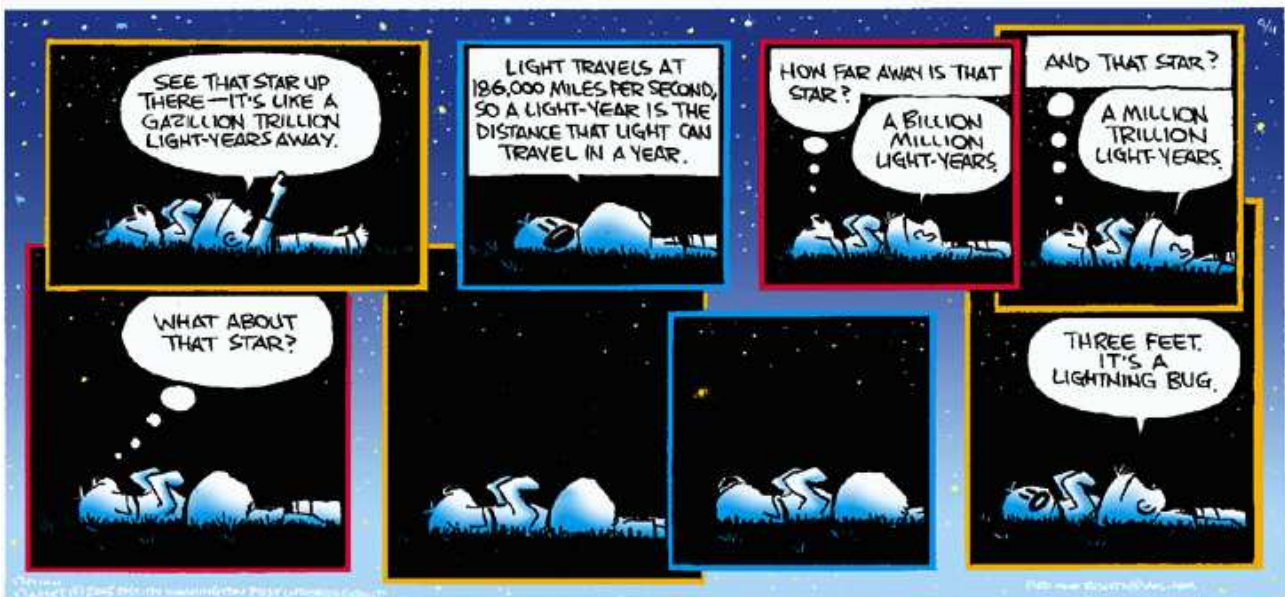


Photometry – I. “All sky”

Dave Kilkenny

Red and Rover by Brian Basset



Contents

1	Review of basic definitions	3
1.1	Apparent Magnitude	3
1.2	Absolute Magnitude	3
1.3	Bolometric Magnitude and effective temperature	3
2	Elements of “all sky” photometry	4
2.1	Atmospheric Extinction	4
2.2	Standard stars	10
2.3	Colour Equations and Zero Points	10
3	Some Photometric systems	15
3.1	Filters	15
3.2	UBVRI	18
3.3	JHKLMN	19
3.4	Strömngren uvby and $H\beta$ photometry	20
3.5	SDSS - $u'g'r'i'z'$	23
4	Applications	26
4.1	Visual	26
4.2	Photographic surveys	26
4.3	Calibrations	26
4.4	Interstellar reddening	27
4.5	Metallicity	30
4.6	Cluster sequence fitting	31
4.7	Variability	35

1 Review of basic definitions

1.1 Apparent Magnitude

Recall that the magnitude scale is defined by:

$$m_1 - m_2 = -2.5 \log \frac{I_1}{I_2}$$

or:

$$m = -2.5 \log I + \text{constant}$$

which is sometimes called **Pogson's formula**.

1.2 Absolute Magnitude

Applying the inverse-square law for the propagation of light and setting the standard distance to be 10pc gives the expression relating apparent and absolute magnitude to distance (d):

$$m - M = 5 \log d - 5$$

or parallax (π):

$$m - M = -5 \log \pi - 5$$

The absolute magnitude of a star is the apparent magnitude the star would have if seen from a standard distance of 10 parsecs.

1.3 Bolometric Magnitude and effective temperature

The **Bolometric magnitude** of a star is the magnitude integrated over all wavelengths.

$$M_{bol} = M_V + BC$$

where the subscript V refers to the “ V ” or “Visual” passband (a yellow-green filter) of the UBV or $UBVRI$ photometric systems, and BC is the **bolometric correction** which is determined as a function of temperature or spectral type.

2 Elements of “all sky” photometry

We have seen that we can define magnitude using the Pogson formula

$$m = -2.5 \log I + \text{constant}$$

In the case of a photomultiplier system, for a given passband, we will have :

- A measure of the star (plus background “sky”) and a measure of the sky, as a number of “counts” in a time interval, t .
- Both can be converted to a (counts/sec) rate and corrected for the pulse-coincidence loss.
- Then (star + sky) – (sky) gives a star count which can be converted to a magnitude.

And for a CCD :

- The CCD frame can be corrected using the “flat fields”
- The data for a given star (or many stars) can be extracted using
 - a profile-fitting technique – where the brighter stars on the frame are used to determine a mean profile which can then be fitted to all stars
 - an aperture technique – analogous to the photoelectric method – can be used by extracting all data within a given area and then subtracting an equal area of carefully chosen “sky”
- The extracted number can be converted to magnitude

Before we can go further, however, there is one important correction that must be made – certainly in the case of “all sky” photometry and usually for other techniques as well. This is:

2.1 Atmospheric Extinction

This involves a complicated set of processes which we can generally deal with adequately in a simplified way. But first ...

The atmosphere affects incident starlight in a number of ways. As well as the effects of scintillation, “seeing” and so on, there is:

- **Absorption by molecules.** This is particularly troublesome in the infrared, but also affects the red end of the visible spectrum.
- **Scattering.** This is complex and involves:
 - **Rayleigh scattering** – scattering by molecules – which is approximately proportional to $\lambda^{-4.1}$
 - **Aerosol scattering** – includes scattering by dust (meteoritic and volcanic), pollution, organic products from trees and plants, and even sea salt from breaking waves. Scattering from small *spherical* particles (comparable size with the wavelength of light) is known as **Mie scattering** and is roughly proportional to λ^{-1} . Scattering from large particles is almost independent of wavelength.

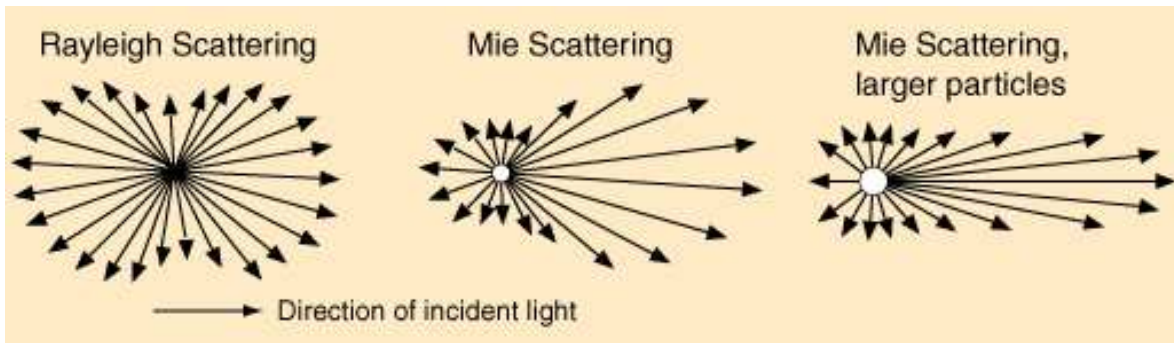


Figure 1: Schematic representation of Rayleigh and Mie scattering.

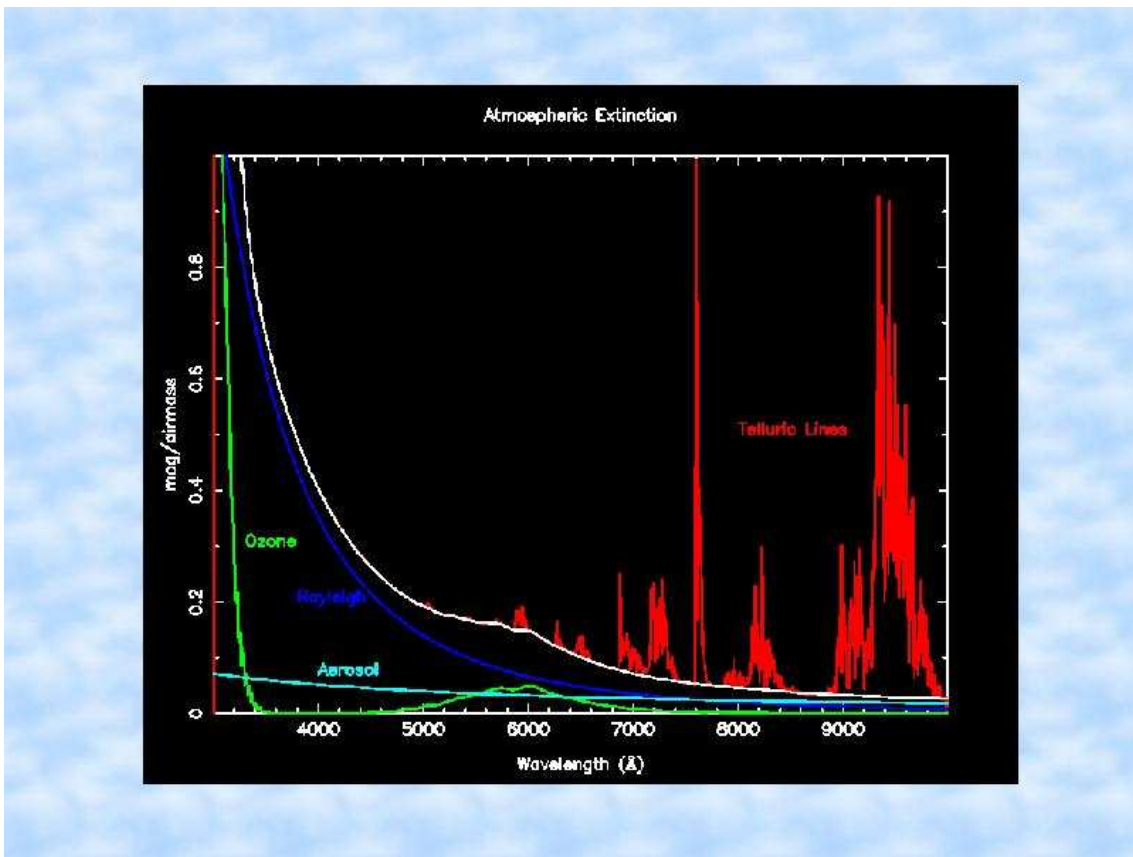


Figure 2: The components of the atmospheric extinction.

The wavelength dependencies noted above mean that the extinction increases rapidly as we go towards the ultraviolet – and luckily for us – with ozone absorption cutting off most of the light shortwards of about 3200\AA (320nm). It also accounts for the fact that the Sun looks redder as it sets and that the sky is blue.

If a light beam of intensity I passes through a thickness of material dX , with an opacity, τ (related to the absorption). The extinction loss – the amount absorbed/scattered (dI) is:

$$dI = -I \tau dX$$

so

$$\frac{dI}{I} = -\tau dX$$

and integrating

$$\log I - \log I_0 = -\tau X$$

where I_0 is the incident beam (outside the atmosphere) and I is the final beam (at the telescope). So the observed magnitude is:

$$m = -2.5 \log I = -2.5 \log I_0 + 2.5 \tau X$$

so

$$m = m_0 + k X$$

where the k is called the extinction coefficient – and is, of course, wavelength dependent – and X is the total path length traversed by the light.

Observing a star as it rises or sets and plotting the uncorrected magnitude, m , against air mass will give a (more-or-less) straight line graph with a gradient equal to the extinction coefficient and an intercept on the magnitude axis (at $X=0$) equal to the “outside atmosphere” magnitude of the star. We can thus correct all data to effectively zero extinction – though it might sometimes be preferable to correct all data to, say, 1 air mass.

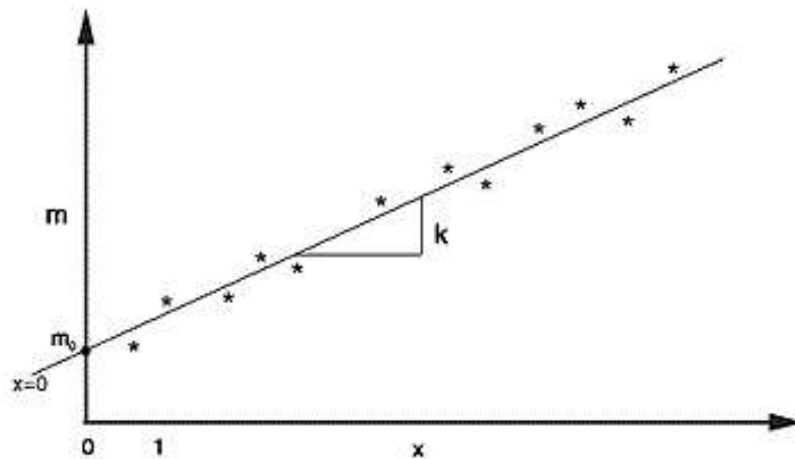


Figure 3: A **Bouguer line** – a plot of observed magnitude (m) vs. airmass (X).

From the figure, which assumes a plane-parallel atmosphere, it can be seen that the path length light travels through the atmosphere is $h \sec z$ where h is the zenithal path length and z is the zenith angle or **zenith distance**, easily calculated for any instant from the location of the observer, the (α, δ) of the star and the local sidereal time.

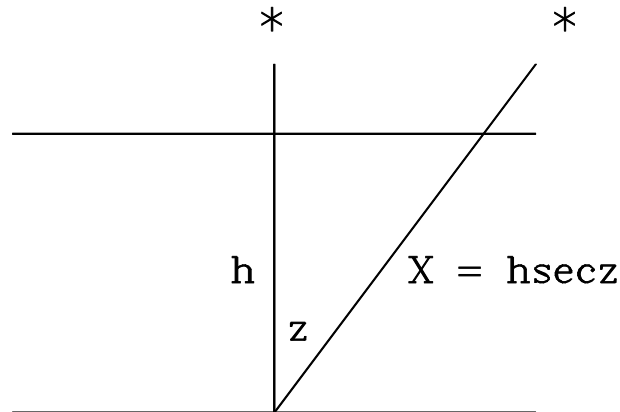


Figure 4: Schematic for the first order determination of air mass, X .

If we say that the zenithal path length, h , is **one air mass** then $\sec z$ is a good approximation for X for zenith angles smaller than about 60° (2 air masses – or twice the zenithal path).

The extinction coefficient, k , is usually given in magnitudes per air mass, so the above equation is very simply applied to correct any observations.

Note that the above derivations make a number of assumptions:

- That the atmosphere is plane-parallel – when it is a spherical shell.

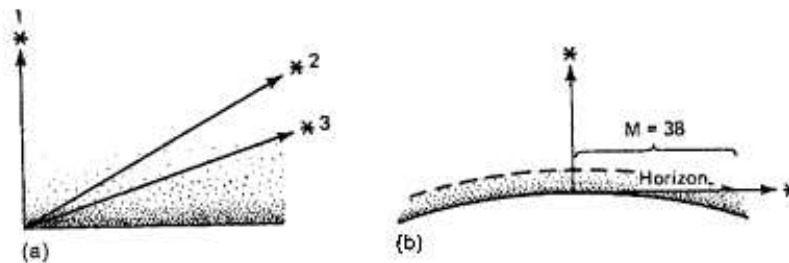


Figure 5: Schematic for the (a) plane-parallel and (b) spherical effect of the atmosphere/air mass.

- That the path through the atmosphere is straight – when it is *slightly* curved by refraction.
- That the atmosphere is uniform and unchanging (and you know that's not true !)
- That we are dealing with monochromatic light when, in fact, we are often dealing with passbands of the order of 1000\AA (100 nm).

These potential problems can be dealt with practically, for example:

- More exact equations can be used. for example, Hardie gives:

$$X = \sec z - 0.0018167 (\sec z - 1) - 0.002875 (\sec z - 1)^2 - 0.0008083 (\sec z - 1)^3$$

and Young & Irvine propose:

$$X = \sec z (1 - 0.0012(\sec^2 z - 1))$$

- We can apply small “drift” corrections to the data from a given night – effectively correcting for *small* extinction variations by allowing the zero-point of the data to change
- We might need to use “colour” terms (equivalent to allowing for different stellar temperatures) by for example re-writing the extinction correction as:

$$m = m_0 + k X + k' (\text{colour})X$$

and the k' is a second-order coefficient and “colour” is some temperature-related quantity such as the $(B - V)$ colour.

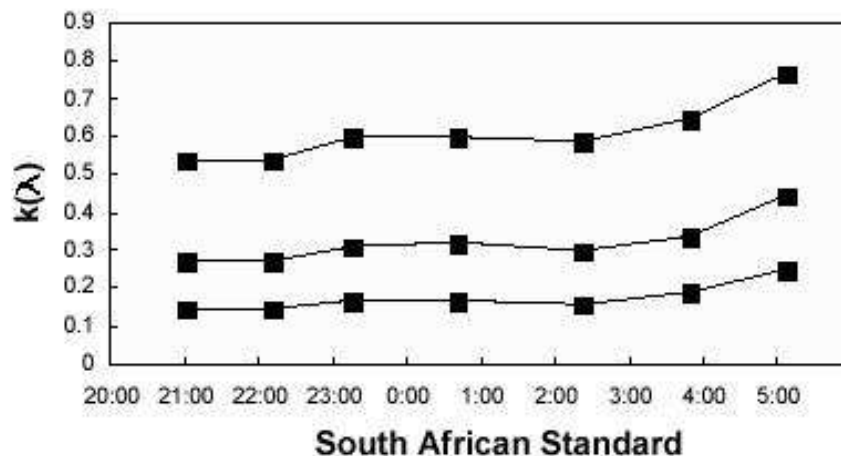


Fig. 2. Variations in the total k_U , k_B and k_V extinction coefficients (from top to bottom) during the night 29–30 September 1997.

Figure 6: An unusual night at Sutherland – attributed to pyrogenic aerosols (Winkler).

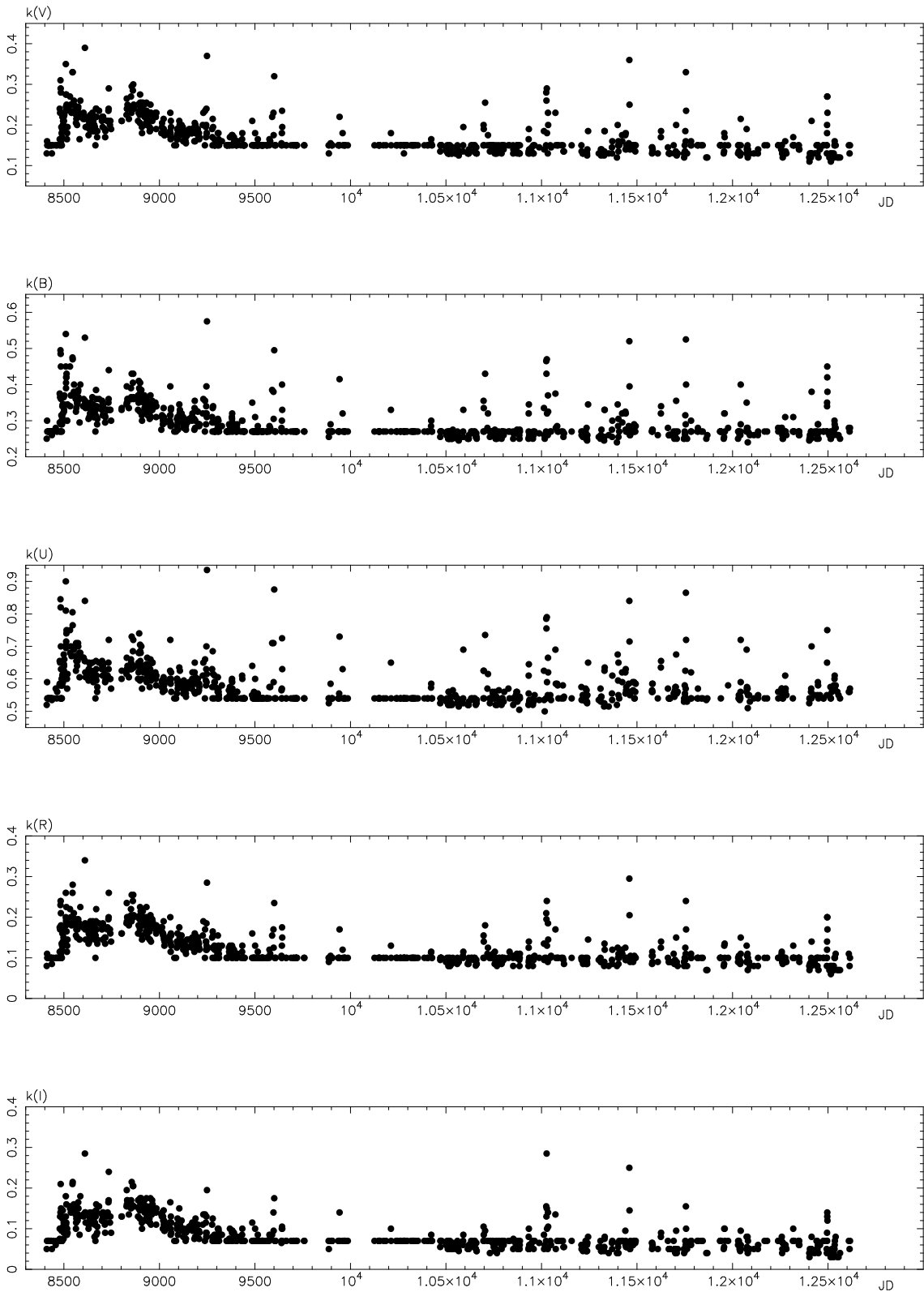


Figure 7: Extinction coefficients (k_U, k_B, k_V, k_R, k_I) from June 1991 to the present as measured on the SAO 0.5m telescope. The enhanced extinction at all wavelengths around JD 8500–9100 was due to volcanic ash aerosols from the Mt Pinatubo (Philippine Islands) explosion.

2.2 Standard stars

So far, we have looked at the concept of magnitude in a rather generalised way. When we come to actually make photometric measurements, we usually need to use a specified system. This is because different observers will often use different sized telescopes at different sites (which means different altitudes and therefore somewhat different extinction coefficients) and even with filters cut from the same piece of glass, you cannot guarantee that they will be identical. After a few years, you might not even be able to get the same glass that was used previously. Detectors are also not really uniform; CCDs are much more red-sensitive than photomultipliers and different types (of either) might have significantly different responses as a function of wavelength.

At the same time, when we measure a star from Sutherland, we want to get the same answer that would be obtained by someone working in Chile, for example. Otherwise, the whole procedure is pointless.

To get around these problems, we use systems of **standard stars** – stars which have been observed many times, have been carefully transformed on to a homogeneous and reproducible system and from which variable stars have been eliminated (as far as possible).

The main sources of standard stars for probably the most used photometric system in the southern hemisphere (the $UBV(RI)_C$ system) are the “E region” standards set up largely by Alan Cousins at SAAO (see also www.sao.ac.za/news/awjc.html) and the “equatorial” standards set up by Arlo Landolt at CTIO. The Cousins stars (all near declination -45°) really define the southern system; the Landolt standards are very useful because they are available from both hemispheres and because they include numbers of fainter stars, and also stars close together on the sky – both of which make them more suitable for CCD operations.

Standard stars can of course be used to determine extinction coefficients as described above. They are also used to determine the so-called “colour equations” and zero-points which enable us to transform data to a standard homogeneous system and thus compare data from different sources.

2.3 Colour Equations and Zero Points

We illustrate the fundamentals of colour equations using the UBV filters from the $UBVRI$ system. Transmission curves from typical UBV filters are shown in the figure.

When photographic plates came into widespread use, it was quickly realised that a magnitude derived from a visual observation and a photographic observation might have systematic differences dependent on the colour of the star – because the response of the photographic plate and the eye are not the same.

A so-called “Colour index” could be formed, where:

$$C.I. = m_{pg} - m_{vis}$$

This was defined to be zero for A0 stars (T_{eff} around 10000K), so hotter stars would have a negative colour index and cooler stars a positive colour index.

The colour index is – or can be – a temperature indicator. And the idea of colour indices – or “colours” – has carried over into modern methods. In the UBV system, for example, we typically

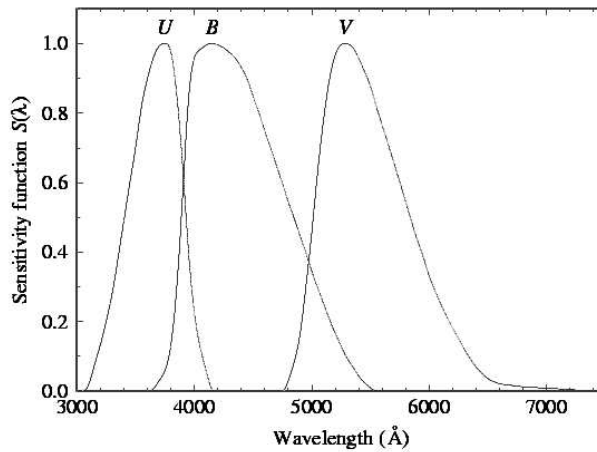


Figure 8: *UBV* filter passbands (Ultraviolet, Blue, Visual).

reduce the data as V , $(B - V)$ and $(U - B)$, where the magnitude V now replaces the old “ m ” or “ m_{vis} ” with conspicuously more accuracy and $(B - V)$, for example, is a useful temperature indicator. $(B - V)$ and $(U - B)$ together can be used to correct for the effects of interstellar reddening – the way interstellar dust affects starlight.

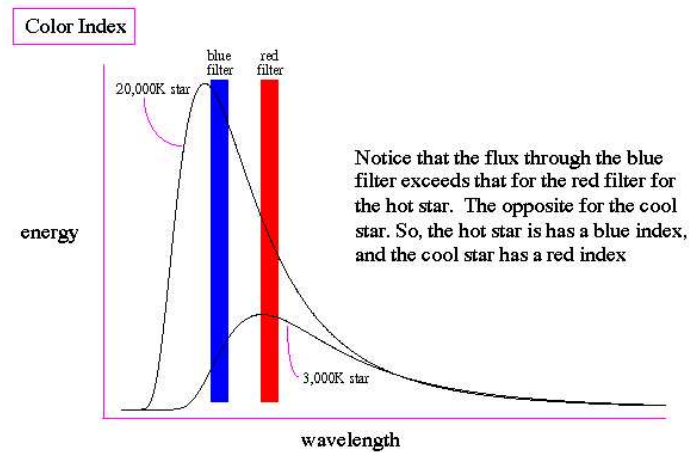


Figure 9: Schematic illustration of “colour index”.

Given that we can calculate magnitudes (as described in 5.3) and correct them for atmospheric extinction, we can then measure a large number of standard stars, determine a mean zero point and plot the residuals (standard values – observed) against colour.

This results in plots similar to the following figure:

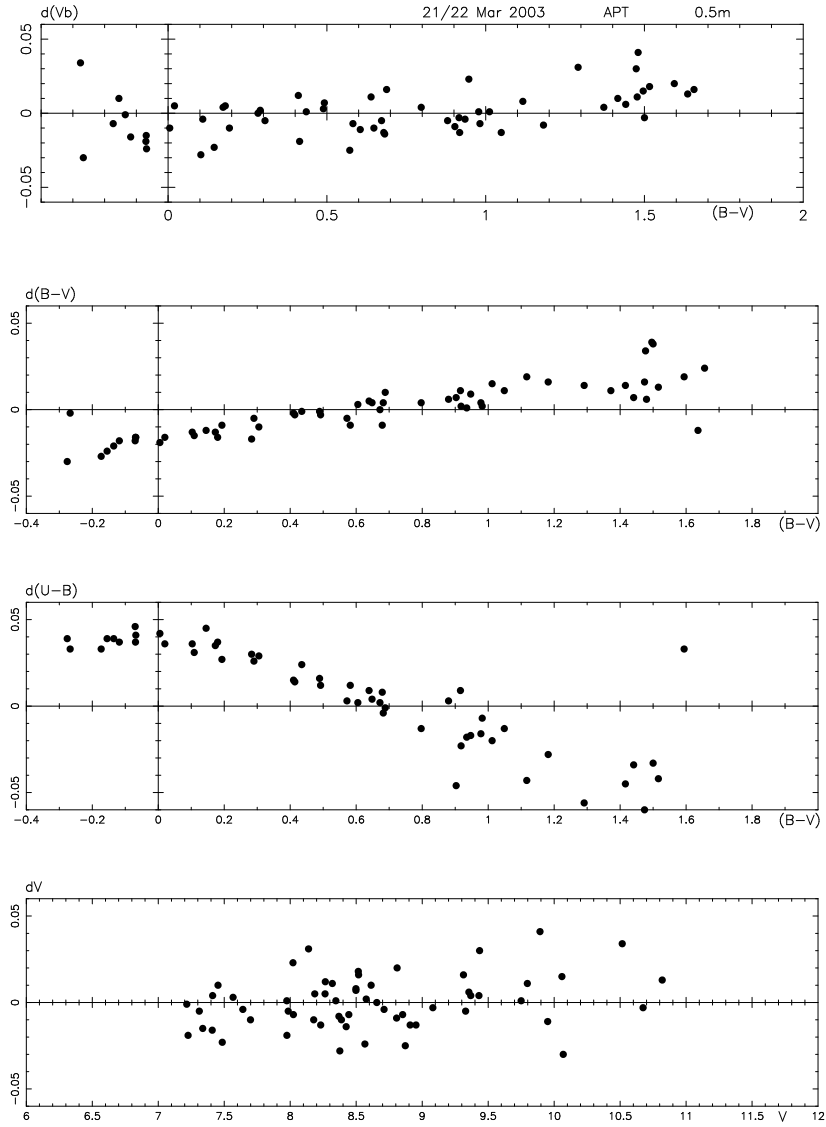


Figure 10: Residuals (standard values – observed) for a single night of Automatic Photometric Telescope (APT) observations.

With solutions:

$$\begin{aligned}
 V &= V_i + 20.309 \\
 (B - V) &= (B - V)_i - 0.159 \\
 (U - B) &= (U - B)_i - 1.314
 \end{aligned}$$

where the “i” subscripts indicate the instrumental data and the unsubscripted are (nominally) the standard values. We can then correct for the slopes or colour terms:

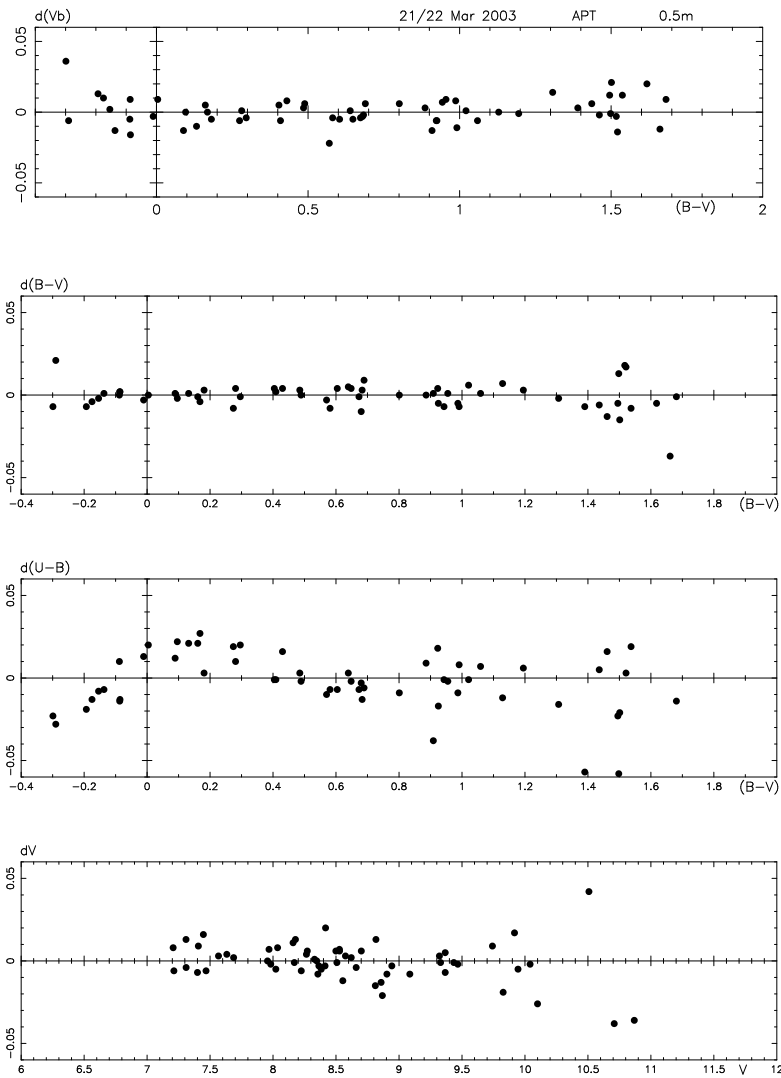


Figure 11: Residuals after scale factors applied.

With solutions:

$$V = V_i - 0.021 (B - V) + 20.311$$

$$(B - V) = 1.025 (B - V)_i - 0.180$$

$$(U - B) = 0.960 (U - B)_i - 1.269$$

which clearly leaves “non-linear” residuals. A time-dependent correction of -0.080 magnitudes/day has been applied to the V data.

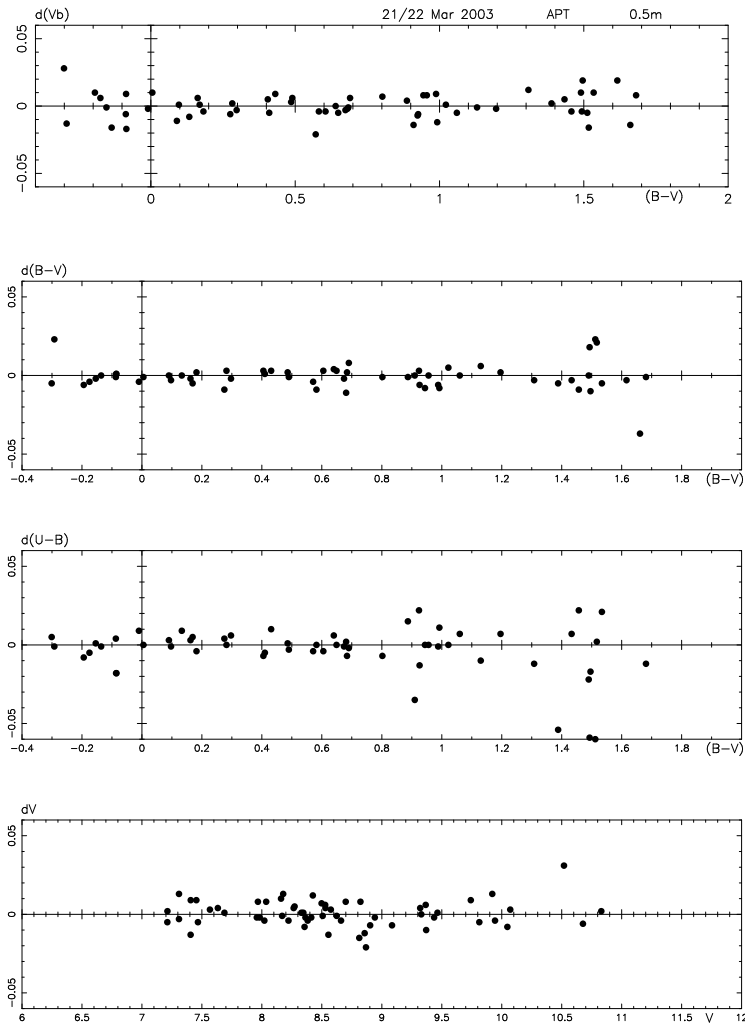


Figure 12: As for the previous plot but with non-linear corrections applied.

The colour equations thus derived can be applied to all subsequent observations. Normally, the colour equations are good for several months unless something in the system changes, and normally we observe a few standards during each night to determine the zero-points, correct for any time-dependent drift and check the extinction coefficients k_V , k_B , and so on.

Note that there is no obvious magnitude dependence in the V which indicates that we have the pulse-coincidence correction about right. (We could in fact now re-reduce the data with different values of ρ as a way of determining the correction for pulse coincidence loss).

3 Some Photometric systems

There are **many** different photometric systems, the Asiago Database on photometric systems lists over 200 ! (see <http://ulisse.pd.astro.it/Astro/ADPS/>). Early photometric systems – mainly broad-band – were used in a very general way; many systems were devised for rather specific applications (eg for the investigation of specific types of star such as hot stars, cool stars, etc). Some systems arose from existing systems but with slight modifications or an additional filter added, again to meet rather specific needs. Some systems arose because the observers just couldn't transform accurately to the standard system and simply adopted the best they could do !

A few common systems are described below, but first a few words on:

3.1 Filters

An optical filter can be a simple piece of glass or a complex compound structure. Its purpose is to allow certain wavelengths of light to pass through it while reflecting or absorbing other frequencies. Some common terminology:

- A **transmission curve (or response curve)** shows the percentage of light transmitted as a function of wavelength.
- The **pass band** is the range of wavelengths transmitted.
- The **peak wavelength** is the wavelength of maximum transmission.
- The **central wavelength** is the mid-point of the maximum and minimum wavelengths transmitted.
- The **effective wavelength** is the “centre” of the filter weighted by the intensity $S(\lambda)$, defined as:

$$\lambda_{\text{eff}} = \frac{\int_0^{\infty} \lambda S(\lambda) d\lambda}{\int_0^{\infty} S(\lambda) d\lambda}$$

- **Long pass (or “cut on”) filters** only transmit light redwards of a limit
- **Short pass (or “cut off”) filters** only transmit light bluewards of a limit. (For both long and short pass filters, this may only apply in a limited region (e.g the visible).
- **Dichroic filters** are interference filters which reflect certain wavelengths and transmit others.
- **Beamsplitters** reflect and transmit the same wavelengths (more-or-less). The simplest example would be a piece of glass at 45° to a light beam; most of the light will go straight through, but a fraction will be reflected at 90° to the original beam.

There are two main types of filter in common use in astronomical photometry:

- simple **glass filters** – these are usually rather broad band filters, typically $\sim 1000\text{\AA}$ - and are often used in combinations to produce the required pass-band.

Note, for example, the UG11 filter which might be used in the ultraviolet. It has an ultraviolet component and an red/infrared component – sometimes called the “red leak”. Early photomultipliers were blue sensitive, so that they had essentially no response redwards of

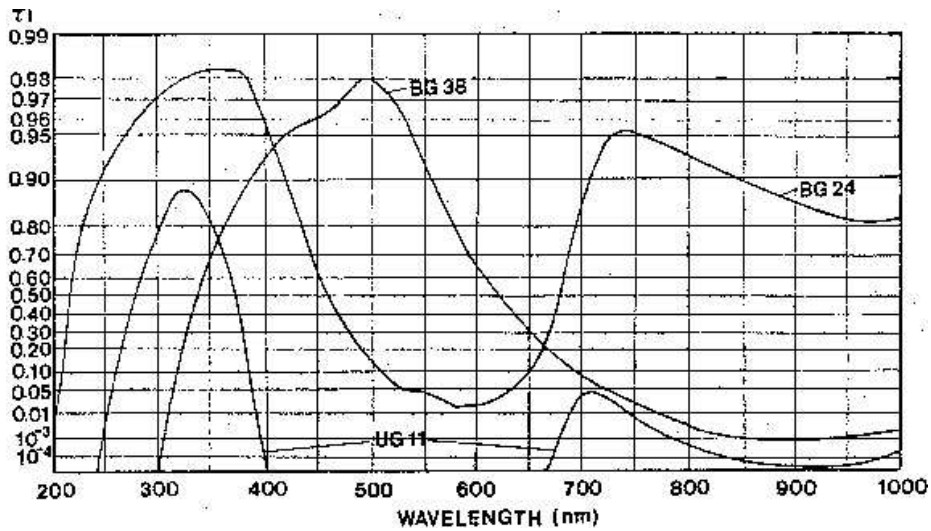


Figure 13: Transmission curves for some Schott glass filters.

about 6500\AA . Red-sensitive photomultipliers (and CCD detectors) are responsive well into the “red leak” region, making it difficult, if not impossible to reproduce a standard system based on a blue sensitive detector. We could add a red cut-off filter – a CuSO_4 filter, for example, will not transmit light redwards of about 7000\AA ($=700\text{nm}$); BG39 glass does much the same thing.

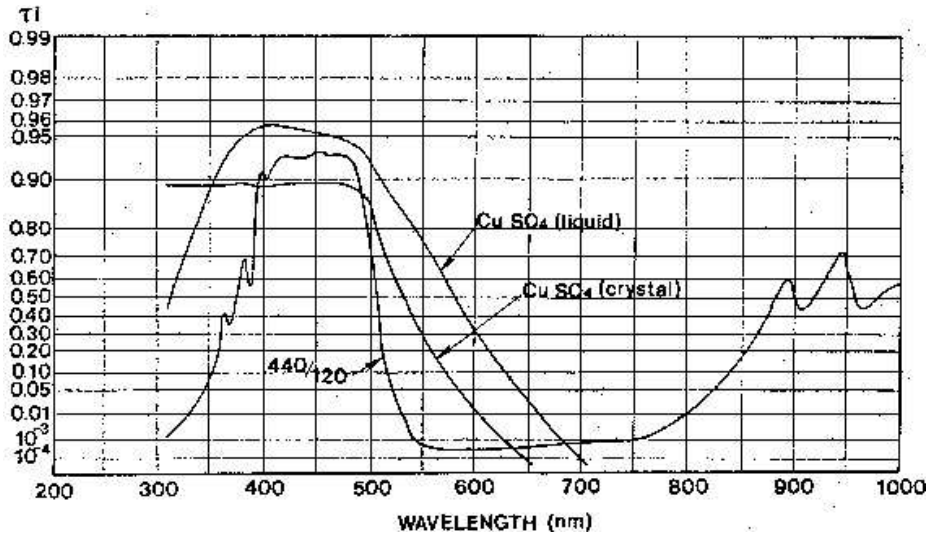


Figure 14: Transmission curves for CuSO_4 filters.

Note also, that the UG11 filter extends to well below the atmospheric cut-off ($\sim 3200\text{\AA}$) which means that the atmosphere determines the blue cut-off of the filter and might therefore be variable (!) – not a desirable characteristic.

- **interference filters** – these are complex filters, and usually consist of several layers of **thin films**, vacuum deposited on to an optically flat glass or quartz substrate; the principle is the same as the Fabry-Perot interferometer. The thin films are made of various dielectric materials – zinc sulphide, zinc selenide and sodium aluminium fluoride (cryolite) have been commonly used, but these tend to be hygroscopic and interference filters are usually sealed between two glass or quartz surfaces. Even so, there is a tendency for these filters to deteriorate from the edges inwards.

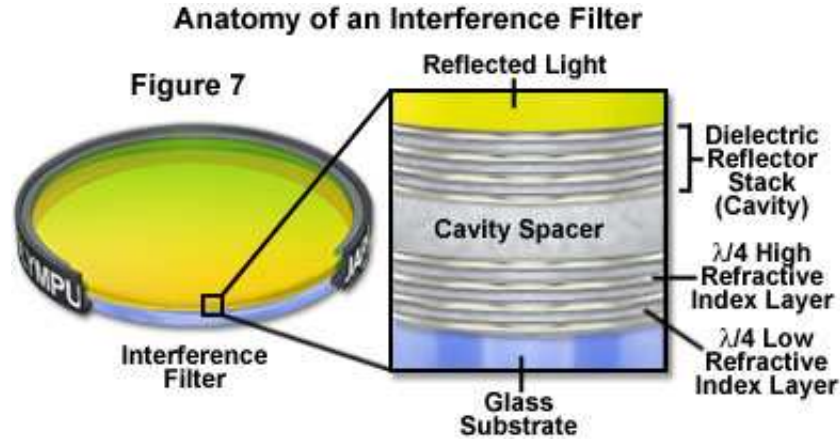


Figure 15: Schematic of an interference filter.

The use of metal oxides (**hard coatings**) gets around the environmental problems to some extent, and interference filters can be made with very precisely defined band-passes (and even multiple pass-bands) and sharp cut-offs.

There are some disadvantages:

- narrow pass-band filters tend not to have very high transmission (only $\sim 50\%$ at peak is not unusual) which is a disadvantage in faint object work,
- interference filters are sensitive to the angle of incidence of the light - the filter pass-band shifting to shorter wavelengths as the filter is tilted. Some care has to be taken to keep the filter normal to the incident light. Actually, this property has been utilised to “scan” spectral features by tilting the filter in a controlled way.

Other types of filter include:

- Chemical suspensions (e.g. CuSO_4 solution)
- Crystals (e.g. CuSO_4 - which is a good longpass filter)
- Various dyes in gelatin (on a substrate of some kind)

3.2 UBVR

The UBV system was one of the first well-defined broad band photometric systems to be introduced (in about 1953) after photoelectric detectors (photomultipliers) began to be used commonly. The system was introduced by H.L. Johnson and W.W. Morgan and is usually referred to as the “Johnson” system. Somewhat later, Johnson added the R and I (red and “infrared”) filters and this system has persisted to the present – and enjoyed widespread use – AND been very successful – even though the filter system itself is not optimally designed. For example, the U lies across the Balmer discontinuity which complicates things like transformation (colour equation) determinations and the correction of atmospheric extinction effects.

The V magnitude is similar to the old m_{pg} and the $(B - V)$ resembles the old Colour Index = $m_{pg} - m_{vis}$. The system colours are set at zero for an unreddened AOV star, so that OB stars have negative colours (unless significantly reddened by interstellar dust). We shall look at some applications of UBV photometry in the later examples.

In the southern hemisphere, a great deal of work over several decades by Alan Cousins (SAAO) resulted in the establishment of a very sound set of standard stars – the “E-region standards” – but for $UBV(RI)_C$ photometry – where the “C” subscript refers to “Cape” or “Cousins” photometry (because Cousins used somewhat different filters to Johnson for R and I – filters sometimes referred to as “Kron” or “Kron-Cousins” filters. **So you have to make sure you know which RI or VRI system you are dealing with.** The northern and southern UBV systems are essentially identical (for most practical purposes).

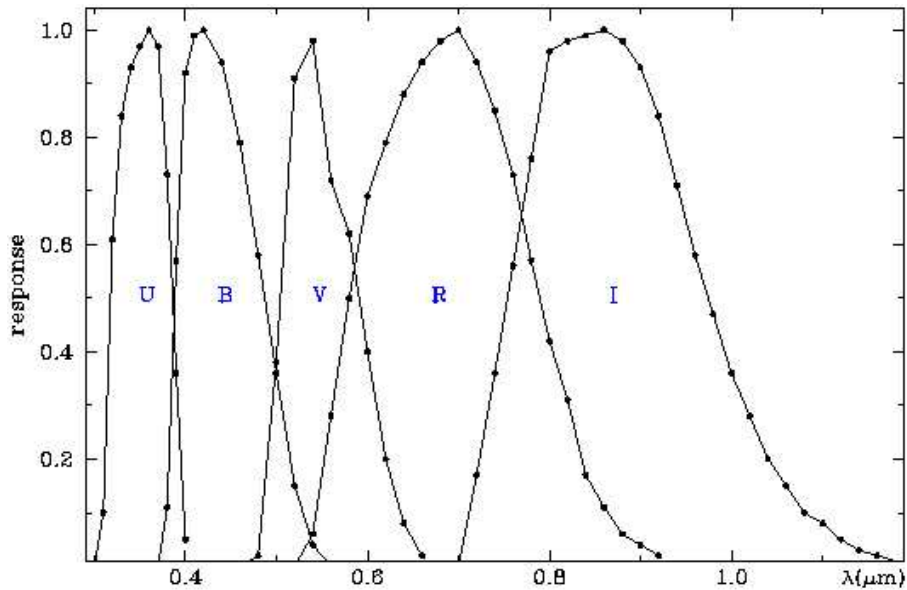


Figure 16: $UBVRI$ filter passbands.

3.3 JHKLMN

A later addition to the *UBVRI* filter system was the use of filters at wavelengths beyond a micron ($10000\text{\AA} = 1000\text{nm}$) – the “near” infrared colours (Johnson, *AstrophysJ* 141, 923, 1965). The set *UBVRIJHKLMN* is sometimes called the **Arizona system**

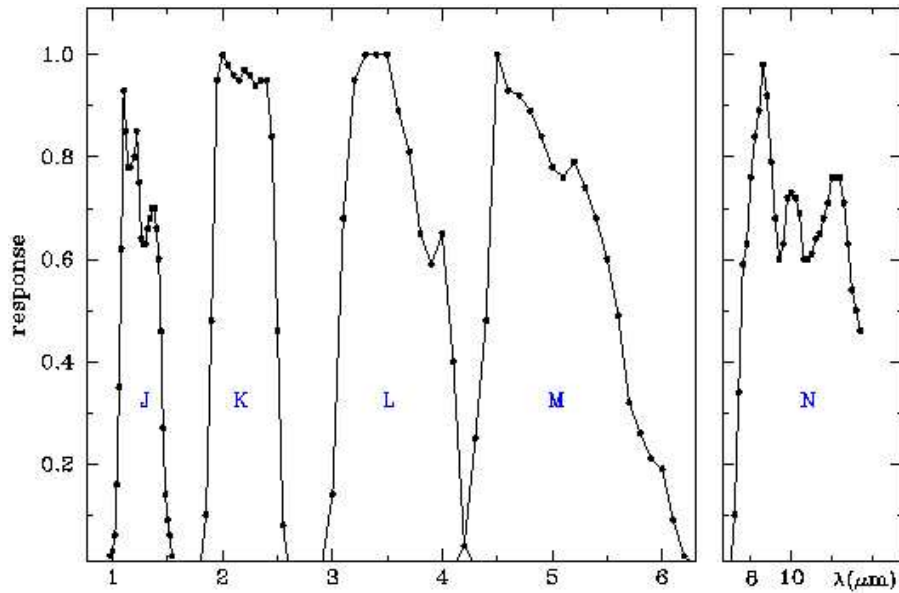


Figure 17: *JHKLMN* filters compared to late-type (very cool) M dwarf star.

We don't look at these in any detail, since they will be discussed in the lecture on infrared techniques (by Ian Glass). Because the techniques for infrared photometry are generally so different from those used for optical photometry, the infrared broad band colours are normally measured separately from the *UBVRI*. The *JHKL...* filters are very useful in the study of cool objects – for example late-type (very cool) M dwarf stars (which have *very* little flux at optical wavelengths; the flux peak is redwards of $1\mu\text{m}$) and circumstellar dust.

3.4 Strömgren $uvby$ and $H\beta$ photometry

The $UBVRI\dots$ and similar systems are usually called **broad-band** systems. There is no specific definition of what constitutes broad or narrow pass-bands, but typically these might be **broad** – with filters having FWHM around 1000\AA or more; **intermediate** – filters a few hundred \AA wide; and **narrow** – filters less than about 200\AA wide.

The Strömgren $uvby$ photometric system (Vistas in Astronomy 2, 1337, 1956) is an intermediate-band system and the allied $H\beta$ photometry is a narrow-band system. The filter transmissions are shown in the figure.

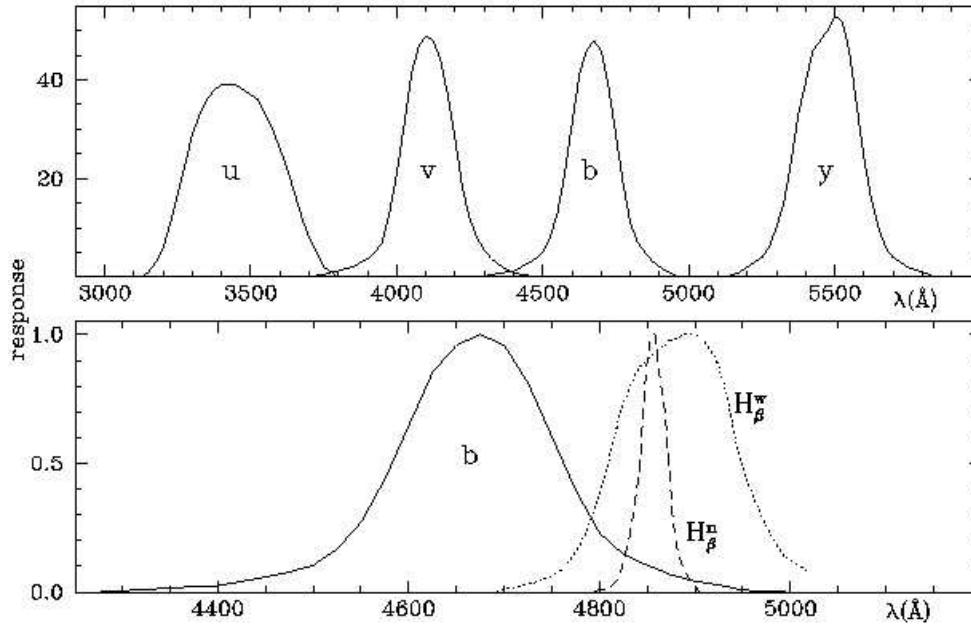


Figure 18: Strömgren $uvby$ and $H\beta$ filter transmissions.

Now the “ v ” filter is a “violet” filter (rather than the Johnson “ V ” or “visual” filter) and the Strömgren y is similar to (in fact is usually forced to be) the Johnson V . The narrower filters tend to reduce transformation problems (because the narrower filters are not as sensitive to atmospheric effects, for example) although the system itself is usually restricted to (and was defined to investigate) the earlier-type stars – say O to about G2. The narrower filters also sample the spectrum more precisely.

The colour indices usually formed are:

- $(b - y)$ – which is a **temperature index**, similar to $(B - V)$;
- $m_1 = (v - b) - (b - y)$ – which is a **metallicity index** which measures the strength of metal lines around $H\delta$, relative to the spectral continuum;
- $c_1 = (u - v) - (v - b)$ – which is a **Balmer discontinuity index**, which is a measure of **temperature** in the OB stars and **luminosity** in AFG stars.

Two two-color diagrams are usually formed:

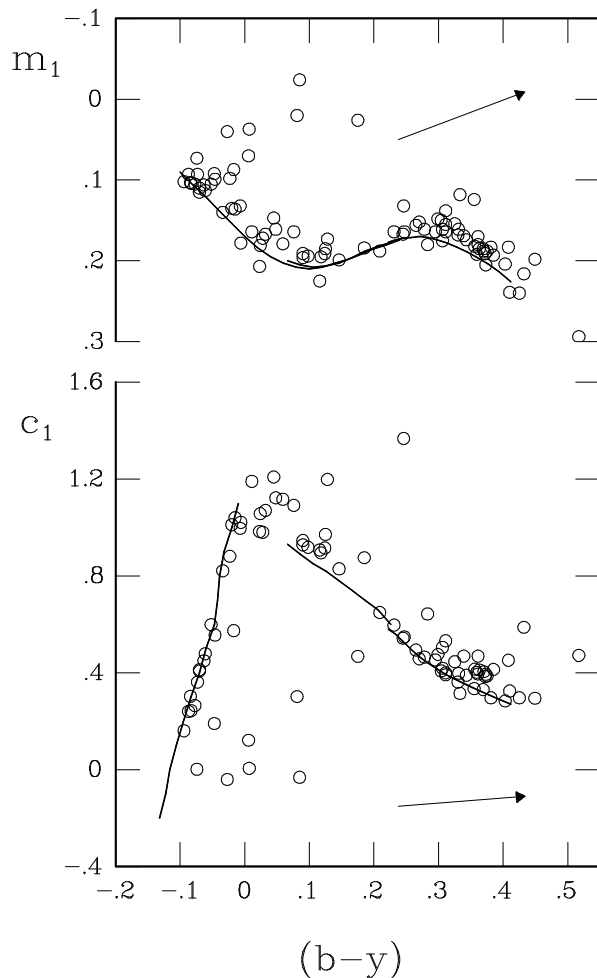


Figure 19: Strömrgren *wby* two-colour diagrams. The solid lines are the **intrinsic colour lines** – the loci of ‘normal’ stars with no interstellar reddening. The arrows indicate the direction of the effect of interstellar reddening in each diagram and the points indicate the Strömrgren indices for some standard stars.

The *wby* measurements are often combined with $H\beta$ photometry. This comprises two filters, both centred on the Balmer line of Hydrogen at 4861\AA . One filter has a pass-band (FWHM) of about 150\AA and the other has a pass-band of about 30\AA . These are usually called $H\beta(\text{wide})$ and $H\beta(\text{narrow})$, or $H\beta(W)$ and $H\beta(N)$ or some such.

Because the narrow filter is affected much more by the strength of the $H\beta$ line in the stellar spectra than the wide filter, the quantity:

$$\beta = -2.5 \log \frac{H\beta(W)}{H\beta(N)} = H\beta(W) - H\beta(N) \text{ (in magnitudes)}$$

is a measure of the **equivalent width** or strength of the Balmer line – and therefore a measure of luminosity in OB stars and temperature in AFG stars. Notice how this complements the c_1 index which works the other way round – so β/c_1 give us a measure of luminosity and temperature for the OBAF(G) stars.

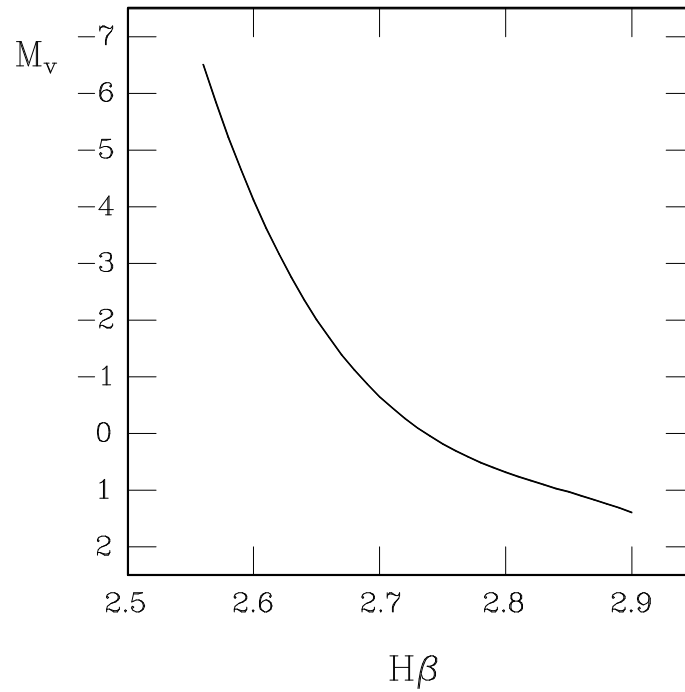


Figure 20: Calibration of the $H\beta$ index in terms of absolute magnitude for Zero Age Main Sequence (ZAMS) O and B stars (Crawford, AstrJ 83, 48, 1978).

Recall that:

$$m - M = 5 \log d - 5$$

alternatively:

$$V - M_v = 5 \log d - 5$$

and since β photometry can give us a measure of M_v and either Johnson or Strömngren photometry will give us a V (or y), we can get stellar distances – assuming we can correct for interstellar reddening, which will be described later.

3.5 SDSS - $u'g'r'i'z'$

The Sloan Digital Sky Survey (SDSS – see www.sdss.org) is a custom-built system for surveying a quarter of the sky down to rather faint limiting magnitude. It will measure the shape, brightness and colours of hundreds of millions of objects and will carry out follow-up spectroscopic measurements of a million galaxies and a hundred thousand quasars.



Figure 21: The Sloan Digital Sky Survey (SDSS) site at Apache Point, New Mexico. The 2.5m survey telescope is on the left.

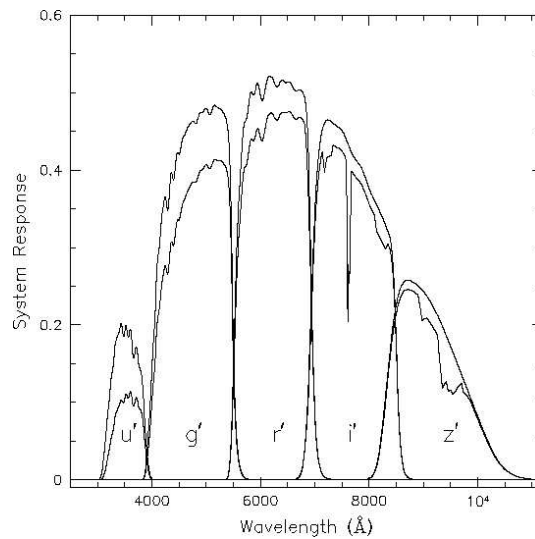


Figure 22: The SDSS filters – $u'g'r'i'z'$ which cover the spectrum from the atmospheric UV cut-off to the IR sensitivity limit for silicon CCDs. The lower curves are the filter transmissions plus 1.2 airmasses of atmosphere.

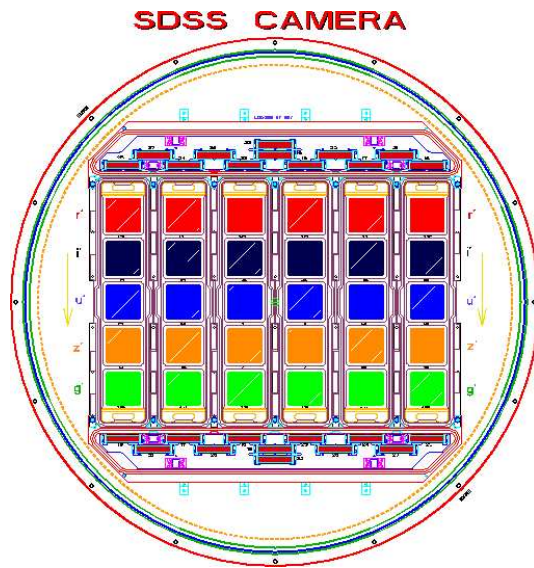


Figure 23: Schematic of the arrangement of the thirty (2048 x 2048 pixel) imaging CCDs operated in “drift scan” mode plus twenty-four (2048 x 400 pixel) CCDs for astrometry.

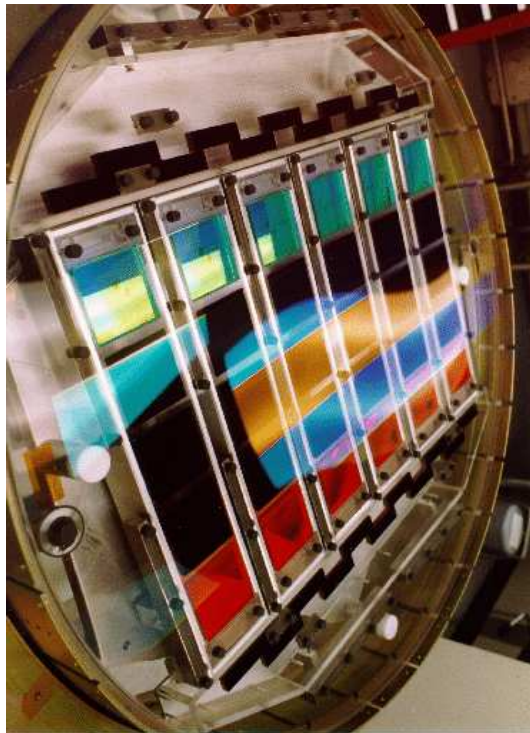


Figure 24: Photograph of the actual camera array. Filters from top to bottom are $gzurig'$, z' , u' , r' , i' .

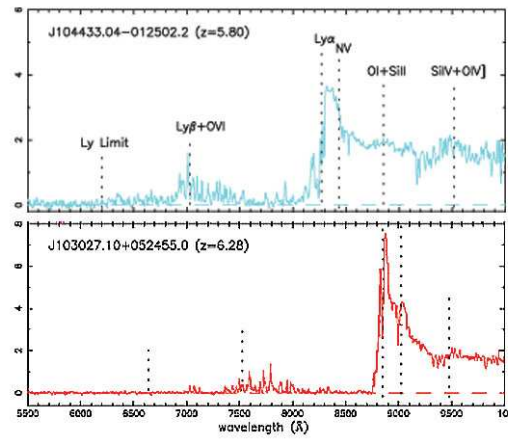


Figure 25: Two very high red-shift quasars from the SDSS.

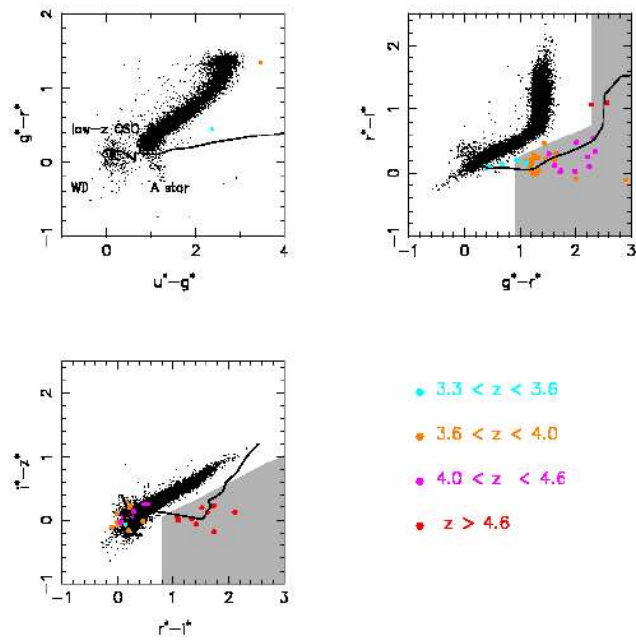


Figure 26: Colour selection of high redshift quasars using SDSS photometry.

4 Applications

4.1 Visual

Curiously, visual (eye + telescope) observations are still useful. Many amateurs regularly observe objects which are likely to undergo rapid and unpredictable light variations – brightening (eg cataclysmic variables) or fading (eg R Coronae Borealis variables). Early warning of these occurrences is often very valuable. The Reverend Evans in Australia had discovered more supernovae in external galaxies than any other person or group (at least until the inception of automated supernova searches) by memorising hundreds of galaxy fields and checking them all regularly.

4.2 Photographic surveys

Although photographic work is becoming less useful (particularly in light of such surveys as the SDSS), still some useful work is being done, because of the wide-angle of view of Schmidt telescopes. The Edinburgh-Cape faint blue object survey was based on photographic material, as was the recent galactic plane $H\alpha$ survey.

4.3 Calibrations

A plot of luminosity against temperature for “normal” stars looks something like:

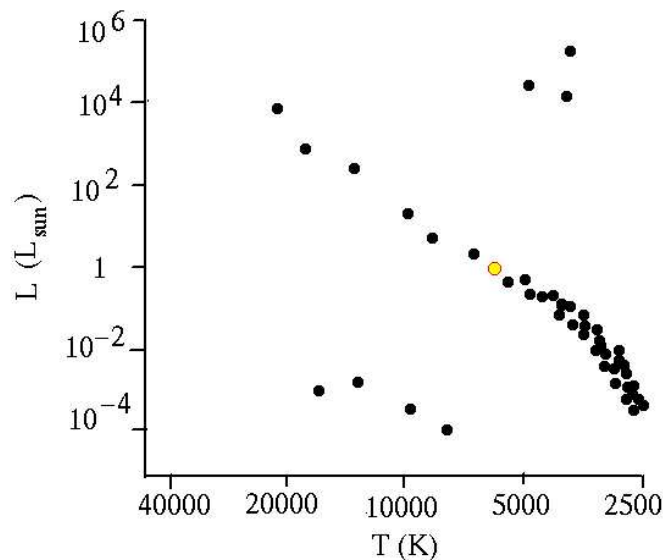


Figure 27: Schematic Hertzsprung–Russell diagram

This – or some variant of it (eg M_v vs T_{eff}) – is usually called a **Hertzsprung–Russell diagram** – or **HR diagram** and is very commonly used in stellar astronomy. If we look at plots of – say – M_v vs spectral type or a colour such as $(B - V)$ or $(V - I)$, we see a similar appearance: This is because – as we have seen – colour is strongly correlated with temperature. It is therefore very handy to have colours *calibrated* in terms of temperature, because it’s a lot easier to measure

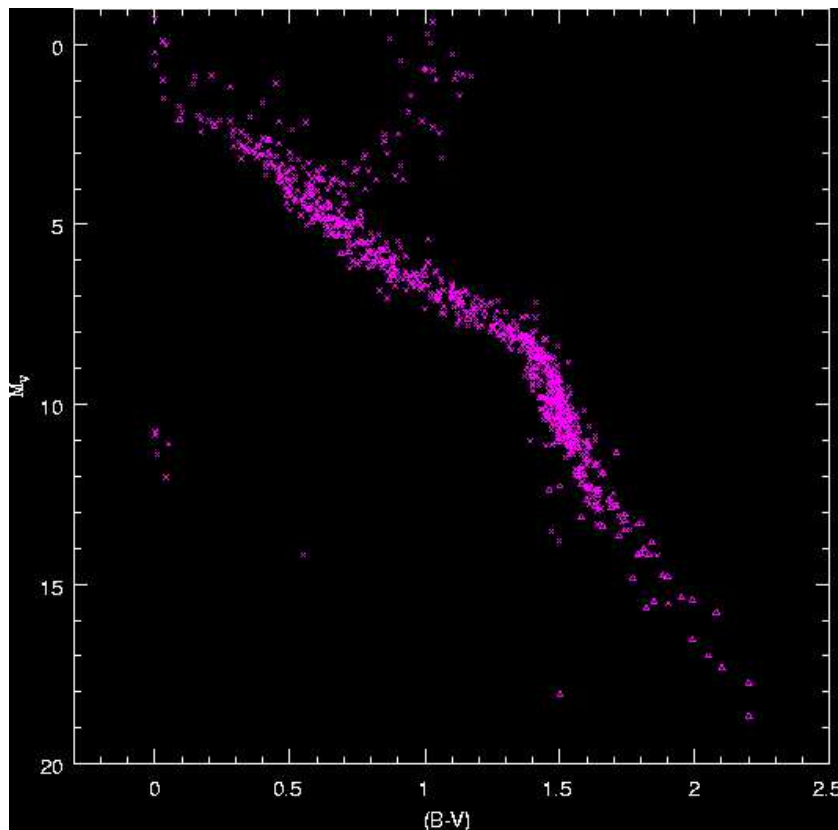


Figure 28: Stars with high-quality *Hipparcos* distances (absolute magnitudes plotted against $(B - V)$ color).

colour than temperature (by, for example, spectral analysis) **and** of course, with a CCD, we can measure colours for thousands of stars at the same time.

4.4 Interstellar reddening

If we plot, say, $(U - B)$ against $(B - V)$ – a colour–colour diagram – we can see that, like the HR diagram, most of the stars lie on a fairly well defined locus – the **intrinsic colour** line.

Unless stars are rather nearby, almost all of them have colours which are affected by interstellar dust – in a very similar way to the extinction of starlight by the Earth’s atmosphere. Because of the sizes of the dust particles, the “reddening” of starlight by interstellar dust is roughly proportional to λ^{-1} – at least in the visible spectrum.

This diagram gives us a nice empirical way of determining the effects of interstellar dust quantitatively. We can:

- determine the intrinsic – or unreddened – colours of stars by using the colours of nearby stars, or by plotting lots of stars and determining the “blue-most” envelope of the data.
- measure lots of stars which we know to have the same spectral type and therefore temperature – and therefore colours – and plot the observed colours in a two–colour diagram and determine the gradient of the “reddening line”.
- try to relate this to the total extinction.

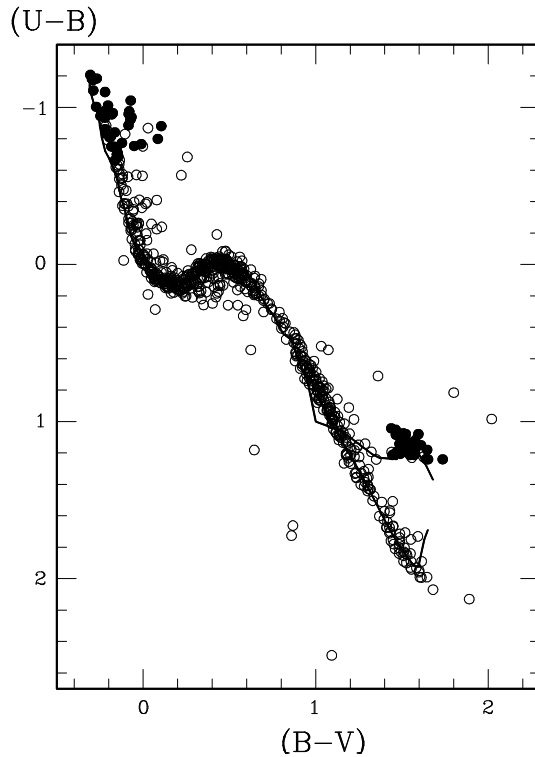


Figure 29: $(U - B)/(B - V)$ diagram for E-region and other nearby standard stars. (Note that $(U - B)$ would not be a good temperature indicator – *in the absence of other information* – because at certain values of $(U - B)$ it is multi-valued.)

If we write, for example, the “colour excess”:

$$E_{(B-V)} = (B - V) - (B - V)_0$$

where $(B - V)$ is the observed quantity and $(B - V)_0$ is the intrinsic or unreddened quantity, then it turns out that:

$$\frac{E_{(U-B)}}{E_{(B-V)}} = 0.72 + 0.05 (B - V)$$

Couple of notes:

- In actual fact, there is some evidence that the gradient (0.72) is slightly dependent on spectral type (or colour, though this is a small effect).
- It is common, if spectral types are available, to assume an intrinsic colour based on the spectral type. I prefer not to do this generally because:
 - it assumes that the spectral types are without error and forgets that any given spectral type is a “box” covering a range of values.
 - given that, it effectively lets you assume different reddening laws for different stars – which might be unphysical.
 - the photometry is probably significantly more accurate than the spectral type.

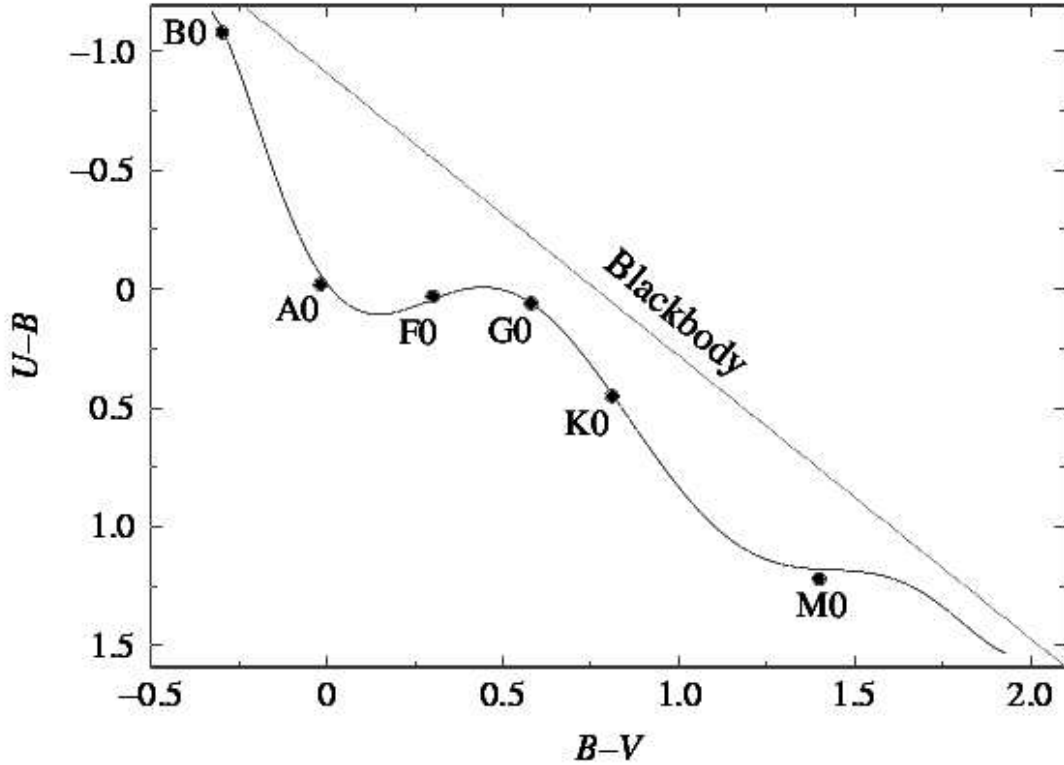


Figure 30: Schematic two-colour diagram. Note that the differences between the intrinsic colour line and the perfect radiator (black body) are caused by absorption in the stellar atmospheres. Around A0, the Balmer series lines are at their strongest, so the Balmer discontinuity (the Balmer series limit strongly affects U). For later-type stars, the Balmer lines weaken, but molecular absorption gets stronger.

Another way of looking at this is to say that we can calculate a “reddening-free parameter” – usually written:

$$Q = (U - B) - 0.72 (B - V)$$

where Q can be calibrated against temperature in the same way as intrinsic colours.

By comparing spectrophotometry of reddened and unreddened stars, it is possible to determine the **ratio of total to selective absorption** – in other words the relation between the observable E_{B-V} and the total interstellar extinction for a star. This turns out to be simply:

$$A_V = R_V E_{(B-V)}$$

where R_V is close to 3.05 in many (most) galactic regions, but can be significantly higher in dense dust clouds, such as star-forming regions. Generally, it is fairly simple to correct data for the effects of interstellar reddening – and to measure the reddening quantitatively.

Note that similar quantities can be derived for other colours than $(B - V)$ and in other photometric systems.

4.5 Metallicity

Astronomers tend to refer to everything except Hydrogen and Helium as “**metals**”. This is not terribly surprising, even in stars like the Sun metals constitute only about 2% of the mass; older stars might be considerably weaker in metals.

Many metallic absorption lines are present in the ultraviolet, so that in (for example) the UBV colours, a **metal-weak** star would be bluer than a metal-rich star, due to less **line blanketing** in the ultraviolet. In particular the $(U - B)$ colour will be more negative for the more metal-weak star.

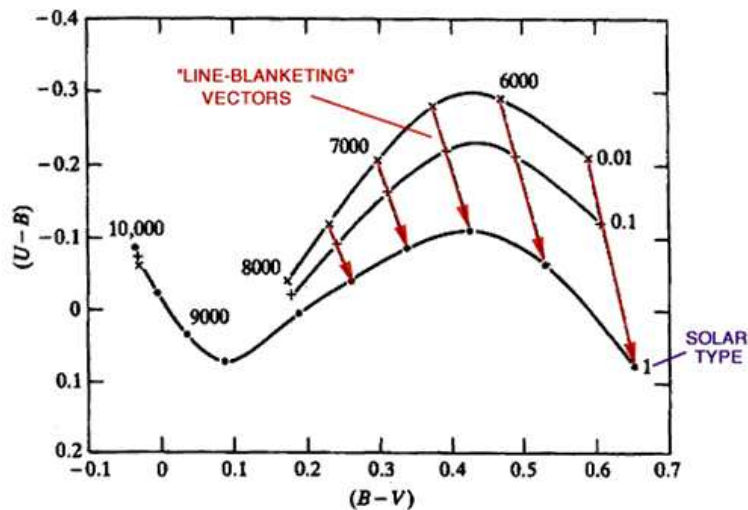


Figure 31: Line “blanketing” effects in the UBV system. U is affected most, B a little, so $(U - B)$ is affected more than $(B - V)$.

Iron is commonly used as a measure of metallicity, though in detailed analyses, a wide range of metals might be measured. Metallicity is often written:

$$[\text{Fe}/\text{H}] = \log[n(\text{Fe})/n(\text{H})]_{\text{Star}} - \log[n(\text{Fe})/n(\text{H})]_{\text{Sun}}$$

or, for example:

$$[\text{Ca}/\text{H}] = \log[n(\text{Ca})/n(\text{H})]_{\text{Star}} - \log[n(\text{Ca})/n(\text{H})]_{\text{Sun}}$$

Metal-weak stars – such as those in globular clusters – might have $[\text{Fe}/\text{H}]$ as low as -2, or a metallicity relative to the Sun of 0.01.

Other metallicity indicators in “designed” narrower band systems – such as m_1 in the Strömgren system – are more accurate, but spectral analysis is required for the most accurate results.

4.6 Cluster sequence fitting

Galactic open and globular clusters are important (amongst other reasons) for our understanding of stellar evolution. Because all the stars are at (almost) the same distance and because we (think we) can assume they all formed at much the same time (in astronomical terms, at least) and from material of much the same composition (same metal abundance), they provide natural laboratories for stellar evolution.



Figure 32: The “Jewel Box” cluster.

If we can measure magnitudes and colours for many stars in a cluster (such as the one shown) and can correct these for interstellar reddening as described above, we can get a good distance determination for the cluster by making its main sequence fit the main sequence of clusters nearer the Sun for which we have accurate distances (such as the Pleiades and Hyades).

This not only allows us to determine absolute magnitudes for very hot, massive stars – including Cepheid variables, which are needed to determine extra-galactic distance scales – but also allows the cluster age to be determined by comparing the cluster “**turn-off**” point to theoretical models – or by the fitting of “**isochrones**” – lines of constant time – to the observed details of the cluster. Conversely, the cluster data provide tests of stellar evolution theory.

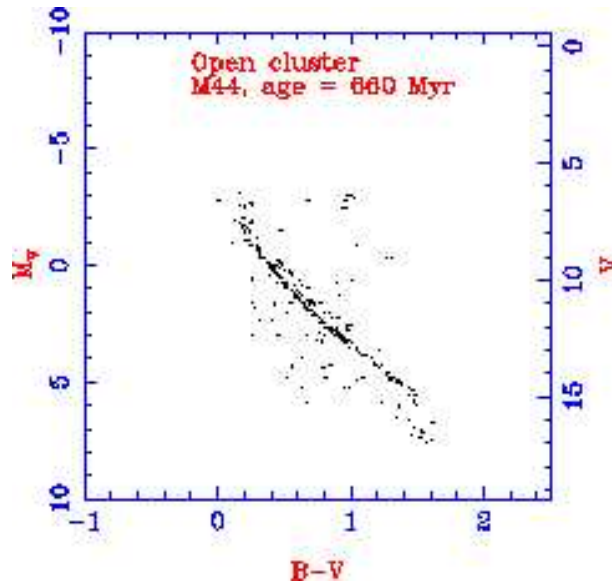


Figure 33: Open cluster M44 with age estimated from main sequence turn off.

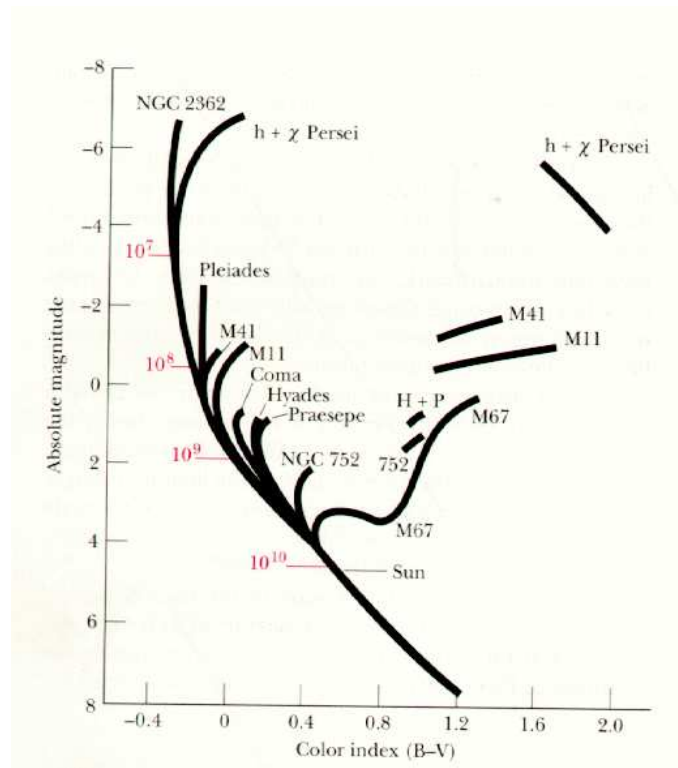


Figure 34: Composite diagram of several open clusters from the very young (NGC2362) to the very old (M67). Note that some of the clusters have stars evolved into the red giant/supergiant region. Ages of main sequence turn-off are indicated.

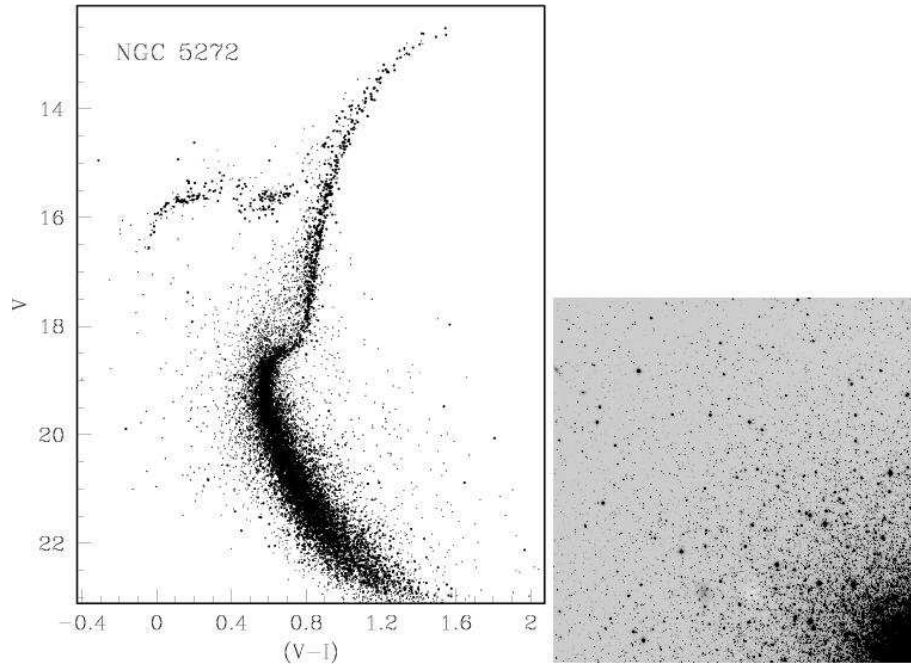


Figure 35: Frame of part of the globular cluster NGC5272 and colour-magnitude diagram showing main sequence, giant branch and horizontal branch.

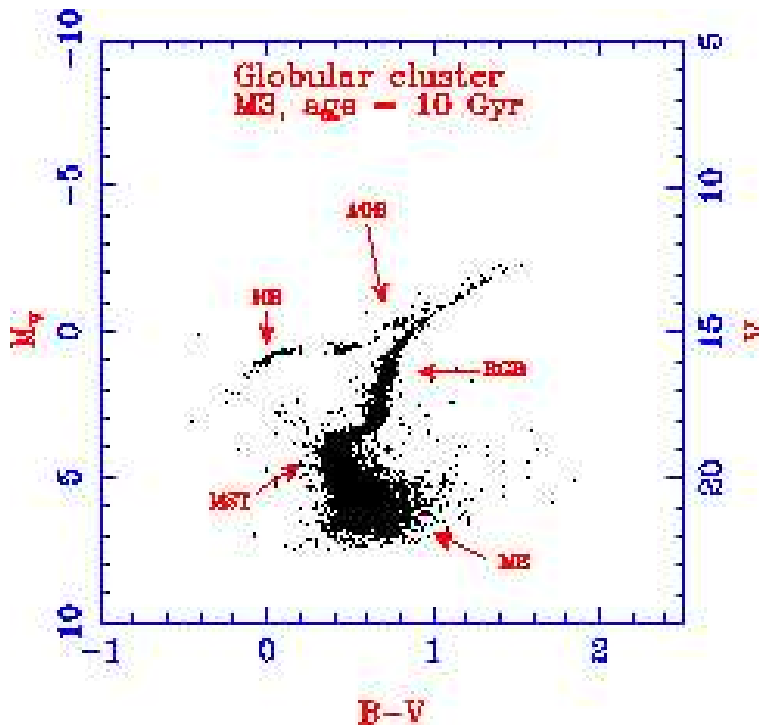


Figure 36: Colour-magnitude diagram of 10 billion year old globular cluster M3 indicating main sequence, main sequence turn off, red giant branch, horizontal branch and asymptotic giant branch.

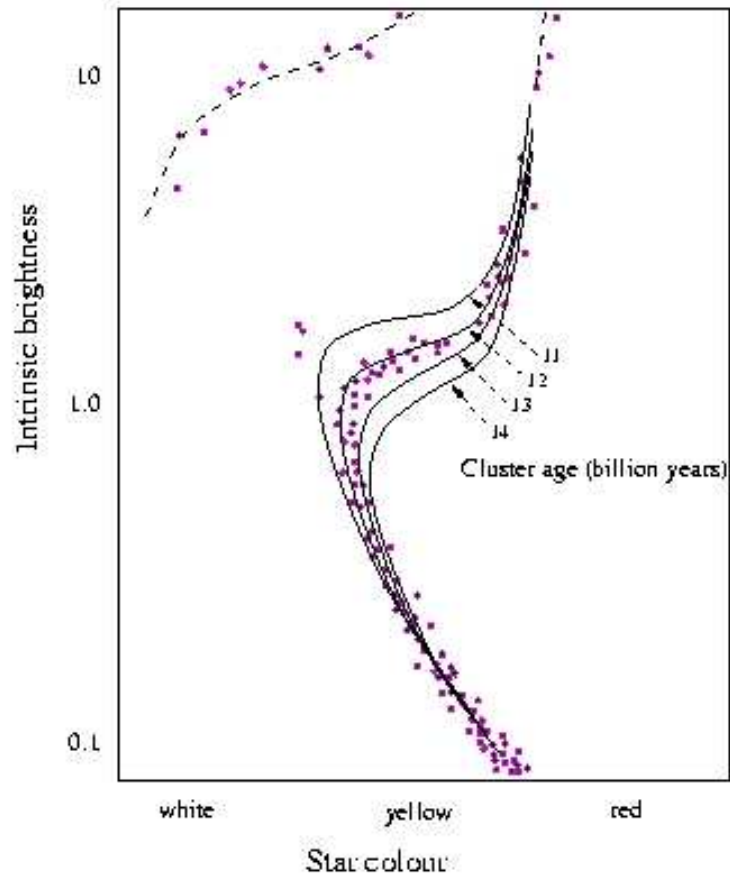


Figure 37: Schematic globular cluster colour-magnitude diagram and computed isochrones.

4.7 Variability

With “all sky” photometry, we can also regularly measure numbers of long-term variable stars with low frequency.

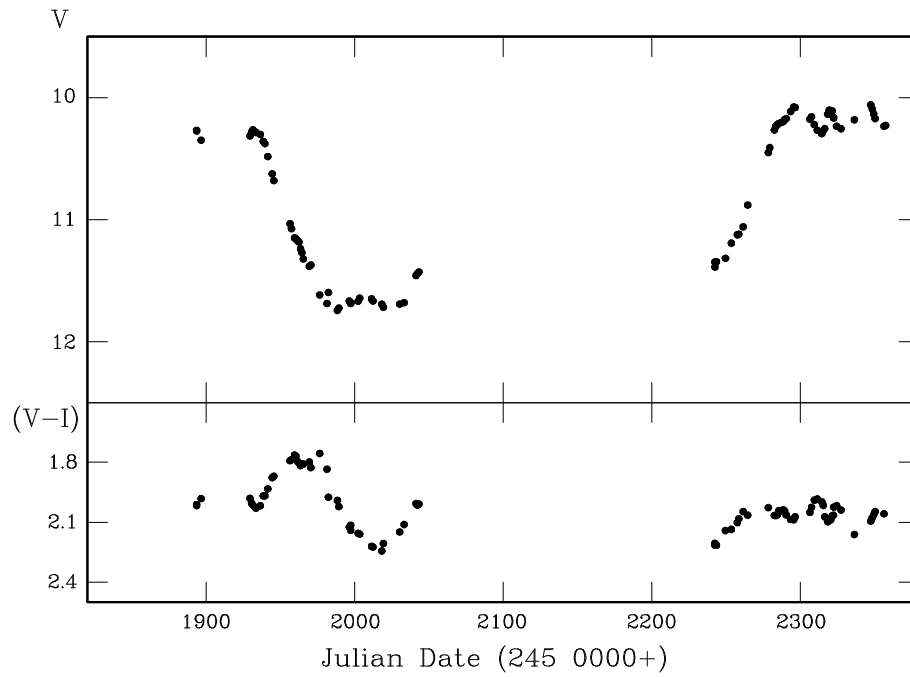


Figure 38: APT data for two seasons for an IRAS source.