2003; Goudfrooij \textit{et al.} 2007) and ACS (Riess 2003; Muchler & Sirianni 2005). The degradation of the CCD detectors due to CTE loss affect the precision of stellar photometry and astrometry on HST science programs (Riess & Mack 2004; Kozhurina-Platais \textit{et al.} 2007; Kozhurina-Platais \textit{et al.} 2008; Chiaberge \textit{et al.} 2009, Anderson & Bedin, 2010). Therefore it is important to monitor WFC3 UVIS CTE degradation over time.

In this paper we report CTE measurements for the new instrument Wide Field Camera 3 (WFC3), a fourth - generation imaging instrument which was installed in HST during Servicing Mission 4 in May 2009. A popular technique of CTE measurement is the Extended Pixel Edge Response (EPER) described by Janesick (2001). EPER measures the excess charge in the CCD over-scan pixels (extended pixel region), which appears as an exponentially decreasing tail moving across the over-scan pixels away from the science pixels. The EPER technique has been successfully used for the ACS Wide Field Camera (WFC) and High Resolution Camera (HRC) to measure and monitor CTE (Muchler & Sirianni, 2005). The same technique was used by Robberto (2007) to measure CTE for the WFC3 UVIS-2 detector during the ground-based ambient test campaign in April 2007, and by Kozhurina-Platais \textit{et al.} 2009, using observations for the flight detector UVIS-1 taken during the ground-based Thermal Vacuum 3 (TV3) test on April 10, 2008.

Here, we present the results of one year of CTE measurements based on data acquired during the monthly internal EPER observations (CAL–11924, PI Kozhurina-Platais) with the internal TUNGSTEN lamp.

2. Observations and Analysis

The calibration program of CTE measurements consists of internal TUNGSTEN lamp flat field observations of short exposures through three filters, F390M, F390W and F438W, in order obtain a large range of illumination levels, namely 200, 400, 800, 1600 and 5000
electrons. The observation of flat fields with different levels of illumination will be used to monitor the CTE detector with time, by measuring the signal profiles into the trailing over-scan region (EPER), as it described by V.Kozhurina-Platais et al. 2009. Table 1 lists one set of the typical monthly observations with TUNGSTEN lamp, the filters and the exposure times used to generate the average signal levels in the flat fields images.

<table>
<thead>
<tr>
<th>Image ID</th>
<th>Filters</th>
<th>Exp. Time (sec)</th>
<th>Intended Level (e⁻)</th>
<th>Measured Level (e⁻)</th>
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</thead>
<tbody>
<tr>
<td>ibc620dmq</td>
<td>DARK</td>
<td>0.5</td>
<td>-</td>
<td>-</td>
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<tr>
<td>ibc621zeq</td>
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<td>0.5</td>
<td>-</td>
<td>-</td>
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<td>200</td>
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<td>400</td>
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<tr>
<td>ibc621zjq</td>
<td>F438W</td>
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</tr>
</tbody>
</table>

The WFC3 flight software does not allow for a true bias (with exposure time equal 0) to be taken in EPER mode, therefore the dark exposures are used as a bias. Thus, the first two images (ibc620dmq and ibc621zeq) listed in Table 1 are dark exposures with 0.5sec, were used as bias reference images.

Before describing the data reduction techniques, we outline here the WFC3 CCD detector format and its structure as they relate to EPER measurements. As described in the WFC3 Instrument Handbook for Cycle 17 (Bond et al. 2007), each of the default readout format of the UVIS CCD detectors is 4096 × 2051 science pixels. Both CCD chips have 25 extra columns on each side of the detectors, which form the serial physical over-scan regions. Each CCD chip also has 30 × 2 virtual serial over-scan areas in the center of each row. Next to the inter-chip gap there are 19 rows of parallel overscan for each CCD chip. Figure 1 shows the format of a raw image obtained with the nominal full-chip four-amplifier readout.

Figure 1: Schematic illustration of the standard WFC3 UVIS image from Figure 6.1 of the WFC3 Instrument handbook (Bond et al. 2007). The enclosed pale yellow boxes mark the CCD image area (science pixels). The bright yellow color shows serial physical over-scan and the pink color indicates serial virtual and parallel virtual over-scan.
For EPER measurements, readouts have been modified such that the over-scan regions contain up to 300 pixels of serial over-scan and 300 pixels of parallel over-scan for each CCD chip. Figure 2 shows the format of an EPER sub-array exposure. As can be seen, the CCD image area associated with each amplifier (indicated in yellow) is only 1635 × 1770 pixels. Pixels which are not read out are shown in white.

Figure 2: Schematic illustration of the EPER WFC3/UVIS sub-array exposures from Roberto (2007, Figure 2). The yellow color indicates the CCD image area while the white color indicates areas that are not readout. The blue vertical stripes show the area of the parallel virtual over-scan and the grey color denotes areas of serial virtual over-scan.

Following the EPER technique described by Janesick (2001), we measure the amount of deferred charge in the extended pixel region. Several lines are averaged to reduce noise and improve the signal-to-noise ratio in the extended pixel region. Then, CTE is calculated from the ratio of the total deferred charge in the extended pixel region $S_D$ (in e$^-$) over the charge level of the last column of trailing over-scan region $S_{LC}$ (in e$^-$), multiplied by $N_p$ which is the number of signal transfers to the CCD register ($N_p = 2051$ for WFC3), namely (Janesick, 2001, Eq. 5.21):

$$ CTE_{EPER} = 1.0 - \frac{S_D}{S_{LC}} \times N_p $$

(1)

It is well-known that CTE depends on signal level, and approaches the ideal of 1.0 at high signals, similar to ACS/WFC & HRC (Mutchler & Sirianni 2005) and STIS (Goudfrooij et al. 2006). This asymptotic distribution of CTE vs. signal level has a power-law relationship, which can be defined in the following form:

$$ CTE = 1 - m \times (S_{LC})^{\rho} $$

(2)

where $\rho$ and $m$ are free parameters and $S_{LC}$ is the signal level of the last columns in electrons. The numerical implementation of the power-law fitting was realized by employing...
a non-linear least-square fit using the IDL library by C. Markwardt (2006). Figure 3 shows over more than one year of the CTE measurements and uncertainties for each individual exposure as a function of the signal level in the last column of the trailing region \(S_{LC}\). Uncertainties were calculated by dividing the standard deviation of \(S_D\) by the denominator in Eq.(1). The over-plotted red line is the best power–law fit.

On the other hand, CTE can be specified in terms of charge transfer inefficiency (CTI), which is the fraction of charge left behind in a single pixel transfer i.e. \(CTI = (1 - CTE)\). Then, Eq.2 can be rewritten as a linear equation of the \(\log\)–\(\log\) of the parameters \(m\) and \(\rho\), namely:

\[
\log(CTI) = \log(m) + \rho \times \log(S_{LC})
\] (3)

Figure 4 (top panel) shows the \(CTI\) calculated from Eq.(2) for each amplifier as function of signal level \(S_{LC}\) (in e\(^-\)), where different symbols show the \(CTI\) measurements for all four amplifiers - the asterisk symbols are for amplifier A, diamonds for amplifier B, triangles and squares are for amplifiers C & D respectively. The red symbols are the averaged \(CTI\) of the four amplifiers and the solid red line is the best linear fit of Eq.(3).
Figure 4: The top panel shows the log–log plot of CTI vs. signal level $S_{LC}$, the solid red lines are the best linear fit. The middle panel shows $\rho$ parameter vs. time (Modified Julian Date MJD), over-plotted red line is the best linear fit. The bottom panel shows $m$ as function of MJD with over-plotted red line as the best linear fit.
The middle panel of Figure 4 shows that the parameter \( \rho \) (the slope of the log-log CTI) vs. signal level which has been relatively constant over time for WFC3, at the level of \( \sim -0.6 \). The value of \( \rho \) has been time-independent in other HST CCDs as well, implying that this parameter may be a measure of detector characteristic(s) intrinsic to the device, i.e., established during the manufacturing process. For comparison, the value of \( \rho \) was measured at -0.61 and -0.85 for ACS/WFC and HRC, respectively (Mutchler & Sirianni, 2005) and -0.82 for STIS (Goudfrooij et al., 2006); there is no value for WFPC2 as those chips were not able to take images in EPER-format.

The bottom panel of Figure 4 shows \( m \) parameter (intercept of the log-log CTI) vs. time. As it expected, CTI is increasing over time as the CTE decreases.

3. Comparison with ACS

The WFC3/UVIS CTE measurements have been compared with the ACS/WFC CTE measurements. Figure 5 shows the WFC3 UVIS parallel EPER CTE measurements at several signal levels over time and illustrates the linear dependence of any given signal level vs. time. The measured UVIS CTE is 0.99998 at a signal level of 160e\(^-\) after \( \sim 400 \) days on board HST. This is similar to the ACS/WFC value (see Figure 7 in Mutchler & Siriani, 2005) of 0.99996 at a signal level of 185e\(^-\) \( \sim 400 \) days after the ACS installation. WFC3/UVIS and ACS/WFC EPER CTE measurements are also similar at higher signal levels for example, the WFC3/UVIS CTE is 0.999993 at \( \sim 1700 \)e\(^-\) after \( \sim 400 \) days, consistent with the ACS/WFC value of 0.99999 measured at \( \sim 1620 \)e\(^-\) after \( \sim 400 \) days. These comparisons indicate that the on-orbit degradation in WFC3/UVIS is similar to that in ACS/WFC.

![Figure 5: WFC3 UVIS parallel EPER measurements at selected signal levels vs. time (MJD). The best linear fit for different signal levels is over-plotted by different color lines.](image-url)
4. Conclusion

The EPER method, which assesses the CTE level utilizing the trailing over-scan in EPER-format images, measures CCD CTE losses via internal exposures. The advantage of this method is that it does not require the use of valuable external observing time and it is an effective, low-impact way monitor the CTE over time. The internal calibration of CTE however, do not lead directly to correction in stellar photometry due to CTE induced flux losses or correction in stellar astrometry due to CTE induced centroid shift. For development of such a correction would require external observations of stellar cluster, for example.

The WFC3/UVIS CTE degradation will continue to be monitored with both internal and external observations. Mitigation of CTI can be achieved by invoking the WFC3 charge injection capability. Analogous to a pref-lash mode, the charge is injected before the science observation, albeit electronically rather than optically, in order to temporarily fill the traps causing the CTI. There will be a noise penalty although and it is expected to be significantly lower (15e- for 10,000 e- injection; Giavalisco 2003) than photon shot noise. The charge injection capability is expected to be made available as an observing option in 2011 or 2012.

WFC3/UVIS EPER measurements spanning over more than 1 year have been presented here. These results have shown, that the CTE at low and high signal levels are similar to the corresponding ACS/WFC EPER measurements during its first year of its operation.

Acknowledgments. We are grateful to A. Martel and P. McCullough for their keen interest in the early study of WFC3/UVIS EPER assessments during the ground-based thermal vacuum campaign in April 2008.

References

Dither and drizzle strategies for Wide Field Camera 3

Max Mutchler
Space Telescope Science Institute
3700 San Martin Drive, Baltimore, Maryland 21218

Abstract. Hubble’s 20th anniversary observation of Herbig-Haro object HH 901 in the Carina Nebula is used to illustrate observing strategies and corresponding data reduction methods for the new Wide Field Camera 3 (WFC3), which was installed during Servicing Mission 4 in May 2009. The key issues for obtaining optimal results with offline Multidrizzle processing of WFC3 data sets are presented. These pragmatic instructions in “cookbook” format are designed to help new WFC3 users quickly obtain good results with similar data sets.

1. Introduction

MultiDrizzle provides a powerful algorithm for combining and cleaning multiple images. It gives the user great control in the application of distortion corrections, exclusion and rejection of bad detector pixels and cosmic rays, enhancing the resolution of subsampled data, and assembling large mosaics.

The MultiDrizzle Handbook (Fruhchter & Sosey, 2009), and the WFC3 Data Handbook (Kim Quijano, 2009) provide detailed descriptions of the philosophy and mechanics of dithering and drizzling. This brief “cookbook” is intended as a quick-start guide that should help new WFC3 users generate reasonable first-pass results.

Observing strategies and data reduction methods for the WFC3 ultraviolet-visible (UVIS) channel inherit much from the Advanced Camera for Surveys (ACS), since they have similar features. The infrared (IR) channel of WFC3 is an extraordinary new capability for Hubble, and there is less conventional wisdom established for its usage. Therefore, the HH 901 example below utilizes a mosaic data set for the F128N filter of the IR channel (HST observing program 12050, PI Mario Livio). A complete and well-documented set of drizzled output images, prepared by the author for all UVIS and IR filters, are available as High Level Science Products (HLSP) in the Multimission Archive at STScI (MAST):

http://archive.stsci.edu/prepds/carina

2. Observing strategy: dithers and mosaics

Since drizzling strategy is largely dictated by the observing strategy employed, in designing a Phase II program it is important to choose a dither or mosaic pointing pattern which will achieve optimal results within the allocated number of orbits. For small-scale dithers, tradeoffs must often be made between optimizing for artifact rejection (e.g. detector features and cosmic rays) or subsampling. Pre-defined convenience patterns are described in the WFC3 Instrument Handbook (Dressel, 2010), and can be selected in APT, but users can also modify patterns (with caution!) or use POS TARGs for greater freedom in shifting and sequencing exposures, to minimize overhead time. Large-scale mosaic shifts can also be defined, but typically require multiple visits (guide star pairs) to execute, which then requires careful image registration.
The most common UVIS dither pattern spans the gap between the two CCDs, and the new camera is still relatively free of the effects of space radiation damage (hot pixels, bad columns, CTE loss). Documented ACS/WFC drizzling strategies can be readily applied to WFC3/UVIS images.

The primary IR detector features to consider in designing a pointing pattern are the "death star", dust motes, and the "wagon wheel" features (seen in Figure 1, and described elsewhere in these Proceedings). The pointing strategy will determine whether these features can be easily rejected during drizzling. Persistence from previous IR exposures can also be present, and along with other dark-imprint features, it is not rejected by MultiDrizzle (which is optimized to reject bright artifacts like cosmic rays and hot pixels). However, offline steps can be taken to mask the dark-imprint features, which can then be included with the masks created by MultiDrizzle. Well-sampled data can also be drizzled to smaller pixel scales, and made into large mosaics by applying shifts measured offline, as illustrated below.

3. MultiDrizzle: pipeline versus offline

At present, the automatic drizzling of WFC3 images in the pipeline does not make use of the latest distortion corrections, which continue to be refined. The user must download the latest distortion tables (IDCTAB) from the WFC3 website, and specify them in their input image headers to apply the best correction. The MultiDrizzle table (MDRIZTAB) sets parameters for pipeline drizzling, but the user can often produce more optimal results offline by setting parameters tailored to the characteristics of a specific data set, as described here.

Data quality flags are generated by steps within the WFC3 calibration pipeline (calwf3), and are stored in the DQ arrays of the flt images used as input for MultiDrizzle. All of the flagged pixels would be excluded during drizzling, but by setting bits (summing any flag types to be included), the user can control which pixels are excluded. For example, the user may set bits=4864 to include pipeline-rejected pixels (flag=4096), saturated pixels (flag=256), and dust motes (flag=512) for an IR data set that did not employ a dither pattern large enough to span them.

Pipeline-drizzled drz images should be considered quick-look products at best, but basic WFC3 pipeline calibrations (bias, dark, flat) are maturing, and "on-the-fly" (OTFR) pipeline-calibrated archival flt images from the HST archive (MAST) form good input for running MultiDrizzle offline. Subsampled data sets can be drizzled offline to smaller pixel scales (using a smaller drop size, or pixfrac), to fully extract all the spatial information contained in them. Mosaic data sets must be drizzled offline to align and combine images from many different visits (various pointings or different epochs).

4. Getting started: input images and distortion tables

The sample commands below are shown at a PyRAF prompt (so IRAF commands can also be used), but MultiDrizzle can also be run with python syntax (see the MultiDrizzle Handbook section 5.4). De-archive a set of calibrated flt images of uniform data quality: typically using the same instrument modes and filter, with exposure times within 20% of each other. Make an input list in your working directory:

```bash
--> ls i*flt.fits > list_flt_f128n
```
Download the latest distortion reference table, currently available from the WFC3 website, and specify its path name in your image headers (IDCTAB keyword). You may wish to make your own local \texttt{iref} directory and collect reference files there:

\begin{verbatim}
http://www.stsci.edu/hst/wfc3/lbn_archive/
---> mkdir iref
---> set iref = "/grp/hst/cdbs/iref/"  (STScI iref, or make your own)
---> hedit i*flt.ft[0] IDCTAB "iref$t20100519_ir_idc.fits"
\end{verbatim}

\texttt{MultiDrizzle} and the core \texttt{drizzle} task (Fruchter & Hook, 2002) are available in the STSDAS \texttt{dither} package. Load \texttt{multidrizzle} and set parameters as suggested below. Parameters can be saved in a "par" file to record and reuse them later, or a command script can be used to quickly make parameter changes and iteratively reprocess images. The initial run produces the \texttt{single\_sci} images needed to measure shifts.

### 5. Image registration and shift file

\texttt{MultiDrizzle} uses the \textbf{World Coordinate System} (WCS) information in each image header to align the images. But any data set including images taken in different visits or epochs will have small misalignments between visits, due to errors in the cataloged positions of the different guide stars used. Objects in the overlaps between different pointings can be used to measure "delta" shifts and rotations, or update the World Coordinate System in each image, to register the images for drizzle combination. Initial shifts can be quickly measured by visually selecting a few ideal objects (e.g. stars which are unsaturated and uncontaminated by cosmic rays) in the undistorted \texttt{single\_sci} images produced by the \texttt{driz\_separate} step. The results must be put into a \texttt{shiftfile} with the following format:

\begin{verbatim}
# frame: output
# refimage: ibdz21mcq_wcs.fits
# form: delta
# units: pixels
ibdz21mcq_flt.fits  0.00  0.00  0.0  1.0
ibdz21mhq_flt.fits  0.02  0.01  0.0  1.0
ibdz21mq_flt.fits  -0.11 -0.04  0.0  1.0
ibdz21msq_flt.fits  -0.06  0.00  0.0  1.0
ibdz22h9q_flt.fits  -1.76  4.49  0.0  1.0
ibdz22hq_flt.fits   -1.76  4.51  0.0  1.0
ibdz22hiq_flt.fits  -1.78  4.49  0.0  1.0
ibdz22hnq_flt.fits  -1.79  4.49  0.0  1.0
ibdz22hq_flt.fits   -1.79  4.49  0.0  1.0
ibdz23hqq_flt.fits  -1.77  3.11  0.0  1.0
ibdz23hqq_flt.fits  -1.75  3.09  0.0  1.0
ibdz23q_flt.fits    -1.98  2.92  0.0  1.0
ibdz23i6q_flt.fits  -1.94  2.99  0.0  1.0
ibdz25wdq_flt.fits  1.72  12.22  0.0  1.0
ibdz26q_flt.fits    1.71  12.20  0.0  1.0
ibdz25wmq_flt.fits  1.70  12.16  0.0  1.0
ibdz25wrq_flt.fits  1.77  12.22  0.0  1.0
\end{verbatim}

After applying the initial x,y shifts (in the first two columns above), tweakshifts or geomap can be run on the \texttt{single\_sci} images to further refine the shifts, and also solve for small rotational or scale offsets (the 3rd and 4th columns in the shift file).
6. Key drizzling parameters

The following are a key subset of MultiDrizzle parameters optimized for an IR mosaic dataset, with output at both the native detector pixel scale, and an enhanced scale:

```
-> epar multidrizzle
multidrizzle.mdriztab = no # ignore pipeline parameters
multidrizzle.clean = no # keep intermediate files for inspection
multidrizzle.ra = '161.02219' # convenient to fix the central coordinates
multidrizzle.dec = '-59.49465'
multidrizzle.build = no # separate SCI and WHT output files
multidrizzle.shiftfile = 'shifts_f128n.txt' # apply your shifts
multidrizzle.static = no # sometimes flags real features
multidrizzle.skysub = no # handle sky subtraction offline?
multidrizzle.driz_separate = yes
multidrizzle.driz_sep_outnx = 2806 # 6 arcminutes on each side
multidrizzle.driz_sep_outny = 2806
multidrizzle.driz_sep_kernel = 'square' # better than turbo
multidrizzle.driz_sep_scale = 0.1283 # specify explicitly
multidrizzle.driz_sep_pixfrac = 1.0
multidrizzle.driz_sep_rot = 0.0 # rotate North up, East left
multidrizzle.driz_sep_fillval = 9999.9 # arbitrary high fill value
multidrizzle.driz_sep_bits = 4352 # 4864 to include dust motes
multidrizzle.median = yes
multidrizzle.combine_type = 'median'
multidrizzle.combine_nlow = 0 # zero helps suppress motes, persistence
multidrizzle.combine_nhigh = 0 # zero helps suppress motes, persistence
multidrizzle.combine_lthresh = -8.8 # exclude fill values
multidrizzle.combine_hthresh = 8888.8 # exclude fill values
multidrizzle.blot = yes
multidrizzle.driz_cr = yes
multidrizzle.driz_cr_snr = '10.0 8.0' # set very high for IR, no CRs
multidrizzle.driz_cr_scale = '1.2 0.7'
multidrizzle.driz_combine = yes
multidrizzle.final_wht_type = 'EXP'
multidrizzle.final_outnx = 2806
multidrizzle.final_outny = 2806
multidrizzle.final_kernel = 'square' # try different kernels
multidrizzle.final_wt_scl = 'exptime'
multidrizzle.final_scale = 0.1283
multidrizzle.final_pixfrac = 1.0
multidrizzle.final_rot = 0.0 # rotate North up, East left
multidrizzle.final_bits = 4352 # same as single drizzle
multidrizzle.crbit = 0 # don’t record rejections in flt DQ arays
```
Drizzle to native scale, and make a scaled sum image (Figure 1) for diagnostic purposes:

```plaintext
--> multidrizzle input='@list_flt_f128n' output='hh901_wfc3_f128n'

--> epar imcombine
imcombine.combine = 'sum'
imcombine.reject = 'none'
imcombine.lthreshold = -8.8  # exclude fill values
imcombine.hthreshold = 8888.8  # exclude fill values
--> imcombine *single_sci.fits temp_sumN.fits

--> imcalc temp_sumN.fits hh901_wfc3_f128n_drz_sum.fits "im1/4"
```

Although this observing program did not employ an optimally subsampling dither pattern, much of the nebula resides in the overlap areas with extra sampling, so the following command is an example of drizzling to enhance resolution. The smaller pixel scale is 62% of the detector pixel scale, and happens to be conveniently twice the UVIS scale. Also note the smaller pixfrac “drop size” and alternate kernel:

```plaintext
--> multidrizzle input='@list_flt_f128n' output='hh901x_wfc3_f128n'
final_scale=0.07920 final_pixfrac=0.8 final_kernel='gaussian'
```

7. Inspecting output and iteration

To visually verify the quality of your drizzled output, blink your `drz_sci`, `drz_weight`, and `drz_sum` images (Figure 1) to look for signs of bad rejections (too much or too little), misregistration, or other signs that some parameters may need to be adjusted. The median image produced by MultiDrizzle is also worth inspecting: it is used to identify cosmic rays and other bad pixels to be rejected, so it should look almost as good as the final drizzled image. The rms of the exposure weight map (`drz_weight`) should typically be under 30% of the mean, to ensure photometric consistency throughout the image. Several iterations of trial and error are typically required to arrive at optimal results.

References

Figure 1: F128N sum image of HH 901, which shows the relationships between IR detector features, pointing overlaps, and the target. The lower mosaic pointing is at a noticeably different orientation than the other three, because that observation initially failed, and had to be repeated several weeks later. Blinking this sum image against the clean drizzled output and the corresponding exposure weight map is a good initial diagnostic inspection.
New NICMOS Flat-fields

Deepashri Thatte and Tomas Dahlen

Space Telescope Science Institute, Baltimore, MD 21218

Abstract. The A-to-D conversion, or gain, of the NICMOS detectors depends on a combination of detector temperature and bias voltage. The response of the three NICMOS detectors changed slightly due to changes in the A-to-D conversion over both the era before the installation of the NICMOS Cooling System (NCS) 1997-1998 and after the installation 2002 -2008. This change can been expressed as a change in an effective temperature (or equivalently A-to-D conversion) referred to as biastemp, which is calculated directly from the bias level of the science exposures. Here we discuss the creation of new flat-field reference files used to correct for changes in the structure of flat-fields caused by the change in A-to-D conversion. These reference files (*_tdf.fits) populate the keyword TDFFILE in the headers of pre-NCS data retrieved from the OTFR after January 23, 2009 and post-NCS data retrieved after November 19, 2008 and are used by the FLATCORR step in calnica version 4.4 and later. Each *_tdf.fits file consists of five different extensions each valid at a different biastemp range. During calibration, the pipeline calculates the biastemp of the science image and then uses the flat-field extension with the best matching biastemp. For consistency check we also created epoch dependent flat-fields for post-NCS data and compared them with the biastemp dependent flat-fields.

1. Introduction

The detector quantum efficiency (DQE) of NICMOS detectors varies both on pixel-to-pixel scale and on large scale across the chip. This causes the science data to have large spatial variations even for a case with uniform illumination. Before 2008 static flat-fields were used to correct pixel-to-pixel and large scale DQE variations by multiplying the science data by inverse flat-field image normalized to unity. One flat-field was used for each camera and filter combination.

The response of the three NICMOS detectors changed slightly due to (1) variation in detector temperature in the pre-NCS era, 1997-1998 and (2) changes in the A-to-D conversion due to a varying bias voltage after the installation of NCS in 2002. Both (1) and (2) can be expressed as a change in an effective temperature (or equivalently A-to-D conversion) referred to as biastemp, which is calculated directly from the bias level of the science exposures. This variation was noticed in 2008-2009. Variation in biastemp affects the pixel response (hotter pixels are relatively more affected than cooler) causing biastemp dependent changes in the structure of flat-fields even after normalizing to unity.

Throughout Cycle 7 and 7N (pre-NCS era, operating temperature 62K) and in Cycle 11 and beyond (post-NCS era, operating temperature 77K) the variation was approximately 2K as measured by biastemp. To correct for biastemp dependent changes in the structure of flat fields, we have created new flats-fields for all filters in NIC1, NIC2 and NIC3. These reference files each consist of five different flat-field extensions, each valid at a different biastemp range. Calnica, version 4.4 and later now uses these reference files.
to choose the flat-field extension appropriate for the biastemp of the file that is being reduced. The effect of this on the photometry is at most 1% and depends on the position on the detectors. Users who use data retrieved earlier than 2008 and for whom this effect is significant, should re-calibrate the data with calnica version 4.4 (or later) and the latest reference files or retrieve the data again from the archive using the OTFR.

In this report we discuss the creation of new temperature dependent flat-field reference files for all three NICMOS cameras, for both the pre-NCS and post-NCS eras. The term temperature refers to the bias-derived temperature (biastemp) calculated by PyRAF task CalTempFromBias. This temperature populates the keyword TFTEMP in primary headers. Note again that this is not a true "physical" temperature, but an "effective temperature" used to quantify changes in the detector A-to-D conversion. For more details about the CalTempFromBias code and biastemp measurements please refer to Dahlen et al. 2009 NICMOS ISR-2009-002, Pirzkal et al. 2009 NICMOS ISR-2009-007 and NICMOS Data Handbook version 8.0.

2. Data

Table 1 shows the data that was used to create the new flat-fields.

Table 1: Program information table for pre and post NCS data used to make temperature dependent flat-fields.

<table>
<thead>
<tr>
<th>Program ID</th>
<th>Program Title</th>
<th>Program PI</th>
<th>Cameras</th>
<th>Date Range</th>
</tr>
</thead>
<tbody>
<tr>
<td>7689</td>
<td>NIC2 Narrow Band Filter Pointed Super Flats</td>
<td>A. Schultz</td>
<td>1, 2</td>
<td>02/26/1998 - 02/27/1998</td>
</tr>
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<td>7690</td>
<td>NICMOS Pointed Super Flats</td>
<td>A. Schultz</td>
<td>1, 2</td>
<td>07/08/1997 - 11/15/1998</td>
</tr>
<tr>
<td>7957</td>
<td>Lamp Flats II: NICMOS Pointed Flats</td>
<td>A. Schultz</td>
<td>1, 2, 3</td>
<td>08/03/1998 - 08/05/1998</td>
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<td>8083</td>
<td>FLATS: Warming Up - continuation</td>
<td>D. Calzetti</td>
<td>1, 2, 3</td>
<td>12/14/1998 - 01/10/1999</td>
</tr>
<tr>
<td>8974</td>
<td>NICMOS Flats and temperature dependence of the DQE</td>
<td>T. Boeker</td>
<td>1, 2, 3</td>
<td>4/20/2002 - 5/7/2002</td>
</tr>
<tr>
<td>8985</td>
<td>NICMOS Internal Flats</td>
<td>A. Schultz</td>
<td>1, 2, 3</td>
<td>5/13/2002 - 5/19/2002</td>
</tr>
<tr>
<td>9326</td>
<td>NICMOS Cycle 10 Early Calibration Monitor</td>
<td>A. Schultz</td>
<td>1, 2, 3</td>
<td>5/30/2002 - 9/18/2002</td>
</tr>
<tr>
<td>9327</td>
<td>NICMOS Flats: narrow filters for NIC1 + NIC2, NIC3 in parallel</td>
<td>S. Arribas</td>
<td>1, 2, 3</td>
<td>7/26/2002 - 7/28/2002</td>
</tr>
<tr>
<td>9557</td>
<td>NICMOS flats: Camera 3 narrow filters and grisms</td>
<td>S. Arribas</td>
<td>1, 2, 3</td>
<td>6/2/2002</td>
</tr>
<tr>
<td>9640</td>
<td>Flats Stability</td>
<td>A. Schultz</td>
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<td>9/10/2002 - 9/9/2003</td>
</tr>
<tr>
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<td>A. Schultz</td>
<td>1, 2, 3</td>
<td>10/15/03 - 9/4/2004</td>
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<td>10379</td>
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<td>A. Schultz</td>
<td>1, 2, 3</td>
<td>12/30/2004 - 8/2/2005</td>
</tr>
<tr>
<td>10728</td>
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<td>A. Schultz</td>
<td>1, 2, 3</td>
<td>10/28/2005 - 8/6/2006</td>
</tr>
<tr>
<td>11016</td>
<td>NICMOS Flats: narrow and broad filters for NIC1 + NIC2, NIC3 in parallel</td>
<td>N. Pirzkal</td>
<td>1, 2, 3</td>
<td>10/12/2006 - 09/22/2007</td>
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<td>11059</td>
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<td>11352</td>
<td>NICMOS Cycle 16 Time</td>
<td>T. Dahlen</td>
<td>1, 2, 3</td>
<td>06/14/2008 - 7/25/2008</td>
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</table>

- Pre-NCS: The data obtained during the NICMOS flat-field calibration programs during 1997-1998 were used to generate static pre-NCS flat-fields. Data from program 8083 were used to calculate the pixel-to-pixel response as a function of temperature.
• Post-NCS:-The data obtained during the NICMOS flat-field monitoring from 2002 to 2008 were used to regenerate biastemp dependent flat-fields. These programs represent the full span of post-NCS flat monitor programs.

3. Creating Flat-field images

3.1. Pre-NCS flat-fields

These flat-fields are created as follows.

1. First we determine the temperature of the existing static pre-NCS flat-field images by calculating the mean temperature of the individual exposures that were used to produce the flat-fields. We use the bias-derived temperature given by the TEBTEMP keyword in the image header of each file.

![Graph](image)

Figure 1: Response in electrons/s as a function of temperature, normalized at T=63K. Curves are shown for wave-lengths 1.1, 1.6 and 2.2 micron. Shorter wave-lengths have steeper slopes.

2. We thereafter calculate the pixel-to-pixel response as a function of temperature and wave-length using flat-field images in F110W and F160W (and also including F222M for NIC3) taken during the end-of-life warm-up monitoring in January 1999. Adding to these a few images taken at lower temperatures earlier in the pre-NCS era, the available flat-fields span over the temperature range 55-70K. The first fit of the pixel-to-pixel response in electrons/s vs. temperature for each pixel is made at the particular wave-lengths of the available filters using an IDL script makefits written by Eddie Bergeron. Another script makewavefits is thereafter used to make wave-length fits to the output from above. The end products from these scripts are arrays with coefficients relating the response in electrons/s of each camera with temperature and wave-length. New flat-fields at desired temperature and wavelength are created using these relations and existing static flat-fields. An example of the response on
electrons/s as a function of temperature for NIC3 at wave-lengths 1.1, 1.6, and 2.2 microns is shown in Figure 1. The symbols are measured points and the lines are fits to these points. Shorter wave-lengths have steeper temperature gradient. Curves are normalized to unity at 63 K.

For each filter and camera, flats are calculated at the following five different output temperatures: 61.2 K, 61.6 K, 62.0 K, 62.4 K, and 62.8 K. This approximately divides the temperature range used in the pre-NCS era in five equally spaced bins. For consistency with previously created flat-fields, each flat-field is normalized to unity in $[*,36:256]$ for NIC1 and NIC2 and in $[*,56:256]$ for NIC3. The flat-fields are thereafter inverted. Each flat-field, having an extension *.df.fits, consists of a stack of five flats for the above mentioned temperatures. When applying flat-field correction, the biastemp, TFBTEMP, is used to choose the flat closest in temperature, which is thereafter used for the flat-field correction.

3.2. Post-NCS flat-fields

![Figure 2: Temperature from bias for NIC1, NIC2 and NIC3 data](attachment:flat-fields.png)
In Figure 2 we show the temperature variation of individual flat-field images taken during the NICMOS flat-field monitoring programs 2002-2008. From the figure it is clear that there is a strong evolution in temperature-from-bias with time and we therefore should expect that the shape of the post-NCS flat-fields will also change over this period. We therefore create new flat-fields also for the post-NCS era, where each flat-field consists of a data-cube with five separate flat-fields extensions, each valid for a specific biastemp range.

To do this we first create single exposure flat-field images by subtracting pairs of calibration images taken first with the NICMOS internal flat-field lamp on then off. This is done for all filters and all cameras using the available monitoring data taken 2002-2008. In total there are about 800-900 single exposure images per camera divided among the different filters. We thereafter use the temperature (TFBTEMP) of these images to divide the flat-fields into five temperature bins so that approximately the same number of exposures falls in each bin for each camera. The limits for the temperature bins for each camera are given in Section 5.

We finally combine the single exposure images in each temperature bin to create a median flat-field image. Each image is thereafter normalized to unity (using [*,36:256] for NIC1 and NIC2 and [*,56:256] for NIC3) and inverted.

For filters monitored regularly, like F110W, F160W and F222M, there are enough single exposure flat-fields to create final images in all five temperature bins. For some less observed filters, mostly narrow band filters, images are not available in all five temperature bins. In those cases final images are generated by interpolating. We also generate a pedigree file that includes information on all individual images going into each final flat-field together with S/N for all the combined images. This pedigree file is available on the NICMOS web page, www.stsci.edu/hst/nicmos.

4. Epoch Flat-fields

Figure 2 shows that the biastemp changes close to linearly with time and an alternative approach to create flat-fields would therefore be to divide the flat-fields in temporal bins as discussed in Dahlen et al. 2007 NICMOS ISR-2007-002. For consistency check epoch dependent flat-fields were also created using a method similar to that of biastemp dependent flat-fields. Instead of temperatures the flat-fields were divided into bins for different epochs.

4.1. Comparison of Epoch and Temperature flat-fields

To quantify the difference between the epoch and temperature flat-fields, the signal-to-noise ratio for flat-fields in all temperature bins was compared with signal-to-noise ratio for flat-fields in all epoch bins (not including bins for which interpolated flat-fields were created). Table 2 shows the median S/N for biastemp dependent and epoch dependent flat-fields for all three NICMOS cameras.

The median of the S/N for biastemp dependent flat-fields is higher than for the epoch dependent flat-fields. Note also that the total number of different temperature images created is larger than the number of epoch images. This means that on average fewer single flat-field images were used to produce the temperature dependent flat-fields. Despite that, the S/N for the temperature flat-fields are higher, assuring us that the temperature dependent flat-fields are a better choice to use.
Table 2: Comparison of post-NCS epoch and temperature flat-fields.

<table>
<thead>
<tr>
<th>Flat-field Type</th>
<th>Number of Images</th>
<th>Median S/N</th>
</tr>
</thead>
<tbody>
<tr>
<td>NIC1 temperature flat-fields</td>
<td>64</td>
<td>733.510</td>
</tr>
<tr>
<td>NIC1 epoch flat-fields</td>
<td>52</td>
<td>722.210</td>
</tr>
<tr>
<td>NIC2 temperature flat-fields</td>
<td>67</td>
<td>728.600</td>
</tr>
<tr>
<td>NIC2 epoch flat-fields</td>
<td>57</td>
<td>713.710</td>
</tr>
<tr>
<td>NIC3 temperature flat-fields</td>
<td>63</td>
<td>823.690</td>
</tr>
<tr>
<td>NIC3 epoch flat-fields</td>
<td>55</td>
<td>772.180</td>
</tr>
</tbody>
</table>

5. Structure of new flat-field reference files

Each flat-field image (*_tdf.fits) consists of five individual images (imsets) each having five extensions: SCI, ERR, DQ, SAMP, and TIME. The *_tdf.fits images therefore consist of 25 extensions with the SCI images, i.e., the flat-fields, in extensions 1, 6, 11, 16, and 21. Keywords TFBLOW and TFBHIGH show the temperature range for each extension. The full temperature range covered by all five bins is given by the TFBLOW and TFBHIGH keyword in the primary header (EXT=0). The extension for which TFBLOW < TFBTEMP < TFBHIGH, where TFBTEMP is the temperature from science data header, is used for the flat-fielding.

The temperature ranges (in Kelvin) are:

Pre-NCS NIC1, NIC2, NIC3 flat-fields
extension 0: TFBLOW=52.0, TFBHIGH=70.0
extension 1: TFBLOW=52.0, TFBHIGH=61.4
extension 6: TFBLOW=61.4, TFBHIGH=61.8
extension 11: TFBLOW=61.8, TFBHIGH=62.2
extension 16: TFBLOW=62.2, TFBHIGH=62.6
extension 21: TFBLOW=62.6, TFBHIGH=70.0

Post-NCS flat-fields
NIC1
extension 0: TFBLOW=64.95, TFBHIGH=86.85
extension 1: TFBLOW=64.95, TFBHIGH=74.95
extension 6: TFBLOW=74.95, TFBHIGH=75.55
extension 11: TFBLOW=75.55, TFBHIGH=76.25
extension 16: TFBLOW=76.25, TFBHIGH=76.85
extension 21: TFBLOW=76.85, TFBHIGH=86.85

NIC2
extension 0: TFBLOW=64.65, TFBHIGH=86.80
extension 1: TFBLOW=64.65, TFBHIGH=74.65
extension 6: TFBLOW=74.65, TFBHIGH=75.30
extension 11: TFBLOW=75.30, TFBHIGH=76.15
extension 16: TFBLOW=76.15, TFBHIGH=76.80
extension 21: TFBLOW=76.80, TFBHIGH=86.80

NIC3
extension 0: TFBLOW=65.60, TFBHIGH=87.05
extension 1: TBFLOW=65.60, TFBHIGH=75.60
extension 6: TBFLOW=75.60, TFBHIGH=76.08
extension 11: TBFLOW=76.08, TFBHIGH=76.56
extension 16: TBFLOW=76.56, TFBHIGH=77.05
extension 21: TBFLOW=77.05, TFBHIGH=87.05

Figure 3: Panels 1 to 5 show the ratio images obtained by dividing a flat corresponding to each temperature bin in the biastemp dependent flat by the static flat for pre-NCS NIC3, F110W data.

Figure 4: Panels 1 to 5 show the ratio images obtained by dividing a flat corresponding to each temperature bin in the temperature dependent flat by the static flat for post-NCS NIC3, F110W data.

Figures 3 and 4 show the ratio images obtained by dividing each SCI extension in NIC3, F110W temperature dependent flat by the SCI extension in the static flat file, for pre and post-NCS data respectively. The temperature goes up as we go from panel 1 (top left) to panel 5 (bottom right). Panel 3 in figure 3 and panel 4 in figure 4 are relatively flat due to the fact that the temperature range of SCI[11] image is close to the temperature of the static flat file. From both the figures we can conclude that the biastemp dependent
flat-fields do have some structure as compared to the static flat-field and the structure of flat field does depend upon temperature. The evolution of the pre-NCS flats is due to a true change in detector temperature, while the change in post-NCS flats is due to a drift in bias voltage.

6. FLATCORR

The new flat-fields populate the header keyword TDFFILE. This reference file contains the flat-field image for a given detector and filter (or polarizer) combination. As explained above this file contains five imsets, each with a flat-field image valid for a particular temperature range. In the FLATCORR step, calnica reads the bias-derived temperature from the TFBTEMP keyword in the science data header and selects the appropriate imset from the TDFFILE. The imset used is written to the TDFGROUP keyword in the header of the calibrated (*_cal.fits) files. See figure 5

Figure 5: Use of new Flat-fields by calnica.

The new flat fields were delivered to the archive on November 19, 2008 (updated NIC3 flat-fields delivered on May 27, 2009) for Cycle 11 and later observations and on January 23, 2009 for data taken during Cycles 7 and 7N. There is no TDFFILE listed in the header for the data retrieved prior to these dates. In this case, the FLATFILE reference file which is a single static (non-temperature dependent) flat-field image is used instead. If both TDFFILE and FLATFILE reference files are given in the header, calnica chooses the TDFFILE. FLATFILE is used instead of TDFFILE in the cases where TFBTEMP keyword is missing or TFBTEMP lies outside the range given by TFBHIGH and TFBLOW keywords in the primary header. Error estimates and DQ flags contained in the TDFFILE/FLATFILE are propagated into the processed images. The difference between the new flat-fields and the static flat-fields is small but it can affect the photometry on a 1% level.
7. Summary

Data from NICMOS calibration programs during 1997-1998 and 2002-2008 are used to generate new pre-NCS and post-NCS flat-fields. Variations in the A-to-D conversion in the NICMOS detectors, as quantified by the bias-derived effective temperature (given by header keyword TFBTEMP), affects the pixel-to-pixel response and the structure of the flat fields for all detectors. The new flat-fields (*_tdf.fits) correct for this effect in the FLATCORR step of calnica version 4.4 and later. The *_tdf.fits files populate the keyword TDFFILE in the headers of data retrieved after November 19, 2008 for post-NCS flat-fields and January 23, 2009 for pre-NCS flat-fields.

Acknowledgments. We would like to thank L. Bergeron for providing the scripts to create pre-NCS temperature dependent flat-fields.

References

Dahlen, T., Barker, E., Bergeron, E., & Smith, D. NICMOS ISR-2009-002, STScI, Temperature Dependent Dark Reference Files: Linear Dark and Amplifier Glow Components
New Bad Pixel Mask Reference Files for the Post-NCS Era

Elizabeth A. Barker and Tomas Dahlen

*Space Telescope Science Institute, Baltimore, MD 21218*

**Abstract.** The last determined bad pixel masks for the three NICMOS cameras were made in September 2002. Those masks were created using data from calibration programs following the installation of the NCS and are therefore based on the relatively limited data set available at the time. Since then, the NICMOS calibration monitoring programs have regularly obtained calibration images of both flat-fields and darks, images used to create the mask reference files. With numerous images taken during a long base-line (2002-2008), this data set allows us to create high signal-to-noise reference files, as well as investigate any temporal dependence of the mask files. Here we describe the creation of new mask files based on this extended data set and compare the new masks with the previous versions. The new masks created contain a higher number of bad pixels compared to the old versions, while the number of pixels thought to be affected by “grot” is lower.

1. Introduction

The NICMOS detectors are regularly calibrated through the use of flat-field and dark reference images. Studying these references images provides a tool to understand the behavior of individual pixels, as well as any evolution with time. The NICMOS data quality (DQ) extension of calibrated images (*ima.fits or *cal.fits) contains specific information regarding problematic pixels, which should be considered for exclusion when science images are combined or dithered. This DQ extension identifies several different types of problems that may affect pixels, including both static flag values, which do not depend on the observation, and dynamic flag values, which are set based on the observation or the subsequent calibration. The static flags for “bad” pixels and “grot” pixels are identified with separate DQ values (DQ=32 and DQ=16, respectively). Each NICMOS camera has two unique static masks, one for use with pre-NCS data and one for use with post-NCS data. The relevant mask for each observation is identified in the header keyword MASKFILE of the science images and has the *msk.fits extension.

A proper removal of bad pixels from NICMOS images is important for securing the data quality of reduced images. Initial bad pixel masks were created during Systems Level Thermal Vacuum ground testing prior to the installation of NIMCOS on HST. Newer bad pixel masks for pre-NCS data were created in 2002, based on in-flight data taken in 1997-8 (Sosey 2002, NICMOS ISR 2002-001), however these masks were not delivered to the CDBS database until 2009.

In 2002, the NICMOS Cooling System (NCS) was installed on HST and connected to NICMOS during Service Mission 3b. The NCS consists of a cryocooler and re-enabled NICMOS for science observations, but at a higher nominal operating temperature. After the installation of the NCS, masks for post-NCS data were created in September 2002 using in-flight data.

The dark reference frames measure the accumulation of signal due to dark current, per pixel, during observations. Studying the variation of dark current pixel-by-pixel is necessary to understand the inherent variations across the detectors. Since the dark images measure
New Bad Pixel Mask Reference Files for the Post-NCS Era

the signal accumulated without any external illumination, they are an excellent device to
detect defective pixels with deviant behavior. Hot pixels have excessive charge and cold
have little or no charge, even with significant exposure times. To study the bad pixels of the
detectors, we have analyzed dark reference files taken during the post-NCS era of 2002-2008.

An additional type of bad pixel exists on the NICMOS detectors, often referred to as
grot-affected pixels. As a result of the deformation of the NICMOS instrument, we believe
paint flecks (ranging in size from a fraction of a pixel to a few pixels) were deposited on
the detectors of all three cameras. They are apparent in all images where an external
source of illumination is used. Grot-affected pixels appear as pixels with lower signal
response compared to neighboring pixels. Since flat-field images have a high and uniform
illumination, they are well suited for detecting grot pixels. Thus, we have analyzed the flat-
field calibration images to study the behavior of grot pixels since SM3b and have determined
a new set of grot pixels appropriate for post-NCS observations. Since these data have been
taken during a large part of the post-NCS era, investigating changes of the bad pixels with
time is now also possible.

2. Data and Reduction

In order to determine hot and cold pixels, we use data taken during the dark monitoring
programs 2002-2008 (programs 9321, 9636, 9993, 10380, 10723, 11057, 11318) together with
data taken during the extended darks program (11330), which started early 2008. The
number of exposures for each different SAMP SEQUENCE per NICMOS camera is shown
in Table 1. We use SPARS64 with NSAMP of 20, 24 and 26, giving exposure times of over
1000s. All used the BLANK filter position, as is typical for dark observations.

<table>
<thead>
<tr>
<th>SAMP_SEQ</th>
<th>NSAMP</th>
<th>Exp Time</th>
<th>NIC1</th>
<th>NIC2</th>
<th>NIC3</th>
</tr>
</thead>
<tbody>
<tr>
<td>SPARS64</td>
<td>20</td>
<td>1088s</td>
<td>290</td>
<td>290</td>
<td>290</td>
</tr>
<tr>
<td>SPARS64</td>
<td>24</td>
<td>1344s</td>
<td>477</td>
<td>478</td>
<td>478</td>
</tr>
<tr>
<td>SPARS64</td>
<td>26</td>
<td>1472s</td>
<td>30</td>
<td>30</td>
<td>30</td>
</tr>
<tr>
<td>Total Images</td>
<td></td>
<td></td>
<td>797</td>
<td>798</td>
<td>798</td>
</tr>
</tbody>
</table>

Table 1: Number of dark images for the read out modes for the three NICMOS
detectors.

The grot-affected pixels were analyzed by examining flat-field data from 2002 to 2008
(programs 8974, 8985, 9326, 9327, 9640, 9996, 10379, 10728, 11016, 11059, 11321) with the
F160W filter, which is well-sampled including more than 100 individual images per camera.

3. Bad (DQ=32) Pixels

In order to identify both hot and cold (collectively “bad”) pixels, we examine long exposure
dark images with exposure times greater than 1000s (see Table 1). Long exposures make the
identification of both types of bad pixels easier to identify, when compared to neighboring
pixels and expected values for dark current. “Hot” pixels are defined to be those that
exhibit excessive charge when compared to surrounding pixels. “Cold” pixels are those that
have extremely low or near zero response or dark current (also known as “dead” pixels).

For consistency and continuity, we utilize the same method of identifying bad pixels as
that given in NICMOS ISR 2002-001 (Sosey 2002). Specifically:

1. Each dark reference image is CR cleaned.
2. A composite median dark image is created from the CR-cleaned darks.
3. A smoothed composite image is created from the composite median dark image.

4. The smoothed composite is subtracted from the composite median dark image.

5. The subtracted composite image is rescaled to units of its RMS.

6. Bad pixels are identified in the subtracted composite image as those pixels outside of $5\sigma$.

### 3.1. Time Dependence of Bad Pixels

Since we are not only interested in knowing if any new pixels have changed into bad pixels, we also look for changes in time. This is now possible, since we have nearly-continuous dark calibration images between 2002 and 2008 (post-NCS era). Table 2 shows the number of bad pixels, per camera, found in the old post-NCS masks from 2002, as well as the number of additional bad pixels found each year compared to the old mask.

<table>
<thead>
<tr>
<th></th>
<th>NIC 1</th>
<th>NIC 2</th>
<th>NIC 3</th>
</tr>
</thead>
<tbody>
<tr>
<td>Old DQ=32</td>
<td>193 (0.29%)</td>
<td>656 (1.0%)</td>
<td>446 (0.68%)</td>
</tr>
<tr>
<td>2002</td>
<td>46</td>
<td>17</td>
<td>16</td>
</tr>
<tr>
<td>2003</td>
<td>49</td>
<td>21</td>
<td>21</td>
</tr>
<tr>
<td>2004</td>
<td>70</td>
<td>28</td>
<td>28</td>
</tr>
<tr>
<td>2005</td>
<td>69</td>
<td>24</td>
<td>30</td>
</tr>
<tr>
<td>2006</td>
<td>70</td>
<td>24</td>
<td>29</td>
</tr>
<tr>
<td>2007</td>
<td>66</td>
<td>29</td>
<td>34</td>
</tr>
<tr>
<td>2008</td>
<td>70</td>
<td>31</td>
<td>36</td>
</tr>
<tr>
<td>All new</td>
<td>88 (+0.13%)</td>
<td>40 (+0.06%)</td>
<td>42 (+0.06%)</td>
</tr>
</tbody>
</table>

**Table 2:** Number of additional bad pixels per year per camera. “Old DQ=32” represents the number pixels in the 2002 post-NCS bad pixel masks. “All new” represents the number of completely new bad pixels over the whole time range 2002-2008.

The number of bad pixels in the old post-NCS bad pixel masks from 2002 are listed for each camera. For each year in the post-NCS era, we have tabulated the number of additional bad pixels found in this investigation. On the last line, we list the total number of different pixels that are flagged as bad during at least one year. This can also be seen in Figure 1.

Note that all “new” bad pixels we find are hot and there are no new “dead” pixels found compared to the 2002 post-NCS bad pixel masks. There could still be new cold pixels that are not selected by the above criteria, however such pixels will be selected as grot pixels in the next step if the decrease in response of the pixel is significant.

Figure 1 shows the number of bad pixels in the previous bad pixel masks, for each camera, as a black dot in 2002. Shown in red dots are the total number of bad pixels found in this investigation. All of these bad pixels, both hot and cold, are flagged with DQ=32 in the *msk.fits* files, as well as in the calibrated *ima.fits* and *cal.fits* files.

As a primary approach, we create new bad pixel masks by adding all new pixels (bottom row of Table 2) to the existing masks. However, since we know that the number of bad pixels is changing with time, it is worthwhile to study how many pixels can be “saved” as scientifically useful pixels by having multiple bad pixel masks, one per year, for instance. If the number is significant, such an approach may be justifiable.

To examine the number of pixels that can be returned to scientific use during at least one year by creating a bad pixel mask per year, we plot in Figure 2 the RMS of all pixels...
Figure 1: Number of new bad pixels by year and camera. The lower black dot for 2002 represents the old mask. The remaining dots show the total number of bad pixels per year, including all old pixels and additional new pixels for each year.

Figure 2: The dark current of the darks (in RMS units) versus year, by camera, for all additional bad pixels found in this investigation. The dotted line represents the $5\sigma$ cutoff.
that are selected as hot (RMS > 5σ) during at least one of the seven years. Good pixels that suddenly turn bad are characterized by initially having a low RMS (<1σ), where after they turn bad and show RMS > 5σ. Similarly, there are pixels that start out bad (RMS > 5σ), but with time drop in RMS. If the number of these pixels had been high, then the multiple mask approach would be justifiable. However, studying Figure 2 shows that only a handful of pixels have low RMS (~1σ) during part of the whole period. This small number of pixels does not justify having multiple masks. Instead, we create new bad pixel masks including all pixels that have RMS > 5σ during at least one epoch.

Figure 3 shows the old bad pixel mask (top row) and the new bad pixel mask (bottom row) for each camera.

![Figure 3: The old (top row) and new (bottom row) bad pixel masks (DQ=32 only) for each camera.](image)

4. Grot (DQ=16) Pixels

Grot is believed to be flecks of anti-reflective paint deposited on the detectors when the expansion of the solid nitrogen caused the optical baffles to scrape against each other (see NICMOS ISR 99-008 for a complete discussion). Grot produces small areas of reduced sensitivity, ranging in size from 25µm to 100µm. Since NICMOS pixels are 40µm on each side, grot can affect less than one pixel up to several pixels by a single fleck. The largest
example of a single grot region is called the “battleship” feature in NIC1 (see Figure 4), which affects approximately 35 pixels.

![Flat-field Image and Bad Pixel Mask](image)

Figure 4: The largest grot feature, known as the “battleship”, can be seen in NIC1.

<table>
<thead>
<tr>
<th>Camera</th>
<th>Existing</th>
<th>New</th>
</tr>
</thead>
<tbody>
<tr>
<td>NIC1</td>
<td>180</td>
<td>170</td>
</tr>
<tr>
<td>NIC2</td>
<td>243</td>
<td>123</td>
</tr>
<tr>
<td>NIC3</td>
<td>249</td>
<td>113</td>
</tr>
</tbody>
</table>

Table 3: The number of old and new grot-affected pixels in each camera.

Since grot consists of a physical substance on the detectors, it affects the incoming light landing on the detectors. Since flat-field exposures are uniformly illuminated and have high counts over the entire detector, we can use them to estimate the response of grot-affected pixels. The last study of grot-affected pixels was done in 2003 (Schultz et al, NICMOS ISR 2003-003) as part of the flat-field monitoring of stability in NICMOS.

In a manner similar to that used in 2003, we utilized the post-NCS flat-field observations:

1. A well-sampled (F160W) flat-field, based on 2002-2008 data, is re-inverted\(^1\).
2. The re-inverted flat is smoothed.
3. The smoothed flat is subtracted from the re-inverted flat-field.
4. Grot pixels are defined as pixels deviating more than $4\sigma$ in the subtracted image, excluding bad (DQ=32) pixels.

A smoothed image was created by replacing each pixel value with the mean value within a radius of 10 pixels after excluding flagged pixels and pixels deviating by more than $3\sigma$ from the mean. We find that the new data results in fewer grot-affected pixels compared to the existing grot mask made in 2002, as can be seen in Table 3.

---

\(^1\) This is because NICMOS flat-fields are, by construction, inverted images, where a high pixel value means a low response.
Due to the large difference in the number of old and new grot pixels, it is useful to examine the relative Detector Quantum Efficiency (DQE) response of the pixels that are not being flagged in the new analysis, but were previously flagged. The relative DQE response can be determined by dividing the flat-field image by the smoothed flat-field image. In the resultant ratio image, a “normal” pixel (not grot-affected) has a response equal to unity.

Figure 5: The number of pixels by relative DQE response of pixels, per camera. The left panel (blue) shows the old grot mask. The center panel (red) shows the new selected grot pixels. The right panel (green) shows the “extra” grot pixels found in the old mask, but not in the new mask.

The left panel of Figure 5 shows the relative response of the pixels in the old grot mask, while the middle panel shows the new selection. The analysis of the new data selects grot-affected pixels as those with a response ≤0.8. The pixels that are flagged as grot-affected in the old masks and not flagged as grot in the new analysis (“extra” grot) are shown in the right panel of the figure. It is clear that these pixels have close to normal responses (0.8–1.2).

In order to further examine the nature of the “extra” grot pixels and see if we can safely recover these pixels for science observations, we compare the RMS of these pixels to the RMS of normal non-grot pixels, by camera. The RMS is calculated pixel by pixel from the stack of individual flat-field observations used to create the combined flat-field used here. If, for example, grot has moved between epochs, then it is expected that the RMS would be higher for pixels affected by the moved grot.

We see in Figure 6 that the RMS of the “extra” grot pixels (large green dots) are similar to those of non-flagged, “normal” pixels (small black dots). This, along with the fact that the “missing grot” pixels have relative DQE responses in the normal range, provides confidence that these pixels have proper sensitivities for science observations. Therefore, we should be able to recover these pixels for science and set their DQ flags back to zero, or “deflag” them.
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Figure 6: RMS per pixel versus relative DQE response. Small black dots are non-flagged, “normal” pixels and large green dots are “extra” grot pixels.

We also want to know if there is a significant number of grot pixels that change with time. In Table 4, we list the number of grot-affected pixels in the old grot masks, the new grot masks, during 2002, and during 2008. We see that all three cameras show fewer grot-affected pixels in the new masks, but the numbers are not significantly different between 2002 and 2008. We therefore do not create time-dependent grot masks, but incorporate all data between 2002 and 2008 into single masks, one for each camera.

<table>
<thead>
<tr>
<th></th>
<th>Old</th>
<th>New</th>
<th>2002</th>
<th>2008</th>
</tr>
</thead>
<tbody>
<tr>
<td>NIC1</td>
<td>180</td>
<td>170</td>
<td>172</td>
<td>174</td>
</tr>
<tr>
<td>NIC2</td>
<td>243</td>
<td>123</td>
<td>119</td>
<td>139</td>
</tr>
<tr>
<td>NIC3</td>
<td>249</td>
<td>113</td>
<td>113</td>
<td>112</td>
</tr>
</tbody>
</table>

Table 4: The number of grot-affected pixels with time.

As a final selection of grot-affected pixels, we include all “$4\sigma$ pixels” from the above discussion as grot pixels (except those that have a response $>0.8$ and a normal RMS, which are instead addressed by the flat-fields). This selection somewhat reduces the number of grot pixels compared to that given in Tables 3 and 4. The final number of grot pixels for the three cameras are given in Table 5. These pixels will be marked with DQ=16 in the same mask files as the bad pixels (DQ=32). The difference between the old and new grot masks can be seen clearly in Figure 7 below, with the old masks shown across the top and the new grot masks shown along the bottom. The reduced number of grot pixels is most apparent in NIC3.
5. New Static DQ Masks for NICMOS

The NICMOS bad pixel masks for each camera were generated by placing the DQ flag values in the DQ extension of the MASKFILE reference files. Table 5 lists the number of pixels for both the old and new masks, with grot and bad pixels distinguished.

<table>
<thead>
<tr>
<th></th>
<th>Grot, DQ=16</th>
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<th>Bad, DQ=32</th>
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<tbody>
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<td></td>
<td>Old</td>
<td>New</td>
<td>Old</td>
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<td>656</td>
</tr>
<tr>
<td>NIC 3</td>
<td>249</td>
<td>96</td>
<td>446</td>
</tr>
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</table>

*Table 5:* The final, new static DQ masks for all three NICMOS cameras.

We also present the new final, complete, static DQ masks for each camera in Figure 8 below. We can visually see the reduction in grot-flagged pixels, notably in NIC2, but especially in NIC3.

These new mask files have been delivered to CDBS and are available to the pipeline for calibration and to users via the NICMOS Reference Files web page at http://www.stsci.edu/hst/observatory/cdbs/SIfileInfo/NICMOS/refstablequeryindex.
New Bad Pixel Mask Reference Files for the Post-NCS Era

6. Conclusions

We have used NICMOS SPARS64 images to create new “hot+cold” pixel masks and flat-field calibration image to determine current grot-affected pixels, for calibrating NICMOS images. We have shown that there is only a weak temporal dependence of the number of flagged pixels and have therefore created only a single MASKFILE for each camera for the post-NCS era. Compared to the previous mask from 2002, these new masks contain a slightly higher number of bad pixels (DQ=32), while the number of grot pixels (DQ=16) has decreased. As always, well-dithered observations is the recommended way to handle bad pixels of all types and for most types of science.

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Sosey, M. 2002, NICMOS ISR 2002-001

Figure 8: Old and new final DQ masks for each camera. Yellow pixels are bad pixels, purple pixels are grot-affected pixels.
The Detection and Removal of Large-scale Detector Background Structures in NICMOS Observations

Lawrence Berkeley National Laboratory

C. Lidman
Australian Astronomical Observatory

K. S. Dawson
University of Utah

L. E. Bergeron, S. Deustua, A. S. Fruchter
Space Telescope Science Institute

K. Barbary, H. K. Fakhouri, J. Meyers, D. Rubin
University of California, Berkeley

Abstract. After applying the standard corrections for well-studied NICMOS anomalies, significant large-scale spatial background variation remains. We report on the detections of a sky-dependent fringe pattern in the F110W filter, and a time-dependent residual corner glow in the calibrated NICMOS deep science images. We also describe methods developed to further correct these anomalies. A model of the background structure is derived from the algebraic manipulation of stacked science images and consists of the following two components. The first component is constant, dominated by a residual glow as high as 20 DN at the corners and by residual flat and persistence structures at the center. The second component, which scales with sky level, displays a clear fringe pattern with 10% variation for F110W images. However this pattern is not detected for F160W images. Using these model components to correct for the anomalies significantly improves the cosmetic appearance of NICMOS images and reduces the magnitude scatter in the photometry of distant galaxies by 20%.

1. Introduction

The standard calibration pipeline CALNICA and additional STSDAS software have been excellent tools for observers, removing many of the well-studied NICMOS anomalies. Despite the aid of these tools, large-scale spatial background structures are still quite visible in the final calibrated deep images. These background structures cause variation as high as 60% in counts from the center to the corners. This makes the precision study of faint sources, which have counts near the sky level, exceptionally difficult. In this paper, we outline novel methods for deriving model images of the observed background structures and for removing them from individual images. We also present a technique for recovering the flux from the erratic middle column.
The data used for this study are deep NICMOS supernova follow up observations from the HST Cluster Supernova Survey (PI: Perlmutter) and the GOODS Transient Search (PI: Riess). The data are taken with NIC2 camera and F110W/F160W passbands between June, 2002 and August, 2008, after the installation of the NICMOS Cooling System. The data set contains 385 exposures in F110W and 188 exposures in F160W. The exposures are deep, typically $\sim 1000$ seconds of exposure time. They are also sparsely populated with distant faint sources, making them ideal for studying spatial background structures.

The science exposures are calibrated with the latest CALNICA pipeline (version 4.1.1; Dahlen et al. 2008) and STSDAS software PEDSKY. Exposures taken within one orbit of the last exit from the NICMOS SAA contour are put through the STSDAS software SAACLEAN, and corrections are made when needed. Four of our exposures suffer from significant bright earth persistence (BEP; Riess & Bergeron 2008), and are corrected using the STSDAS software NICREM_PERSIST.

2. Principal component analysis

Principal component analysis (PCA) is a powerful statistical technique which reduces the dimensionality and identifies patterns in a set of multi-dimensional data. PCA has been employed for analyzing astronomical data in a wide variety of applications (e.g., Suzuki 2006; Blanton & Roweis 2007). Here, we use PCA to characterize the intensity of the corner glow (Section 3.) and to reconstruct the erratic middle column (Section 5.).

The simplifying “bra ket” notation is adopted here for the formulation of PCA. For an $n$-dimensional data set $|f_i\rangle$ and its mean properties $|\mu\rangle$, the principal components (PCs) $|\xi_j\rangle$ are simply the eigenvectors of the covariance matrix for $|f_i - \mu\rangle$. The eigenvalue accounts for the variance of the data in the direction of the associated PC. When the PCs are ranked by their eigenvalues, the first PC points in the direction of maximum variance in the $n$-dimensional data space. The rest of the PCs are orthogonal vectors to the first: $\langle \xi_i | \xi_j \rangle = \delta_{ij}$.

The projection $p_{ij}$ of $|f_i\rangle$ onto a PC $|\xi_j\rangle$ is computed as: $p_{ij} = \langle f_i - \mu | \xi_j \rangle$. The $n$-dimension data set can then be represented as the sum of all of its PCs:

$$|f_i\rangle = |\mu\rangle + \sum_{j=1}^{n} p_{ij} |\xi_j\rangle. \quad (1)$$

If the first few PCs constitute a large fraction of the overall variations, then they alone can provide decent reconstructions of the original data.

Here, we obtain the eigenvectors and eigenvalues of the covariance matrix using internal IDL routines which utilize Householder reductions and the QL method (Press et al. 1992). For the case where there is missing data (Section 5.), we adopt the expectation maximum (EM; Roweis 1998) algorithm to obtain the PCs. The EM algorithm naturally accommodates missing data and allows for the extraction of only the first few leading PCs without having to diagonalize the entire covariance matrix.

3. Time dependence of residual corner glow

The most prominent large-scale background is the corner structure which resembles amplifier glow and linear dark current. Without understanding the origin of this structure, it is here loosely termed “residual corner glow.” PCA (Section 2.) is performed on the calibrated images to gain a better understanding of the shape and variation of this structure.

PCA is performed on calibrated and sky-subtracted images in absolute counts, with the center and sources masked out. The first PC (or eigen-image) has the shape of the corner glow and accounts for over half of the variation (Figure 1). The corner glow is thus not only prominent, but also varies between exposures. The eigen-images of the F110W
and F160W data sets show very similar characteristics, consistent with the orientation at the detector.

![Figure 1](image-url)

**Figure 1:** First PC eigen-image of the residual corner glow. The left and right panels show the eigen-images for the F110W and F160W data sets, respectively.

The projections of individual images onto the first PC eigen-image are used here as a measure of the corner glow intensity (Figure 2). They reveal that the intensity of the corner glow varies with time: starts high, decays exponentially, resets every orbit, but never goes to zero. This behavior may be the result of larger variation in detector temperature that cannot be accounted for by the bias temperature. At the beginning of the orbit, the detector is slightly warmer after spending the occultation in auto-flush. The high corner glow intensity can then be attributed to higher linear dark current, higher amplifier glow, and even persistence from the amplifier glow.

The intensity and the time evolution of the corner glow are quite uniform. For several consecutive exposures taken within an orbit, the behavior of exponential time decay is clear. However, most images in this data set have exposure times much larger than the e-folding time of the exponential decay. The intensity then appears binary in nature as shown in Figure 2. This allows the images to be segregated into two groups of approximately constant corner glow. With this simplification, we can devise a simple model of the background structure that can be extracted with algebraic manipulation of stacked images.

4. Deriving model images

The background structures are modeled as two components: 1) component $G$ which is constant in absolute counts, and 2) component $F$ which scales with the sky level. Component $G$ is designed to capture the residual corner glow (Section 3.). Segregating images according to the residual corner glow justifies the use of a constant component. The models are derived for each filter (F110W and F160W) and for each glow group (high and low). The sky-subtracted calibrated science image $I$ in DN/s is then modeled as the sum of the contributions from the two background model components, $G$ and $F$, in the following two ways:

$$t(I + s) = stF + G + tI_0,$$

$$t(I + s) = t(I_0 + s)F + G,$$

where $I_0$ denotes the desired zero-background image $I_0$ in DN/s, $s$ denotes the sky level in DN/s, and $t$ denotes the exposure time in seconds. To measure the sky level, we fit a Gaussian to the count histogram of the source-masked image before sky subtraction. The
The Detection and Removal of Large-scale Detector Background Structures in NICMOS Observations

Figure 2: Time sequence of the residual corner glow intensity. The projections on the first PC are shown in the number of standard deviations away from the mean, and are used as a measure of the corner glow intensity. Each horizontal mark represents one exposure and its exposure time. Consecutive observations are connected with a vertical line. The left and right panels show the time sequence for the F110W and F160W data sets, respectively. Histograms are plotted to show the distribution of intensities. Note the uniformity of the intensity and that none of the exposures reaches zero.

Gaussian is truncated on the high side to avoid including the corner glow. The resulting sky levels agree well with the PEDSKY output of SKYVAL.

The model components, $F$ and $G$, can be solved algebraically using stacked images, where the over-line average symbol denotes median values or median stacked images:

$$F = \frac{t(I+s) - (I+s)}{st}$$

$$G = t(I + s) - stF.$$  \hspace{1cm} (3)

Both cases in Equation 2 yield the same solution for model components $G$ and $F$. The resulting models (Figure 3) show that we have successfully separated out two unique background structures for components $G$ and $F$ from the stacked images. The models derived from the high and low glow level sets of images show the same structures. They differ only in the intensities of the $G$ component and the signal-to-noise. The models derived from the F110W and F160W filters show identical structures for the $G$ component. However, fringe pattern is detected in F110W and not in F160W (Figure 3).

Not surprisingly, the corner structure of the $G$ component largely resembles amplifier glow and linear dark current (panel $a$ of Figure 4), as it is the main background structure we are attempting to extract (Section 3.). The nonzero component $G$ suggests that the bias temperature used in the current NICMOS reduction pipeline cannot account for the observed variation. The residual flat structure in component $G$ also points to a temperature effect (panel $b$ of Figure 4). The center negative structure resembles the structures in BEP (panel $c$ of Figure 4). Riess & Bergeron (2008) reported significant BEP in 5% of the NICMOS exposures in the SHOES program. Here, for the fact that the structure survives median stacking and that it is visible in the inspection of individual images, we conclude that residual BEP-like structure affects every exposure.

The $F$ component of F110W shows very clear fringe pattern, but none is detected in that of F160W. The pattern is well described by a simple model of concentric ellipses (panel $d$ of Figure 4). The ellipses are centered approximately at [128, 28], tilted at $\sim \frac{\pi}{3}$.
with respect to the x axis, have minor radii in multiples of \(\sim 20\) pixels and major-to-minor semi-axes ratio of \(\sim 1.5\). The fringe pattern resembles neither the reference dark, flat (panel e of Figure 4) nor BEP (panel f of Figure 4). Fringe patterns in NICMOS have previously been reported in NIC1 narrow-band earth flats (Gilmore et al. 1998), but have never been reported for a wide filter. Gilmore et al. (1998) reported the effect to be particularly apparent in the bluest filters, F095N, F097N and F113N, and not detectable in the redder filters, F164N and F166N. These wavelength regions are consistent with our observation of fringing in F110W and non-detection in F160W, and suggest that only the bluest photons are affected by this anomaly.

It is unclear whether the observed fringe pattern in F110W is a result of true detector-induced fringing, although the characteristics of this structure seem to point in that direction. The fact that the structure does not resemble the reference files rules out a simple bias or color temperature effect. The ubiquity of the effect makes hypotheses such as stray light from optics and persistence less likely. The fact that the fringe structure scales with the sky and only affects the bluest photons gives clues to the source of the fringing photons. It is most likely zodiacal light which is blue at these wavelengths (Aldering 2002), although there are indications that bright earth limb could also be a factor. The interference rings are quite widely spaced; a mere sub-milliradian wedge in a detector layer can potentially produce such an effect. A model of the NIC2 detector layers with descriptions of index of refraction, transmission coefficient and thickness is required to confirm whether zodiacal light can indeed produce the observed pattern.

Figure 3: Model components for the background structure (available for download at http://supernova.lbl.gov/~hsiao/nicmos/). The left and right panels show model components from the F110W and F160W image sets, respectively. The top and bottom panels show component \(G\) which is constant in absolute counts and component \(F\) which scales with sky level, respectively.
Figure 4: Comparisons between F110W model background structures and known NICMOS structures. The top panels compare component $G$ (panel a) to reference flat (panel b) and BEP (panel c). The contour lines of component $G$ are drawn at $-4$ (yellow) and $-0.5$ (green). The bottom panels compare component $F$ (panel d) to reference flat (panel e) and BEP (panel f). The contour lines of component $F$ are best-fitting concentric ellipses at the scale of 1.

5. Reconstructing erratic middle rows and columns

The erratic middle row and column contain the first pixels read out in each quadrant. This anomaly is suspected to be the result of the steep and highly non-linear detector shading corrections (Thatte et al. 2009), making these pixels especially susceptible to changes in the detector environment. Nonetheless, there is no reason to believe that these pixels are less sensitive. In this section, we outline methods for recovering the flux in these pixels. An image with corrected middle row and column yields substantial improvements in the extraction of astronomical sources and in the construction of model images (Section 4.).

The erratic middle row and column in calibrated science images have drastically different characteristics (Figure 5). The noise characteristics of the middle row and neighboring rows are identical. The residual corner glow, although weak in this region, is preserved in the middle row. The application of a simple constant offset determined from source-masked middle and neighboring rows will recover the flux.

The behavior of the middle column is more complex. Here, we introduce a novel method for reconstructing the middle column using PCA with missing data. PCA is performed here to determine the bases for the variations in the middle columns (Section 2.). Specifically, we adopt the algorithm of EM PCA (Roweis 1998). The EM PCA naturally accommodates missing data, which in this case are masked-out astronomical sources. Warm and hot pixels are also identified and masked using multiple iterations of the EM algorithm.

The first two principal component eigenvectors derived are plotted in the right panel of Figure 5. The first eigenvector describes the broad overall shape of the middle column and accounts for most of the variation (80.2%). Interestingly, a middle column with shorter exposure time tends to have a stronger curve, but the shape does not correlate with the residual glow level (the time sequence in the orbit; Section 3.). The correlation perhaps reflects the state of the detector at the time of read out. The second eigenvector describes the coherent modulating signal which produces the larger-than-normal noise levels.
6. Post-processing procedures and improvements

We summarize here the steps for deriving the two-component background model and for correcting individual exposures:

1. Construct masks from source extraction, data quality image and bad pixel map
2. Reconstruct erratic middle column and row of individual exposures
3. Segregate exposures using the intensity of residual corner glow
4. Derive model components from algebraic manipulations of stacked images
5. Fit model components to individual masked images

The construction of model images, recovery of the middle rows and columns, and the model fitting of individual exposures all require masks which exclude astronomical sources and defective pixels, while astronomical source extraction performs best on background-removed images. The above procedures thus need to be iterated. Here we use the SExtractor program (Bertin & Arnouts 1996) to build the catalog of all the sources for every image.
Figure 6: An example for the reconstruction of the middle column using PCA. Panel a plots the PCA reconstructed model in red, and the middle column before and after the correction in blue and black, respectively. Panels b and c show the images before and after the correction, respectively. The correction successfully recovers the galaxy flux.

and use the measured Petrosian radii to build the source mask. Also note that while the reconstructed middle row has the corner glow structure preserved, the reconstructed middle column does not (Section 5.). The iterations thus also help obtain the characteristics of the corner structure for the extractions of gapless background models.

The two cases in Equation 2 can be written in terms of the desired zero-background image $I_0$ as follows:

$$I_0 = I - \left[ s(F - 1) + \frac{G}{t} \right],$$

$$I_0 = \frac{1}{F} \left( I - \left[ s(F - 1) + \frac{G}{t} \right] \right).$$

(4)

After obtaining the model images (Section 4.), applying the corrections is trivial and the bulk of the background structure can be removed using Equation 4 with the sky value $s$ and exposure time $t$ for the image in question. Alternatively, since there is a spread in the intensities in the corner structures (Figure 2), fitting model components to individual masked images with variable sky value and exposure time can potentially improve the removal of the structures. Here we fit the model components using three free parameters: a fiducial sky value $s'$, a fiducial exposure time $t'$, and a constant offset. Non-linear least-squares IDL fitter MPFIT (Markwardt 2009), which adopts the efficient Levenberg-Marquardt technique, is used here to perform the fits. A typical example is illustrated in Figure 7.

The background-removed images have very obvious cosmetic improvement from the original calibrated images. To obtain a quantitative gauge of the improvement, we measure the dispersions in the fluxes of faint galaxies. Galaxies which appear in more than three images are included. The photometry adopted uses six-pixel radius apertures and local background subtraction. Median absolute deviation is adopted to provide a robust measure
Figure 7: An example of the removal of background structures. Panels a and b show the example image before and after the removal of background structures, respectively. Panels c and d show the image before background removal with contours from the background model components G and F over-plotted, respectively. The contours for component G are −3 (yellow) and 3 (green) in DN. The contours for component F are concentric ellipses fitted at the scale of unity.

Figure 8: Improvements (reductions) in the magnitude dispersions for faint galaxies with respect to their average AB magnitude. The left and right panels plot the improvement for the F110W and F160W data sets, respectively. The first case of Equation 4 is applied here as an example. The dotted curves mark the expected magnitude dispersion from Poisson noise. The histograms and the best fit Gaussian functions show the distributions for both passbands shifted toward approximately 20% improvements.
of dispersion. Removing the background structures with the methods described here yields on average a 20% reduction of magnitude dispersion in the photometry of faint galaxies (Figure 8). The two methods of applying the corrections for component $F$ from Equation 4 show similar improvements and no clear advantage for one over the other.

Acknowledgments. We are grateful to Anton Koekemoer, Tomas Dahlen, Deepashri Thatte, Glenn Schneider for valuable discussions. This work has been supported by the Office of Science, U.S. Department of Energy, through contract DE-AC02-05CH11231 and in part by NASA through grants associated with HST-GO-10496 and HST-GO-11799.

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Calibrating the HST Focal Plane post SM4

Colin Cox and Matthew Lallo
Space Telescope Science Institute, Baltimore, MD 21218

1. Introduction

Knowing the location of the science instrument apertures within the HST focal plane to within Guidestar Catalog 2 errors (~0.25 arcsec) is valuable for two primary reasons. 1.) It improves pointing and makes target acquisitions more expeditious for a small aperture instrument like COS, and 2.) It improves the astrometry metadata which aids in archival and multi-mission object matching.

Figure 1: HST field of view showing the instrument apertures superimposed over the M35 open cluster in the December orientation.
Calibration Proposal 11878 contains several visits observing the open cluster M35. Within each visit exposures are made using several of the HST instruments with only small changes in pointing so that the same pair of guide stars may be used throughout. An astrometric catalog of M35 (McNamara & Sekiguchi 1986) as re-reduced by Welter & Kimmer (1994) based on HST FGS observations, is used for these calibrations. It contains positions and proper motions for the cluster members measured well enough so that a final positional accuracy of 30 mas is achieved for an individual star. By taking guide stars and targets from this catalog, the positions of apertures relative to the FGSs are determined to within 50 mas and possibly better for the larger apertures when several stars may be used. In these cases an approximate measure of the aperture orientation can also be derived. The figure above illustrates the pointing in one visit for the WFC3 exposure. The STIS and ACS exposures are performed with exactly the same pointing. A small shift is needed to place the 2.5 arc-second diameter COS aperture on a suitable star. (The COS aperture is too small to show on the figure. HRC and NICMOS apertures are shown but neither were operable during Cycle 17.) The position of the two guide stars are indicated and connected by the red line.

2. Measurements

Measurements were made in December 2009, March 2010 (COS only) and April 2010. Information was also collected from each instrument’s alignment measurements following the May 2009 Servicing Mission. These alignment measurements utilized more stars and telescope positions to accurately measure scales and rotations. The analysis method consists of first obtaining accurate celestial coordinates of the measured stars from the catalog. Next we calculate the corresponding positions in the (V2,V3) coordinate system describing the HST focal plane. This requires knowledge of the telescope pointing which is in turn derived from the guide star positions taken from the same catalog and a detailed model of the FGS positions and distortions. This calculation is carried out by the HST Sensors Calibration (SAC) group at GSFC. For the same set of stars we measure the centroids in each instrument’s (x,y) pixel coordinate system. Each instrument has its own distortion solution which converts from (x,y) to (V2,V3) measured in arc-seconds. We apply this to the measured centroids to find another set of (V2,V3) coordinates which we compare with those derived from the sky positions and the HST pointing solution. Normally the two sets agree to within less than one arc-second. We rely on the distortion solutions for the overall shape and scale factors and find a linear transformation involving only shifts and rotations to match the two sets of (V2,V3) positions. From this we establish new reference positions and sometimes orientations. Normally the derived orientation change is smaller than the estimated error on the measurement and no change is made. The derived shifts imply new values for the constant terms of the distortion solutions. These are applied to the instrument distortion calibration (IDC) tables which are used for detailed image analysis. If the shifts are not significantly less than the GSC2 accuracy of 0.25 arc-seconds then they are also applied to the Science Instrument Aperture Files (SIAF) which support telescope pointing proposal preparation. Measurements of the COS instrument are the most critical because of its small aperture size and expected drift as new instrument’s optical benches tend to outgas. The method for measuring the COS position is simpler, using just the target acquisition observation. In the first phase the telescope points the V2V3 position given by the current SIAF to the RA and Dec of the star. This typically placed the aperture center a short distance from the star. Subsequently the acquisition software moves the star to the exact aperture center. Again using the GSFC calculations we determine the true V2V3 position matching this pointing. If this is significantly different, more than about 0.1 arc-seconds from the stored values, they will be updated.
3. Outcome

Over the past year we have seen changes of up to 0.5 arc-seconds as illustrated. The largest step of about 2 arc-seconds is between the initial pre-flight modeled position and the first in-flight measurement. Updates to the SIAF were made following the July and December 2009 and the March 2010 measurements, but not after the April 2010 measurements. Deceleration of the aperture-position trend is clear from the diagram and consistent with previous experience with newly flown Science Instruments.

![Diagram of WFC3 UVIS and COS Aperture Moves](image)

Figure 2: Calculated and measured positions of WFC3 and COS reference positions over a period of almost one year. The measured positions, which include initial alignment values from July 2009, do not vary by more than 0.4 arc-seconds.

Like COS, the WFC3 camera was installed in May 2009 and so received extra attention to measure its placement. The first measured position was about 3 arc-seconds away from the pre-flight modeled position and required a rotation correction of 0.075 degrees. For the large UVIS aperture such deviations normally presented no risk of missing a target, but would affect the astrometry of the science products. The positions of the IR channel moved in a similar manner and are not shown here. SIAF updates were made following
the July and December measurements. The IDC tables were updated on these occasions and also in April 2010. The newly resurrected instruments, ACS and STIS, experienced the physical disturbances of SM4 crew interactions, which included (among other things) the cutting of grates and snapping of handrails. Despite this, both instrument’s post-SM4 position determinations were found within 0.2 arc-seconds of their SIAF ground system values, which were last updated in 2002 (ACS) and 1996 (STIS). No SIAF changes are deemed necessary at this time.

References

HST Cycle 19 Exposure Time Calculators

R. I. Diaz, I. Busko, P. Greenfield, T. Miller, M. Sienkiewicz, and M. Sosey

*Space Telescope Science Institute, 3700 San Martin Dr., Baltimore MD 21030*

V. G. Laidler

*Computer Sciences Corporation, Space Telescope Science Institute, 3700 San Martin Dr., Baltimore MD 21030*

**Abstract.** The Exposure Time Calculator (ETC) is a web-based application to assist users in calculating the exposure time needed for their HST observations, or the Signal-to-Noise Ratio (SNR) that can be achieved with a given HST observing time. These quantities are key for the preparation of proposals and observations during Phase I and Phase II of the proposing cycle and therefore must be sufficiently accurate for each of the supported observing modes of all the HST instruments. Developing a general tool that shares commonality among the different instruments is complicated, not only from the point of view of attaining accuracy of the calculations but also regarding reliability, portability, and maintainability. We are currently in the final stages of the development of the new ETC for Cycle 19 in Python. This new version improves these qualities and meets the needed level of reliability for the ETC. It also provides a basis for JWST ETCs which are in preliminary development, with a limited-functionality prototype planned to be available in Spring 2011. This paper shows the new tool, its improvements over the previous ETC and the current status of the new version.

1. **Introduction**

The ETC is a web-based application developed to assist HST observers in preparing their observations with any of the supported observing modes for HST. The main purpose of the ETC is to calculate the exposure time needed for observations requiring a particular SNR, and also to ascertain the SNR for simulated observations for a specified exposure time. Currently the ETC provides support for all active HST instruments: ACS, COS, NICMOS, STIS, and WFC3.

The ETC relies heavily on *pysynphot*. *Pysynphot* is a synthetic photometry package that performs photometry by computing detected photons as a function of wavelength (and, for spectrographs, detector bin) using a specified input spectrum, known throughput, and detector quantum efficiency curves (see Figure 1). It uses the throughput tables maintained as part of the Calibration Database System (CDBS) at STScI to compute instrument throughput and efficiency. This information is then used by the ETC to further evaluate the feasibility of the observations for the selected mode and the selected target. The results of the calculations are displayed to a web output form along with any relevant information or flags needed on the planning of the observations (see Figure 2).
The ETC provides also the result of the calculations in the form of tables and plots that could be useful to the user. The results are saved in a database and the associated files are retained for future reference. This allows contact scientist to retrieve any observers calculation in case of later questions or problems. It also allows the STScI ETC helpdesk to determine what happened in cases where the observer encountered errors.

Other ETC capabilities include:

- The option to upload an user-input spectra for the ETC calculation. The input spectra can be provided via an ASCII or a FITS file
- Calculation of optimal SNR as would be obtained by PSF fitting
- Inclusion of instrument-specific effects like dark current and read noise
- Bright limit check for those modes where there are safety concerns.

2. ETC for Cycle 19

For Cycle 19, the ETC software is being re-written in Python. This new version looks to maintain the current functionality of the ETC while being flexible enough to accomodate
quick updates to the software and the data; which sometimes is needed for the accurate planning of HST observations. The new ETC will also provide improved performance, reliability and testing. In addition, the computation engine will be readily adaptable to other telescopes, in particular JWST.

2.1. The Requirements

Previous experiences managing the Java version of the ETC as well as requests from the HST Instrument Teams and external ETC users were considered carefully when deciding on the configuration and design for the new system. The list of requirements that the new software were to satisfy are:

- **The new system can be installed in one step.** A build of a specific Compute Engine and Web Interface defined by a specific configuration file using a specific (and unique) set of uniform resource locators (URLs). The configuration of an installation includes version numbers of all subsystems, including CDBS.

- **Ability to support multiple installations on the same computer.** Different versions of the ETC can be run concurrently in the same computer.

- **Consistent test scheme and nightly regression testing.** All the instruments have a consistent set of test defined following the same input and reporting format. These test should be run every day in order to identify any problems with the system.

- **Separation of web and computational functionality.**
• **Ability to script ETC calculations from Python without a web server.** Astronomers with Python experience can use the calculation engine directly for batch processing or science exploration.

• **Handle failover and load balancing.** Using standard Apache/Database server schemes it should be possible to handle failover and load balancing.

• **Simple database structure.**

• **Concentrate instrument information in one place.** Instrument data should be stored and tested in accordance to the the CDBS quality and testing standards.

• **Same look and feel of previous ETCs**

• **Offer interactive form features that weren’t working well in the old ETC.**

### 2.2. The Improvements

The functionality and look from the previous ETC is preserved in ETC 19.0. However, when necessary, changes have been made in order to optimize the tool and handling of the code as well to improve the rendering of information to the users.

What are the technological changes?

• The underlying technology (web infrastructure, database, plotting, and so on) has been implemented using several tools according to the part of the system they will manage. These include:
  
  – **Python**, the underlying script language used by pyetc  
  – **Django**, the web framework used by pyetc  
  – **MySQL**, the Database system that will store information on each of the ETC calls  
  – **matplotlib**, Used to provide plots of the throughput curves, source spectra, and calculated quantities.  
  – **PyFITS**, used to read FITS files  
  – **distutils**, provides support for building and installing additional modules into a Python installation  
  – **nose**, provides an alternate test discovery and running process for unittest  
  – **Pandokia**, a test management and reporting system.

• Version control for both software and data (instrument and model) has been improved

• Configuration and installation procedures have been simplified and standardized

• The underlying database schema has been simplified

• Testing is being significantly expanded so we find problems before users encounter them. New tests cover more of input parameter space as well as better end-to-end testing of the system.

What are the other improvements from previous ETC versions?

• Web pages (HTML) have been simplified & brought into closer compliance with W3C standards (see Figure 1 and 4)

• Internal handling of previous calculations has been simplified
PyETC is inherently parallel (see Figure 3). It can service multiple requests concurrently on a single computer, taking full advantage of multiple CPUs, if available. If necessary, standard web server load-balancing systems can distribute the workload across multiple machines to service a greater workload.

This architecture offers improved reliability and performance.

The computation engine is readily adaptable to other telescopes.

The computation engine is being developed to support direct interactive or batch use as well as the web application, and will be publicly released and available to Python-savvy astronomers.

3. Development Status

Nearly all the basic computational functionality has been implemented, as well as elements of the web interface including the plotting functionality, for all supported instruments and observing modes (ACS, STIS, WFC3, and COS). The new version is being commissioned by comparing its computational results to the previous version, with a maximum acceptable discrepancy of 5% (smaller for some values) except for known/expected differences between the two versions. More than 14000 tests, including old and new tests, are running nightly in an automatic testing and reporting system.

Currently we are completing unimplemented functionality (primarily in the web interface) and working on testing the user interface and load balancing capabilities. As development progresses and we look to resolve test differences, we keep modifying the code and data, as well as adding new test to improve the parameter space coverage. In order to remove redundancy, distributed logic, and dependencies, and make the code more readable and maintainable, we are also working on refactoring the code from the original set.

We are also implementing configuration control for ETC instrument, configuration, and model data via the CDBS system, and implementing the needed tools for the Bright Object Table generation. Before the start of the Cycle 19th call for proposals, and as time allows, we will continue working on other enhancements for the ETCs.
Figure 4: A section of the input ETC form for COS Spectroscopic modes.

We are also drafting the User and Developer documentation and starting to implement support for JWST instruments. The JWST ETCs are in preliminary development, with a limited-functionality prototype to be available in Spring 2011.

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Laidler et al. 2010, “Pysynphot Tutorial” (Baltimore, STScI; http://stsdas.stsci.edu/pysynphot/)
UBVRI-ZY and ugriz zeropoints from 20 calspec standards

Andrew Pickles
Las Cumbres Observatory Global Telescope, Santa Barbara, CA

Abstract. Synthesized UBVRI-ZY and ugriz zeropoints are presented, calibrated on 20 calspec standards with standard photometry and covering a wide range of color, from DA white dwarfs through G/K stars to VB8 (M7V). The zeropoint dispersions are 0.027, 0.020, 0.008, 0.014 and 0.016 mag respectively in UBVRI, where the $UMBVRCI$ filter system profiles minimize these dispersions. $V(Vega)$ is found to be $0.021 \pm 0.008$ mag from an average of 17 calspec standards, with Vega colors which are close to zero. Zeropoints for u & u' are found to be $-0.03 \pm 0.02$ mag; those for griz and g'r'i'z' are $0.00 \pm 0.02$ mag.

1. Filters and zeropoints based on spectrophotometric standards

Mean fluxes and synthetic magnitudes in the system bandpasses of all the filters have been computed for up to 20 standards which have both accurate digital spectra available from calspec1, and measured photometry from the literature.

The filter-bands discussed here are illustrated in figure 1. Synthesized magnitudes of flux calibrated spectra have been determined using these filter system bandpasses and the procedures described in the Synphot User’s Guide2

\[
mag_{\nu}(b) = mag_{\lambda}(b) - 5 \times \log_{10}\lambda_{pivot} + 18.692 - ZeroPoint_{\nu}
\]

where the numeric constant, as derived in Sirianni et al (2005), has the advantage of bringing $mag_{\nu}$ close to the AB79 system of Oke & Gunn (1983). The bandpass zeropoints are to be determined.

The approach is similar to that presented by Holberg & Bergeron (2006) but deliberately attempts to calibrate a large number of filters with standards covering a wide range of color and spectral type. The small zeropoint dispersions achieved here demonstrate the validity of this approach over a wide variety of spectral types and colors. They reinforce the advances that have been made in accurate flux calibration by Bohlin (1996), Colina & Bohlin (1997), Bohlin et al (2001), & Bohlin (2010).

2. Calspec standards

There are 13 standards with catalog photometry and STIS_NIC_003 calibrated spectrophotometry covering 0.1 to 2.5 µ, which include the latest HST calspec calibration enhancements and the 2010 corrections to STIS gain settings. The spectra are mainly of white dwarfs, but include four G dwarfs and VB8(=LHS 429, a late M dwarf) so provide significant color range: $-0.3 < V - I < 4.6$.

1http://www.stsci.edu/hst/observatory/cdbs/calspec.html
2http://www.stsci.edu/resources/software_hardware/stsdas/synphot
Figure 1: The filter system bandpasses discussed here are illustrated. In the first panel the Tycho2 $B_T V_T$ filter system transmissions are taken from Maiz Apellániz (2006), and overplotted (dotted lines) in the third panel for convenient reference. The 2MASS $J_2 H_2 K_2/S$ system total response curves are from the IPAC website; the $JHK_{MKO}$ filter transmission curves (dotted) are from Tokunaga et al (2002). In the second panel the Landolt $U_L B_L V_L R_L I_L$ transmission curves for the filters alone are shown (dotted, Landolt 1992a); their normalized system response including typical atmosphere and detector QE are shown (solid, Cohen 2003a). In the third panel the synthetic $U_M B_M V_M$ filters are from Maiz Apellániz (2006) ($V_M$ plotted as dashed line); the $V_C R_C I_C$ synthetic curves are from Bessell (1979) and the $Z_V Y_V$ curves are from the UKIRT/VISTA collaboration. An $R_p$ “Mould-type” R-system response is also illustrated. In the bottom panel the Sloan system bandpasses are shown for both the unprimed survey imaging bandpasses (dotted lines, where the filter interference red edges move bluer in the vacuum of the survey camera) and primed bandpasses used in air for PT and Southern standard observations ($u'g'r'i'z'$ – solid lines); they are from the JHU skyservice and the USNO FNAL websites.
Figure 2: The bandpasses, Vega magnitudes (colored symbols) and ±2σ zeropoint dispersions (bars) from table 1 are over-plotted on the STIS.005 calspec/Kurucz spectrum of Vega, which defines the nominal 0-mag definition for filters on the vega system. The horizontal line at \(3.631 \times 10^{-20}\) erg cm\(^{-2}\) sec\(^{-1}\) Hz\(^{-1}\) (3631 Jy) illustrates the AB=0 mag reference.

There are two calspec K giants (KF08T3, KF06T1) with low reddening which were observed with NICMOS to provide IRAC calibration (Reach et al 2005). There is little optical standard photometry for these two K giants, so they are only included in the 2MASS \(JHK_{2/S}\) zeropoint averages, but optical colors estimated from their types are shown to be consistent with the derived optical zeropoints.

There are three additional calspec white dwarfs with coverage to \(\sim 1\)\(\mu\) (G93-48, GD 50, Feige 34), and two subdwarfs (G158-100, BD +26 2606) observed by Oke (1990) and Oke & Gunn (1983) for which fairly extensive photometric data are available in the literature. The latter 5 spectra calibrated from the uv to \(\sim 1\)\(\mu\), were extended to 2.5\(\mu\) for illustrative purposes. Their synthetic infrared magnitudes are computed and shown for comparison, but only their optical zeropoints are combined in the averages.

2.1. Standard Catalog Magnitudes

The matching “catalog” photometry for calspec standards comes from i) the Tycho2 catalog for \(B_TV_T\) (Høg et al 2000), ii) the 2MASS catalog for \(J_2H_2K_{2/S}\) (Cutri 1998), iii) \(UBVRI\) data for GD 71 (DA1), G93-48 (DA3) & GD 50 (DA2) from Landolt (2009), and for G191-B2B (DA0), BD +17 4708 (sdF8), BD +26 2606 (sdF), AGK +81 266 (sdO), GRW +70 5824 (DA3), LDS 749B (DBQ4), Feige 110 (DOP), Feige 34 (DA) & G158-100 (sdG) from Landolt & Uomoto (2007), iv) from Bessell (1991) for optical and infrared photometry of VB8 (M7V), v) from the UKIRT standards listed on the JAC/UKIRT website\(^3\) for \(JHK_{MKO}\) and their WFCAM \(Z_VY_V\) data, and vi) from Wegner (1983), Lacomb & Fontaine (1981) and

\(^3\)http://www.jach.hawaii.edu/UKIRT/astronomy/calib/phot_cal/
Hauck & Mermilliod (1998) for Stromgren standards (white dwarfs). The above photometric data is on the “Vega” system where the magnitudes of Vega are nominally zero in all bands.

Sloan catalog (vii) data on the AB system are from Smith et al (2002) for primary standards, from the SDSS IPT catalog Davenport et al (2007) and Southern SDSS standards Smith et al (2005), from the SDSS DR7 database\(^4\) for LDS 749B, VB8, GD 50, and in 2 cases (G191-B2B & GD 71) from Holberg & Bergeron (2006) for \(u'g'r'i'z'\) magnitudes.

Additional \(UBVRI\) and \(ugriz\) photometric data for GD 153 (DA0) are from Holberg & Bergeron (2006). Additional \(BV\) data for P041C, P177D, P330E (G0 V), KF08T3 (K0.5 III) and KF06T1 (K1.5 III) are from the calspec website.

For each standard and each filter, magnitudes have been computed via equation 1, and the zeropoints calculated to match synthetic to observed magnitudes.

3. Zeropoint means and dispersions

Complete tables are presented in Pickles & Depagne (2010). Table 1 summarizes the derived zeropoint means and dispersions for each filter, adopted zeropoints, synthesized magnitudes and fluxes of the STIS\(005\) calspec/Kurucz spectrum of Vega, and derived values of \(< F(\nu) >\) for 0-magnitude in all filters. The number of standards included in each filter zeropoint calculation ranges from 5 for \(V\) to 17 for \(B\) and \(V\).

The filter bands and derived magnitudes and zero-point dispersions for Vega from table 1 are illustrated in figure 2, where the STIS\(005\) calspec spectrum illustrates the nominal 0-mag definition for filters on the vega system, and the horizontal line at 3631 Jy illustrates the AB=0 mag reference. The error bars show the synthesized magnitudes with \(\pm 2\sigma\) zeropoint dispersion errors.

The 2MASS and SDSS coordinates of VB8 differ by 7.4-arcsec, corresponding to the large proper motion (-0.77, -0.87 arcsec/year) for VB8 between the epochs of the two surveys. Good VB8 zeropoint fits are obtained for \(BVRI\), \(J_2H_2K_2/S\) and \(HK_{MKO}\) and \(grz\), but not for \(J_{MKO}\), \(u\) or \(i\) bands. There is no U-band photometry for VB8.

4. Choice of Landolt synthetic filter bandpasses

Several possible synthetic system bandpass profiles were tested for \(UBVRI\), and an empirical choice made of the best system bandpass(es) that minimize zeropoint scatter in the fitted mean zeropoints for each filter, over the full color range.

Table 1 lists zeropoints for \(UBVRI\) using both Landolt (system) filter response functions convolved with a typical atmosphere and detector response Cohen et al (2003a), and synthetic system bandpasses for \(UMBMV\) from Maiz Apellániz (2006) and \(VCRCIC\) from Bessell (1979).

It may seem that the Landolt system response curves would provide the optimum synthetic matches to Landolt photometry, but this is not the case, as shown by the results in table 1.

Both Landolt and Kron-Cousins measurements seek to emulate the original Johnson system for UBV. Both apply color-calibrations to their instrumental magnitudes, to bring them into correspondence with catalog values extending back several decades. These steps are summarized in Landolt (2007) for example, and previously in Landolt (1983, 1992a, 1992b) to illustrate how equipment changes over time have required slightly different color corrections to maintain integrity with the original system definition. These calibration steps are further reviewed in Sung & Bessell (2000).

\(^4\)http://www.sdss.org/dr7/
Figure 3: Zeropoints are shown for $U_L B_L V_L R_L I_L$ (left) and $U_M B_M V_C R_C I_C$ (right) as a function of standard $V - I$ color. Vertical error bars indicate the average and $\pm 1\sigma$ dispersion in each panel. Crosses indicate White Dwarfs, Pentagons indicate dwarfs. Three point stars indicate the two K giants, which are shown but not included in these zeropoint averages or dispersion as they lack accurate UBVRI photometry. The reddest point corresponding to the M7V VB8 lacks U-band data. The selected UBVRI system profiles on the right minimize zeropoint scatter and trend with color.

It is possible to measure synthetic magnitudes in the Landolt bandpasses and apply corrections to achieve “standard” values. But it is clearly preferable to find the synthetic bandpasses that best match the spectrophotometric data with zeropoints, and no color term(s). For UBVRI, the results in table 1 and figure 3 indicate that these are best provided by the $U_M B_M V_C R_C I_C$ bandpasses. In figure 3 and subsequently, the ordinate scale is set so that vertically higher zeropoints result in magnitudes which are larger (fainter).

The zeropoint dispersions are 0.027, 0.020, 0.008, 0.014 and 0.016 mag for $U_M B_M V_C R_C I_C$ respectively in table 1, for the full color range from White Dwarfs to VB8 ($-0.3 < V - I < 4.6$). The effective wavelengths vary from 355 to 371 nm, 432 to 472 nm, 544 to 558 nm, 640 to 738 nm and 785 to 805 nm for $U_M B_M V_C R_C I_C$ respectively, between White Dwarfs and VB8. Figure 3 illustrates that the selected bands show much less zeropoint dispersion than do $U_L B_L V_L R_L I_L$, with negligible trend with color.

This is not a criticism of Landolt system response curves, which enable accurate photometry with appropriate calibration and color corrections, but indicates that the selected $U_M B_M V_C R_C I_C$ system profiles are best for deriving synthetic spectrophotometry of flux calibrated spectra to properly match Landolt photometry.

The zeropoint dispersions for $U_3 B_3$ from Azusienis & Straizys (1969) and Buser (1978) in table 1 are worse, at 0.050 and 0.024 mag respectively, and thus support the “inflected” $U_M$ transmission curve as a better U-band measure for stars of varying type & color. The dispersion for $V_M$ is marginally worse than for $V_C$, where $V_C$ has a slightly more elongated red tail than $V_M$.

The zeropoint dispersions for $UBVRI$ and $ugriz$ (primed and unprimed) in table 1, typically closer to 0.02 mag than 0.01 mag, indicate both the accuracy and the current limitations of comparing synthetic photometry of well calibrated spectra with good standard photometry. Tighter fits can be obtained by restricting the selection of comparison stars,
for instance to only WD standards. But such zeropoints are then a function of color, and lead to errors for other stellar types much larger than the nominal dispersion.

It is gratifying that these comparisons match so well over a large range of color and standard types, confirming the increasing correspondence of spectrophotometric to photometric standards, and setting the basis for accurately calibrated magnitudes of an extended library of synthetic spectra.

4.1. Other synthetic filter zeropoints on the Vega system

The transmission functions for $B_T V_T$ have been taken from Maiz Apellániz (2006). The zeropoint dispersions for $B_T V_T$ are 0.045 and 0.020 mag respectively, which is acceptable given the typical errors in the photometric values for fainter stars, and for several standards included here. There are a total of 7 values covering a color range from white dwarfs to G dwarfs for $B_T$, but only 5 with accurate $V_T$ information, with HD209458 (G0 V – out of planet occultation) being the reddest comparison standard for $B_T V_T$.

The filter transmission functions for $Z_V Y_V$ have been taken from the Vista website\(^5\), where the subscript “V” is used to refer to both the VISTA/UKIDSS consortium and the fact that these are *vega* based magnitudes. The WFCAM detector QE is not included but, unlike a CCD, is roughly flat over these wavelengths. The dispersions for $Z_V Y_V$ zeropoints, compared to only 5 and 4 UKIRT standards measured with the WFCAM filters, are 0.038 and 0.031 mag respectively. The $Y_V$ zeropoint for G158-100 is suspect as its spectrum is not well defined at 1$\mu$m. $Z_V Y_V$ zeropoint determinations may improve as more photometric standards in common with spectrophotometric standards are measured. The zeropoint results and color ranges for $B_T V_T Z_V Y_V$ are illustrated in figure 4.

The transmission functions of the 2MASS filters, including detector and typical atmosphere, have been taken from the IPAC website\(^6\). The zeropoint dispersions for $J_2 H_2 K_2/3$, \(^5\)http://www.vista.ac.uk/Files/filters
\(^6\)http://www.ipac.caltech.edu/2mass/overview/about2mass.html
Figure 5: Zeropoints are shown for $J_2H_2K_{2/S}$ (left) and $JHK_{MKO}$ (right) as a function of standard $V-I$ color, with averages and dispersions. Symbols as before. The KIII points (triangles) have been included for $JHK_{2/S}$ The zeropoints for $JHK_{MKO}$ show some trend with color, but the $J_{MKO}$ value for VB8 has been excluded due to uncertainty over its catalog magnitude.

indicated graphically in figure 5, are about 0.02 mag, corresponding to the typical (low) 2MASS errors for these stars.

The transmission functions for the JHK filters on the Mauna Kea Observatory (MKO) system have been taken from Tokunaga, Simons & Vacca (2002). The zeropoint dispersions for $JHK_{MKO}$ are 0.02, 0.02 & 0.04 mag respectively, for relatively few standards, and they do show trends with color. The $J_{MKO}$ zeropoint for VB8 is indicated in figure 5 but, due to uncertainty with its catalog magnitude, it is not included in the $J_{MKO}$ zeropoint average.

The transmission functions for the Stromgren filters (not illustrated) are from Maiz Apellániz (2006). The zeropoints derived in table 1 are $-0.290 \pm 0.035$, $-0.316 \pm 0.012$, $-0.181 \pm 0.024$ and $-0.041 \pm 0.030$ for Stromgren $uvby$ respectively. These are averaged over 7 calspec white dwarfs for which Stromgren photometry was found in the literature.

The Stromgren zeropoints listed in table 1 lead to Vega magnitudes of 1.431, 0.189, 0.029 and 0.046 in Stromgren $uvby$ respectively, compared to values of 1.432, 0.179, 0.018 and 0.014 derived in Maiz Apellániz (2007).

5. Sloan filters on the AB system

The transmission functions of the unprimed $ugriz$ filters used in the imaging camera, including typical atmosphere and detector response, have been taken from the skyservice website\textsuperscript{7} at Johns Hopkins University (JHU). Those for the primed $u’g’r’i’z’$ filters used for standard observations have been taken from the United States Naval Observatory (USNO) website\textsuperscript{8}, for 1.3 airmasses.

The SDSS survey u-band data for VB8 appears to be too bright, likely because of the known red leak. The DR7 i-band magnitude is close to the r-band magnitude, and appears too faint for such a red star. The $u$ and $i$ band data for VB8 has been omitted from the zeropoint averages.

\textsuperscript{7}http://skyservice.pha.jhu.edu

\textsuperscript{8}http://www-star.fnal.gov/ugriz/Filters/response.html
UBVRI-ZY and ugriz zeropoints from 20 calspec standards

Figure 6: Zeropoints are shown for $u'g'r'i'z'$ and ugriz as a function of standard $V - I$ color, with averages and dispersions. VB8 is present in the grz plots, but has been omitted in the ui plots. The K III points are missing as there are no ugriz data for them. Symbols as before.

In order to compute zeropoints for the imaging survey (unprimed) Sloan bandpasses ugriz, conversion relations between the primed and unprimed Sloan system values were computed by comparing synthetic magnitudes of digital library spectra from Pickles (1998).

With the exception of two zeropoint adjustments for $g$ and $r$, these relations are identical to those listed on the SDSS site\(^9\), and are summarized here.

- $u = u' + 0.060 \times (g' - r' - 0.53)$ \(\sigma = 0.011\)
- $g = 0.037 + g' + 0.060 \times (g' - r' - 0.53)$ \(\sigma = 0.012\)
- $r = 0.010 + r' + 0.035 \times (r' - i' - 0.21)$ \(\sigma = 0.007\)
- $i = i' + 0.041 \times (r' - i' - 0.21)$ \(\sigma = 0.015\)
- $z = z' \quad (\sigma = 0.003)$

These relations were used to convert SDSS standard values on the primed system to unprimed values (and vice versa when only DR7 data was available, eg. for LDS 749B, P177D, P330E & GD 50), then compared with synthetic magnitudes computed in the unprimed bandpasses to derive their zeropoints and dispersions.

The zeropoint dispersions for $u'g'r'i'z'$ and ugriz listed in table 1 are all about 0.02 mag, for up to 14 standards covering a wide color range.

There is little evidence for non-zero zeropoints in griz; they are zero to within the measured dispersions, and are adopted zero in table 1. Zeropoints for both $u'$ and $u$ are $-0.03 \pm 0.02$ mag. The zeropoint results for $u'g'r'i'z'$ and ugriz are illustrated in figure 6.

\(^9\)http://www.sdss.org/dr5/algorithms/jeg_photometric_eq_dr1.html
6. V-band magnitude of Vega

The $V_C$ zeropoint derivation here, based on 17 standards (including VB8), results in a Landolt V magnitude of $V_C = 0.021 \pm 0.008$ for the STIS_005 calspec spectrum of Vega, with colors of $U-B = -0.007 \pm 0.031$, $B-V = -0.009 \pm 0.022$, $V-R = -0.002 \pm 0.016$ and $V-I = -0.010 \pm 0.018$. These values, based on 13–17 standards, are within quoted errors of those derived in Maiz Apellániz (2007).

Note that using the Cohen et al (2003a) filter definition results in a zeropoint using all 17 standards which leads to $V_L(Vega) = 0.013 \pm 0.021$, or alternatively $V_L(Vega) = 0.024 \pm 0.002$ when using a zeropoint based solely on the bluest three DA white dwarfs. The latter value is close to the recently adopted value of Landolt $V(Vega) = 0.025$ mag. However the $V_L$ zeropoint is clearly a function of color, as seen in figure 3. For the appropriate $V-I = -0.01$ color for Vega, the color-corrected zeropoint ($ZP_{VL} = -0.022$), leads to $V_L(Vega) = 0.021$, i.e. the same as the $V_C(Vega)$ derivation above — when color effects are taken into account.

We argue that our $V_C(Vega)$ magnitude of $0.021 \pm 0.008$ represents a realistic mean value and error for the Landolt V magnitude of the STIS_005 spectrum of Vega. The zeropoint dispersion can probably be improved by standard photometric measurements of more red spectrophotometric standards, including the IRAC K giants, which are shown in figure 3 but not used, as they lack optical standard photometry.

The further question of the small differences between Landolt, Kron-Cousins (SAAO) and Johnson V magnitudes are discussed in Landolt (1983, 1992a) and Menzies et al (1991) and summarized again in Sung & Bessell (2000), but is sidestepped here. Here we compare $UBVRI$ synthetic spectrophotometry of calspec standards to available published photometry on the Landolt system.

7. Vega and Zero magnitude fluxes

The final column of table 1 show the fluxes measured on the STIS_005 spectrum of Vega, the derived magnitudes of Vega with the adopted zeropoints, and the inferred fluxes (\(F(\nu)\) in Jy) for a zero-magnitude star (i.e. zero vega mag for $B_T V_T, UBVRI, Z_Y Y_V, JHK$, Stromgren uvby; and zero AB mag for Sloan u′g′r′i′z′ and ugriz filters).

The zero magnitude fluxes can be compared with other values given for instance by Bessell (1979) Bessell & Brett (1998), Cohen et al (2003b), Hewett et al (2006), and the IPAC\footnote{http://www.ipac.caltech.edu/2mass/releases/allsky/doc/sec6_4a.html}, SPITZER\footnote{http://casa.colorado.edu/~ginsbura/filtersets.htm} and Gemini\footnote{http://www.gemini.edu/?q=node/11119} websites. The largest discrepancy is for $U$, where the 0-mag U-band flux is about 4% less than the IPAC value for example, compared to our measured sigma of 2.7%. The zero magnitude fluxes for $griz$ and $g′r′i′z′$ are close to 3631 Jy, and close to 3680 Jy for $u$ and $u′$.

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Table 1: ZeroPoints, Zero-Mag fluxes and Selected System/Filter bandpasses

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Recovery of Very Low Signal-to-Noise Transiting Lightcurves Using Self-coherence

Ingo Waldmann

*University College London*

**Abstract.**

The detection and characterisation of telluric planets from space and ground facilities is one of the main goals of future exoplanet research, thus the importance of being able to fully recover and understand data in low signal-to-noise (SNR) conditions is crucial. Methods such as “self-coherence” allow for the retrieval of signals buried far below the noise level. In particular, at SNRs \( \leq 1 \) the noise may suppress high-order coefficients in the eclipse signal’s Fourier series. As a consequence the shape of the retrieved transit-curve becomes inherently linked with noise, as well as the overall morphology of the timeseries. Understanding the details of this process is key to fully recovering the shape and depth of the original signal. In this paper we have used Monte-Carlo simulations of very low SNR data, to investigate morphological changes of the original eclipse at low-SNRs. Furthermore, we demonstrate the full recovery of the eclipse transit-depth down to SNRs of 0.15 per spectral channel using calibration techniques described here. A clear understanding of methods such as self-coherence is essential for probing ever weaker atmospheric signatures such as the ones observed for extrasolar planets.
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