Astrophysics 3, Semester 1, 2011–12

Observational Astronomy

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1 Quantitative Measures of Light

For the quantitative description of the radiation field from astronomical objects, and its measurement, we use the following quantities:

Flux Density: this is the radiation energy received per unit time, per unit area (normal to the propagation direction of the radiation) per unit frequency (or wavelength) range. For most astronomical observations, 'per unit frequency' is used, and the flux density, f_ν, is therefore measured in units of $\mathrm{W\,m^{-2}\,Hz^{-1}}$. A commonly–used unit for measurement of flux density is the **Jansky**. $1\,\mathrm{Jansky} = 10^{-26}\mathrm{W\,m^{-2}\,Hz^{-1}}$.

Wavelength is used instead of frequency as a measure of bandwidth when there are practical reasons for doing so, for example in spectrographs which produce a nearly linear wavelength scale. We then use f_λ, measured in units of $\mathrm{W\,m^{-3}}$. The two expressions for flux density are related as:

$$f_\lambda = \frac{d|\nu|}{d|\lambda|}\, f_\nu = \frac{c}{\lambda^2} f_\nu. \tag{1}$$

The quantities λf_λ and νf_ν have the units of W $\mathrm{m^{-2}}$, and are often used when the spectral energy distribution is discussed in log space of wavelengths or frequencies.

Flux: this is the integrated flux density within a given range of wavelengths or frequencies:

$$F = \int_{\nu_1}^{\nu_2} f_\nu d\nu; \qquad F = \int_{\lambda_1}^{\lambda_2} f_\lambda d\lambda; \tag{2}$$

Surface brightness: this is the flux density received per angular area of sky, $B_\nu = f_\nu/\Omega$, in units of $\mathrm{W\,m^{-2}Hz^{-1}sr^{-1}}$ (where sr is for steradian. Ignoring cosmological effects of curved space, *surface brightness is conserved with distance.*

Specific Intensity: also with units of $\mathrm{W\,m^{-2}Hz^{-1}sr^{-1}}$, this is a similar quantity to surface brightness, but specific intensity is measured from the source frame (flux density *emitted* per solid angle), rather than from the observer's frame (brightness being the flux density *received* per solid angle). Specific intensity is also conserved with distance. Written as I_ν, it is related to flux density by

$$f_\nu = \int_\Omega I_\nu \cos\theta d\Omega. \tag{3}$$

Luminosity: this is the total power (radiation energy per unit time) emitted by an astrophysical source. *Monochromatic luminosity* is luminosity per wavelength or frequency unit. The luminosity is related to the flux (and monochromatic luminosity to flux density) by the distance to the source, d as:

$$L = 4\pi d^2 f; \qquad L_\nu = 4\pi d^2 f_\nu \tag{4}$$

2 Measuring distances

A difficult question in astronomy is how one measures the distance to an object. The most direct way to do this is by means of *parallax*. The position of a (nearby) source is followed during the course of a year, and changes relative to the more distant sources due to the rotation of the earth around the sun. The distance corresponding to a parallax of 1 arcsec is defined to be 1 parsec (pc), which is 3.08×10^{16}m, or 3.26 light-years. For example, Vega, the brightest star in the northern sky, has a parallax of 0.123 arcsec, so its distance is 8.13 parsec.

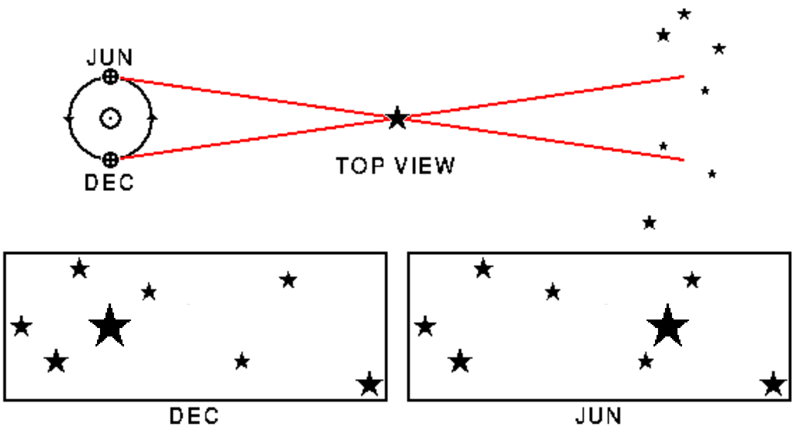

Figure 1: Concept of parallax

This technique for measuring distances only works for nearby stars, but can be used to calibrate a 'distance ladder' to more distant objects. For example, in 1912 it was discovered that a particular type of variable star known as Cepheid Variables show a very tight relationship between their variability period and their luminosity. This relationship was calibrated using nearby Cepheid stars, for which the distance could be determined, but can then be applied to more distant Cepheids: by measuring the variability period of these, the luminosity can be determined, and comparing this with the measured flux density then provides the distance. This method was used by Edwin Hubble in 1925 to demonstrate for the first time that the Andromeda galaxy was extragalactic, rather than being a distant cloud of gas within our own galaxy.

3 Magnitudes

In practice, we measure flux densities or brightnesses within some certain frequency (or wavelength) range. The total energy measured is then the integral of the source flux times some frequency-dependent effective filter response. This last quantity includes all the factors that modify the energy arriving at the top of the Earth's atmosphere. The main factor is the instrumental filter, but atmospheric absorption and frequency-dependent sensitivity of the detector also matter.

A notion of relative brightness can still be maintained by using the instrumental output, O, which is what leads to the notion of magnitude. The definition of magnitude is one of astronomy's unfortunate pieces of historical baggage. Around 150 BC, Hipparcos catalogued $\sim$1000 visible stars in 6 categories of apparent brightness, and in 1856 Pogson suggested they be called magnitudes, with the brightest stars to the eye being 'first magnitude', and fainter stars having larger magnitude. Most human senses are logarithmic, and it turned out that a difference of 1 magnitude corresponds roughly to a factor 2.5 difference in flux density, and so the magnitude system was defined as

$$\boxed{m = -2.5\ \log_{10}\left(\frac{\mathrm{O}_{\mathrm{object}}}{\mathrm{O}_0}\right).} \tag{5}$$

A magnitude system is then defined by the total effective filter and by the **zero point** (the object of zero magnitude that produces output O_0). The magnitude is often written in terms of flux as

$$\boxed{m = m_0 - 2.5\ \log_{10} f.} \tag{6}$$

In the commonly–used 'Vega–magnitude' system, Vega is defined to have zero magnitudes in all bands/filters. This is a simple system, but makes conversions from magnitudes to Jy very messy, because of the spectral properties of Vega: the zero-point has to be determined for every filter. In the alternative 'AB–magnitude' system, the zero point is defined to be 3631 Jy in all bands (the flux density of Vega in the middle of the optical atmospheric window).

Table 1: Parameters for common filter systems

Filter	λ_{eff}/nm	$\Delta\lambda$/nm
U	360	65
B	430	100
V	550	80
R	680	95
I	900	230
J	1220	150
H	1630	170
K	2190	190
L	3450	280

Once we have the apparent magnitude, this can be converted to the magnitude form of intrinsic luminosity: the **absolute magnitude**. This is the apparent magnitude that would be observed if the source lay at a distance of 10 pc:

$$\boxed{M = m - 5\log_{10}\left(\frac{d}{10\mathrm{pc}}\right),} \tag{7}$$

where d is the distance to the source. The quantity $m - M = 5\log_{10}(d/10\mathrm{pc})$ is also called the *distance modulus*.

4 Black Body Radiation

Black body radiation is a special example of a brightness or specific intensity which applies to the thermal radiation from a black body, an object with absorptive power of unity. The brightness of a black body is given by the *Planck function*:

$$B_\nu(T) = \frac{2h\nu^3}{c^2}\frac{1}{e^{h\nu/(kT)} - 1}, \tag{8}$$

$$B_\lambda(T) = \frac{2hc^2}{\lambda^5}\frac{1}{e^{hc/(\lambda kT)} - 1}, \tag{9}$$

where h is one of two incarnations of *Planck's constant*: $h = 2\pi\hbar$, where $\hbar = 1.055 \times 10^{-34}$ Js. Some key things to note about Black Body radiation are:

(1) The form of the function is uniquely determined by the temperature T.

(2) The black–body spectrum peaks at a wavelength $\lambda_{\rm peak}/\mu{\rm m} \simeq 2900\,{\rm K}/T$.

(3) Many astronomical objects turn out to be in good enough thermal equilibrium that they closely resemble black bodies.

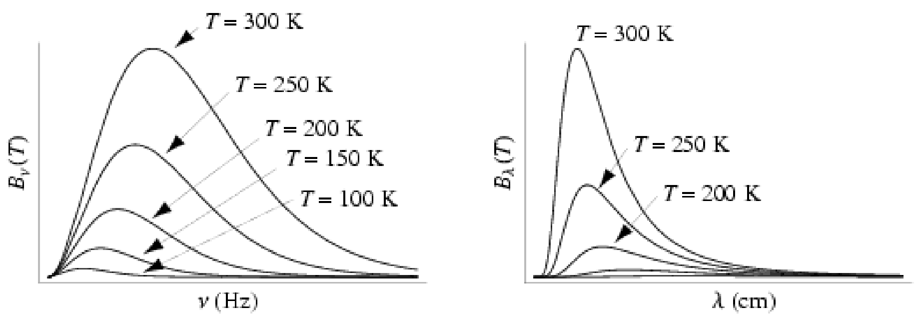

Figure 2: Spectrum of black body radiation

The amount of energy emitted per unit time per unit surface area of a black body is σT^4 where T is the temperature of the black body, and $\sigma = 5.67 \times 10^{-8}\ {\rm W\,m^{-2}K^{-4}}$ is the *Stefan-Boltzmann constant.* Although astronomical objects are not perfect black bodies, they are close enough that it is possible to define an **effective temperature**, such that a perfect black body of this temperature would radiate the same total luminosity as the real star:

$$\boxed{L = 4\pi R^2\,\sigma\,T_{\rm eff}^4,} \tag{10}$$

where R is the radius of the star. The effective temperature of the Sun is about 5770K.

The effective temperature is difficult to determine observationally, because one needs to know the total output over all frequencies. Therefore source temperatures are often estimate by other means:

Wien's Displacement Law: As mentioned above, differentiating the Planck function shows that it peaks at $\lambda_{\rm peak}/\mu{\rm m} \simeq 2900\,{\rm K}/T$. If one can determine the wavelength at which the observed spectral energy distribution (SED) peaks, this therefore yields an estimate of T. For the Sun, $\lambda_{\rm max} \simeq 0.5\mu$m. For a very cool star $\lambda_{\rm max} \simeq 1\mu$m. For our environment $\lambda_{\rm max} \simeq 10\mu$m. For the cosmic background $\lambda_{\rm max} \simeq 1$ mm.

Colour temperature: One cannot always guarantee to be able to locate the peak of an SED. However, one can estimate T by determining the local gradient (colour) of the SED and comparing that with a black body. It is useful for stars in the visible region, but not in the radio or infrared. In the visible we can assume the Wien approximation:

$$B_\nu d\nu = \frac{2h\nu^3}{c^2} e^{-h\nu/kT} d\nu. \tag{11}$$

There are other methods to estimate temperature which do not depend upon the black body spectrum. For example, the temperature of stars can often be estimated from the observed absorption lines from their stellar atmospheres, the ratio of which depends on the population of the atomic energy levels, which in turn depends on temperature via the Boltzmann equation.

5 Stellar Classification

5.1 Spectral types

The spectra of stars are close to black body in shape, since they are optically very 'thick': an individual photon takes many years ($\sim 3 \times 10^4$) to random walk to the surface of the sun from its centre due to repeated scattering by electrons, and so there is plenty of opportunity for extensive exchange of energy between photons and matter, leading to thermal equilibrium locally.

The spectra of stars also show large numbers of absorption lines (e.g. see Figure 3). These arise largely in the thin surface layers comprising the atmospheres of the stars. These atomic absorption lines come in families, of which the most familiar is the spectrum of Hydrogen:

$$h\nu = E_n - E_m = 13.6(1/n^2 - 1/m^2) \text{ eV} \tag{12}$$

(where photon energy $h\nu$ is approximately 1.25 electron volts (eV) at $\lambda = 1\mu$m). The most prominent lines in the optical are the *Balmer series*, which have $n = 2$ as their lowest level. The hydrogen series produce some of the strongest absorption lines in the hottest stars, with temperatures $8000 < T_{\rm eff} < 40000$ K, but absorption lines due to a wide range of other chemical elements are also found: Si, N, Ca, Fe, etc. The coolest stars, with $T_{\rm eff} < 4000$ K, even show absorption bands from molecules like TiO, CN, CH, and others.

In the 1880s, Pickering and Cannon developed a scheme for classifying stars into different spectral classes (the 'Harvard' types) based on the strongest absorption lines apparent in the star's spectrum. This is the classification scheme commonly used today. Unfortunately, their logical classification (ABCDE. . . , with 10 sub-classes 0. . . 9) was based on the relative strength of hydrogen absorption lines, which do not vary monotonically with temperature. It was Annie Jump Cannon who argued in the early 20th Century for a physical classification in order of temperature, but unfortunately the old letters were retained. The order of the sequence is therefore:

O B A F G K M.

This is often remembered by the mnemonic

Oh, Be A Fine Girl/Guy, Kiss Me

Alternatives include 'Only Bored Astronomers Find Gratification Knowing Mnemonics'.

The Sun has spectral type G2 ($T \approx 6000$K). Vega is a type A0 star ($T \approx 10000$K).

Table 2: Spectral properties of stars

O	T>20000 K	Once ionised He (HeII) lines visible; H lines weak; higher ionisation species (CIII etc) present; strong UV continuum
B	T=20-10000 K	Neutral Helium dominate (He I); HeII lines absent
A	T=10-7000 K	Neutral Hydrogen lines dominate (Balmer lines).
F	T=7-6000 K	H lines weaker; lines from neutral and once-ionised metals
G	T=6-5000 K	Solar type spectra; absorption lines of neutral metallic atoms and ions (e.g. once ionised calcium) grow in strength
K	T=5000-3500 K	Metal lines dominate; molecular bands (CH, CN) getting stronger
M	T=3500-2500 K	Strong molecular bands, particularly TiO; very red continuum

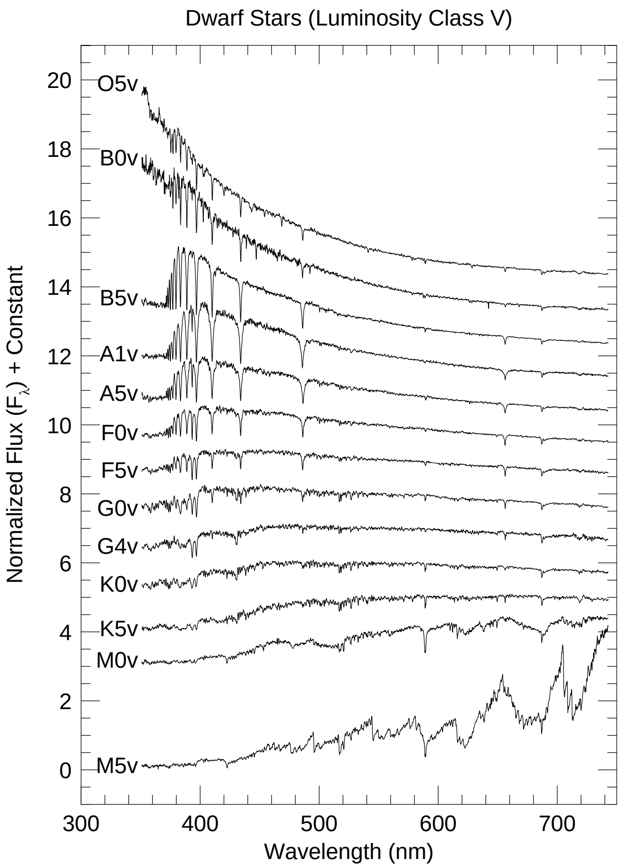

Figure 3: Illustrating how the spectra of main-sequence stars change systematically from types O5v through M5v ('early' to 'late') spectral types (from Jacoby et al. 1984, ApJS, 56, 257).

Table 2 provides details of the absorption line spectra of the different stellar types, and Figure 3 illustrates their spectra. Notice that the *colour* of stars within the visible spectrum gets bluer as the temperature goes up, as expected due to the nearly black-body nature of the emission. This means that the colour of a star can also be used estimate the temperature. Stellar colours most commonly involve the UBV system developed by Johnson and Morgan in 1951. A colour such as $B-V$ is simply the difference of the magnitudes measured in these two wavebands, and so larger colours indicating redder spectra. Since, by definition, Vega has $B-V=0$, and $U-B=0$, we would expect any A0 star to have zero colour, and would expect *hotter* stars (*i.e.* O and B stars) to have negative colours, and *cooler* stars (F $\rightarrow$ M) to have increasingly positive colours. This is broadly true, although the relation isn't monotonic for $U - B$ due to, for example, the Balmer discontinuity at about 400nm.

5.2 The Main Sequence

One of the most important relationships in all of stellar astronomy is that between the absolute magnitude of a star and its spectral type. It was noticed in 1905 by Hertzsprung and independently in 1913 by Russell that most stars (including our sun) fall on a narrow locus in the $M - Sp$ plane (M = absolute magnitude; Sp = spectral type). This locus is called the **Main Sequence**. The plot of absolute stellar magnitudes against stellar type is credited to both and is known as the **Hertzsprung-Russell diagram** (or often just as the 'HR diagram').

5.3 MK Classification

It was noticed by Hertzsprung that although some stars deviated from the Main Sequence, these were often of the same spectral type, as determined by their absorption lines, as other stars on the Main Sequence. Since stars are good blackbodies, it follows that more luminous stars of the same effective temperature must have larger surface areas, and thus must be bigger. In the 1940s and 1950s, Morgan and Keenan generalised the classification of stars by dividing them into luminosity classes, as detailed in Table 5.3. Main Sequence stars are known as dwarf stars, whilst the more luminous stars comprise different classes of giants.

Table 3: The MK classification scheme for stars

I	Supergiants
II	Bright Giants
III	Giants
IV	Subgiants
V	Main Sequence (dwarfs)
VI	Subdwarfs

In practice, this classification uses the shape of spectral lines to measure surface gravities of stars. The gravitational acceleration on the surface of a giant star is much lower than for a dwarf; the lower gravity means that gas pressures and densities are lower, so that spectral lines are narrower in these stars. Classifying a star according to both its luminosity class and its spectral type forms the basis of the MK Classification scheme of stars, which is central to stellar astronomy.

5.4 Colour–magnitude diagrams

Instead of plotting absolute magnitude against spectral type, it is often convenient to plot it against colour, since both of these may be measured directly from stellar photometry with no need for doing spectroscopy. An example HR diagram is shown in Figure 4.

The colour-magnitude diagram shown in Figure 4 fails to do justice to the precise physical correlations seen in systems of stars. The main sequence is blurred because the parallax distances are not perfectly accurate, and because the location of the main sequence depends on metallicity (more absorption lines in high-metal stars). Moreover, the Solar neighbourhood contains a mixture of stars of different ages, and age is a critical feature of the colour–magnitude diagram. This is illustrated in Figure 5, which shows the colour–magnitude diagram for a globular cluster. Here, all

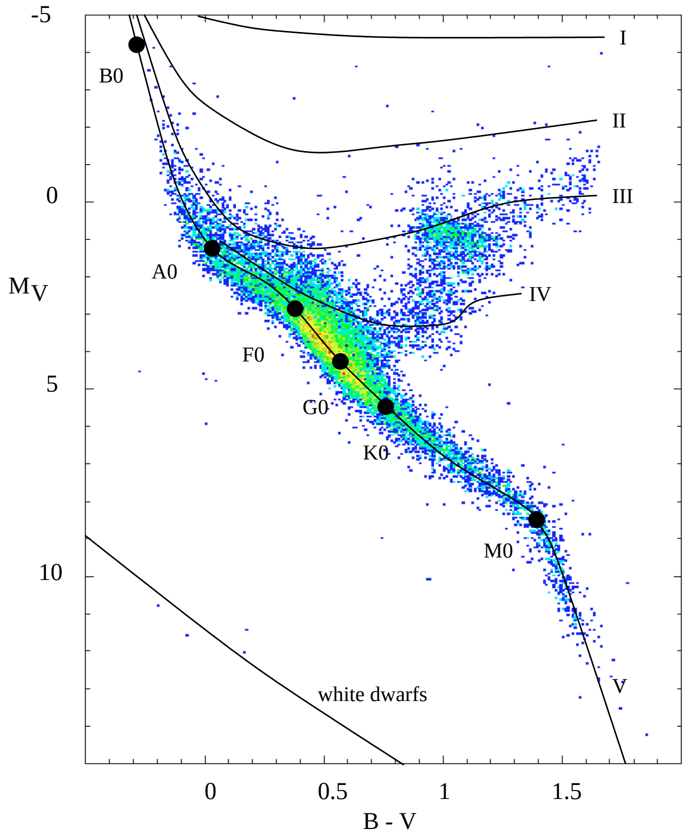

Figure 4: The colour-magnitude diagram (the observational version of the Hertzsprung–Russell diagram) for the Solar neighbourhood, showing 16631 stars from the Hipparcos catalogue with parallax distances known to better than 10%. The loci of the different luminosity classes are shown, as well as the locations of the different main-sequence spectral types.

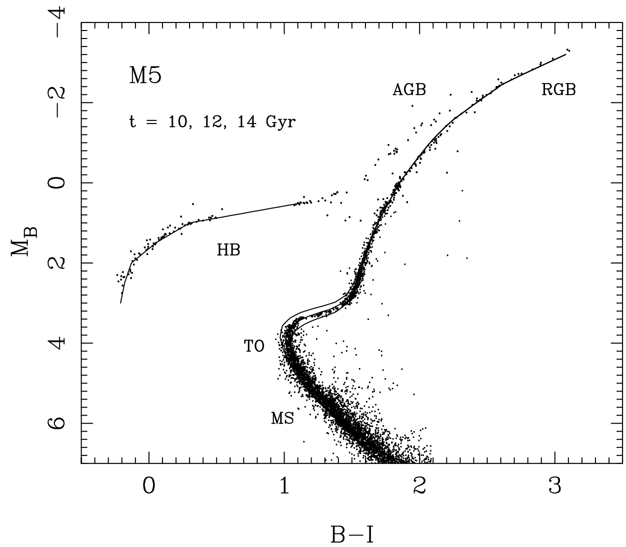

Figure 5: A colour–magnitude plot for stars in the globular cluster M5. This illustrates well the main features of stellar evolution: the main sequence (MS) and its turn-off point (TO), followed by the red giant branch (RGB), horizontal branch (HB) and asymptotic giant branch (AGB). The data are well-fitted by a theoretical isochrone of age 12 Gyr.

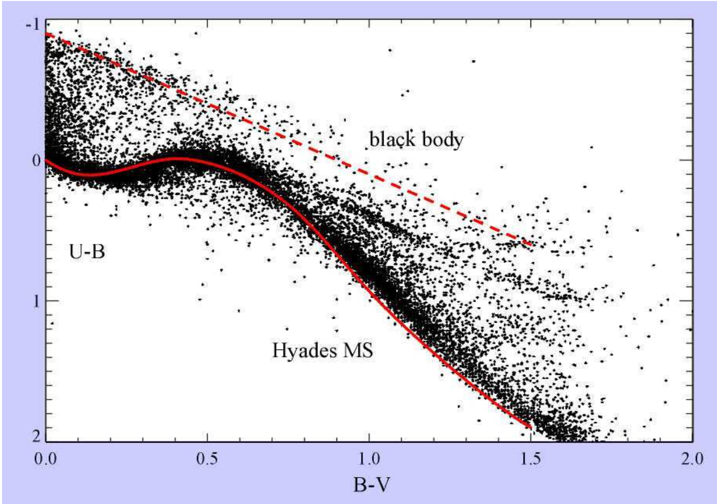

Figure 6: A $U - B$ vs $B - V$ colour diagram for the Hyades cluster. The dashed line indicates the locus for a perfect black body, whilst the S-shaped grey line shows the expected locus of the main sequence.

the stars are at the same distance, age and metallicity, so we can see how well defined the stellar loci are, illustrating not only the main sequence, but the giant branches. The form of this diagram is due to stars at the top left of the main sequence having used up a large amount of the 'fuel' in their cores, which causes them to move off the main sequence. They become giant stars with lower surface temperatures, but hugely greater sizes and hence greater luminosities.

The arguments above suggest that the colours of main sequence stars are well-defined, such that they can be used directly to classify the stellar types. Equivalently, in a two–colour plot of $U - B$ against $B - V$, the stars should follow one fixed curve. However, the actual plot with observed data look like Figure 6. The reason why this diagram looks so messy is that interstellar dust grains absorb the light from a star on the way to the Earth. The amount of absorption depends on wavelengths: within the optical/infrared range, the amount is larger at shorter wavelengths (varying roughly as $\sigma(\lambda) \propto \lambda^{-1}$), so the colour of the stars gets *reddened*. Therefore this effect is called **interstellar reddening**. We'll discuss this in more depth in the section of the course on interstellar nebulae.

6 Astronomical Observations

6.1 Co-ordinate systems

To locate an object on the sky, astronomers need to define two–dimensional co-ordinate systems within the celestial sphere. These co-ordinate systems follow the same principal as the definition of the geographic co-ordinates latitude and longitude on the earth. They have: (i) a principal axis about which the systems rotates [the earth's rotation axis for geographic co-ordinates]; (ii) perpendicular to this, a great circle which is the principal reference circle along which one co-ordinate is measured [the equator, around which longitude is defined, for geographic co-ordinates]; and (iii) an infinite number of secondary reference circles that are great circles perpendicular to the principal reference circle, which meet at the poles of the principal axis [lines of constant longitude for geographic co-ordinates], and along which the second co-ordinate [latitude] is measured.

The most important co-ordinate system in astronomy is the **Equatorial** co-ordinate system, which is universally used for extra-galactic objects, and often used for galactic targets as well. It uses

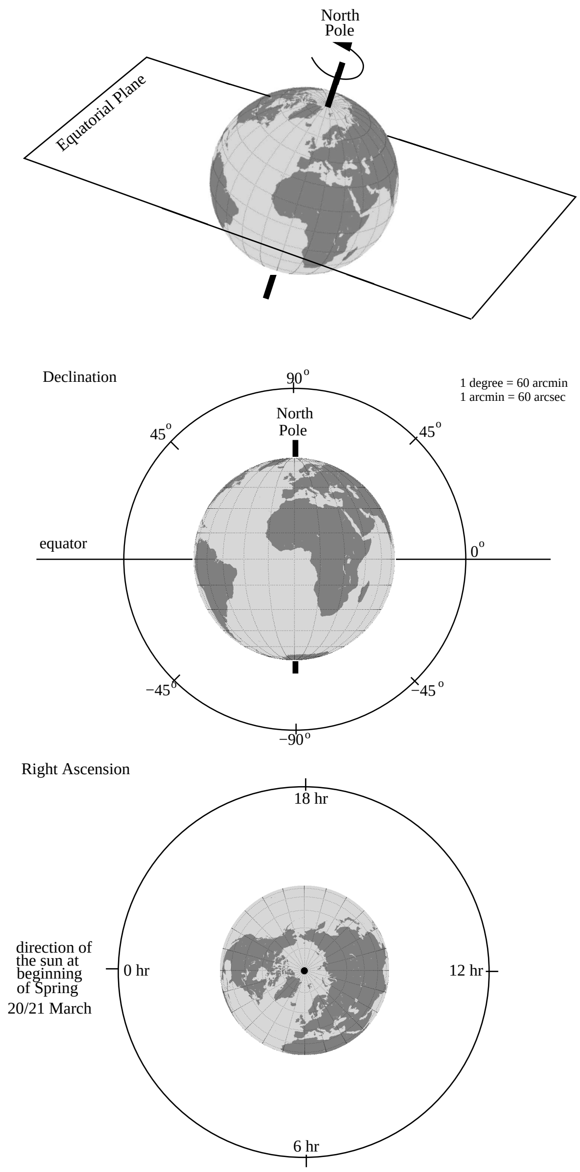

Figure 7: Equatorial coordinates - right ascension and declination.

the rotational plane of the earth as its reference. The co-ordinates are **Right Ascension (RA)**, defined around that rotational plane, and usually measured in hours (from 0 to 24, increasing to the east), and **Declination (Dec)** ranging from 0 at the equator to ±90 degrees at the north and south poles (see Figure 7). We need to define the zero point of Right Ascension (i.e. 0h direction) in some way. The definition involves the **ecliptic**, which is the apparent path of the Sun's motion on the celestial sphere as seen from Earth. The ecliptic plane is tilted 23.5° with respect to the plane of the celestial equator since the Earth's spin axis is tilted 23.5° with respect to its orbit around the sun. The celestial equator therefore intersects the ecliptic at two points, which are called *equinoxes*. The direction of the right ascension 0h is defined as the direction of the Sun at the northern vernal (spring) equinox (around 21 March).

The exact rotational plane of the earth changes, mainly due to a wobble with a period of 26,000 years called **precession**. Therefore the coordinates have to be specified with the date which those positions refer to. The positions usually refer to the coordinates with the equinox direction of B1950 (the year/date 1950.0 in Besselian epoch) or J2000 (the year/date 2000.0 in Julian epoch).

If the object in question has a proper motion, you also need to specify the exact observing date (epoch) for the positions you are talking about.

6.2 Where and when to observe?

The Sun travels on the celestial sphere from RA=0h at around 21 Mar, to RA=6h around 21 Jun, RA=12h at around 21 Sep, to RA=18h at around 21 Dec, and then comes back to RA=0h again in Mar. For optical observations, we of course need to avoid the Sun. Therefore, the best observing season should be when the RA of the Sun is different from that of the object by about 12 hours.

Another astronomical way of saying this is that the best time is when the **Local Sidereal Time (LST)** at midnight for the telescope site is close to the RA of the object. LST is the RA of the objects transiting the Meridian of the observer at a given location. So the LST at midnight for that location is essentially the RA which is 12h different from the RA of the Sun.

The declination of the object determines the suitable latitude range of the observing sites. For example, Mauna Kea in Hawaii, which is one of the main observing sites, is at latitude $\approx 20°$. Objects with declination less than about -70° are invisible from there, and we usually don't want to observe objects with declination less than about -30°, since they are always quite far from zenith.