What Properties of Electromagnetic Radiation Can We Measure?

- Specific flux = Intensity (in ergs or photons) per unit area (or solid angle), time, wavelength (or frequency), e.g., $f_\lambda = 10^{-15}$ erg/cm$^2$/s/Å - a good spectroscopic unit
- It is usually integrated over some finite bandpass (as in photometry) or a spectral resolution element or a line
- It can be distributed on the sky (surface photometry, e.g., galaxies), or changing in time (variable sources)
- You can also measure the polarization parameters (photometry $\rightarrow$ polarimetry, spectroscopy $\rightarrow$ spectro-polarimetry); common in radio astronomy

Measuring Flux = Energy/(unit time)/(unit area)

Real detectors are sensitive over a finite range of $\lambda$ (or $\nu$). Fluxes are always measured over some finite bandpass.

Total energy flux:

$$ F = \int F_\nu(\nu) d\nu $$

Integral of $f_\nu$ over all frequencies

Units: erg s$^{-1}$ cm$^{-2}$ Hz$^{-1}$

A standard unit for specific flux (initially in radio, but now more common):

1 Jansky (Jy) = $10^{-23}$ erg s$^{-1}$ cm$^{-2}$ Hz$^{-1}$

$f_\nu$ is often called the flux density - to get the power, one integrates it over the bandwith, and multiplies by the area

(From P. Armitage)
Fluxes and Magnitudes

For historical reasons, fluxes in the optical and IR are measured in magnitudes: \( m = -2.5 \log_{10} F + \text{constant} \)

If \( F \) is the total flux, then \( m \) is the bolometric magnitude. Usually instead consider a finite bandpass, e.g., \( V \) band.

\[
m_V = -2.5 \log_{10} F + \text{constant}
\]

flux integrated over the range of wavelengths for this band

(From P. Armitage)

There are way, way too many photometric systems out there …

(Bandpass curves from Fukugita et al. 1995, PASP, 107, 945)
Using Magnitudes

Consider two stars, one of which is a hundred times fainter than the other in some waveband (say V).

\[ m_1 = -2.5 \log F_1 + \text{constant} \]
\[ m_2 = -2.5 \log(0.01 F_1) + \text{constant} \]
\[ = -2.5 \log(0.01) - 2.5 \log F_1 + \text{constant} \]
\[ = 5 - 2.5 \log F_1 + \text{constant} \]
\[ = 5 + m_1 \]

Source that is 100 times fainter in flux is five magnitudes fainter (larger number).

Faintest objects detectable with HST have magnitudes of \(~28\) in R/I bands. The sun has \(m_V = -26.75\) mag

(From P. Armitage)

Magnitude Zero Points

Alas, for the standard UBVRIJKL... system (and many others) the magnitude zero-point in any band is determined by the spectrum of Vega \(\neq \text{const!}\)

Vega calibration \(m = 0\): at \(\lambda = 5556: f_\lambda = 3.39 \times 10^{-6}\ \text{erg/cm}^2/\text{s/Å}\)

\(f_\lambda = 3.50 \times 10^{-20}\ \text{erg/cm}^2/\text{s/Hz}\)

\(N_\lambda = 948\ \text{photons/cm}^2/\text{s/Å}\)

A more logical system is AB magnitudes:

\[ AB_v = -2.5 \log f_v \ [\text{cgs}] - 48.60 \]

Magnitudes, A Formal Definition

\[ m = -2.5 \left[ \log \int d\lambda R(\lambda) f_\lambda - \log \int d\lambda R(\lambda) f_\lambda (\alpha \ \text{Lyrae}) \right] \]

\(c\), e.g.,

\[ U = -2.5 \log \int d\lambda R_u(\lambda) f_\lambda - 14.08 + c_U, \]

\[ B = -2.5 \log \int d\lambda R_B(\lambda) f_\lambda - 13.00 + c_B, \]

\[ V = -2.5 \log \int d\lambda R_V(\lambda) f_\lambda - 13.76 + c_V, \]

where the peak of the response function is normalized to unity, and \(c\) represents the magnitude of \(\alpha\ \text{Lyrae}\); \(c_U = 0.02, c_B = c_V = 0.03 \) (Johnson and Morgan 1953).

(From Fukugita et al. 1995)

Photometric Zero-Points (Visible)

<table>
<thead>
<tr>
<th>bandpass system</th>
<th>band</th>
<th>ref</th>
<th>(\lambda_{\text{eff}}) (Å)</th>
<th>FWHM (Å)</th>
<th>(\lambda_{\text{Vega}}) (Å)</th>
<th>(f_{\lambda_{\text{Vega}}}) ((\times 10^{-20}\ \text{erg/cm}^2/\text{s}))</th>
<th>(c(f_{\lambda_{\text{Vega}}})) ((\times 10^{-20}\ \text{erg/cm}^2/\text{Hz}))</th>
</tr>
</thead>
<tbody>
<tr>
<td>Johnson-Morgan</td>
<td>U5</td>
<td>Bauer 78</td>
<td>3652</td>
<td>526</td>
<td>3709</td>
<td>4.28</td>
<td>2617</td>
</tr>
<tr>
<td></td>
<td>B5</td>
<td>AS69</td>
<td>4448</td>
<td>1008</td>
<td>4383</td>
<td>6.19</td>
<td>4363</td>
</tr>
<tr>
<td></td>
<td>V</td>
<td>AS69</td>
<td>5555</td>
<td>822</td>
<td>5439</td>
<td>3.60</td>
<td>5437</td>
</tr>
<tr>
<td>Cousins</td>
<td>R1</td>
<td>Bessell 90</td>
<td>6588</td>
<td>1568</td>
<td>6410</td>
<td>2.15</td>
<td>6415</td>
</tr>
<tr>
<td></td>
<td>R2</td>
<td>Bessell 90</td>
<td>8050</td>
<td>1542</td>
<td>7977</td>
<td>1.11</td>
<td>7980</td>
</tr>
<tr>
<td>Johnson</td>
<td>R1</td>
<td>6930</td>
<td>3606</td>
<td>6688</td>
<td>1.87</td>
<td>6693</td>
<td></td>
</tr>
<tr>
<td></td>
<td>R2</td>
<td>8785</td>
<td>1706</td>
<td>8571</td>
<td>0.912</td>
<td>8545</td>
<td></td>
</tr>
<tr>
<td>SDSS</td>
<td>u'</td>
<td>3585</td>
<td>556</td>
<td>3594</td>
<td>3.67</td>
<td>3530</td>
<td></td>
</tr>
<tr>
<td></td>
<td>g'</td>
<td>4858</td>
<td>1207</td>
<td>4765</td>
<td>5.11</td>
<td>4748</td>
<td></td>
</tr>
<tr>
<td></td>
<td>r'</td>
<td>8290</td>
<td>1368</td>
<td>6285</td>
<td>2.40</td>
<td>6210</td>
<td></td>
</tr>
<tr>
<td></td>
<td>i'</td>
<td>7706</td>
<td>1547</td>
<td>7617</td>
<td>1.28</td>
<td>7623</td>
<td></td>
</tr>
<tr>
<td></td>
<td>z'</td>
<td>9222</td>
<td>1530</td>
<td>9123</td>
<td>0.783</td>
<td>9098</td>
<td></td>
</tr>
<tr>
<td>Tphot-Gunn</td>
<td>u</td>
<td>3536</td>
<td>412</td>
<td>3542</td>
<td>3.33</td>
<td>3519</td>
<td></td>
</tr>
<tr>
<td></td>
<td>g</td>
<td>4927</td>
<td>709</td>
<td>4888</td>
<td>4.84</td>
<td>4888</td>
<td></td>
</tr>
<tr>
<td></td>
<td>r</td>
<td>6538</td>
<td>893</td>
<td>6496</td>
<td>2.09</td>
<td>6498</td>
<td></td>
</tr>
</tbody>
</table>
(From Fukugita et al. 1995)
Defining effective wavelengths (and the corresponding bandpass averaged fluxes)

$$\lambda_{\text{eff}} = \frac{\int d\lambda \lambda R(\lambda)}{\int d\lambda R(\lambda)},$$

$$f^\text{eff}_{\lambda}(\alpha \text{ Lyr}) = \frac{\int d\lambda f_{\lambda}(\alpha \text{ Lyr}) R(\lambda)}{\int d\lambda R(\lambda)},$$

$$\lambda_{\text{eff}}(\alpha \text{ Lyr}) = \frac{\int d\lambda \lambda f_{\lambda}(\alpha \text{ Lyr}) R(\lambda)}{\int d\lambda f_{\lambda}(\alpha \text{ Lyr}) R(\lambda)},$$

$$f^\text{eff}_{\nu}(\alpha \text{ Lyr}) = \frac{\int d\nu f_{\nu}(\alpha \text{ Lyr}) R(\nu)}{\int d\nu R(\nu)},$$

$$\nu_{\text{eff}}(\alpha \text{ Lyr}) = \frac{\int d\nu \nu f_{\nu}(\alpha \text{ Lyr}) R(\nu)}{\int d\nu f_{\nu}(\alpha \text{ Lyr}) R(\nu)},$$

where $f_{\nu} = \lambda^2 f_{\lambda}/c$ and $R_{\nu} = R_{\lambda}$.

The Infrared Photometric Bands...

... where the atmospheric transmission windows are

![Atmospheric transmission graph](image)

### Infrared Bandpasses

<table>
<thead>
<tr>
<th>Filter name</th>
<th>$\lambda_{\text{iso}}$ ($\mu$m)</th>
<th>$\Delta \lambda$ ($\mu$m)</th>
<th>$F_{\lambda}$ (W m$^{-2}$ $\mu$m$^{-1}$)</th>
<th>$F_{\nu}$ (Jy)</th>
<th>$N_{\nu}$ (photons s$^{-1}$ m$^{-2}$ $\mu$m$^{-1}$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>V</td>
<td>0.5556$^d$</td>
<td>...</td>
<td>$3.44 \times 10^{-8}$</td>
<td>3540</td>
<td>$9.60 \times 10^{10}$</td>
</tr>
<tr>
<td>J</td>
<td>1.215</td>
<td>0.26</td>
<td>$3.31 \times 10^{-9}$</td>
<td>1630</td>
<td>$2.02 \times 10^{10}$</td>
</tr>
<tr>
<td>H</td>
<td>1.654</td>
<td>0.29</td>
<td>$1.15 \times 10^{-9}$</td>
<td>1050</td>
<td>$9.56 \times 10^{9}$</td>
</tr>
<tr>
<td>K$_s$</td>
<td>2.157</td>
<td>0.32</td>
<td>$4.30 \times 10^{-10}$</td>
<td>667</td>
<td>$4.66 \times 10^{9}$</td>
</tr>
<tr>
<td>K</td>
<td>2.179</td>
<td>0.41</td>
<td>$4.14 \times 10^{-10}$</td>
<td>655</td>
<td>$4.53 \times 10^{9}$</td>
</tr>
<tr>
<td>L</td>
<td>3.547</td>
<td>0.57</td>
<td>$6.59 \times 10^{-11}$</td>
<td>276</td>
<td>$1.17 \times 10^{9}$</td>
</tr>
<tr>
<td>L'</td>
<td>3.761</td>
<td>0.65</td>
<td>$5.26 \times 10^{-11}$</td>
<td>248</td>
<td>$9.94 \times 10^{8}$</td>
</tr>
<tr>
<td>M</td>
<td>4.769</td>
<td>0.45</td>
<td>$2.11 \times 10^{-11}$</td>
<td>160</td>
<td>$5.06 \times 10^{8}$</td>
</tr>
<tr>
<td>8.7</td>
<td>8.756</td>
<td>1.2</td>
<td>$1.96 \times 10^{-12}$</td>
<td>50.0</td>
<td>$8.62 \times 10^{7}$</td>
</tr>
<tr>
<td>N</td>
<td>10.472</td>
<td>5.19</td>
<td>$9.63 \times 10^{-13}$</td>
<td>35.2</td>
<td>$5.07 \times 10^{7}$</td>
</tr>
<tr>
<td>11.7</td>
<td>11.653</td>
<td>1.2</td>
<td>$6.31 \times 10^{-13}$</td>
<td>28.6</td>
<td>$3.69 \times 10^{7}$</td>
</tr>
<tr>
<td>Q</td>
<td>20.130</td>
<td>7.8</td>
<td>$7.18 \times 10^{-14}$</td>
<td>9.70</td>
<td>$7.26 \times 10^{6}$</td>
</tr>
</tbody>
</table>

Notes:

1. In $\mu$m
2. $F_{\lambda}(10^{-15}$ W cm$^{-2}$ $\mu$m$^{-1}$), $F_{\nu}(10^{-30}$ W cm$^{-2}$ Hz$^{-1}$) for a 0.03 magnitude star from Drelling and Bell, and Bell Vega models for adopted passbands.
3. Mag = - 2.5 log$_{10} F_{\nu}$ - 66.08 - ZP
**Colors From Magnitudes**

The color of an object is defined as the difference in the magnitude in each of two bandpasses: e.g. the \((B-V)\) color is: \(B-V = m_B - m_V\)

Stars radiate roughly as blackbodies, so the color reflects surface temperature.

Vega has \(T = 9500\) K, by definition color is zero.

(From P. Armitage)

---

**Apparent vs. Absolute Magnitudes**

The absolute magnitude is defined as the apparent mag. a source would have if it were at a distance of 10 pc:

\[
M = m + 5 - 5 \log d/\text{pc}
\]

It is a measure of the *luminosity* in some waveband.

For Sun: \(M_\odot = 5.47, M_\odot = 4.82, M_\odot = 4.74\)

Difference between the apparent magnitude \(m\) and the absolute magnitude \(M\) (any band) is a *measure of the distance* to the source

\[
m - M = 5 \log_{10} \left( \frac{d}{10 \text{ pc}} \right)
\]

Distance modulus

(From P. Armitage)
Why Do We Need This Mess?

Relative measurements are generally easier and more robust than the absolute ones; and often that is enough.

An example: the Color-Magnitude Diagram
The quantitative operational framework for studies of stellar physics, evolution, populations, distances …

Signal-to-Noise (S/N)

- Signal = \( R_s \cdot t \) (detected rate in e-/second)
- Consider the case where we count all the detected e- in a circular aperture with radius \( r \).

The Concept of Signal-to-Noise (S/N) or: How good is that measurement really?

- \( S/N = \text{signal/error} \) (If the noise is Gaussian, we speak of 3-\( \sigma \), 5- \( \sigma \), … detections. This translates into a probability that the detection is spurious.)
- For a counting process (e.g., photons), error = \( \sqrt{n} \), and thus \( S/N = n / \sqrt{n} = \sqrt{n} \) (“Poissonian noise”). This is the minimum possible error; there may be other sources of error (e.g., from the detector itself)
- If a source is seen against some back(fore)ground, then
  \[ \sigma^2_{\text{total}} = \sigma^2_{\text{signal}} + \sigma^2_{\text{background}} + \sigma^2_{\text{other}} \]

- Noise Sources:
  \[ \sqrt{R_s \cdot t} \quad \Rightarrow \quad \text{shot noise from source} \]
  \[ \sqrt{R_{\text{sky}} \cdot t \cdot \pi r^2} \quad \Rightarrow \quad \text{shot noise from sky in aperture} \]
  \[ \sqrt{RN^2 \cdot \pi r^2} \quad \Rightarrow \quad \text{readout noise in aperture} \]
  \[ \sqrt{[RN^2 + (0.5 \times \text{gain})^2] \cdot \pi r^2} \quad \Rightarrow \quad \text{more general RN} \]
  \[ \sqrt{\text{Dark} \cdot t \cdot \pi r^2} \quad \Rightarrow \quad \text{shot noise in dark current in aperture} \]

\( R_s = \text{e}^-/\text{sec} \) from the source
\( R_{\text{sky}} = \text{e}^-/\text{sec/pixel} \) from the sky
\( RN = \text{read noise (as if RN}^2 \text{ e}^- \text{ had been detected)} \)
\( \text{Dark} = \text{e}^-/\text{second/pixel} \)
S/N for object measured in aperture with radius \( r \): \( n_{\text{pix}}=\pi r^2 \)

\[
S/N = \left[ R \cdot t + R_{\text{sky}} \cdot t \cdot n_{\text{pix}} + \left( RN + \frac{\text{gain}}{2} \right) \cdot n_{\text{pix}} + \text{Dark} \cdot t \cdot n_{\text{pix}} \right]^{\frac{1}{2}}
\]

Noise from the dark current in aperture

Noise from sky electrons in aperture

Read noise in aperture

\[ \sqrt{(R \cdot t)^2} \]

All the noise terms added in quadrature

Note: always calculate in e-

\[ \text{Area of 1 pixel} = (\text{Scale})^2 \]

\section*{Sky Background}

Signal from the sky background is present in every pixel of the aperture. Because each instrument generally has a different pixel scale, the sky brightness is usually tabulated for a site in units of mag/arcsecond\(^2\).

\[
\text{(mag/}\text{arcsec}^2) \]

\begin{table}

<table>
<thead>
<tr>
<th>Linear age (days)</th>
<th>U</th>
<th>B</th>
<th>V</th>
<th>R</th>
<th>I</th>
</tr>
</thead>
<tbody>
<tr>
<td>0</td>
<td>22.0</td>
<td>22.7</td>
<td>21.8</td>
<td>20.9</td>
<td>19.9</td>
</tr>
<tr>
<td>3</td>
<td>21.5</td>
<td>22.4</td>
<td>21.7</td>
<td>20.8</td>
<td>19.9</td>
</tr>
<tr>
<td>7</td>
<td>19.9</td>
<td>21.6</td>
<td>21.4</td>
<td>20.6</td>
<td>19.7</td>
</tr>
<tr>
<td>10</td>
<td>18.5</td>
<td>20.7</td>
<td>20.7</td>
<td>20.3</td>
<td>19.5</td>
</tr>
<tr>
<td>14</td>
<td>17.0</td>
<td>19.5</td>
<td>20.0</td>
<td>19.9</td>
<td>19.2</td>
</tr>
</tbody>
</table>

Scale = \( \sigma/\text{pix} \)

\[
\text{Area of 1 pixel} = (\text{Scale})^2
\]

this is the ratio of flux/pix to flux/\( \text{arcsec}^2 \)

In magnitudes:

\[
I_{\text{pix}} = I \cdot \text{Scale}^2
\]

\[ \Rightarrow \text{Intensity} (e^-/\text{sec}) \]

\[ -2.5 \log(I_{\text{pix}}) = -2.5[\log(I_0) + \log(\text{Scale}^2)] \]

\[ m_{\text{pix}} = m_0 - 2.5\log(\text{Scale}^2) \]

(for LRIS - R : add 3.303 mag)

and

\[
R_{\text{sky}} (m_{\text{pix}}) = R(m = 20) \times 10^{0.4 - m_{\text{pix}}}
\]

Example, LRIS in the R-band:

\[
R_{\text{sky}} = 1890 \times 10^{0.4(20-24.21)} = 39.1 \text{ e}^-/\text{pix/sec}
\]

\[
\sqrt{R_{\text{sky}}} = 6.35 \text{e}^-/\text{pix/sec} = \text{RN in just 1 second}
\]

\section*{Side Issue: S/N \( \leftrightarrow \) \( \delta \)mag}

\[
m \pm \delta(m) = c_0 - 2.5\log(S \pm N)
\]

\[
= c_0 - 2.5\log[S(1 \pm \frac{r}{3})]
\]

\[
= c_0 - 2.5\log(S) - 2.5\log(1 \pm \frac{r}{3})
\]

\[
\delta(m) = 2.5\log(1 + \frac{1}{7\pi})
\]

\[
= \frac{2.5}{2.3} \left[ \frac{r}{3} - \frac{1}{2} \left( \frac{r}{3} \right)^2 + \frac{1}{2} \left( \frac{r}{3} \right)^3 - \ldots \right]
\]

\[
= 1.087 \left( \frac{r}{3} \right) \text{ Fractional error}
\]

This is the basis of people referring to \( \pm 0.02 \) mag error as “2%”
S/N - some limiting cases. Let’s assume CCD with Dark=0, well sampled read noise.

\[
R_t = \frac{R_s \cdot t + R_{\text{sky}} \cdot t \cdot n_{\text{pix}} + (RN)^2 \cdot n_{\text{pix}}}{\sqrt{R_s \cdot t + R_{\text{sky}} \cdot t \cdot n_{\text{pix}} + (RN)^2 \cdot n_{\text{pix}}}}^{1/2}
\]

Bright Sources: \((R_s \cdot t)^{1/2}\) dominates noise term

\[
S/N = \frac{R_t}{\sqrt{R_s \cdot t \cdot t}} \propto t^{1/2}
\]

Sky Limited \((\sqrt{R_{\text{sky}} \cdot t} > 3 \times RN)\): \(S/N \propto \frac{R_t}{\sqrt{n_{\text{pix}} R_{\text{sky}} \cdot t}} \propto \sqrt{t}
\]

Note: seeing comes in with \(n_{\text{pix}}\) term

### Photometry With An Imaging Array

- Aperture Photometry: some modern ref’s
  - DaCosta, 1992, ASP Conf Ser 23
  - Stetson, 1990, PASP, 102, 932

### Aperture Photometry

\[
I = \sum_{ij} I_{ij} - n_{\text{pix}} \times \text{sky/pixel}
\]

Total counts in aperture from source

Number of pixels in aperture

Counts in each pixel in aperture

\[
m = c_0 - 2.5 \log(I)
\]
Aperture Photometry

- What do you need?
  - Source center
  - Sky value
  - Aperture radius

Marginal Sums

- With noise and multiple sources you have to decide what is a source and to isolate sources.

Centers

- The usual approach is to use "marginal sums".

\[ \rho_{x_i} = \sum_j I_{ij} \quad \text{Sum along columns} \]

- Find peaks: use \( \partial \rho_x / \partial x \) zeros
- Isolate peaks: use "symmetry cleaning"
  1. Find peak
  2. compare pairs of points equidistant from center
  3. if \( I_{\text{left}} >> I_{\text{right}} \), set \( I_{\text{left}} = I_{\text{right}} \)
- Finding centers: Intensity-weighted centroid

\[
x_{\text{center}} = \frac{\sum_i \rho_{x_i} x_i}{\sum_i \rho_{x_i}}; \quad \sigma^2 = \frac{\sum_i \rho_i x_i^2}{\sum_i \rho_i} - x_i^2
\]
• Alternative for centers: Gaussian fit to $\rho$:

$$\rho_i = \text{background} + h \cdot e^{-\frac{(x_i-x_c)^2}{2\sigma^2}}$$

• DAOPHOT FIND algorithm uses marginal sums in subrasters, symmetry cleaning, reraster and Gaussian fit.

Sky

• To determine the sky, typically use a local annulus, evaluate the distribution of counts in pixels in a way to reject the bias toward higher-than-background values.
• Remember the 3 Ms.

Sky From Minmax Rejection

Because essentially all deviations from the sky are positive counts (stars and galaxies), the mode is the best approximation to the sky.
Aperture size and growth curves

- First, it is VERY hard to measure the total light as some light is scattered to very large radius. Perhaps you have most of the light within this radius.

Radial intensity distribution for a bright, isolated star.

Radius from center in pixels

Radial profile with neighbors

- Neighbors OK in sky annulus (mode), trouble in star apertures

One Approach Is To Use “Growth Curves”

- Idea is to use a small aperture (highest S against background and smaller chance of contamination) for everything and determine a correction to larger radii based on several relatively isolates, relatively bright stars in a frame.

- Note! This assumes a linear response so that all point sources have the same fraction of light within a given radius.

- Stetson, 1990, PASP, 102, 932
Aperture Photometry Summary

1. Identify brightness peaks

2. \[ \sum_{i \text{xy}} (s \cdot \text{aperture area}) \]

   Use small aperture

3. Add in \"aperture correction\" determined from bright, isolated stars

Easy, fast, works well except for the case of overlapping images

Growth Curves

<table>
<thead>
<tr>
<th>Aperture</th>
<th>2</th>
<th>3</th>
<th>4</th>
<th>5</th>
<th>6</th>
<th>7</th>
<th>8</th>
</tr>
</thead>
<tbody>
<tr>
<td>Star#1</td>
<td>0.43</td>
<td>0.31</td>
<td>0.17</td>
<td>0.09</td>
<td>0.05</td>
<td>0.02</td>
<td>0.00</td>
</tr>
<tr>
<td>Star#2</td>
<td>0.42</td>
<td>0.33</td>
<td>0.19</td>
<td>0.08</td>
<td>0.21</td>
<td>0.11</td>
<td>0.04</td>
</tr>
<tr>
<td>Star#3</td>
<td>0.43</td>
<td>0.32</td>
<td>0.18</td>
<td>0.10</td>
<td>0.06</td>
<td>0.02</td>
<td>-0.01</td>
</tr>
<tr>
<td>Star#4</td>
<td>0.44</td>
<td>0.33</td>
<td>0.18</td>
<td>0.22</td>
<td>0.14</td>
<td>0.12</td>
<td>0.14</td>
</tr>
<tr>
<td>Star#5</td>
<td>0.42</td>
<td>0.32</td>
<td>0.18</td>
<td>0.09</td>
<td>0.19</td>
<td>0.21</td>
<td>0.19</td>
</tr>
<tr>
<td>Star#6</td>
<td>0.41</td>
<td>0.33</td>
<td>0.19</td>
<td>0.10</td>
<td>0.05</td>
<td>0.30</td>
<td>0.12</td>
</tr>
<tr>
<td>cMean</td>
<td>-0.430</td>
<td>0.324</td>
<td>0.184</td>
<td>0.094</td>
<td>0.057</td>
<td>0.02</td>
<td>0.00</td>
</tr>
</tbody>
</table>

Sum of these is the total aperture correction to be added to magnitude measured in aperture 1
Crowded-field Photometry

- As was assumed for aperture corrections, all point sources have the same PSF (linear detector)
- PSF fitting allows for an optimal S/N weighting
- Various codes have been written that do:
  1. Automatic star finding
  2. Construction of PSF
  3. Fitting of PSF to (multiple) stars
- Many programs exist: DAOPHOT, ROMAPHOT, DOPHOT, STARMAN, …
- DAOPHOT is perhaps the most useful one: Stetson, 1987, PASP, 99, 191

To construct a PSF:
1. Choose unsaturated, relatively isolated stars
2. If PSF varies over the frame, sample the full field
3. Make 1st iteration of the PSF
4. Subtract psf-star neighbors
5. Make another PSF
- PSF can be represented either as a table of numbers, or as a fitting function (e.g., a Gaussian + power-law or exponential wings, etc.), or as a combination

Photometric Calibration

- The photometric standard systems have tended to be zero-pointed arbitrarily. Vega is the most widely used and was original defined with $V=0$ and all colors = 0.
- The AB scale (Oke, 1974, ApJS, 27, 21) is a physical-unit-based scale with:

$$m(AB) = -2.5 \log(f) - 48.60$$

where $f$ is monochromatic flux is in units of erg/sec/cm²/Hz. Objects with constant flux/unit frequency interval have zero color on this scale.
### Photometric calibration

1. **Instrumental magnitudes**
   
   
   $m = c_0 - 2.5 \log(I \cdot t)$
   
   $= c_0 - 2.5 \log(I) - 2.5 \log(t)$
   
   $m_{\text{instrumental}}$

   - Extinction coefficients:
     - Increase with decreasing wavelength
     - Can vary by 50% over time and by some amount during a night
     - Are measured by observing standards at a range of airmass during the night

### Photometric Calibration

- To convert to a *standard* magnitude you need to observe some standard stars and solve for the constants in an equation like:

  $m_{\text{inst}} = M + c_0 + c_1 X + c_3(\text{color}) + c_3(\text{UT}) + c_4(\text{color})^2 + ...$

  - $m_{\text{inst}}$ (Instr mag)
  - $M$ (Stnd mag)
  - $c_0 + c_1 X + c_3(\text{color}) + c_3(\text{UT}) + c_4(\text{color})^2 + ...$
  - $zpt$
  - $\text{airmass}$
  - $\text{Color term}$

- **Extinction coeff** (mag/airmass)

### Photometric Standards

- Fields containing several well measured stars of similar brightness and a big range in color. The blue stars are the hard ones to find and several fields are center on PG sources.
- Measure the fields over at least the airmass range of your program objects and intersperse standard field observations throughout the night.
Photometric Calibration: Standard Stars

Magnitudes of Vega (or other systems primary flux standards) are transferred to many other, secondary standards. They are observed along with your main science targets, and processed in the same way.

Alas, Even The “Same” Photometric Systems Are Seldom Really The Same …

This Generates Color Terms …

… From mismatches between the effective bandpasses of your filter system and those of the standard system. Objects with different spectral shapes have different offsets:

A photometric system is thus effectively (operationally) defined by a set of standard stars - since the actual bandpasses may not be well known.
Surface Photometry

- The way to quantify the 2-dimensional distribution of light, e.g., in galaxies
- Many references, e.g.,
  - Davis et al., AJ, 90, 1985
- Could fit (or find) isophotes, and the most common procedure is to fit elliptical isophotes.
- Isophotal parameters are: surface brightness itself ($\mu$), $x_{\text{center}}$, $y_{\text{center}}$, ellipticity ($\varepsilon$), position angle (PA), the enclosed magnitude ($m$), and sometimes higher order shape terms, all as functions of radius ($r$) or semi-major axis ($a$).

Start with guesses for $x_c$, $y_c$, $R$, $\varepsilon$ and p.a., then compare the ellipse with real data all along the ellipse (all $\theta$ values).

Fit the $\Delta I - \theta$ plot and iterate on $x_c$, $y_c$, p.a., and $\varepsilon$ to minimize the coefficients in an expression like:

$$I(\theta) = I_0 + A_1 \sin(\theta) + B_1 \cos(\theta) + A_2 \sin(2\theta) + B_2 \cos(2\theta)$$

Changes to $x_c$ and $y_c$ mostly affect $A_1$, $B_1$, p.a.

Good isophote

$$\text{mag} = c_0 - 2.5(\text{cnts}_\text{aper} - \pi r^2 \text{sky})$$

Typically working with much larger apertures
- prone to contamination
- $\varepsilon$ determination even more critical
- often want to know more than total brightness
More specifically, iterate the following:

\[ \Delta(\text{major axis center}) = \frac{-B_1}{I'} \]
\[ \Delta(\text{minor axis center}) = \frac{-A_1(1 - \epsilon)}{I'} \]
\[ \Delta(\epsilon) = \frac{-2B_2(1 - \epsilon)}{a_0I'} \]
\[ \Delta(\text{p.a.}) = \frac{2A_2(1 - \epsilon)}{a_0I'[(1 - \epsilon)^2 - 1]} \]

where:
\[ I' = \frac{\delta I}{\delta R} |_{a_0} \]

- After finding the best-fitting elliptical isophotes, the residuals are often interesting. Fit:

\[ I = I_0 + A_n\sin(n\phi) + B_n\cos(n\phi) \]

already minimized \( n=1 \) and \( n=2 \), \( n=3 \) is usually not significant, but:

- \( B_4 \) is negative for “boxy” isophotes
- \( B_4 \) positive for “disky” isophotes

---

Examples of Surface Photometry of Ellipticals

- Major axis surf. brightness profiles
- Isophotal shape and orientation prof’s

One can also measure the deviations from the elliptical shape (boxy/disky)

---

Disky and Boxy Elliptical Isophotes

- Calculate mean and RMS pixel intensity for annulus, toss any values above mean + nRMS

---

Examples for boxy and disky isophotes from Bender et al. (1988)
From your isophotal fits, you can then construct the best 2-dimensional elliptical model for the light distribution.

... And subtract it from the image to reveal any deviations from the assumed elliptical symmetry.
Panoramic Imaging

- Generally used to study populations of sources (e.g., faint galaxy counts; star clusters; etc.)
- Commonly done in (wide-area) surveys
- Image (pixels) → catalogs of objects and their measured properties
- If done properly, essentially all information is extracted into a more useful form; but not always…
- Key steps:
  1. Object finding (there is always some spatial “filter”)
  2. Object measurements / parametrization
  3. Object classification (e.g., star/galaxy)

- One very common program for panoramic photometry is **Sextractor**
- Not for good a detailed surface photometry, but good for classification and rough photometric and structural parameter derivation for large fields.

Steps:
1. Background map (sky determination)
2. Identification of objects (thresholding)
3. Deblending
4. Photometry
5. Shape analysis

---

Image (De-)Blending and Segmentation

**Thresholding** is an alternative to **peak finding**. Look for contiguous pixels above a threshold value.
- User sets area, threshold value.
- Sometimes combine with a smoothing filter
**Deblending** based on multiple-pass thresholding

---

Sextractor Flowchart
Star-Galaxy Separation

- Galaxies are resolved, stars are not
- All methods use various approaches to comparing the amount of light at large and small radii.

Star-Galaxy Separation

- Important, since for most studies you want either stars (or quasars), or galaxies; and then the depth to which a reliable classification can be done is the effective limiting depth of your catalog - not the detection depth
  - There is generally more to measure for a non-PSF object
- You’d like to have an automated and objective process, with some estimate of the accuracy as a $f(mag)$
  - Generally classification fails at the faint end
- Most methods use some measures of light concentration vs. magnitude (perhaps more than one), and/or some measure of the PSF fit quality (e.g., $\chi^2$)
- For more advanced approaches, use some machine learning method, e.g., neural nets or decision trees

Typical Parameter Space for S/G Classif.

- Lots of talk about neural-net algorithms, but in the end it is a moment analysis.
- “Stellarity”. Typically test it with artificial stars and find it is very good to some limiting magnitude.
More S/G Classification Parameter Spaces: Normalized By The Stellar Locus

Then a set of such parameters can be fed into an automated classifier (ANN, DT, …) which can be trained with a “ground truth” sample.

Automated Star-Galaxy Classification: Artificial Neural Nets (ANN)

Input:
- various image shape parameters.

Output:
- Star, \( p(s) \)
- Galaxy, \( p(g) \)
- Other, \( p(o) \)

(Odewahn et al. 1992)

Automated Star-Galaxy Classification: Decision Trees (DTs)

Automated Star-Galaxy Classification: Unsupervised Classifiers

No training data set - the program decides on the number of classes present in the data, and partitions the data set accordingly.

An example:
- AutoClass (Cheeseman et al.)
  - Uses Bayesian approach in machine learning (ML).

This application from DPOSS (Weir et al. 1995)
Summary of the Key Points

- Photometry = flux measurement over a finite bandpass, could be integral (the entire object) or resolved (surface photometry)
- The arcana of the magnitudes and many different photometric systems …
- The S/N computation - many sources of noise
- Issues in the photometry with an imaging array: object finding and centering, sky determination, aperture photometry, PSF fitting, calibrations …
- Surface photometry and isophote fitting lore …
- Star-galaxy classification in panoramic imaging
- The effects of the seeing: flux measurements and morphology

Seeing - A Key Issue

- The “image size”, given by a convolution of the atmospheric turbulence, and instrument optics
- Affects two important aspects of imaging photometry:
  1. S/N: because a larger image spreads the available signal over more noise-contributing pixels
     - Defines the detection limit
     - Should have pixel scale matched to the expected image quality, to at least Nyquist sampling
  2. Morphological resolution: as high spatial frequency information is lost
     - Defines the star-galaxy classification limit
     - Can be recovered only partly by image deconvolutions, which require a good S/N, sampling, and a well measured PSF
- A tricky issue is how to combine data obtained in different seeing conditions

Example of seeing variations in the ground-based data (from the PQ survey)