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Subject: Photometry

Second Semester

Photometry

Many people are interested in astronomy because it is visually exciting. The many marvelous pictures of celestial objects taken using large telescopes on the ground or in space are certainly the most visible manifestation of modern research astronomy. However, to do real science, one needs far more than pictures. Pictures are needed as a first step in classifying objects based on their appearance (morphology). To proceed past this initial stage of investigation, we need quantitative information- i.e. measurements of the properties of the objects. Observational astronomy becomes science only when we can start to answer questions quantitatively: How far away is that object? How much energy does it emit? How hot is it?

The most fundamental information we can measure about celestial objects past our solar system is the amount of energy, in the form of electromagnetic radiation, that we receive from that object. This quantity we will call the flux. The science of measuring the flux we receive from celestial objects is called photometry. As we will see, photometry usually refers to measurements of flux over broad wavelength bands of radiation. Measurement of flux, when coupled with some estimate of the distance to an object, can give us information on the total energy output of the object (its luminosity), the object's temperature, and the object's size and other physical properties. If we can measure the flux in small wavelength intervals, we start to see that the flux is often quite irregular on small wavelength scales. This is due to the interaction of light with the atoms and molecules in the object. These "bumps and wiggles" in the flux as a function of wavelength are like fingerprints. They can tell us lots about the object– what it is made of, how the object is moving and rotating, the pressure and ionization of the material in the object, etc. The observation of these bumps and wiggles is called spectroscopy. A combination of spectroscopy, meaning good wavelength resolution, and photometry, meaning good flux calibration, is called spectrophotometry. Obviously, there is more information in a spectrophotometric scan o f an object compared with photometry spanning the same wavelength range. Why would one do low wavelength resolution photometry rather than higher resolution spectrophotometry or spectroscopy, given the fact that a spectrum gives much more information than photometry? As we will see, it is much easier to make photometric observations of faint objects than it is to make spectroscopic observations of the same object. With any given telescope, one can always do photometry of much fainter objects than one can do spectroscopy of. On a practical note, the equipment required for CCD imaging photometry is much simpler and cheaper than that needed for spectroscopy. With low cost CCDs now readily available, even small telescopes can do useful photometric observations, particularly monitoring variable objects.

Visible EMR

Almost all astronomical information from beyond the Solar System comes to us from some form of electromagnetic radiation (EMR). (Can you think of any sources of information from beyond the Solar system that do not involve EMR in some form?) We can now detect and study EMR over a range of wavelength or, equivalently, photon energy, covering a range from short wavelength, high photon energy gamma rays to long wavelength, low energy radio photons. Out of all this vast range of wavelengths, our eyes are sensitive to a tiny slice of wavelengths- roughly from 4500 to 6500 $\overset{\circ}{A}$. The range of wavelengths our eyes are sensitive to is called the visible wavelength range. We will define a wavelength region reaching somewhat shorter (to about 3200 $\overset{\circ}{A}$) to somewhat longer (about 10,000 $\overset{\circ}{A}$) than the visible as the optical part of the spectrum. (Note: Physicists measure optical wavelengths in nanometers (ηm). Astronomers tend to use Angstroms. 1 $\overset{\circ}{A}$ = 10^{-10}m = 0.1 ηm. Thus, a physicist would say the optical region is from 320 to 1000 ηm.) All EMR comes in discrete lumps called photons. A photon has a definite energy and frequency or wavelength. The relation between photon energy (E_{ph}) and photon frequency (ν) is given by:

$$E_{\mathrm{ph}} = h\nu$$

or, since c= λν

$$E_{\mathrm{ph}} = \frac{\mathrm{hc}}{\lambda}$$

where h is Planck's constant and λ is the wavelength, and c is the speed of light. The energy of visible photons is around a few eV (electron volts). 1 eV = 1.602E−19 Joules or 1.602E−12 erg).

we have direct sensory experience with this region. Today, virtually no research level astronomical observations are made with the human eye as the primary detecting device. However, the fact that we see in visible light has driven a vast technological effort over the past century or two to develop devices - photographic emulsions, photomultipliers, video cameras, various solid state images that detect and record visible light. The second overriding reason to study optical light is that the Earths atmosphere is at least partially transparent to this region of the spectrum- otherwise you couldn't see the stars at night (or the Sun during the day)!

Much of the EMR spectrum is blocked by the atmosphere, and can only be studied using telescopes placed above the atmosphere. Only in the optical and radio regions of the spectrum are there large atmospheric windows – portions of the EMR spectrum for which the atmosphere is at least partially transparent- which allow us to study the universe. Study of wavelengths that don't penetrate the atmosphere using telescopes and detectors out in space- which we will call space astronomy - is an extremely important part of modern astronomy which has fantastically enriched our view of the universe over the past few decades. However, space astronomy is very expensive and difficult to carry out.
In purely astronomical terms, the optical portion of the spectrum is important because most stars and galaxies emit a significant fraction of their energy in this part of the spectrum. (This is not true for objects significantly colder than stars - e.g. planets, interstellar dust and molecular clouds, which emit in the infrared or at longer wavelengths - or significantly hotter- e. g. ionized gas clouds or neutron stars, which emit in the ultraviolet and x-ray regions of the spectrum. Now, you may ask yourself why I included planets along with dust clouds in the above sentence. The reason is that the bright visible light you are seeing from planets is reflected sunlight and not light emitted by planets itself). Another reason the optical region is important is that many molecules and atoms have diagnostic electronic transitions in the optical wavelength region.

Imaging, Spectrophotometry and Photometry

The goal of the observational astronomer is to make measurements of the EMR from celestial objects with as much detail, or finest resolution, possible. There are several different types of detail that we want to observe. These include angular detail, wavelength detail, and time detail.
The perfect astronomical observing system would tell us the amount of radiation, as a function of wavelength, from the entire sky in arbitrarily small angular slices. Such a system does not exist! We are always limited in angular and wavelength coverage, and limited in resolution in angle and wavelength. If we want good information about the wavelength distribution of EMR from an object (spectroscopy or spectrophotometry) we may have to give up angular detail. If we want good angular resolution over a wide area of sky (imaging) we usually have to give up wavelength resolution or coverage. The ideal goal of spectrophotometry is to obtain the spectral energy distribution (SED) of celestial objects, or how the energy from the object is distributed in wavelength. We want to

measure the amount of power received by an observer outside the Earth's atmosphere, or energy per second, per unit area, per unit wavelength or frequency interval. Units of spectral flux (in cgs) look like:

$$f_\lambda = \text{erg s}{-}1\text{cm}^{-2}\text{A}^{\circ -1}$$

if we measure per unit wavelength interval, or

$$f_\nu = \text{erg s}{-}1\text{cm}^{-2}\text{HZ}^{-1}$$

To get true spectrophotometry, we must use some sort of dispersing element (diffraction grating or prism) that spreads the light out in wavelength, so that we can measure the amount of light in small wavelength intervals. Now, this obviously dilutes the light. Thus, compared to imaging, spectrophotometry requires a larger telescope or is limited to relatively bright objects. Spectrophotometry also requires a spectrograph, a piece of equipment to spread out the light. Good research grade spectrographs are complicated and expensive pieces of equipment.

Instead of using a dispersing element to define which wavelengths we are measuring, we can use filters that pass only certain wavelengths of light. If we put a filter in front of a CCD camera, we obtain an image using just the wavelengths passed by the filter. We do not spread out the light in wavelength. If we use a filter with a large bandpass (broadband filter), then we have much more light in the image than in a single wavelength interval in spectrophotometry. Thus, a given telescope can measure the brightness of an object through a filter to far fainter limits than the same telescope could do spectrophotometry, at the trade off, of course, of less information on the distribution of flux with wavelength. Filters typically have resolutions (here $\Delta\lambda$ is the full width at half maximum or FWHM of the filter bandpass) of $\lambda / \Delta\lambda$ of 5 to 20 or so. Filters will be discussed in more detail in a later chapter. Thus you can think of filter photometry as very low resolution spectrophotometry. We sometimes take images with no filter. In this case, the wavelengths imaged are set by the detector wavelength sensitivity, the atmosphere transmission, and the transmission and reflectivity of the optics in the telescope. If we image without a filter we get no information about the color or SED of objects. Another problem with using no filter is that the wavelength range imaged is very large, and atmospheric refraction (discussed later) can degrade the image quality. Filter photometry, or just photometry, is easier to do than spectrophotometry, as the equipment required is just a gizmo for holding filters in front of the detector and a detector (which is now usually a CCD

camera). A substantial fraction of time on optical research telescopes around the world is devoted to CCD photometry.
So let's say you want to know the spectral flux of a certain star in at a particular wavelength, with a wavelength region defined by a filter. How does one go about doing this? Well, you might think you point the telescope at the star, measure the number of counts (think of counts as photons for now) that the detector measures per second, then find the energy of the counts detected (from their average wavelength), and then figure out the energy received from the star. Well, that’s a start, but as we will see it’s hard, if not impossible, to go directly from the counts in the detector to a precise spectral flux! The first obvious complication is that our detector does not detect every single photon, so we must correct the measured counts for this to get photons. If you measure the same star with the same detector but a bigger telescope, you will get more photons per unit time. Obviously, the flux of the star cannot depend on which telescope we use to measure it! Dealing with various telescope sizes sounds simple- simply divide by the collecting area of the telescope. Well, what is the collecting area of the telescope? For a refractor its just the area of the lens, but for a mirror, you must take into account not only the area of the mirror, but also the light lost due to the fact that the secondary mirror and its support structure blocks some of the light. That’s not all you have to worry about- telescope mirrors are exposed to the outside air. They get covered with dust, and the occasional bird droppings and insect infestations. The aluminum coating that provides the reflectivity (coated over the glass that holds the optical figure) gets corroded by chemicals in the air and loses reflectivity over time (and even freshly coated aluminum does not have 100% reflectivity). The aluminum has a reflectivity that varies somewhat with wavelength. Any glass in the system through which light passes (glass covering over the CCD or, for some telescopes, correctors or reimaging optics) absorbs some light, always a different amount at each wavelength. How the heck can we hope to measure the amount of light blocked by dust or the reflectivity and transmission of the optics in our telescope? Even if we could, we still have to worry about the effects of the Earth’s atmosphere. The atmosphere absorbs some fraction of the light from all celestial objects. As we will see later, the amount of light absorbed is different for different wavelengths, and also changes with time. The dimming of light in its passage through the atmosphere is called atmospheric extinction. Reading the above list of things that mess up the flux we measure from a star, you might think it impossible to get the accurate spectral flux from any star. Well, it is extremely difficult, but not impossible to get. So, how do we actually measure the spectral flux of a star? The key idea is that we measure the flux of the object that we want to know about and also

measure the flux of a set of stars (called standard stars) whose spectral flux has been carefully measured.
So, how does this help? By measuring our object and then measuring the standard star, we can get the flux of our star as a fraction of the standard star flux.

The Magnitude and Color System

Magnitudes

Optical astronomers almost always use something called the (astronomical) magnitude system to talk about several different kinds of measurements, such as the observed brightnesses (power fluxes, or energy received per unit time per unit area) of stars and the luminosity (total power output in EMR) of stars. The historical roots of the magnitude system go way back to the first star catalog, compiled by a Greek named Hipparchus some 2200 years ago. Hipparchus divided the stars into six brightness classes, and he called the stars that appeared brightest (to the naked eye, of course, there being no telescopes back then) first magnitude stars, and the faintest visible stars the sixth magnitude stars. Much later, when astronomers were able to make more exact measurements of the brightnesses of stars, they found that the Hipparchus magnitude scale was roughly logarithmic. That is, each magnitude step corresponded to a fixed brightness ratio or factor. The first magnitude stars are roughly 2.5 times as bright as the second magnitude stars, the second magnitude are roughly 2.5 times as bright as the third magnitude stars etc.
we have two stars, with flux f_1 and f_2. We can define the magnitude difference between the stars as:

$$\mathrm{m1} - \mathrm{m2} = -2.5\log_{10}(f_1/f_2) \qquad (1)$$

Clearly, if the flux ratio is 100, the magnitude difference is 5. Equation (1) is the fundamental equation needed to define and deal with magnitudes.
Note that we can rearrange the equation to give the flux ratio if the magnitude difference is known:

$$f_1/f_2 = 10^{-0.4(\mathrm{m1}-\mathrm{m2})} \qquad (2)$$

The most common use for magnitudes is for expressing the apparent brightness of stars. To give a definite number for a magnitude of a star (instead of just the magnitude difference between pairs of stars), we must pick a starting place, or zero point, for the magnitude system. we pick the star Vega, and say it has magnitude of 0.00. Then the magnitude of any other star is simply related to the flux ratio of that star and Vega as follows:

$$m1 = -2.5\log_{10} (f_1/f_{Vega}) \qquad (3)$$

Apparent Magnitudes

As early as the second century B. C., Hipparchos divided the visible stars into six classes according to their apparent brightness. The first class contained the brightest stars and the sixth the faintest ones still visible to the naked eye. The response of the human eye to the brightness of light is not linear. If the flux densities of three stars are in the proportion 1:10:100, the brightness difference of the first and second star seems to be equal to the difference of the second and third star. Equal brightness ratios correspond to equal apparent brightness differences: the human perception of brightness is logarithmic. The rather vague classification of Hipparchos was replaced in 1856 by Norman R. Pogson. The new, more accurate classification followed the old one as closely as possible, resulting in another of those illogical definitions typical of astronomy. Since a star of the first class is about one hundred times brighter than a star of the sixth class, Pogson defined the ratio of the brightnesses of classes *n* and *n*+1 as $\sqrt[5]{100}$= 2.512. The brightness class or *magnitude* can be defined accurately in terms of the observed flux density F ([F] =W m^{-2}). We decide that the magnitude 0 corresponds to some preselected flux density $F0$. All other magnitudes are then defined by the equation

$$m = -2.5 \log \frac{f}{f_\circ} \qquad (4)$$

Magnitudes are dimensionless quantities, but to remind us that a certain value is a magnitude, we can write it, for example, as 5mag or 5^{m}.

Absolute Magnitudes

Thus far we have discussed only apparent magnitudes. They do not tell us anything about the true brightness of stars, since the distances differ. A quantity measuring the intrinsic brightness of a star is the *absolute magnitude*. It is defined as the apparent magnitude at a distance of 10 parsecs from the star . We shall now derive an equation which relates the apparent magnitude *m*, the absolute magnitude *M* and the distance *r*. Because the flux emanating from a star into a solid angle *ω* has, at a distance *r*, spread over an area *ωr*2, the flux density is inversely proportional to the distance squared. Therefore the ratio of the flux density at a distance *r, F(r),* to the flux density at a distance of 10 parsecs, *F(*10*),* is

$$\frac{f_r}{f_{10}} = \left(\frac{10\ \mathrm{pc}}{r}\right)^2 \qquad (5)$$

Thus the difference of magnitudes at *r* and 10 pc, or the *distance modulus m−M*, is

$$m - M = -2.5\log\left(\frac{10\ \mathrm{pc}}{r}\right)^2 \qquad (6)$$

or

$$m - M = 5\log\left(\frac{r}{10\mathrm{pc}}\right) \qquad (7)$$

Bolometric Magnitudes

The flux of any object varies with wavelength. To measure all the EMR from a body, we would have to observe at all wavelengths of EMR, from gamma rays to the longest radio waves. Quantities integrated over all wavelengths (or at least over all wavelengths where the object emits significant radiation) are called bolometric quantities, e.g. the bolometric luminosity of the Sun is the total power put out by the Sun in all wavelengths of EMR. Bolometric magnitudes are difficult to actually measure. The object must be observed with a number of different telescopes and detectors- e.g. ground based telescopes for the optical portion of the spectrum, satellite telescopes for the ultraviolet and x rays, which don’t penetrate the atmosphere, ground or space telescopes for the infrared, space telescopes for the very short radio (mm and sub mm range) and ground based radio telescopes for the longer radio waves. The wavelength of peak emission is of course related to the effective temperature of the star by Stefan- Boltzmann law. The wavelength of the peak flux, for most stars, is in or near the visible region of the spectrum, Fortunately, most stars emit the vast majority of their total power within a reasonable interval in wavelength around the wavelength of their peak emission. This is less true for some other objects, for example quasars and other active galactic nuclei, which can emit significant energy over a very wide range of wavelengths.

Bolometric quantities are important to the theorist, as they represent the total amount of power output from an astronomical object. However, observations must be limited to certain wavelength regions, either by the atmosphere, or by the detectors used. The optical region is that region limited by the atmosphere on the short wavelength. Within the optical region, we usually further limit the wavelengths observed by use of filters. Filters are optical components that only allow certain wavelengths to pass through them.

Colors

Although filters will be discussed in more detail later, let us introduce one filter system so that we can discuss the idea of colors. One widely used filter system in the optical region of the spectrum is called the UBV system. The letters correspond to different filters: U for ultraviolet, B for blue, and V for visual. The central wavelengths of the filters are roughly: U - 3600 A°; B - 4400 A°: V - 5500 A°. The pass band, or wavelength range passed by each filter, is roughly 1000 A° for each filter in the broadband UBV system - e.g. the B filter passes only light from about 3900 A° to 4900 A°. e define magnitudes in each filter- e.g. mV (or sometimes just V) is the magnitude in the V filter, for instance. The color of an object related to the variation of flux with wavelength. Using broadband filters (like UBV) we define the color index as the difference between the magnitudes in 2 colors, e.g.

$$\mathrm{B} - \mathrm{V} = \mathrm{m_B} - \mathrm{m_V} \qquad (8)$$

defines the B − V color index.

What does B − V tell us about the color of an object? From the basic equation defining magnitudes equation (1) we see that a magnitude difference corresponds to a flux ratio. The ratio is the flux at B relative to the flux at V, of the same object, instead of different objects.

$$\mathrm{B} - \mathrm{V} = \mathrm{m_B} - \mathrm{m_V} = -2.5\log(\mathrm{f_B}/\mathrm{f_V}) + \text{constant} \qquad (9)$$

where f_B is the flux averaged over the B filter and f_V is the flux averaged over the V filter. The "constant" appears in the above equation because of the way we define the zero point of the color system. You might think that if B−V = 0.00, then $f_B = f_V$. However, this is not how the color system is defined. Historically, astronomers picked a set of stars of spectral class A (including Vega) and defined the average color of these stars to have all colors equal to 0.00. For an A star, f_B is not equal to f_V, so that a non-zero constant is needed in equation (5) to make the color come out to 0.00. Thus, the B−V color of Vega is 0.00, pretty much "by definition". The B−V color of the Sun, redder than Vega, is about 0.63. The B − V colors of the hottest (bluest) stars are about −0.3.

Telescopes

Job of the Telescope

Ground based astronomical research telescopes around the world represent a capital investment of several billion dollars. A single space telescope, the Hubble Space Telescope, had a price tag of about 3 billion dollars, with an annual operating budget large enough to build a large ground based telescope each year.
Why do we bother to build telescopes and equip them with detectors? Why not just use our eyes to study the heavens? Telescopes (1) collect more light than the unaided eye (2) have increased angular resolution, or ability to see fine detail, than the unaided eye, and (3) telescopes (or more specifically, detectors attached to telescopes) allow us to study wavelengths not visible to the unaided eye and (4) detectors allow a permanent record. For collecting light, the bigger the telescope, the better! More light allows us to see and study fainter objects, or make more accurate measurements on bright objects. Bigger telescopes also have better angular resolution, allowing finer detail to be seen, although the full resolving power of research telescopes is usually not attained due to the deleterious effects of the Earth's atmosphere, which smears out the light from celestial objects.

Image Formation

A telescope forms an image in the focal plane. The simplest telescope is simply a convex lens. This forms an image as indicated in Figure (1) . If we put a small magnifying glass (usually called an eyepiece in astronomical terminology) near the focal plane and examine the image with our eye, then we have a simple visual telescope. If we instead put some sort of detector, or device to record the image (such as a piece of film or a CCD), in the focal plane, that is also a telescope.

Types of Telescopes

Telescopes can be divided into refracting, reflecting, or catadioptric. Refractors use a lens (transmissive optical element, meaning light travels through the element) as the primary light gathering element. Reflectors use a mirror (a reflective optical element, meaning light bounces off the surface and does not travel through the element) as the primary light gathering element. Catadioptric telescopes use both transmissive element(s) and reflective element(s) as part of their primary light gathering element.

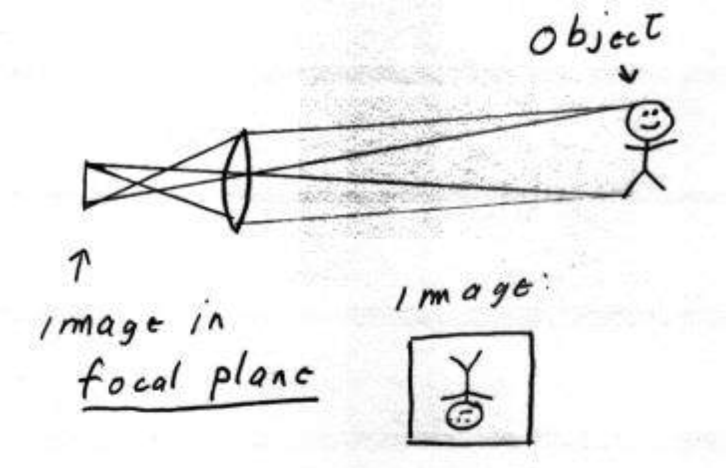


Figure (1) Image formation by a simple lens. The lines show the paths of only a few rays from the object. Note that the rays that pass through the center of the lens are unbent, while those passing through the top of the lens are bent downward, and those hitting the bottom of the lens are bent upwards, resulting in an upside down image in the focal plane

The heyday of the refractor among large research telescopes has long since passed. The largest refractors, built in the late 19th and early 20th century, include the Lick 36 inch and Yerkes 40 inch, (the measurement being the diameter of the main lens). Larger refractors than these have never been built, due to a number of factors. First, since the light must pass through the lens, it must be supported only along the edge of the glass. A large lens can flex as the angle between the lens and the pull of gravity changes, distorting the figure of the lens. Refractors suffer from chromatic aberration, meaning that light of different wavelengths come to slightly different focus.

Today, the majority of amateur and all large research telescopes use a mirror as their primary light collector and so are reflectors. A glass substrate is used to hold the optical figure, while the reflectivity comes from a thin layer of aluminum deposited on the front of the mirror. Because light does not pass through the mirror, it can be supported from the rear, so that glass flex does not limit the size of mirrors the way it does lenses. Most large reflectors use a single large piece of glass for their primary mirror (monolithic mirror), although several important telescopes have used a segmented primary, where the primary is actually composed of a number of separate pieces of glass. The reflective coating (usually aluminum) on mirrors reflects all wavelengths the same, so reflectors do not suffer from chromatic aberration.

Figures (2) and (3) show schematic configurations for several common telescope types. The Newtonian uses a parabolic primary, with a flat diagonal mirror to move the focal plane to the side of the telescope tube. This is a common "home made" telescope type. It suffers from limited field, due to off- axis aberrations meaning image quality deteriorates away from the center of the optical field. Several types of telescopes use 2 curved mirrors, a concave primary and a convex secondary. The convex secondary partially counteracts the converging beam from the primary, making an effective focal length much larger than the focal length of the primary mirror Figure (4). A classical Cassegrain system has a parabolic primary and a hyperbolic secondary.

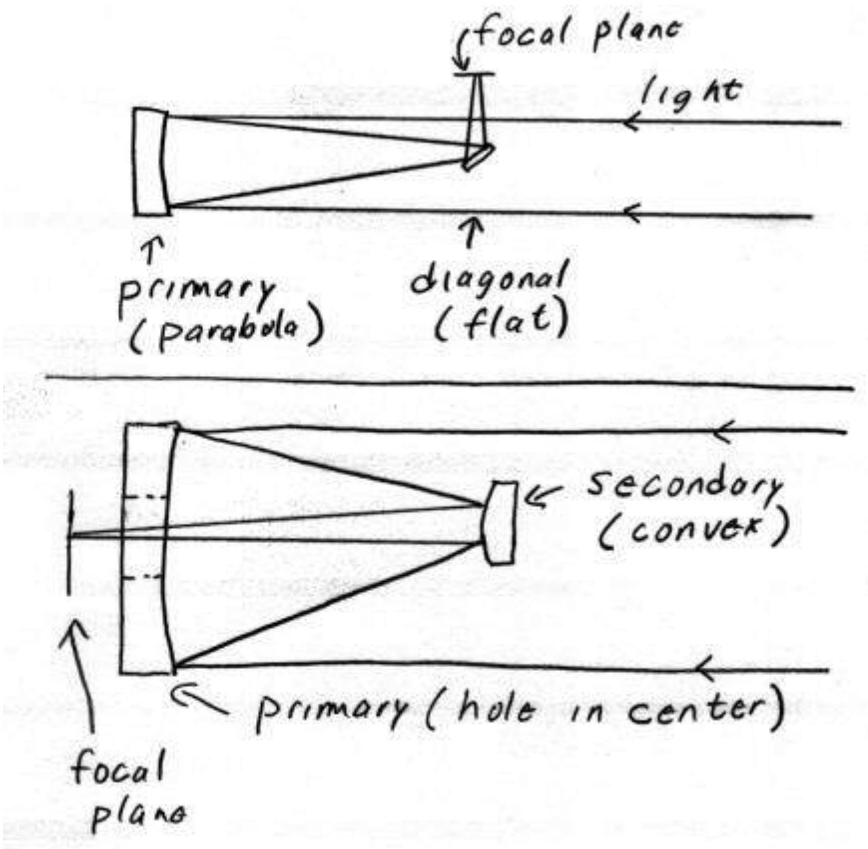


Figure (2) Top: Newtonian ; bottom: Cassegrain. In these types of systems, the diagonal or secondary is usually held in place by 4 vanes attached to the inside of the telescope structure. Diffraction effects from light passing by these vanes are responsible for the familiar lines seen emanating from bright stars on many images of the sky. These lines are called “diffraction spikes”.

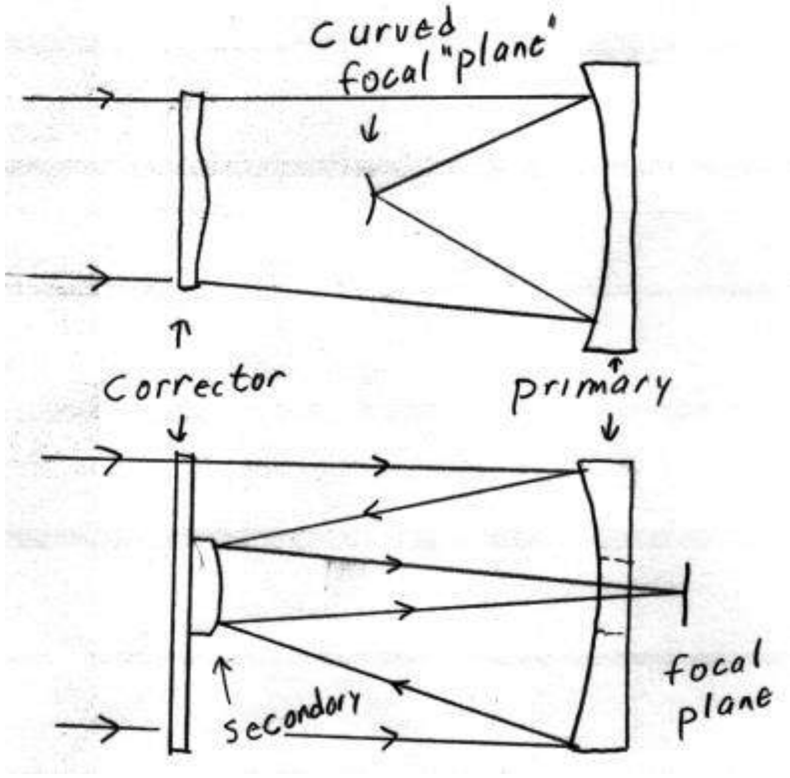
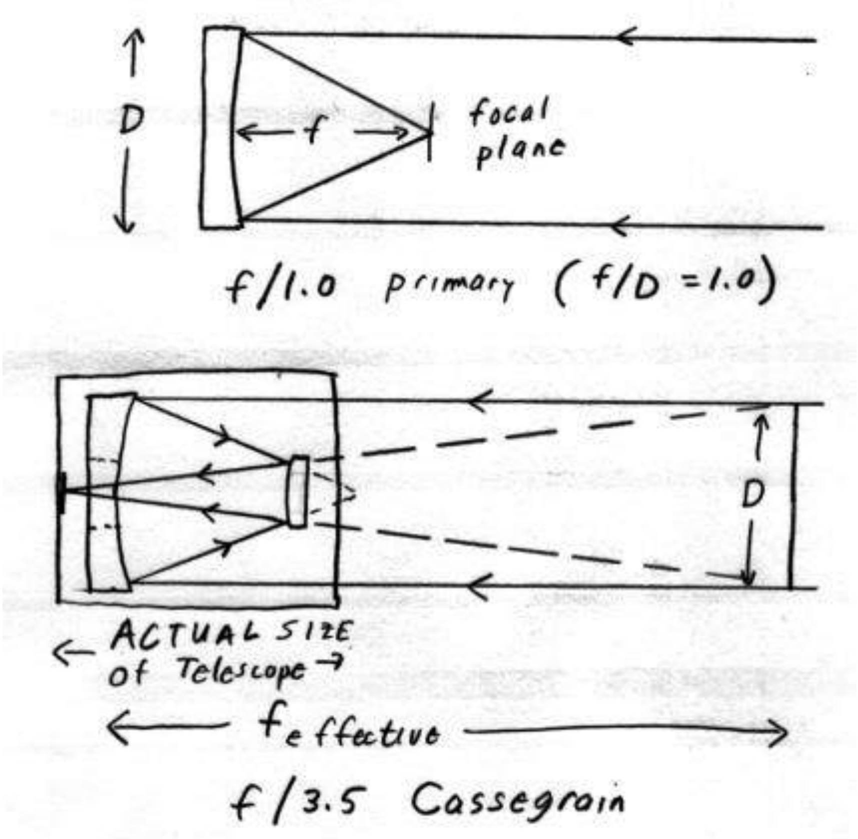


Figure (3) Top: Classic Schmidt camera; bottom: Schmidt- Cassegrain (SCT) configuration. In the SCT the secondary is usually mounted on the corrector plate, so there are no diffraction spikes in images taken with these telescopes.

good images over large fields of a degree or more, a Schmidt camera is often used. This uses a spherical primary. Of course, a spherical mirror suffers from spherical aberration, because rays hitting the central part of the mirror come to a different focus than rays hitting the outer parts of the mirror. In a classic Schmidt camera Figure (3) a weak transmissive corrector is used, which is figured so as to cancel the spherical aberration of the primary. This gives good images over fields of many degrees, but at the expense of a curved focal "plane".

Many amateur telescopes use a catadioptric optical configuration called a Schmidt- Cassegrain system. This combines a weak (almost flat) transmissive corrector plate with a spherical primary mirror and an ellipsoidal secondary. It is relatively easy to make large spherical mirrors, so this configuration has become very popular among amateur telescopes (SCTs= Schmidt- Cassegrain telescopes).



Focal length and f-ratio

Imaging systems are characterized by their focal lengths, f. Focal length is easy to understand for a simple system such as a refractor- it is just the distance from the lens to the image plane, as shown in Figure (1).
Another useful number characterizing a telescope is the f-ratio. The f-ratio is defined as f/D, where D is the diameter of the primary mirror or lens. The focal length sets the size of the image, while the diameter of the primary of course controls the amount of light in the image. Systems with a low f-ratio have a relatively large amount of light in their images, compared to the size of the image, and so are called fast systems, while large f-ratio systems are called slow systems.
The mapping between angles in the sky and linear distance in the image plane is set by the focal length of the system. Consider 2 points of light separated by an angle θ on the sky. The linear distance s between the points in the image is given by

$$s = f\,\theta \qquad (6)$$

provided θ is measured in radians and is reasonably small, as is almost always the case for astronomical telescopes and CCD systems.

Traditionally, the mapping between angle on sky and distance in the focal plane is given by the inverse plate scale, measured in units such as arcsec/mm or arcmin/mm. It is easy to see by equation 5.1 that the scale S (in arcsec/mm) and the focal length f (in mm) are related by:

$$s = \frac{206265}{f} \qquad (7)$$

You should recognize the number 206265 as the number of arcsec in a radian.

Field of View and Sky Coverage

The field of view (FOV) is the sky area covered by an image taken with a telescope. The FOV depends on both the focal length of the telescope and the area of the imaging detector. With single CCD detectors, the angular area covered tends to be smaller with larger telescopes, as larger telescopes usually mean longer focal lengths. To overcome small fields covered by CCDs at large telescopes, astronomers are building cameras with multiple CCDs in the same plane. These cameras are very expensive, and require financial and engineering resources that only the largest observatories can muster.

Angular Resolution

A point source is one that has no angular extent. Although real stars have a finite angular extent, they are, for practical purposes with optical telescopes, point sources. So, is the image of a star a point? No. First, the atmosphere smears out the light from a point source, a very important and deleterious process called seeing. However, even if we could put our telescope outside the atmosphere (`a la Hubble) a real telescope does not focus a point source into a point image, but rather a "bulls eye" pattern called an Airy disk. The reason for this is that light acts as a wave, and waves from different parts of the mirror interfere with each other in such a way as to result in the Airy disk pattern. The Airy disk has a central peak, then a series of dark and light annuli. Figure 5 shows a drawing of an Airy disk, and a cross section of the brightness along a diameter.

The angular size of the Airy disk pattern on the sky is set only by the diameter (D) of the primary mirror or lens, not by its focal length. The linear size (in the focal plane) of the Airy disk in the image plane is set by the angular size (set by D) and the image scale, set by the focal length.

The angular radius (in radians) of the first dark ring is given by

$$\theta = \frac{1.22\lambda}{D} \qquad (8)$$

where λ is the wavelength of the radiation. Note that the larger the telescope primary, the smaller the angular size of the image of a point source. The angle above is traditionally called the Dawes limit, or the diffraction limit. To first order, two point sources with an angular separation larger than the Dawes limit are resolvable, while two point sources closer together than the Dawes limit would be seen as one point and would not be resolvable. In practice, at least in the optical with most telescopes, angular resolving power is set by the seeing or smearing by atmosphere (lots more about this later!), and the Dawes limit plays no role. However, this does not mean that the Dawes limit is not important. For example, the Hubble Space Telescope angular resolution is essentially set by the Dawes limit. Of course, the Dawes limit assumes the optics are figured properly. When Hubble was first used, the optics suffered from spherical aberration. The phrase "in the absence of additional sources of image degradation" turns out to be a crucial one. The Earth's atmosphere smears the light from stars, a process called seeing. Instead of the image of a point source being an Airy pattern, it is a fuzzy blob with a quasi- Gaussian profile. The angular size of the blob is set by the atmosphere, and not by the telescope.

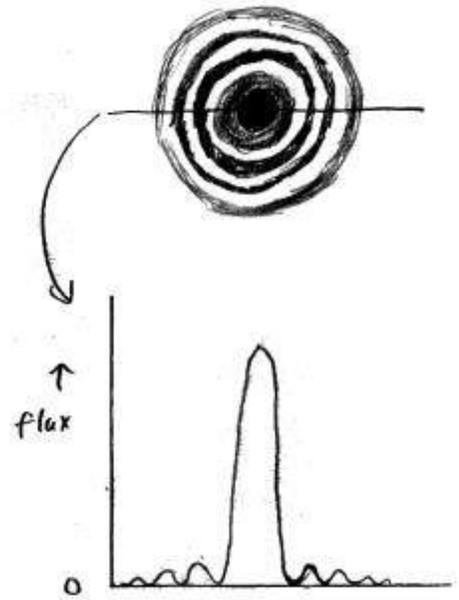


Figure (5): Top: Negative gray scale image of Airy PSF; Bottom: Intensity along a cut across Airy PSF

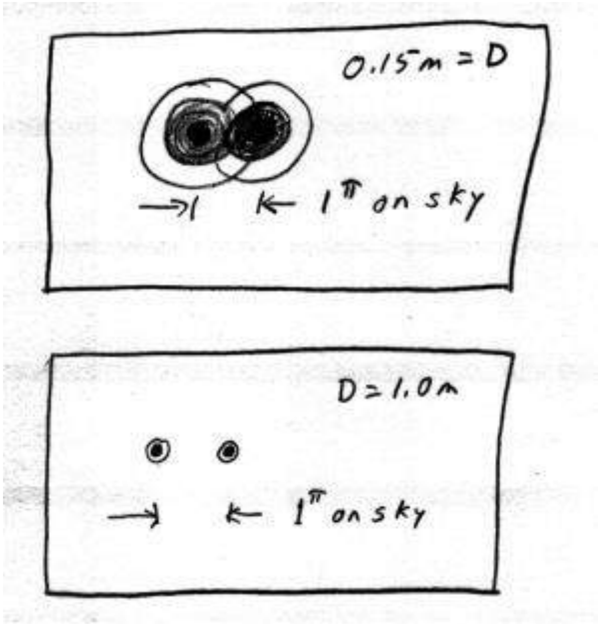


Figure (6): The top panel shows the images from a telescope with D=0.15m. The two stars are barely resolved, because the size of the Airy disks are comparable to the separation. The bottom panel shows the same two stars as seen through a D=1.0 m telescope. The two stars are easily separated, because the angular size of the Airy disk is much smaller than the angular separation of the stars. Note that for almost all optical ground based research telescopes, the resolution is set by seeing, not the Airy pattern.

Telescope Mountings

Most research telescopes are on general purpose mountings which allow them to point at any spot on the sky and track or follow the apparent motion of the stars caused by the rotation of the earth. There are also some special purpose telescopes which can only look at restricted parts of
the sky. One example is a transit telescopes, which only looks on the meridian. The basic general purpose telescope mounting consists of two rotational axes at right angles to each other. In the altaz (altitude-azimuth) mounting, one rotational axis points straight up, and the other axis is horizontal. We can move the telescope in altitude (angle above horizon) and in azimuth (angle relative to north in the plane defined by the horizon). In the equatorial mounting, one axis is tilted to be parallel to the rotational axis of the earth. Both types of mountings have their advantages and disadvantages. An equatorial mounting allows the stars to be tracked by driving only one axis and that at a constant rate of once per sidereal day. In an altaz mount, you must move both axes at the same time to follow the paths of stars across the sky (unless you are at the North or South Pole, where the stars move around the sky at constant altitude above the horizon!) The rates that the two axes are driven are different from each other, and both change with position in the sky. The

optical field for an equatorial mount stays at a constant angle relative to the telescope tube. In an altaz mount, the field rotates relative to the tube.

In an altaz mount, the field rotates relative to the tube. If one took a time exposure using a telescope on an altaz mounting, even if it correctly tracked the stars, the stars would be trailed due to this field rotation. (To understand why field rotation occurs, think about point at the north celestial pole. For an altaz mounting, this would just mean parking the telescope at the point due north, at an altitude above the horizon equal to ones latitude.

The Atmosphere: Bane of the Astronomer

The Earth's atmosphere is, overall, a Good Thing- it provides us with oxygen and destroying things like x-rays and ultraviolet EMR and cosmic rays. But, for ground based astronomers, the atmosphere is nothing but Trouble Problems that you might offhand think are important - clouds and air pollution are not the main source of trouble. We can locate our telescope at a place where clouds are (at least relatively) infrequent, or, if we are stuck someplace, simply wait for clear weather. Nor does atmospheric pollution cause a major problem, as most research observatories are far from civilization. The obvious problems posed by the atmosphere- clouds and pollution- can be largely overcome by telescope location. But, even at 4200 meters above sea level on Mauna Kea on Hawaii in the middle of the Pacific- there are several deleterious effects of the atmosphere. There are five main problems, each of them physically distinct. The problems imposed by the atmosphere are: (1) Limitation to small windows in the EMR spectrum. The Earth's atmosphere allows only a small fraction of all wavelengths of EMR to penetrate. There is an optical window that allows our eyes to see the Sun and stars, and a radio window that allows radio telescopes to observe celestial objects. As we shall see shortly, the atmosphere is not completely transparent even in these windows. (2) "Smearing" or "fuzzing out" of images of celestial objects caused by passage of light through turbulent atmosphere. Astronomers call this seeing. Not only does seeing cause us to lose detail, it also makes it much harder to see and measure the brightness of faint objects. (3) The atmosphere, even under pristine conditions at a site far from city light, glows due to atomic processes in the air. This light emitted by the sky, called sky glow, is a a severe problem when observing faint objects, because the sky glow photons make extra noise which degrades the accuracy of our measurements. Near cities, the situation is much worse, as the atmosphere, besides glowing, also scatters light from artificial sources, making the sky appear even brighter than at pristine sites.

(3) The atmosphere absorbs and scatters some fraction of the light at all optical wavelengths. This causes objects to be dimmer than they would be without the atmosphere. Astronomers call this atmospheric extinction. (4) Except when looking at the zenith, the atmosphere acts as a weak prism, spreading out light in a small spectrum along the line pointing to the zenith. This effect is called atmospheric refraction. This can smear out images, particularly when observing with a filter that covers a wide wavelength range. Refraction also can really mess up spectrophotometry, because the light from different

wavelengths falls on different parts of the detector. Seeing causes us to lose much of the detail that properly designed large telescopes are capable of providing. Seeing and sky glow severely limit the accuracy with which we can measure the light from faint objects. Extinction is by far the least deleterious of these effects Besides these effects, there are other annoyances caused by the atmosphere. Wind shakes our telescopes, degrading image quality; clouds block light a fraction of the time; high humidity, particularly when coupled with dust and pollution, can degrade optical surfaces and rust metal parts. Note that the 5 effects above occur regardless of weather- even the clearest, most cloud free mountaintop hundreds of miles from any city lights has these effects.

Space Astronomy and the Perfect Observing Site

The only way to get totally away from the deleterious effects of the atmosphere is to put your telescope into space. Astronomers, of course, have done this with many telescopes to observe in wavelengths that do not pass through the Earths atmosphere (such as UV, x-rays, rays). The Hubble Space Telescope is the only telescope operating in the the optical window that is in space. The Hubble was put above the atmosphere not only to observe wavelengths that don't pass the atmosphere (ultraviolet), but also to observe in the optical wavelengths above the smearing of images (seeing) caused by the atmosphere. However, putting telescopes in space opens up a whole new set of problems- extreme cost, need to control remotely, inability to easily fix things that break etc. Even in this era of Hubble and other high profile space telescopes, the vast majority of photons from celestial objects are caught with ground based telescopes. This will undoubtedly remain true for many, many decades or centuries into the future, simply due to limited resources. It will always be possible to build bigger telescopes on the ground than in space.

If astronomers had all the resources they wanted to do astronomy, where would we build telescopes? The Moon might be the "perfect" place. The Moon has essentially no atmosphere With much lower gravity than on the Earth, and no wind to shake our telescope structures, telescopes could be much lighter than on Earth . Where on the Moon would be the best place for a telescope? Perhaps on the floor of a crater near one of the poles of the Moon. Such locales would be in perpetual shadow. Without an atmosphere to scatter light, and with the telescope hidden from direct sunlight by the crater walls, the sky would be very dark.

Clouds and Photometric Skies

Back to clouds. Even though clouds are not a fundamental problem, because we can wait them out, they are an annoyance. There are basically two different "modes" of doing photometry. These are sometimes called all sky and differential photometry. In all sky photometry, we must compare the count rates of the object we wish to measure to standard stars in a completely different part of the sky. Obviously, if there are clouds in front of our object and not in front of the standard stars, or vice versa, we will get the wrong answer! All sky photometry thus requires completely cloud free conditions at your observing site. Besides clouds, occasional problems such as lots of dust in the atmosphere can prevent accurate all sky photometry. When conditions are suitable for all sky photometry, we say we have a photometric sky.
differential photometry, can be done, using a CCD camera, from partially cloudy sites. In differential photometry, we compare the brightness of our unknown object, usually some variable object such as a supernova, with the brightness of stars on the same CCD frame. If clouds block some of the light during the exposure, they will dim the light of the stars and the object of interest the same fractional amount (since the objects are close together on the sky), and so the ratio of the fluxes of the two objects will not be affected. We can measure at our leisure on a photometric night (or at a better site) the magnitudes of these secondary standard stars. Using the accurate secondary star magnitudes and the ratio of fluxes observed during non photometric conditions, we can derive accurate magnitudes for the variable object at the time of the non- photometric observation. This is most useful for time- critical observations- e.g. to get a magnitude vs. time plot for a variable object.

Clouds: the Bad and the Ugly

To an astronomer, there are no good clouds, unless maybe when you have been observing for too many long winter nights and need an excuse to get some sleep!) Clouds come in a wide variety of optical thicknesses, or opacities. In some ways, the really optically thick ones (when you look up
and can't see any stars, or when its raining or snowing) are not as bad as those that block only a fraction of light. If the clouds are so thick that you can't see any stars, its time to do something else without guilt. Worse is when there are thin clouds around, or when there are thick clouds around, with some "clear" patches. Conditions with thin clouds or scattered thick clouds can be useful for doing differential photometry with a CCD.

However, it is not possible to do all sky photometry with clouds around. If you are at a really dark site, and the Moon is not up, it is surprisingly difficult, if not impossible, to detect the presence of thin clouds by simply looking at the sky. Near cities, with lots of artificial light around, you can usually detect thin (or thick) clouds because of the light they reflect from artificial sources, rather from any dimming of starlight they might cause.
Many people don't believe the assertion above, that your eye cannot detect thin clouds at a dark place. This is primarily because they have little experience seeing the sky from a really dark place.

Seeing

Without the atmosphere, light rays from a distant star would arrive at our telescope parallel to each other, and the telescope would focus those rays to a small spot (but not exactly a point, due to diffraction effects (Airy disk), as discussed earlier). However, the passage of the light rays through the last few kilometers of their journey in the Earth's atmosphere scrambles the rays slightly and makes them no longer exactly parallel. The direction of the rays is being continuously changed by a slight amount, as the rays traverse the turbulent atmosphere, resulting in an image of a star that appears as a wandering blob of light rather than a nice sharp unwavering diffraction pattern.
Seeing is the term astronomer's use for this smearing and shimmering of light from celestial objects due to its passage through the Earth's atmosphere. Seeing makes the images of stars appear much larger than the limit set by diffraction, and makes the images of extended objects (e.g. planets) appear fuzzy. Seeing, to astronomers, refers to this image smearing. Seeing does not refer to the loss of light, but only to the loss of detail caused by scrambling of light rays. The atmosphere also does cause light to be dimmed somewhat, a process called atmospheric extinction. The process of seeing is exactly the same physical process that you are familiar with when you see far away objects shimmer if they are viewed through turbulent air, or through air parcels of different temperatures, say air over a hot parking lot or roadway. In this case, the air near the road is heated, expands, and rises, causing temperature variations along the line of sight and turbulence. Light rays are bent by the passage through parcels of air with different temperatures, as the index of refraction of air varies slightly with temperature. Seeing causes the image of a star to be a blob of light, centrally concentrated and fading with angular distance from the center. Astronomers

characterize the seeing by the angular FWHM (full width at half maximum), which is the angular size of the star image at a level of half the peak level. The seeing can regularly be as good as 0.5 arcsec. How can we get the best possible seeing? The higher altitude the telescope, the better the seeing tends to be, simply because there is less air to look through the higher one goes. Ideally, one wants a high mountain that has smooth (laminar) airflow over it. In reality, observatory locations are subject to many constraints, from financial considerations to political and access issues. Over the past decade or so, astronomers have begun to realize that, at least at the best astronomical sites, a significant fraction of the smearing of astronomical images occurs during the last few meters the light travels before detection. For instance:

A- If the mirror is warmer than the air in the dome, there will be turbulence and temperature inhomogeneities as the hot air rises above the mirror.

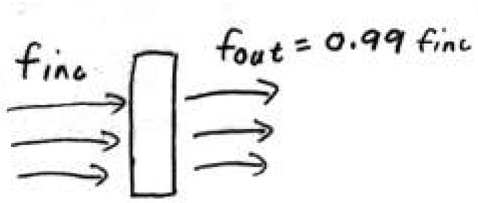
B-Turbulence near the dome slit can be caused if the dome air is warmer than the air surrounding the dome.

Optical Depth and Atmospheric Extinction

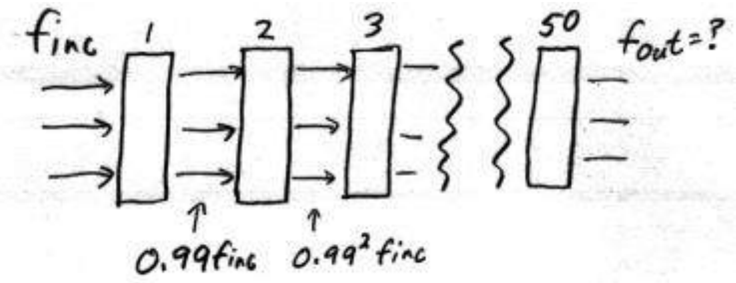
In everyday life, we think of things as transparent or opaque. However, even the clearest looking piece of window glass absorbs some light, so is not totally transparent, and while clear, dry air may look absolutely transparent when looking at something nearby, if you look through many kilometers of the same air, you can see that air is not totally transparent. Observers who must look through the atmosphere to observe celestial objects.

Optical Depth

Radiation transfer is the branch of astrophysics that deals with how electromagnetic radiation (EMR) travels through and interacts with matter. One of the central concepts of radiation transfer is that of optical depth (or optical thickness) which denoted by τ. To understand the basic concept of τ, consider a slab of gas that absorbs a very small fraction of any EMR that falls on it. If we have a beam of EMR of flux f_{inc} (incident flux) and that the output flux (f_{out}) is 0.99 that of f_{inc} (in other words, 1% of incoming light is absorbed) as shown in figure below



This mean that the slab absorb 1% of light incident on it and this value called the optical depth of the slab. What would happen if we have a more opaque slab? In Figure below, we have a slab composed of 50 of the slabs from the previous figure



$$f_{out} = f_{inc}\ (0.99)50 = 0.61 f_{inc} \qquad (1)$$

Or in other way

$$f_{out} = f_{inc}\ (1 - \frac{0.5}{50})^{50} \qquad (2)$$

Equation (2) is similar to exponential function

$$\lim_{n\to\infty}\left(1 + \frac{x}{n}\right)^{n} = e^{-x} \qquad (3)$$

Optical depths simply add linearly. For the configuration of 50 slabs each of which absorbs 0.01 of the incident flux, the optical depth is simply 50 x 0.01 = 0.5. However, the output flux is not half that of the input, but rather

$$e^{-\tau} = e^{-0.5} = 0.6$$

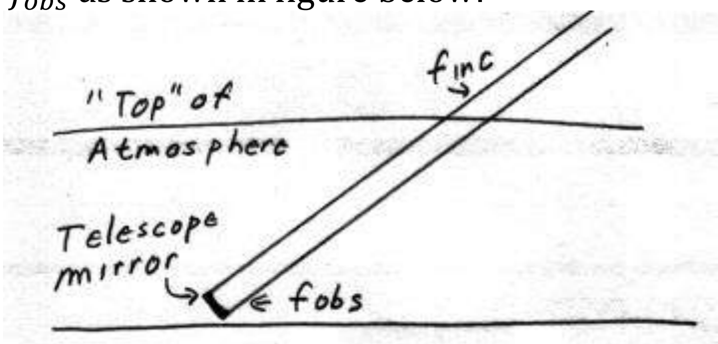
In general, the incident and output flux of a slab of optical depth τ are related by the following equation:

$$f_{out} = f_{inc} e^{-\tau} \qquad (4)$$

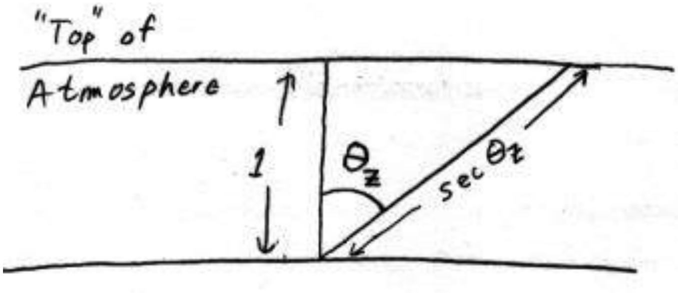
Note that optical depth is a dimensionless (unit less) quantity- it is simply a pure number.

Atmospheric Extinction

You can think of the atmosphere as an absorbing slab. Consider the beam of light from some star that will hit our telescope mirror. Just outside the atmosphere, the beam has flux f_{inc} (for flux incident). At the telescope, the flux of this beam is less due to absorption and scattering of light out of the beam. We will call the observed flux at our telescope f_{obs} as shown in figure below:



Obviously, if our telescope is on the ground, we have to look through some of the Earth's atmosphere to see objects out in space. If we look straight up (at the zenith) we have the minimum possible path length through the atmosphere (for a given altitude of observatory). At an angle θ_z from the zenith (called the zenith angle) the amount of air we look through, relative to that at zenith, is simply given by secant θ_z as shown in figure below:



When we are looking straight up we say we are looking through "1 airmass". At other zenith angles, we look through "secant θ_z airmasses". (NOTE: The secant θ_z formula strictly applies only in an infinite flat slab. Because the atmosphere is curved (due to the curvature of the Earth), airmass is not exactly secant θ_z but the difference between the real airmass and secant θ_z is significant only for lines of sight near the horizon (_ approaching 90 degrees). Also note that the fact that the atmosphere becomes less dense as we go up does not change the secant θ_z formula, as long as the density is the same at different places at the same altitude above sea level, which it is.).
We can express θ_z in term of the observables as follows:

$$\text{Secant } \theta_z = \frac{1}{\sin\lambda \sin\delta + \cos\lambda \cos\delta \cos h}$$

Where λ is the latitude of the observatory δ is the declination of the star and h is the hour angle of the object at the time of the observation.

Night Sky, Bright Sky

Anyone who has "watched the stars come out" as the sky darkens at twilight knows that the darker the sky, the faintest star you can see. Why this is so? The reason is not, as you might first think, that the light from the bright sky somehow blocks the light from the stars, as a cloud would block starlight. Rather, due to the inherent photon nature of light, the light from the sky produces noise that makes it harder to detect the signal from the stars. The star signal must compete with the sky noise to be detected. The brighter the sky, the greater the noise in the sky. The more noise, the harder to detect a given signal. Detection problems, such as how faint a star you can see with your naked eye, or how faint an object you can image with a given telescope, detector, and exposure time, always involve not just the amount of signal from the object you wish to observe, but the ratio of that signal to the noise present, the signal to noise ratio (often written as S/N). The higher the signal to noise ratio, the easier it is to detect the signal from a star. With the idea of signal to noise ratio in mind, it is easy to see that you can detect fainter stars in 2 ways: increase the star signal or decrease the noise. If you wanted to detect fainter stars, you would immediately think of using a larger telescope. With the larger telescope you can see fainter stars because you are increasing the star signal with the larger collecting aperture. (The larger aperture also increases the sky noise, but the star signal increases by a larger factor than the sky noise, so that the all important signal to noise ratio increases.) The other way, not so obvious, to see fainter stars is to decrease the noise. This is the basic idea behind the “stars coming out” at twilight. What is happening as the sky darkens? The signal from any star is constant, but the sky signal, and hence sky noise, is decreasing, so all stars have a signal to noise ratio that increases as the sky darkens. As the sky signal, and hence noise, decrease, fainter and fainter stars have a high enough signal to noise ratio to be detected with your eye. Sky noise, while not the only noise source in astronomical imaging, is usually the dominant such noise source. Once you understand the importance of the signal to noise ratio, you can see that the sky brightness, and hence sky noise, plays a crucial role in determining how faint you can image with a given telescope. What exactly does it mean to measure the sky brightness? When astronomers speak of the sky brightness, they are really referring to the surface brightness of the sky, or the brightness of a small patch of the sky of a given angular extent. All you need to measure the night sky brightness is an image of a star with known magnitude and a dark image of the same exposure time. For comparison with other observing sites, it is preferable to make the image through a standard filter, such as one of the UBVRI filters used by astronomers to measure brightnesses and colors of astronomical objects.

Photon Detectors

Telescopes simply collect photons. To make useful observations requires some sort to detect those photons and make a measurement or record of them.

1- Human Eye

The “natural” detector of visible photons is the retina of the human eye. The human eye is an amazing organ. It does its intended job- provide us with panoramic views over a large solid angle with sub- second time resolution under an enormous range of lighting conditions, from full desert sun to the darkness of the night sky in the same desert at midnight. However, the eye is not a very good photometric device, in the sense of being able to make quantitative measurements of flux. Also, the eye is not an integrating device, that is, the signal does not build up with time. The eye does not provide a permanent record of what it sees. The angular resolution of the human eye is limited by the small size of the lens. Under dark conditions, the lens is about 7 mm in diameter. The theoretical resolution is about 16 arcsec but few people have eyesight that approaches this resolution. The eye is sensitive to only a small slice of EMR wavelengths. The range of wavelength sensitivity of the eye is somewhat dependent on whether the eye is observing in bright light conditions or under very low light level conditions.

2- Photographic Emulsions

The first technological advance in astronomical detectors was the photographic emulsion. Photographic emulsions can make a permanent record of astronomical objects imaged by telescopes. However, photographic emulsions are not very good photometric devices for astronomy because of several drawbacks. Briefly, these are: photographic emulsions only record a small fraction (around 1%) of the photons that hit the emulsion. Because of the analog (rather than digital) nature of the image record on an emulsion, it is difficult to make quantitative measurements of star brightnesses. Photographic emulsions are also nonlinear with input light- if one star is twice as bright as another, it does not produce twice the output on the film. Photographic emulsions are also nonlinear with increasing exposure time- an exposure of 2 minutes does not give twice the output of a one minute exposure. This feature is called reciprocity failure.

3- Modern Detectors - PMT and CCD

Modern detectors are inherently digital. They detect individual photons and output a number which is directly and linearly related to the number of photons that were incident on the detector. The first modern detector in this sense is the photomultiplier tube (or PMT). A PMT consists of an evacuated glass tube, on one end of which is deposited a film of a material (such as indium antimonide) called a photocathode. This material has the property that when it is struck by a photon, an electron is often liberated from the material. Each electron liberated from the cathode is directed away from the cathode by an electric field, and is amplified into a pulse of electrons by a series of metal plates (called dynodes) and an electric field in the tube which accelerates the electrons. Electronics coupled to the PMT counts these pulses. Thus, a single photon hitting the cathode results in an easily counted pulse of many electrons. (Figure (1) shows a diagram of a PMT and Figure (2) represent a figure of a photometer, that holds the PMT and associated equipment onto the telescope.) The drawback of a PMT is that it is essentially a single channel device, meaning there is no positional information in the signal. The output signal does not depend on where on the cathode the photon hit, so we get only a measure of all the light that fell on the photocathode. However, the PMT, unlike a photographic plate, has a digital output, meaning we can easily make quantitative measurements. The fraction of the photons which hit the cathode that are actually detected by the PMT is set by the fraction of photons that hit the cathode that liberate an electron. This fraction is typically 20% or so, so the efficiency of detecting photons is much higher for a PMT than for the best photographic emulsion. The detector of choice for optical astronomy is now the CCD (charge coupled device). The CCD has many advantages- it is a linear, photon counting device which records a large fraction of the photons that fall on it. It is far better than a PMT because it can record a two dimensional image- i. e. there is positional information. CCDs have truly revolutionized astronomy in the last two decades. The next chapter is devoted to further discussion of these amazing devices.

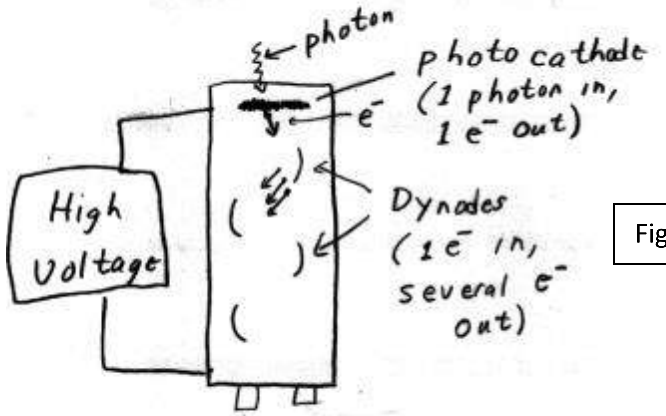


Figure 1

Schematic diagram of a photomultiplier tube (PMT). The high voltage supply creates an electric field that accelerates electrons along the tube. At each dynode, an impinging electron knocks loose several electrons, which are then accelerated towards the next dynode, here each of them knock loose several more electrons. Through this cascade, a single photon hitting the photocathode releases an easily counted pulse of many electrons.

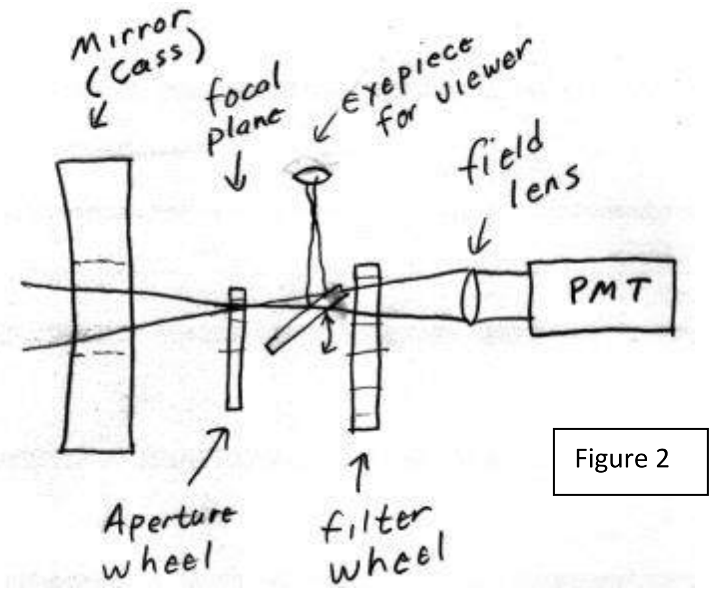


Figure 2

Schematic diagram of a simple photometer. The aperture wheel contains holes of various sizes that define the angular size of the spot in the focal plane which is measured by the photometer. Behind the aperture wheel is a mirror that can be flipped into the light path to direct light to an eyepiece. This “behind the aperture viewer” allows the astronomer to visually center the star in the aperture. The mirror is flipped down to allow light to pass to the detector. The filter wheel holds various filters that define the pass band measured by the PMT. The field lens directs the light to the photocathode of the PMT. The PMT is usually cooled by dry ice (solid carbon dioxide).

CCDs (Charge Coupled Devices)

Basic Concepts

A CCD is a light sensitive silicon “chip” which is electrically divided into a large number of independent pieces called pixels (for “picture elements). Present day CCDs have 512 x 512 (262144) to at least 4096 x 4096 (16,777,216) individual pixels, and are from about 0.5 cm to 10 cm in linear size (typical sizes of each pixel are 10 to 30 micron square). For astronomical use, we use the CCD as a device to measure how much light falls on each pixel. The output is a digital image, consisting of a matrix of numbers, one per pixel, each number being related to the amount of light that falls on that pixel. Of course, one of the beauties of the CCD is that the image, coming out in a digital form, is readily viewed, manipulated, measured, and analyzed with a computer. Research astronomers spend FAR more time sitting in front of computers than anywhere near telescopes!
Several concepts are basic to CCD use as a low light level detector in astronomy.

Quantum efficiency (QE): A CCD detects individual photons, but even the best CCD does not detect every single photon that falls on it. The fraction of photons falling on a CCD that are actually detected by the CCD is called the quantum efficiency (QE), usually expressed as a percentage. QE is a function of wavelength. For optical detection, there are two basic styles of chip: thick chips, in which the light passes through some of the electronic layers of the CCD before hitting the silicon detecting level, and thin chips, in which the silicon layer is mechanically or chemically thinned and the light enters the silicon directly. Thick chips have low QE in the blue, because the electronic layers absorb much of the blue light. Thin chips have better blue QE. Thin chips and thick chips have more similar red QE, but the thin chips usually have higher QE at all wavelengths than the thick chip.
Counts - So, are those numbers that we read out of the CCD the actual number of photons that fell on each pixel? part of the number is an electrical offset called the bias and part may be due to dark current. After we subtract these components, the signal is related to the number of electrons liberated by photons in each pixel. Only a fraction QE of photons generates electrons, so that the number of electrons is: (number of photons) × QE.

Integration time- The CCD (unlike the human eye but like a piece of film) is an integrating device. The signal (electrons knocked loose from the silicon by hitting photons in each pixel) builds up with time. The integration time (or exposure time) is controlled by a mechanical shutter (like in a camera). Except in the case of a very bright star, where the CCD saturates, the signal from a star increases linearly with time, unlike a photographic plate. The CCD does not exhibit reciprocity failure.

Read noise- After an integration (exposure), the CCD must be "read out" to find the signal value at each pixel - because the signal may be as low as a few electrons per pixel, this step involves some very sophisticated amplifiers that are part of the CCD itself ("on chip" amps). Unfortunately, but inevitably, the read out process itself generates some electronic noise. The average noise per pixel is called the read noise. Modern CCDs typically have a read noise of 5 to 20 electrons per pixel per read out (read noise is the same whether exposure is 0.1 sec or 3 hours).

Bias frame- If we simply read out the CCD, without making an integration, (or think of a zero second integration), there will be a signal called the bias signal. (You might think that the bias would be identically zero, but it isn't. Think of it as an electrical offset or background.) A bias frame or zero frame is an image frame taken with an effective exposure time of 0 seconds. This bias frame only contains the inherent system bias as well as the small scale structure in the noise. The number of zero frames required for reduction processes will be depending upon the noise structures present, with about 25 bias frames per observation night being a good average. These zero frames can be combined together to form a master bias frame which can in turn be subtracted from all the data (dark, flat and object) frames to remove the bias

Dark frame- All CCDs have some level of dark current. If we allow the CCD to integrate for some amount of time, WITHOUT any light falling on it, there will be a signal (and more importantly noise associated with that signal) caused by thermal excitation of electrons in the CCD. This is called the dark signal. The dark signal is very sensitive to temperature (lower temperature = lower dark signal), and that is why CCDs used in astronomy are cooled (often to liquid nitrogen temperature). Even with cooling, some CCDs have a non negligible dark current. To remove this dark count, as well as other constant noise patterns present, dark frames are used. A dark frame is an image frame taken with an integration time larger or equal to the longest exposure time used during the observation, and at the same operating temperature as the data frames, with the shutter closed. For increased effectiveness it is recommended that very long integration times be used, with at least 5-10 dark frames per observation night. These dark frames, after bias correction, can be combined together to form a master dark frame which can in turn be subtracted from the flat and object frame. (The dark frame and bias frames are *NOT* the same thing).

Flat frame- All CCDs has non-uniformities. That is, uniformly illuminating the CCD will NOT generate an equal signal in each pixel (even ignoring noise for the moment). Small scale (pixel to pixel) non-uniformities (typically a few percent from one pixel to next) are caused by slight differences in pixel sizes. Larger scale (over large fraction of chip) no uniformities are caused by small variations in the silicon thickness across the chip, non-uniform illumination caused by telescope optics. These can be up to maybe 10% variations over the chip. To correct for these, we want to shine a uniform light on the entire CCD and see what the signal (image) looks like. This frame (called a flat) can then be used to correct for the non-uniformities (we divide our images by the flat).
There are a number of ways of getting flat field frames. You should try several different methods and see which one works best for your telescope, telescope enclosure (or lack thereof), and CCD camera.

1- Twilight Flats

It have been found the twilight flat method to work well, at least for small chips on large telescopes, where the field of view is small. After the Sun sets, we take images of the twilight sky, which should be a reasonably uniform light source. But getting good twilight flats can be hard, particularly if a number of filters are involved. If you are waiting to observe, twilight seems to last too long- if you are trying to get good twilight flats, the sky seems to get dark too quickly, particularly if you have a number of different filters to get flats for! One good practice is to make a number of flat field exposures in each filter, moving the telescope between exposures. Then if stars appear in the frames, you can get rid of them by appropriately scaling and combining the images. To combine the images and remove the stars, we would use a median combine algorithm.

As chips are getting larger, and field of views are also getting larger, we have to worry about the fact that the twilight sky is not uniform in brightness- for example it is obviously brighter in the west, near the setting Sun than overhead. However, by picking the right position in the sky, we can minimize the gradient in the twilight sky brightness.
One common problem with doing twilight flats is the fact that soon after sunset the sky is much brighter in the west near the horizon than elsewhere in the sky. If your telescope has any scattered light problems, then the bright western horizon could easily mess up the twilight flats. One way to eliminate the problem is to observe twilight flats east of the meridian after sunset (or west of the meridian before sunrise!).

2- Dome Flats

Another way to get flat fields is to use a screen in the dome, and illuminate the screen with artificial lights. With effort, this also works well, although there are a number of potential concerns. One is that the electric lights used to illuminate the spot are quite a bit redder than the sky. This concern can be partially addressed by using very hot lamps, or using filters that decrease the red light from the lamps. For exacting work, particularly in the blue, you must also worry about the reflectivity of the screen. A screen may look fine and uniform with your naked eye, but be quite non- uniform at wavelengths less than the eye can see.

Data (object) frame- To take an image of an astronomical object, we point the telescope at the right place in the sky, and open a shutter to allow light to fall on the CCD. We allow the signal to build up (integrate) on the CCD for some length of time (anywhere from 1 second to 1 hour) and then read it out. The exposure time used depends on many things. The basic goal is to get an image of the source with the best signal to noise ratio (S/N) possible in a given amount of available telescope time. Now, since the signal is composed of photons, there is an unavoidable noise associated with photon counting statistics ("root N" noise, where N is the number of photons collected, to be discussed later in detail). There is NO WAY to get rid of this noise. However, by collecting more photons, we can improve the S/N (the signal goes linearly with time, while the noise goes as the square root of time). During the integration, the dark signal is also building up. We also have to worry about other sources of noise- readout, dark, and also the effects of cosmic ray particles, which give a spurious signal. The first thing to insure is that these other sources of noise are much less than the photon noise, so that we are not limiting ourselves unnecessarily. This argues for a long exposure. However, the presence of cosmic rays can argue for several shorter exposures which can then be combined to get rid of the cosmic rays.

** NOTE: the following applies to "professional" CCD systems **

SO the basic steps in observing and reducing a CCD image taken with a telescope are as follows:

Collect a number of bias frames - median combine them to a single low noise bias frame collect a number of dark frames (no light, finite integration time equal to the data frame integration). If dark current is non- negligible, combine dark frames into a single low-noise frame (After subtracting bias frame, of course) Collect a flat frame in each filter- flat frames can be made by pointing the telescope at the twilight sky, or by pointing at the inside of the dome. Bias frames (and dark frames, if the dark current is non- negligible over the time interval covered by the flat exposure) must be subtracted from the flat frame. The signal level in the flat is arbitrary- it is related to how bright the twilight sky was etc- all we need is the information on the differences of the signal across the chip. Thus, we normalize the flat so that the average signal in each pixel is 1.00 (we do this simply by dividing by the average signal). Subtract low - noise bias frame and low noise dark frame from object frame. Then divide this by the normalized flat frame.

$$\text{Reduced frame (output)} = \frac{\text{Raw object frame} - \text{bias frame} - \text{dark frame}}{\text{Flatfield frame}}$$

(For most modern professional grade CCDs the dark current is so low that we can ignore it- We always take dark frames, as one test just to make sure the system is operating properly.)

Data Reduction

Image obtained from a CCD camera is RAW image which needs to be cleaned. The image that has been taken by the telescope, the signal (count) which is measured at each pixel is a combination of the:

1) Astronomical flux from the signal.

2) The instrumental effects (systematic effects introduced by the optics, CCD, etc).

3) The effect of the atmosphere.

The goal of basic CCD data reduction is to remove the instrumental effects so that can be getting the "true" signal from the source. For doing these and many more other tasks, we have Astronomical Image Processing Packages software like IRAF - Image Reduction and Analysis Facility and MIDAS - Munich Image Data Analysis System. The instrumental effects consist of (Bias, Dark current and flat field).

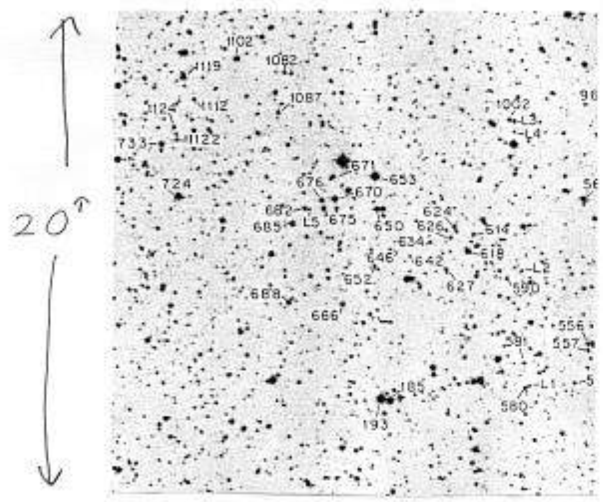
Standard Stars for Photometry

The primary standards for the UBV system are a set of 10 bright, naked eye stars of magnitude 2 to 5. The magnitudes of these stars define the UBV color system. You might think, then, that we should observe one or more of these primary standards along with our objects, so that we can use them to calibrate our photometry. Well, that doesn't work for two reasons: (1) The primary stars are too bright. Modern detectors even on small telescopes simply cannot deal with the flood of photons from naked eye stars! (2) As there are only 10 or so of these stars, they are not always well placed for observation.

Instead of using the primary standards directly, we use a series of secondary standard stars, or just standard stars, whose magnitudes have been carefully measured relative to the primary stars (to oversimplify slightly, a very small telescope is used to observe the primary standards and some stars somewhat fainter than the primary standards, then a larger telescope is used to observe those stars and much fainter ones). For broadband optical work (UBVRI filter system) the standard stars used most frequently today are from the work of the astronomer Arlo Landolt. Landolt has devoted many years to measuring a set of standard star magnitudes. Not the most scientifically glamorous project, but one for which we all give our thanks to Arlo every time we do photometry!

What makes a good standard star? (1) A standard star must not be variable! A variable "standard" star would be like measuring distance with a stretchy rubber ruler! (2) Standard stars must be of a brightness that will not overwhelm the detector and telescope in use, but must be bright enough to give a good S/N in a short exposure. (For very large telescopes, many of the Landolt stars are too bright.). (3) Ideally, a set of stars very close together in the sky will cover a wide range of colors. (4) Standard stars should be located across the sky so that they span a wide range of airmass. (Standards at the north celestial pole would not work for determining extinction!) In practice, the Landolt standard stars are located in the sky at declinations reachable by telescopes in both hemispheres, and are spread out in right ascension, so that there are always some standard stars at reasonable zenith angles. Most Landolt stars are near the celestial equator, and groups of stars are located approximately every 1 hour in right ascension. Selection of standard stars from the Landolt lists must be done carefully to get the best results.

A significant number of the Landolt stars have only been observed a few times, and should not be used as standards. In the Landolt lists, the quantity n indicates the number of observations of each star, and m the number of different nights the star was observed. The error in the mean is also listed. Stars with only a few observations, even if they have a small error, may be variable. The best standard stars combine well-observed stars covering a wide range of colors that fit on a single CCD frame for the telescope and CCD used. Hours spent looking through the Landolt catalog and finding charts (in the daytime, before going to the telescope!) will be repaid by a good selection of standard stars for your particular equipment and scientific program. Ideally, standard stars should be of a brightness to give a well exposed image in an exposure time of something like 10 to 30 seconds. As an example of a standard star field, Figure (1) shows a standard star field that we use a lot at the Steward Observatory 2.3 meter telescope on Kitt Peak. The table of numbers show the V magnitudes and colors of the stars as determined by Landolt.



Star	α(2000)	δ(2000)	V	B-V	U-B	V-R	R-I	V-I	n	m
98 193	06:52:04	-00 27 18	10.030	1.180	1.152	0.615	0.537	1.153	44	35
98 653	06:52:05	-00 18 19	9.539	-0.004	-0.099	0.009	0.008	0.017	65	50
98 650	06:52:05	-00 19 40	12.271	0.157	0.110	0.080	0.086	0.166	31	20
98 652	06:52:05	-00 21 58	14.817	0.611	0.126	0.276	0.339	0.618	2	1
98 666	06:52:10	-00 23 34	12.732	0.164	-0.004	0.091	0.108	0.200	25	14
98 671	06:52:12	-00 18 22	13.385	0.968	0.719	0.575	0.494	1.071	27	15
98 670	06:52:12	-00 19 17	11.930	1.356	1.313	0.723	0.653	1.375	32	19
98 676	06:52:14	-00 19 21	13.068	1.146	0.666	0.683	0.673	1.352	17	8
98 675	06:52:14	-00 19 41	13.398	1.909	1.936	1.082	1.002	2.085	44	21
98 L5	06:52:16	-00 19 39	17.800	1.900	-0.100	3.100	2.600	5.800	6	3

Figure(1)

Top: A portion of the SA98 field finding chart from Landolt article. The field shown is about is 20 x 20 arcmin. Bottom: Small portion of table from Landolt article. Columns are mostly self- explanatory, except for n, which is number of times each star was observed; and m, which is number of different nights each star was observed. The table also shows mean errors for each quantity for each star, but these have been omitted for clarity.

An Overview of Doing Photometry

You want to measure the magnitude and color for some object, say a newly discovered galaxy. How do you make the measurement? Here is an overview of the process.

1- Observe the galaxy and the standard star fields. You must observe one or more standard star fields containing stars of a wide range of colors to enable color transformation equations to be determined. You should observe the standard star field at both low and high airmass to enable determination of the extinction coefficients.
2- Reduce the CCD frames that are correct for bias, flat fielding, and dark current, if needed.
3- Measure the instrumental magnitude for the galaxy and the standard stars.
4- The Earth's atmosphere inevitably absorbs some of the visible light from every celestial object. The amount absorbed depends on atmospheric conditions, wavelength of filter used, and place in the sky where we observe each object. After determining the amount of light absorbed, using the multiple observations of standard stars at different airmasses, we must correct our instrumental magnitudes to what we would have observed outside the atmosphere (or at "zero airmass".)
5- The instrumental magnitudes, corrected for the extinction in the atmosphere, are still specific to our telescope and detector. To convert our numbers to standard system so that we can compare our numbers to those of astronomers around the world, we must derive transformation equations which relate the numbers measured by our setup to the standard system. The transformation equations are derived from the observations of standard stars with known standard magnitudes and colors.

Measuring Instrumental Magnitudes

You can think of a CCD image as simply a two-dimensional set of numbers, one number for each pixel in the image. After the initial reduction steps (bias and dark subtraction, division by a flat field image) the number at each pixel should be linearly related to the number of photons that fell on that pixel. The fact that the CCD image is naturally in a digital form makes computer analysis readily possible, unlike, say an image on a photographic plate, which would have to scanned to turn it into an image that could be dealt with using a computer. (Note: Real CCDs are not linear at all count rates. If we have too many counts in a pixel, the relationship between photons and counts becomes nonlinear. With increasing photons per pixel the CCD will eventually saturate, meaning that more photons will give no more counts. To avoid these problems, we have adjusted our exposure times so that the objects we are interested in do not produce count rates over the linearity limit of the CCD we are using.)
How exactly do we go about measuring the number of counts on our CCD image from a celestial body, say a star? There are several complications that must be understood and dealt with: (1) atmospheric seeing causes the star image to cover a number of pixels, and also causes the shape of the star image to vary with time (from one image to another). (2) the pixels that contain the counts from the star also contain light from the sky foreground (sky glow). The sky signal must be accurately measured and subtracted from the pixels containing the star + sky signal. (3) For many problems (e.g. measuring stars in a rich star cluster) the images overlap, and some way must be found to separate the light from different objects. This problem is often called contamination.

1- Point Spread Function (PSF) and Size of Star Images

For simplicity, let's assume that we have a CCD image of an isolated star, so we don’t have to worry about contamination. One fundamental concept we need to proceed is that of the point spread function (PSF). The PSF function is the shape of the CCD image of a point (unresolved) source of light. (Real stars are not precisely points of light- they must have some finite angular size. However, except for one or two very nearby, very large stars, the angular size of every star in the sky is far smaller than the diffraction limit of our optical telescopes, so we can treat stars as unresolved points). For all research telescopes, the dominant determinant of the PSF is smearing caused by the passage of starlight through the Earth’s turbulent atmosphere. This smearing is called seeing. (In practice, poor telescope focus and tracking errors can also contribute to degradation of the PSF, but these things can be minimized with good observing technique.) What does a PSF look like? Assuming good optics, proper focusing and tracking, the PSF should be circularly symmetric. Assuming circular symmetry, the PSF can be plotted as the flux vs. radius in a star image, as shown in Figure (1). The shape of a real PSF set by seeing is complicated, but can be approximated by a central Gaussian "core" and a large "halo" which can be approximated a power law. The angular size of the PSF can be characterized in several ways. One common measure is the full width at half maximum (FWHM) which is the diameter where the flux falls to half its central value. There is one fundamental fact that you should keep in mind: since the PSF is the shape of a point of light on the CCD, and since all stars are points, then all stars have exactly the same shape and size on the CCD. This statement almost always confuses the novice: don't brighter stars look bigger on images of the sky? They may look bigger, but that is caused by the following effect: On an image of the sky, either printed on paper or viewed on a computer monitor, the darkness of each pixel is related to the intensity in that pixel. Figure (2) shows the intensity along a line through a bright star and a nearby faint star. The shape of the faint and bright star are exactly the same, we are simply looking at a larger diameter at a given intensity for a bright star than for a faint star. Another important point, related to the above, is that the PSF does not have an edge. The intensity of the star fades smoothly to zero with increasing radius, but there is no place that we could call an "edge". You can see this by looking at the image of a bright star on a CCD image.

2- Aperture Correction

The fact that the PSF does not have an edge raises an important question: If we want to measure all the light from a star, how far out in radius do we have to go? (Another way of asking this question is: How big an aperture - in pixels or arcsec - do we want to use when we measure the counts?) One logical answer might be: as big as possible, to get "all" the light from the star. Well, this is not a good answer for several reasons: (1) using a big measurement aperture also means that there will be a lot of sky light contributing to the counts in the aperture containing the star. Now, as we will see, we can subtract off the average sky signal, regardless of aperture size, subtract off the noise associated with the sky signal.

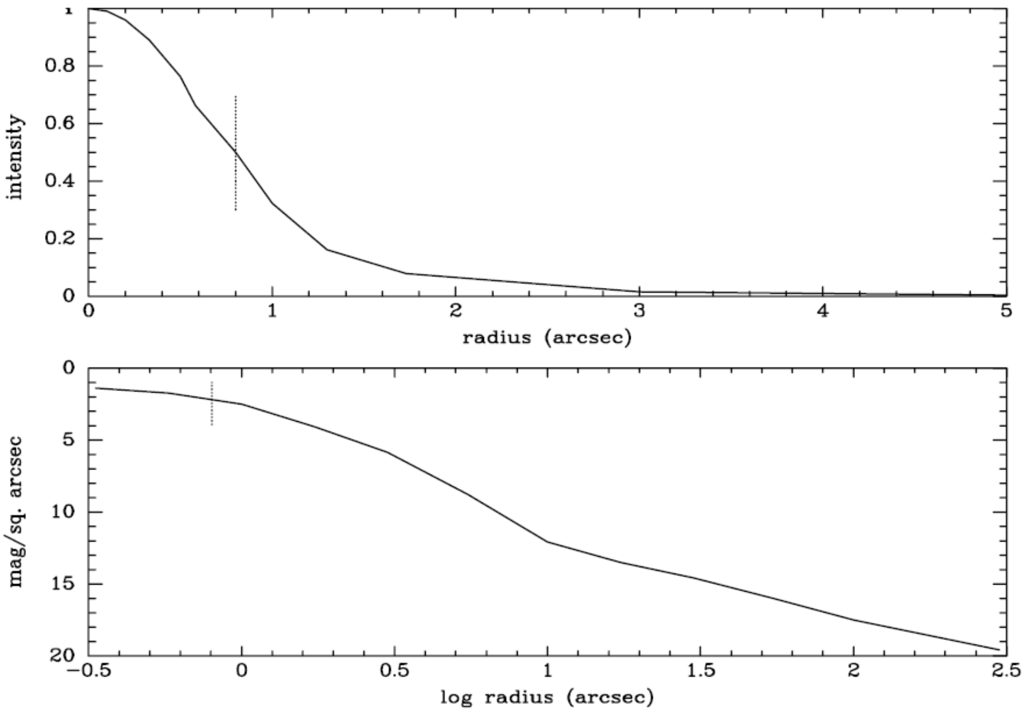


Figure (1)

Typical Stellar Point Spread Function (PSF). This particular PSF is of a star in seeing of about 1.6 arcsec FWHM. Top panel: Radial intensity of light (normalized to center= 1.0) for this PSF. On this plot, the HWHM, the radius at which the PSF drops to half its central or peak value, is marked by the vertical dotted line. Bottom panel: This is a plot of surface brightness (mag/ sq. arcsec) vs. logarithm of radius from the center of the image. The HWHM is again marked by the vertical dotted line. Note that the lower panel shows the PSF to a much larger radius (out to over 300 arcsec) than shown in the top panel (5 arcsec). The logarithmic y axis allows us to see the very faint outer regions of the PSF at large radii that would be invisible in the top panel linear intensity display. Note that the PSF simply fades out- there is no sharp edge where the PSF reaches zero intensity.

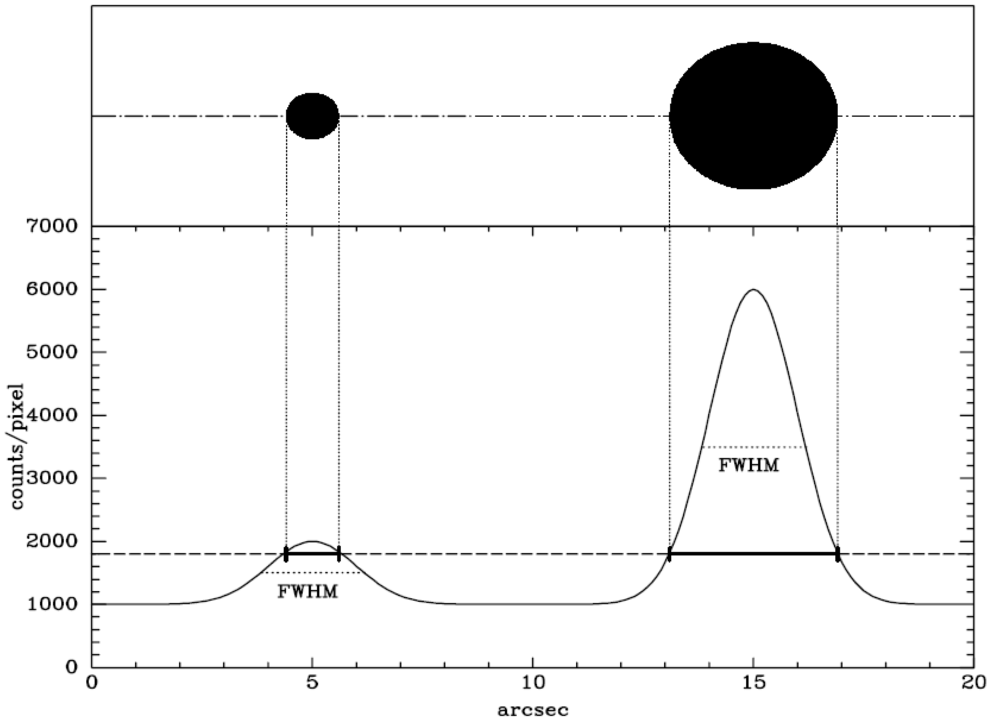


Figure (2)

Why bright stars look bigger than faint stars, even though all stars have same image shape and size. Top: Very schematic image of a faint star and a bright star (bright star 5 times flux of faint star), showing brighter star looking bigger than fainter. This “image” maps all pixels with value less than 1800 to white, and all above 1800 to black. (Ignore the slight ellipticity of the stars.) Bottom: Brightness profile along dot-dash line crossing the centers of the two stars. The shapes of the two stars are exactly the same, the bright star is simply 5 times the intensity above sky at each point relative to the faint star. Brightness along the dashed line at constant level of 1800 counts/ pixel across image shows that, while the bright and faint star have the same shape (same PSF), the bright star looks bigger at each gray level on the image. The dotted line on each star profile marks the FWHM of each star (here the FWHM is about 2.3 arcsec). The solid line intersecting each profile at counts/ pixel = 1800 shows the size of the star on the “image” in the top panel. Note: In a real CCD image and plot, you would be able to see the individual pixels, so that the edge of the star image would be a set of squares, and the intensity profile along a single line would show a “stair step” pattern of individual pixels.

The sky noise in the aperture containing the star will contribute to a larger noise in the measurement of the signal from the star. (2) The bigger the measuring aperture, the more chance that we will have light from objects other than the one we want to measure in the aperture as well as the light from our target object. This is called contamination.

Both effects mentioned above argue for using a small measurement aperture. But, you might well protest, a small aperture will only encompass a fraction of the total light from the star. This is true, but, if the seeing were constant, any aperture would measure the same fraction of light for any star, and when comparing one star with another (which is essentially what photometry is we are comparing unknown stars to standard stars) the effect would cancel out. The problem, of course, is that seeing is not constant. A small aperture might measure 0.5 of the total light from a star on one CCD image, then, if the seeing worsens, the same size aperture might measure only 0.4 of the light from the star on the next CCD image. In practice, we find that seeing affects mostly the inner Gaussian core of the image. Using an aperture 4 to 10 times the diameter of the typical FWHM will get most of the light. In this size aperture, reasonable variations in the seeing will not result in measurable variations in measured counts. However, particularly for faint objects, an aperture, say, 4 times the FWHM will contain a lot of sky signal, which means a lot more noise inevitably associated with the sky signal. Since the signal of a faint object is low, this will result in a low S/N ratio. For bright objects (much brighter than the signal from the sky in the measuring aperture) the sky noise is not much of a problem. This, plus the fact that all stars on the same image have the same PSF, suggest a technique called aperture correction, which greatly helps in obtaining good S/N for faint stars, and also helps greatly in crowded fields. Say we have an image with some faint objects we want to measure and at least one bright star. If we measure the bright object in a small aperture (say radius = 1 FWHM) and also in a bigger aperture which gets “all” the light (say 4 FWHM) we can easily find the ratio of light in the small to large aperture (which we express - of course - as a magnitude difference).

Say we measure an instrumental mag mI(1 FWHM) in the small aperture and mI(4 FWHM) in the large aperture. The aperture correction is defined as:

$$\Delta = mI(4\ \text{FWHM}) - mI(1\ \text{FWHM})$$

(as defined Δ _ is always a negative number- this simply means that there is more light in the larger aperture than in the smaller aperture.) The optimum size of the small aperture has been studied by several authors. For faint objects, where the sky noise dominates, an aperture about as big as the seeing FWHM appears optimum. How do we use the aperture correction? Let's say we want the total counts from a faint star. If we simply measured the star with a large aperture, we would get a poor S/N, because the star signal is very low, except in the center of the image, and the sky noise would result in a low S/N. If we measure with just a small aperture, we will miss a goodly fraction of the light (the light in the outer regions of the star image may be hard to see, because its lost in the sky noise, but the light is there and must be counted for a proper measurement!) However, we can use the aperture correction derived from a bright star on the same image to correct the small aperture measurement of the faint star for light outside the small aperture:

$$\text{total} = mI(1\text{FWHM}) + \Delta$$

Here, "total" is our estimate of the total instrumental magnitude in the faint star, mI(1FWHM) is the measured number of counts in the small aperture for the faint star, and Δ is the aperture correction derived from a bright star in the same frame. How big should the small aperture be? Too small an aperture will result in a poor S/N because too few photons in the star will be detected- too big an aperture will result in a poor S/N due to the inclusion of too much sky noise. There must be an optimum aperture size that gives the maximum S/N. The optimum aperture seems to be achieved when the measurement aperture has a diameter about 1.4 × FWHM of the PSF. At this aperture, the aperture correction is about −0.3 mag. However, the S/N does not appear to be too sensitive to the exact small aperture size.

Uncertainties and Signal to Noise Ratio

All scientific measurements should carry a measurement or estimate of the uncertainty or error of that measurement. If we measure the magnitude of a star a number of times, we will not get exactly the same number each time, due to various sources of noise. To prove, for instance, that a star is variable in brightness, we would have to find a change of magnitude several times larger than the uncertainty in our magnitude measurement. Scientists use the word "error" interchangeably with "uncertainty". In everyday speech, "error" connotes some sort of mistake. That is certainly not the connotation to be attached to the word error as used in the scientific context, where it means uncertainty. The error is due, not to mistakes, but to noise. A crucial concept in photometry is the signal to noise ratio (S/N or sometimes SNR). This is one way of indicating the accuracy of our measurements.

1- Quantitatively investigate the S/N

For a CCD measurement, there are several sources of noise, as discussed in the section on CCDs. These include photon noise, readout noise, dark current noise, and processing noise (noise in flats). However, for most broadband observations, the photon noise of the sky dominates over all other sources of noise. For now, let us assume the only source of noise in our photometry is photon noise.

Because photons obey simple statistics, discussed below, we can make a full quantitative analysis of the S/N of a CCD magnitude measurement, if photon noise is the only important noise source. We can quantitatively investigate the S/N of our photometric measurements using the basic idea of counting statistics of random events. Consider a signal involving discrete elements arriving at random e.g. raindrops falling on a square meter of pavement or photons "falling" on a pixel of a CCD. In a "steady" rain (of raindrops or photons) the number hitting a given area in a given amount of time would not be precisely identical from second to second. Instead the number of raindrops or photons hitting an area per second would vary from second to second. The best way to show such a process is with a histogram graph showing the results of repeated measurements. This is a plot of the number of raindrops or photons hitting the area per interval (on the x axis) vs. the number of times that each number is observed (on the y axis). Note that the quantities on both the x and y axes are integers. A histogram of raindrops or photons in a steady downpour might look something like Figure (3).

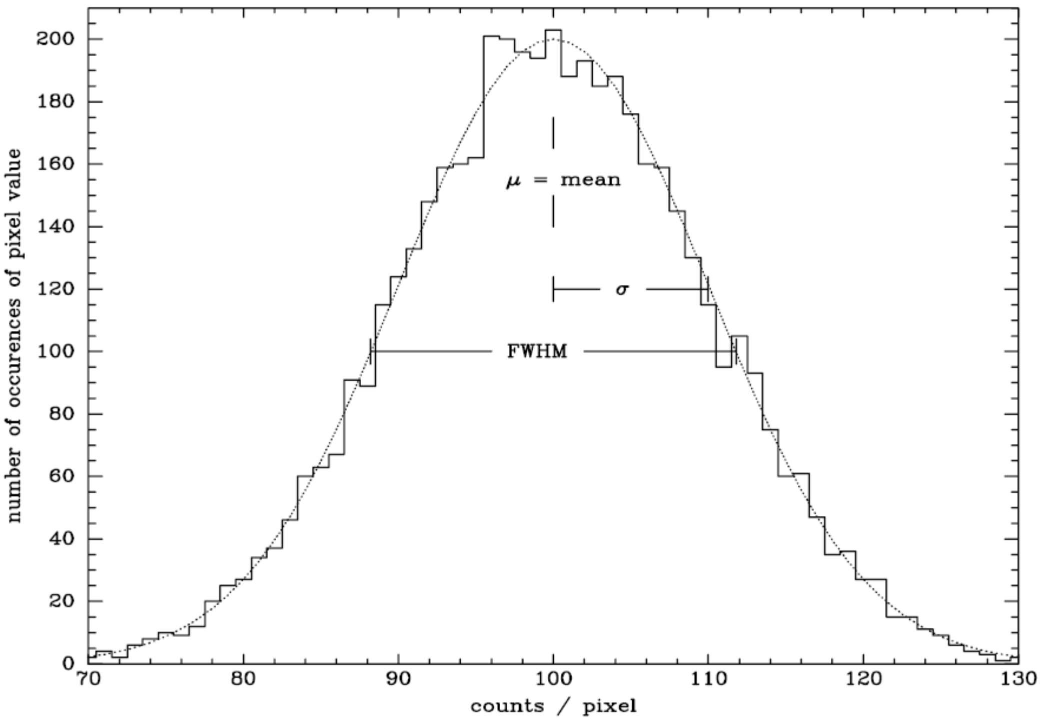


Figure (3)

Histogram of photons falling on the pixels of a CCD. We assume the CCD is uniformly illuminated and that the CCD is “perfect”- i.e. each pixel is exactly the same. In this example the mean number of photons hitting each pixel is 100. The histogram shows (at least schematically) the distribution of the counts measured on a number of pixels of a single exposure of the CCD. The smooth dotted line is the “best fit” Gaussian. If we make repeated measurements of a single pixel, with steady (not varying with time) illumination, we would get a similar distribution. Note that this graph may look like a PSF plot, but this is a histogram, not a plot of something vs. position or time.

The shape of this curve can be approximated by a Gaussian or normal distribution function:

$$P = \frac{1}{\sigma\sqrt{2\pi}} \exp[-\frac{1}{2}\left(\frac{X-\mu}{\sigma}\right)^2]$$

Here μ is the mean or average, σ is the standard deviation. σ is also called the root mean square (or rms) deviation.
For photons, which obey counting statistics, the scatter σ (width of the histogram) would be related to the number of photons counted (n) by the following:

$$\sigma = \sqrt{n}$$

It is important to note that n is simply the number of photons that we count in our observation. It is not the number of photons per second, or per area. If we counted 1000 photons in minute with a large telescope or 1000 photons in a hour with a small telescope, the σ would be the same. Repeated observations would show a distribution of values with a mean of 1000 and a scatter σ of about 32.
(Note: For very low means, less than a few dozen or so, the histogram would not be symmetric because you obviously cannot have negative counts, and you would find a histogram shape approximated by a Poisson distribution, instead of a Gaussian. For all astronomical measurements we will discuss, the count rates will be such that the histograms will be approximated by a Gaussian).
Now we can easily see the relation between signal and S/N. If n photons are counted, the noise is $\sqrt{n}$, so that

$$S/n = \frac{n}{\sqrt{n}} = \sqrt{n}$$

2- Application to Real Astronomical Measurement

So, the noise is just the square root of the photons measured. Note that this applies to photons actually detected. If a million photons hit your detector, but your detector has a QE of 1% so only detects 1% of the incident photons (or 10,000) your S/N is 100 (= $\sqrt{10^4}$), not 1000(= $\sqrt{10^6}$).

If we detect n photons from a star, and those are the only photons we detect, we have indeed measured it with a S/N of “root n”. As usual, that is not the whole story. The problem arises because we cannot measure just the light from the star alone- we also get photons from the sky background. So, we have to measure the star+sky, then measure the sky contribution separately, so that we can subtract the sky contribution to get the star alone. The trouble is, BOTH these measurements carry errors which combine when we try to isolate the counts from just the star. For one measurement, we orient the telescope so that the star is centered in the aperture for the other we move the telescope slightly so that the aperture receives only light from the sky near the star Figure (4).

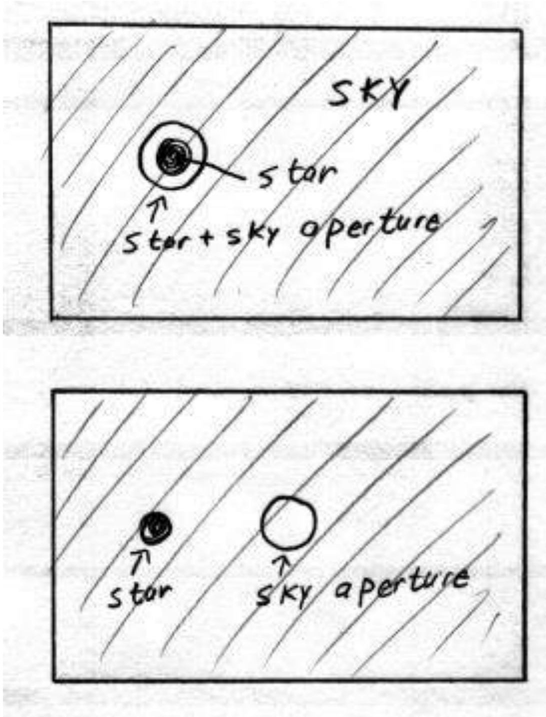


Figure (4)

Sky+star aperture and sky aperture. This is the simple "photometer model" of measuring the sky background, as the star+sky and sky are measured with the same circular aperture.

We can quantitatively analyze the S/N of this observation as follows: the quantity we want to measure is C_{star}, the number of counts from the star alone. However, we can only directly measure the sum of the counts from the star and sky ($C_{star+sky}$) in the star+sky aperture and the counts from the sky (C_{sky}) in the sky aperture.

Obviously, the quantity we want is simply:

$$C_{star} = C_{star+sky} - C_{sky}$$

The noise in star+sky aperture is then $\sqrt{C_{star+sky}}$ and the noise in the sky aperture is $\sqrt{C_{sky}}$.

The noise or uncertainty in the measurement of C_{star}, using the usual rules for propagation of errors, is:

$$N = \sqrt{C_{star} + 2C_{sky}}$$

So that the general equation for the S/N is:

$$^{S}/_{n} = \frac{C_{star}}{\sqrt{C_{star}+2C_{sky}}}$$

We need for the "2" in the above equation because we are forced to observe the sky twice- once along with the star, and once by itself. If the star signal is large compared to the sky signal, then the noise simplifies to $\sqrt{C_{sky}}$, simply "root n". When the star signal is small compared to the sky signal (the usual case for measuring faint stars), the noise is completely dominated by the sky brightness, and can be approximated by $\sqrt{2C_{sky}}$.

To get a better signal to noise, we can always increase the signal (by using a larger telescope or a longer integration time), or, we can try to decrease the noise. One obvious way to decrease the noise is to observe from a darker place. Getting a dark sky is the reason observatories are built far from city lights, and why astronomers prize "dark time", the night time hours during which the Moon is not lighting up the sky.

How does the S/N change with integration time? We can write the signals C_{star}, = t R_{star}, , and C_{sky}= tR_{sky} , where t is the integration time in seconds, and R the count rate in counts s−1. Doing the same for the sky aperture, we can write the S/N as:

$$^{S}/_{n} = \frac{t\,R_{star}}{\sqrt{t\,R_{star} + 2tR_{sky}}} = \sqrt{t}\,\frac{R_{star}}{\sqrt{R_{star} + 2R_{sky}}}$$

so that the dependence of S/N on time is:

$$S/N \propto \sqrt{t}$$

The S/N goes as the square root of the integration time. To improve the S/N by a factor of 2 , we must observe 4 times as long. The longer we observe, the greater the sky signal (and hence sky noise), but the S/N increases because the star signal increases linearly with increased exposure time, while the sky noise increases only as the square root of the exposure time.

Filters

1- Glass and Interference Filters

Filters are used to restrict the wavelengths of EMR that hit the detector. For optical CCDs, there are two main types of filters: colored glass and interference filters. Colored glass filters use chemical dyes to restrict the wavelengths that pass. Most colored glasses are cutoff filters, that is they pass all light above (or below) some particular wavelength. To make a bandpass filter, such as one of the UBV filters, which have both high and low wavelength cutoffs, two glasses are combined. For instance, to make a V filter, which passes light from about 5000 to 6000 °A, we combine a filter which blocks light below 5000 °A, but passes light with higher wavelengths, with a filter that passes light below about 6000 °A, but blocks longer wavelengths. See Figure (5) for a transmission curve of a typical V filter, showing the transmission of the individual glasses and the transmission of the two glasses together. The glasses are usually cemented together with optical cement, which eliminates two air- glass interfaces.
Colored glasses cannot be used to make filters with narrow wavelength bands, because the transition from transmission to blocking provided by the dyes in glass filters is usually one hundred °A or more in width. To make narrow filters, or ones with sharp steep edges to their transmission curves, interference filters are used. These filters are composed of several layers of partially transmitting material separated by certain fractions of a wavelength of the light that the filter is designed to pass. Light of different wavelengths is either reflected by or passed by each layer due to interference effects, which of course depend on wavelength. By using several layers of material, filter makers can make a wide variety of filter widths and central wavelengths. A typical use for an interference filter is to image objects in the light of one particular emission line. Figure (6) shows a typical interference filter set that would be used to image the Hα emission from HII regions in a nearby galaxy. The on band filter passes Hα light, plus continuum light that has wavelength in the range passed by the on band filter. To make a pure emission line image, we would use an off band image, with wavelength above or

below Hα, that does not permit Hα light to pass. The off band image would be used to make an image of the continuum light alone. Then we could use the continuum off band image to estimate what the continuum light in the on band image looked like, and subtract this image from the on band image, leaving a pure emission line image.

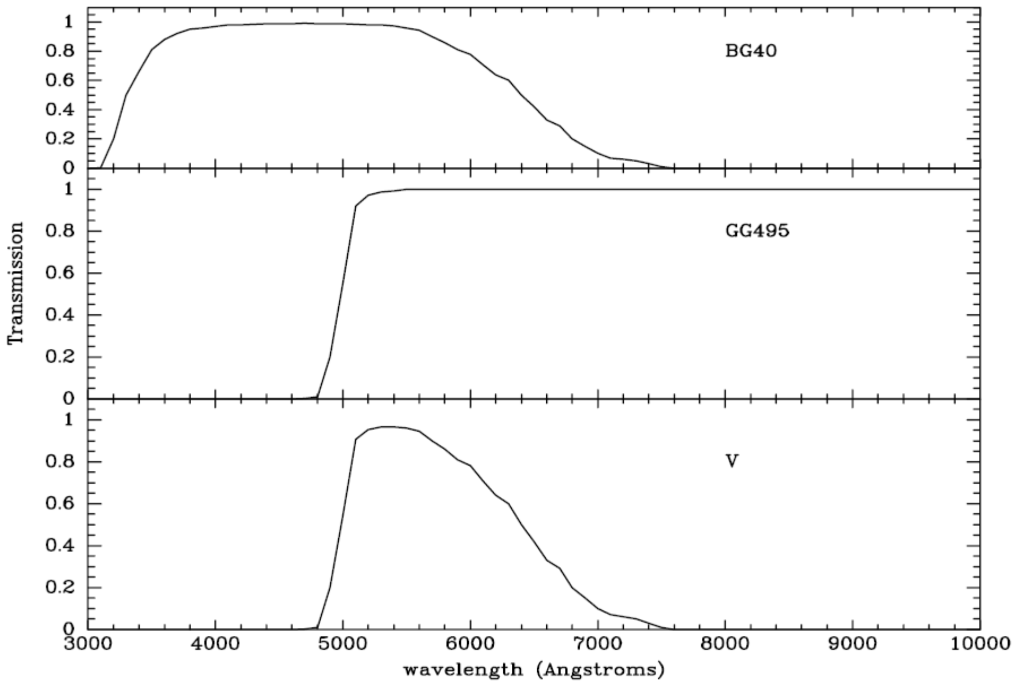


Figure (5)

The transmission of a typical V filter, made up of two pieces of colored glass. Top: the transmission curve of a piece of BG40 glass. Middle: the transmission curve of GG495 glass. Bottom: The transmission curve of the V filter made up of the two pieces of glass together.

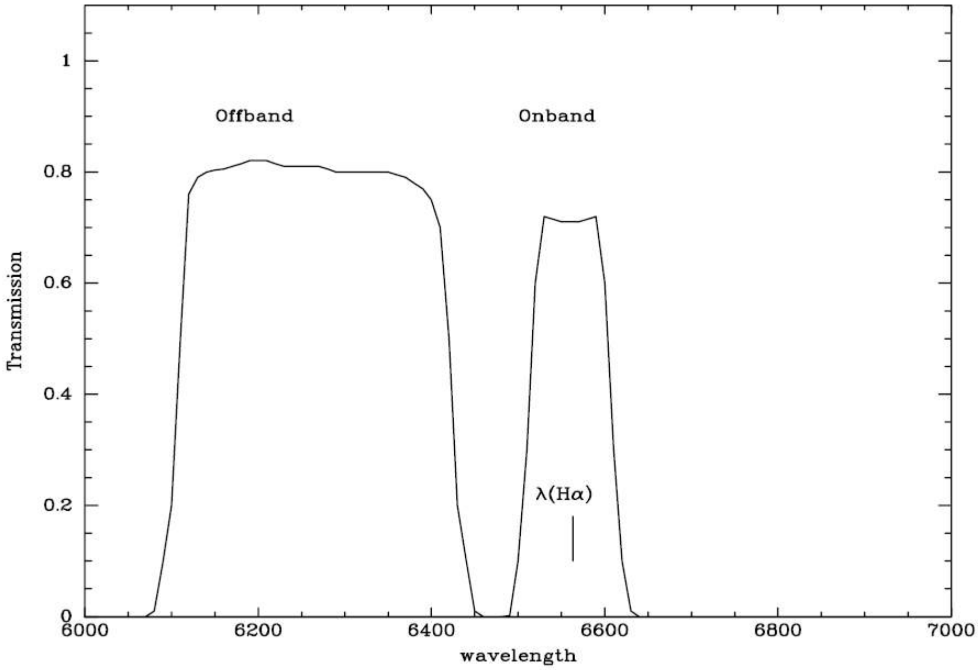


Figure (6)

The pass bands of a typical set of filters used to image Hα radiation. The on band filter passes Hα and continuum, while the off band filter only passes nearby (in wavelength) continuum light. When observing external galaxies, a set of on band filters with different central wavelengths is used for galaxies in different redshift ranges, as the observed wavelength of Hα is different for galaxies of different redshifts. These filters would be interference filters, not colored glass like the V filter in the previous figure, as colored glass filters cannot give such sharp changes of transmission with wavelength.

2- Filter Systems

A filter system is a set of filters of specific central wavelengths and widths, along with standard stars which have carefully measured magnitudes in these filters. Although the UBVRI system(s) (there are several different versions, particularly of the R and I filters) is/are the best known optical system, there are a number of others. Some were specifically designed to solve a particular astrophysical problem, others to mesh with particular detectors. One system that will probably become important in the future is the u'g'r'i'z' that is being used by the Sloan Digital Survey. This CCD survey of a significant fraction of the sky should produce accurate magnitudes and colors for many millions of objects. There are several recipes for making the standard UBVRI filters using various colored glasses.

To accurately mimic the standard pass bands, the glasses must be chosen so that the product of the filter transmission, QE curve of the detector, and atmospheric transmission match the standard pass band. It is impossible to precisely match the standard curve, but that is not necessary as discussed earlier, observations of standard stars are used to calibrate the differences between the standard pass bands and the ones we use.