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ORIGINAL ARTICLE

An Introduction to Observations Relevant to Astrophysical Jets and Nebulae

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Abstract This article reviews the basic physics and jargon associated with astronomical observations of nebulae, with an emphasis on processes relevant to shock waves in astrophysical jets.

Keywords Astronomical observations · Shock waves · Nebulae

1 Motivation

The HEDLA meetings bring together laboratory experimentalists, numerical modeling experts and observational astrophysicists to study how plasmas and fluids behave in a wide variety of conditions in nature. While this synthesis presents unique opportunities for collaborative research, communication between the different disciplines can be problematic, as each field has jargon and conventions that are not immediately transparent to scientists in other fields. As the oldest science, it is perhaps not surprising that astronomical conventions can be particularly arcane, though with a bit of background they quickly become second nature.

This contribution provides a brief overview of the physics, conventions, and nomenclature used when describing and interpreting astronomical images of nebulae, with the objective to make it easier for a non-specialist to understand an observational astronomical talk on this subject. Most astronomical observations are either images or spectra (though some are both!), and in what follows I treat each of these in turn. For more information about physical processes that influence spectra I refer the reader to the classic texts 'Radiative Processes in Astrophysics' by Rybicki and Lightman (1979) for continuum processes, and to 'Astrophysics of Gaseous Nebulae and Active Galactic Nuclei' by Osterbrock (1989) for emission line spectra.

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2 Images

Astronomical images can be spectacularly beautiful, but what they tell us physically about the objects is determined to a large degree by the answers to two questions: (i) *What is the scale/resolution?*, and (ii) *What do the colors represent?*

2.1 Distance scales and units

Astronomers typically use cgs-Gaussian units, but for convenience we like to represent planetary and solar-system scale distances in Astronomical Units (1 AU = 1.495×10^{13} cm), stellar distances in parsecs (1 pc = 3.09×10^{18} cm = 3.26 light years), and masses and luminosities in solar units (1 $M_{\odot}$ = 1.99×10^{33} gm, 1 $L_{\odot} = 3.83 \times 10^{33}$ erg s^{-1}). The AU is the distance from the Earth to the Sun, and a parsec is a typical distance between stars in the solar neighborhood. For reference, the average distance from the Sun to Pluto is about 40 AU, to the nearest star is 1.3 pc, the nearest region of massive star formation (Orion) is 460 pc, the center of our galaxy 8.5 kpc, and the nearest large external galaxy (M31) 0.77 Mpc.

Distances are determined from a variety of methods, but the most direct one, applicable for the closest objects, is to observe the angular shift, known as the parallax, of the object relative to distant background stars as the Earth moves around the Sun (Fig. 1). A parsec is defined by the distance an object would have to be in order to have a parallax of 1 arcsecond. Hence, the number of AU in a pc is 206265, the same as the number of arcseconds in a radian.

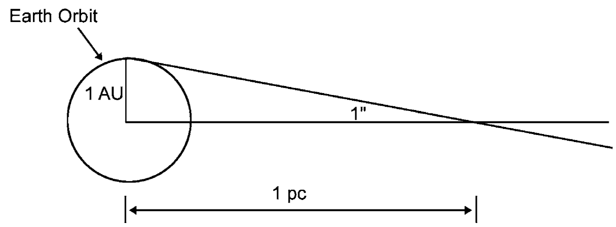

Fig. 1 Definition of a parsec

The theoretical spatial resolution limit of an image is roughly λ/D, where λ is the wavelength of the light and D the diameter of the telescope. However, atmospheric turbulence limits all ground-based optical images to ∼1 arcsecond; observations in the near-infrared fare somewhat better and routinely obtain 0.5 arcseconds at a good site. With the use of adaptive optics to correct for atmospheric distortion one can approach the theoretical limit of λ/D, but such images have very small fields of view at present. A good rule of thumb is that a ground-based image is 1 arcsecond, and a space-based or radio image is λ/D (e.g. Hubble Space Telescope is ∼0.07 arcseconds).

By combining the above considerations it is straightforward to quickly infer distance scales for any astronomical image, provided you know the distance to the source. For example, if you see a ground-based image of the Orion Nebula, then the spatial resolution is $\sim 1''$, which at 460 pc corresponds to 460 AU, or about 6 times the diameter of our solar system. An HST image of a large region of ionized gas at 2 kpc has a resolution of $2000 \times 0.07 = 140$ AU, and so on.

2.2 Colors, magnitudes, photometry, broadband and narrowband images

Modern multicolor astronomical images are created by loading individual images into red, blue, and green channels, each image consisting of continuum and emission lines transmitted by the filter. Often the longest wavelength is put into the red channel, and the shortest into the blue, but one can also put images of the lowest ionization states in red and the highest ionization states in blue, or simply choose whichever channels make the result pleasing aesthetically. Hence, color composites can provide a great deal of information about the physics of the region, or none at all, depending on the composite. It is also possible to create a 'false color' image where the intensities of a single image are assigned a specific color. Such images give no more information than greyscale images or contour plots do, and in this case the colors are completely arbitrary. X-ray and radio continuum images often appear in false color.

The Earth's atmosphere transmits from about 0.35 μm to about 1 μm, and this range is typically broken up into five bandpasses roughly 0.1 μm in width labeled U, B, V, R, and I for ultraviolet, blue, visual, red, and infrared, respectively. The Earth's atmosphere is also transparent in several 'windows' throughout the near- and mid-infrared, including the three near-IR bandpasses J, H, and K at 1.25, 1.65, and 2.2 μm, respectively, and several bandpasses at mid-IR wavelengths that range out to 20 μm. However, ground-based telescopes emit thermal radiation at mid-IR wavelengths, which is why a small space telescope like Spitzer, cooled to cryogenic temperatures, is much more sensitive than are larger telescopes on the ground at these wavelengths.

To quantify brightnesses in these broad bandpass filters, astronomers define a magnitude at each wavelength as Δmag $= 2.5 \log_{10}(F_2/F_1)$, where Δmag is the magnitude difference between objects with fluxes F_1 and F_2. The scale is defined so brighter is smaller, with the Sun having an apparent magnitude of about -26.5 at V, while the stars in the Big Dipper are about $+2$. The zero point for each wavelength is approximately the brightness of the star Vega in the northern summer sky. Color indices are defined as the difference between two magnitudes, e.g. B$-$V, where by convention the bluer filter is first so that redder objects have more positive colors. For objects like stars that radiate nearly like blackbodies, bluer colors mean higher surface temperatures (Fig. 2). In stellar evolution studies one often plots an 'H-R diagram', of either magnitude vs. color, or log(L) vs. log(T), where L is the luminosity (erg/s) and T the surface temperature of the star.

If the object is a nebula that emits primarily line radiation and not continuum, then the apparent brightness simply depends on how many emission lines fall within the bandpass

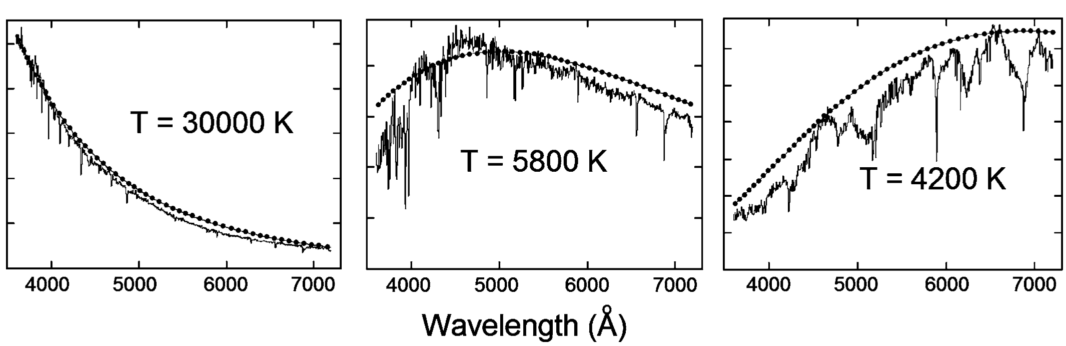

Fig. 2 Examples of three optical spectra of stars that have different surface temperatures. Blackbody fits to the spectra appear as dotted curves

of the filter. Typically one uses a narrowband filters to isolate specific emission lines of interest. Narrowband images also minimize the relative brightness of stars, which, being continuum sources, emit at all wavelengths including those transmitted by narrow band filters.

3 Spectra

3.1 Emission line spectra, densities, temperatures, velocities

The most direct way to determine physical conditions within a nebula is to obtain a spectrum at each point. Many spectrographs disperse the light perpendicular to a long slit, and generate a spectrum at each position along the slit. Mapping the slit across the entire nebula produces a 'data cube' for each emission line, where the counts at a fixed wavelength produce an image at the corresponding radial velocity. Adding all the velocities together gives an emission line image. Instruments such as image slicers and Fabry-Perots give data cubes, as do most molecular line maps obtained with radio telescopes. Data cubes are both images and spectra, and, when combined with proper motion measurements that show how the object moves in the plane of the sky, represent the most thorough observational description of the object possible. Data cubes are time-consuming to obtain and can be challenging to analyze.

The Doppler shift of emission lines in the spectra relative to the velocity of the ambient medium (usually the parent molecular cloud or a protostar in the case of stellar jets) reveals the kinematics of the gas, and the observed emission line width clarifies the gas dynamics such as temperature and turbulence. Because nebulae are optically thin (transparent) to most photons with energies a few eV, the emission lines observed at one position are actually integrals over the entire line of sight through the nebula.

Heating and cooling within nebulae are governed by non-LTE processes. Nebular densities are very low, so nearly all the atoms and ions are in their ground states. Typical nebular densities range 10–10^6 cm^{-3} (note that this number must be multiplied by $\sim 1.67 \times 10^{-24}$ to obtain g cm^{-3} for pure H composition). Nebular gas cools primarily as free electrons collide with atoms and ions and excite them to upper states, which then decay back to the ground level. In this way, kinetic energy of the electrons is converted to light, which escapes, so the nebula cools. Heating is usually accomplished by photoionization or by shock waves. In an 'H II' region, ultraviolet photons from a hot star ionize nebular gas and deposit any energy in excess of the ionization threshold into kinetic energy of the freed electron. Alternatively, shock fronts suddenly decelerate gas and convert a fraction of the bulk kinetic energy into heat.

The kinetic temperature is determined by a balance of heating and cooling, and for an H II region is typically 10^4 K. Temperatures immediately behind shock waves are proportional to the square of the shock velocity, and the temperature then declines in an extended 'cooling zone' as the gas radiates emission lines. Astronomers refer to shocks that cool by emitting photons as 'radiative', a definition which differs from that used in high energy density physics, where radiative shocks are those where radiation is an important component to the total energy content of the gas. In astronomical nomenclature, a nonradiative shock is one where the timescales for cooling are so long that the postshock gas cools by some other means, such as expansion.

It is usually straightforward to measure the electron density and temperature in a nebula. Many of the most abundant ionization states (e.g. O I, O II, O III, N I, N II, C I, S II, S III) have electronic configurations which have 2, 3, or 4 electrons in an outer p-shell. As shown in Fig. 3, p^3 configurations always have a close doublet at a few eV above ground, and transitions from these levels are particularly good for measuring densities. Let us denote level 1 as the ground state, and levels 2 and 3 as a closely spaced doublet excited state. Because levels 2 and 3 have nearly identical excitation energies, their relative density n_2/n_3 is independent of temperature. However, the flux ratio F_{21}/F_{31} of doublet lines to the ground state depends strongly on the density. In the low density limit, every collisional excitation is followed by radiative decay. In this limit, F_{21}/F_{31} equals the ratio at which the two levels are populated from the ground, which is typically the ratio of statistical weights g_2/g_3 ($\nu_{21}/\nu_{31} \sim 1$ for a

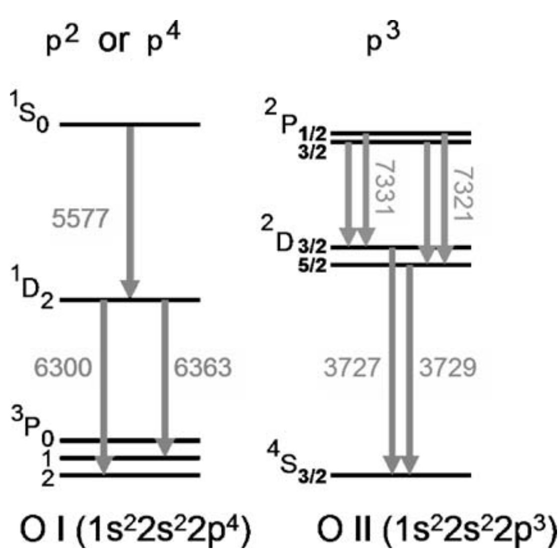

Fig. 3 Left – an energy level diagram for O I, an ion with 4 electrons in an outer p-shell. Line ratios between states highly separated in energy (e.g. λ5577/λ6300) constrain the temperature. Right – the same diagram but for O II, which has 3 electrons in the p-shell. The emission line ratio of λ3727/λ3729 determines the electron density between the low density limit ($\sim$50 cm^{-3}) and the high density limit ($\sim 2 \times 10^4$ cm^{-3}). All the lines depicted in this figure are forbidden transitions (electric quadrupole or magnetic dipole), with wavelengths in Angstroms

closely-spaced doublet). When the density exceeds some critical value where collisional deexcitation rates exceed those for spontaneous decay, n_2/n_3 approaches the LTE value of g_2/g_3, and the flux ratio $F_{21}/F_{31} = (n_2 A_{21} \nu_{21})/(n_3 A_{31} \nu_{31}) \sim (g_2 A_{21})/(g_3 A_{31})$, which differs by a factor of A_{21}/A_{31} from the value in the low density limit. Hence, measuring the line ratio from a closely-spaced doublet gives the electron density if the density lies between the low density limit and high density limit for the transition (see Osterbrock 1989, p134, for examples of diagnostic curves).

In general, an observed flux ratio between two transitions from an ion or atom defines a curve in Ne–Te space. The temperature dependence is more pronounced when the levels are more separated in energy. Line ratios between different ions determine the ionization fraction, and between different elements measure the relative abundances. A good historical example of such an analysis applied to shocks in jets is the classic paper by Brugel et al. (1981).

In shock waves, the gas gradually becomes more neutral as it recombines. A hot star that illuminates a dark cloud of gas and dust sets up a similar stratified ionization structure, where higher ionization states occur on the side of the nebula which faces the star, and more neutral species dominate at greater distances into the cloud where ultraviolet light from the star becomes attenuated. In photoionized regions there is a balance between the ionization rate, which is proportional to $a_\nu F_\nu N_{Xi}$, and the recombination rate αN_{Xi+1}Ne. Here, a_ν is the photoionization cross section, α is the recombination rate coefficient, F_ν is the ionizing flux, Ne, N_{Xi}, and N_{Xi+1} are, respectively, the density of electrons, atoms in ionization state i, and atoms in ionization state $i + 1$. Equating these rates, the ionization parameter N_{Xi+1}/N_{Xi} is proportional to $a_\nu F_\nu$/Ne, which is why lower density nebulae have higher ionization fractions for a given ionizing flux.

3.2 Continuous spectra

Nebulae also emit continuum radiation from a variety of processes. Although less diagnostic than emission lines, continuum radiation also provides information concerning the state of the plasma responsible for the emission. Continuum radiation is characterized by the spectral index α, where $F_\nu \sim \nu^\alpha$. For example, $\alpha = 2$ on the Rayleigh-Jeans portion of a blackbody spectrum. A spectral index is convenient because as long as one observes the source at two different frequencies, it is possible to fit a power law through those two points and obtain a spectral index. Of course, the source is probably not emitting according to a power law, but one can always define $\alpha = \mathrm{d} \ln F_\nu / \mathrm{d} \ln \nu$.

Spectral indices for continuum processes change depending on whether or not the optical depth τ of the source is thin ($\ll$1 or thick $\gg$1) For example, optically thick thermal free-free radiation from a plasma is identical to that from a blackbody, while optically thin free-free has a nearly flat spectral index ($\alpha \sim -0.1$ at optical/IR wavelengths). For synchrotron radiation, which occurs as electrons spiral relativistically around magnetic field lines, optically thick emission at long wavelengths has a spectral index of 5/2, but $\alpha = -(p-1)/2$ at short wavelengths, where p represents the energy spectrum of the electrons, $n(E)\mathrm{d}E \sim E^{-p}$ (cf. Rybicki and Lightman, 1979).

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