

# Deriving chemical abundances in ionised gas nebulae from the optical spectrum

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This is a brief manual of instructions used to calculate electronic densities and temperatures, ionic and total abundances of the main chemical species found in gaseous nebulae showing an emission line spectrum.

Our main objective is to present some of the fittings obtained by means of the IRAF tasks *TEM DEN* and *IONIC* in the STSDAS package. These can then be used inside programs in order to achieve a complete diagnostic for a large number of objects (e.g. Pérez-Montero & Díaz 2005).

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## 1. Introduction

(From Pérez-Montero & Díaz, 2005)

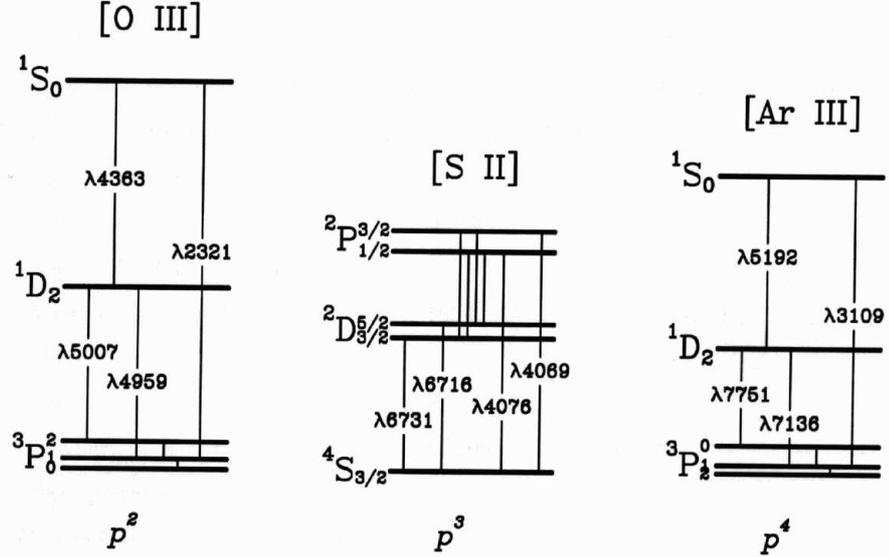
HII regions, from faint ones on the Galactic disk to Giant Extragalactic ones (GEHR) and HII galaxies, constitute the main source of information about the metallicity of all kind of objects, including the most distant ones. Their bright spectra of emission lines can be seen in all those regions that have gone through a recent star formation episode. The analysis of such spectra provides the best method (if not the only one) to determine the chemical abundances of elements like helium, nitrogen, oxygen, neon, argon, sulfur, displaying prominent emission lines in the optical range corresponding to several ionization states. Detailed knowledge of these abundances is essential for fully understanding gaseous and stellar evolution, and has helped us towards the solution of some yet unsolved questions regarding the chemical evolution of galaxies in the local Universe. These questions are becoming more and more relevant for a more distant – hence younger – Universe, as we are able to obtain data on fainter objects.

Recombination lines, weakly dependent on the nebular internal temperature, provide more precise abundance determinations. He abundance, for instance, can be derived with precision better than 5%. Unfortunately, many of the emission lines are collisionally excited and their intensities depend exponentially on the temperature. In principle, the temperature can be determined from appropriate line intensity ratios which however require the detection and measurement of intrinsically faint (or absent) auroral lines. This is in particular the case for regions with high metal content – for which the cooling produced by the metals is very efficient and the lines are no longer detected – and also for distant HII regions and regions with low surface brightness. For these cases, empirical methods based on the strong nebular line intensities are used.

The main sources of uncertainty associated to the derived physical properties are related with the error of the fluxes of the lines. The measure of these fluxes is usually done by means of tasks like *SPLIT* of IRAF, that integrates the intensity of each line over a local continuum. The errors of the fluxes measured in this way can be calculated with the expression:

$$\sigma_l = \sigma_c \cdot \sqrt{N + \frac{EW}{\Delta}}$$

where  $\sigma_l$  is the error of flux of the line,  $\sigma_c$  represents the standard deviation in a box near of the measured emission line and represents the error in the position of the continuum.

FIGURE 1. Energy levels and main transitions for np<sup>2</sup>, np<sup>3</sup> and np<sup>4</sup> type ions.

,  $N$  is the number of pixels in the measure of the flux of the line,  $EW$  is the equivalent width of the line and  $\Delta$  is the dispersion of the wavelength in angstroms per pixel.

Each one of the intensities measured in this way must be reddening corrected. The reddening constant  $c(H\beta)$  can be calculated from the decrement of Balmer of the stronger recombination lines of hydrogen. The measurement of these lines must be done, generally, taking into the account the absorption of the underlying stellar population. The physical law that allows to calculate the reddening constant from this decrement and therefore to correct the other emission lines from reddening takes the form:

$$\frac{I_0(\lambda)}{I_0(H\beta)} = \frac{I(\lambda)}{I(H\beta)} \cdot 10^{-c(H\beta) \cdot [f(\lambda) - f(H\beta)]}$$

where  $I(\lambda)$  is the intensity of the line we want to correct and  $I_0(\lambda)$  is the corrected value in relation to  $H\beta$ ,  $c(H\beta)$  is the reddening constant and  $f(\lambda)$  is the extinction function. The theoretical values of the hydrogen recombination lines that allow to calculate  $C(H\beta)$  are a function of the temperature and density so this imply the necessity of an iterative process to calculate them. This calculation can be done using the tabuled coefficients by Storey & Hummer (1995) and the corresponding program available in the CDS database †. Regarding the extinction function one can assume the law of Miller & Matthews (1972), which normalized to a value of  $R_V = 3.2$ , that is the value for the Galaxy takes the next form:

$$f(\lambda) - f(H\beta) = \begin{cases} 0.477\lambda^{-1} - 1.209, & \lambda^{-1} \leq 2.29 \\ 0.342\lambda^{-1} - 0.645, & \lambda^{-1} > 2.29 \end{cases}$$

† CDS webpage (Centre de données astronomiques de Strasbourg) is <http://cdsweb.u-strasbg.fr/>

where  $\lambda$  is the wavelength of the line we want to correct in micron unities. We can verify that this function takes a value equal to 0 for  $\lambda = 0.4861 \mu$ .

The uncertainty associated with the reddening correction must be taken into the account in the errors of the emission lines and convienly propagated into all the physical magnitudes we calculate. One must take into account, however, that in many cases, calibrations are used obtained from photoionization models for which the uncertainties are difficult to quantify. Even when the electronic temperature has been determined with high precision, there are some problems limiting the confidence on the results attained. Those problems include: (1) the effect of the internal ionization structure on multiple zone models (PMD03); (2) temperature fluctuations across the nebula (Peimbert, 2003); (3) collisional and density effects on the ionic temperatures (Luridiana, Peimbert & Peimbert, 1999); (4) neutral gas zones affecting the determination of ionization correction factors (ICFs) (Peimbert, Peimbert & Luridiana, 2002). ICFs are used to calculate total abundances from the abundances of the observed species

$$\frac{N(X)}{N(H)} = ICF(X_{obs}) \cdot \frac{N(X_{obs})}{N(H^+)}$$

(5) the ionization structure is not adequately described by present models (PMD03); (6) possible photon escape affecting low ionization lines in the outer regions of the nebula (Castellanos, Díaz & Tenorio-Tagle, 2002).

The first three effects can introduce uncertainties regarding the derived O abundance of some 0.2, 0.3 and 0.4 dex respectively, depending on the degree of excitation. The uncertainties associated to the rest of the enumerated problems have not yet been quantified.

## 2. Electron temperature and density

### 2.1. Oxygen

[O III] temperature is calculated from the ratio:

$$R_{O3} = \frac{I(4959\text{\AA}) + I(5007\text{\AA})}{I(4363\text{\AA})}$$

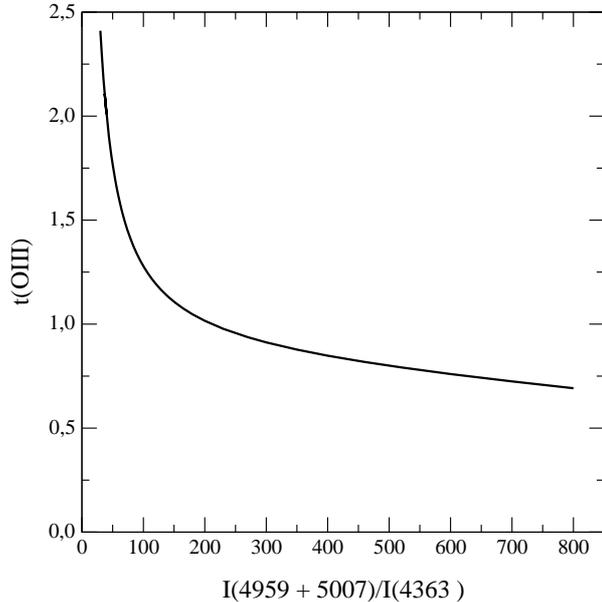
given that, according to Osterbrock (1989), temperature can be obtained from the ratio of collisional transitions that have a similar energy but occupy different levels. The fitting between the ratio and the electronic temperature was obtained using the task *TEM DEN* from the package *STSDAS* inside *IRAF*. This task is based on the program *FIVEL* (De Robertis, Dufour & Hunt 1987; Shaw & Dufour 1995) for a five level atom. In this case, and likewise for the rest of the ions unless stated differently, the transition probability, energy levels and collision coefficients are those given in such program by default. The obtained fitting, as for the rest of the  $np^2$  ions, is independent of the electronic density and is:

$$t([OIII]) = 0.8254 - 0.0002415R_{O3} + \frac{47.77}{R_{O3}}$$

where  $t$  is the temperature in units of  $10^4$  K.

For [O II] the quotient for the electron temperature is calculated from

$$R_{O2} = \frac{I(3726\text{\AA}) + I(3729\text{\AA})}{I(7319\text{\AA}) + I(7330\text{\AA})}$$

FIGURE 2. Relation between  $R_{O3}$  and  $t[\text{O III}]$ 

One has to be careful in this case, because the  $[\text{O II}]$  auroral lines might be contaminated by recombination emission. Such emission however, can be quantified and corrected for. According to Liu et al. (2000), such contribution can be fitted (for  $0.5 \leq t \leq 1.0$ ) by the function:

$$\frac{I_R(7319 + 7330)}{I(H\beta)} = 9.36 \cdot t^{0.44} \cdot \frac{O^{2+}}{H^+}$$

Moreover, the ratio of the  $[\text{O II}]$  lines is strongly dependent on the electron density. Ideally, one should know the  $[\text{O II}]$  density from the ratio  $I(3726\text{\AA})/I(3729\text{\AA})$  but very frequently we lack resolution to separate the doublet in which case one has to resort to the  $[\text{S III}]$  density, also representing the low excitation zone. The fitting obtained is:

$$t([\text{O II}]) = a_0(n) + a_1(n) \cdot R_{O2} + \frac{a_2(n)}{R_{O2}}$$

where the coefficients are respectively:

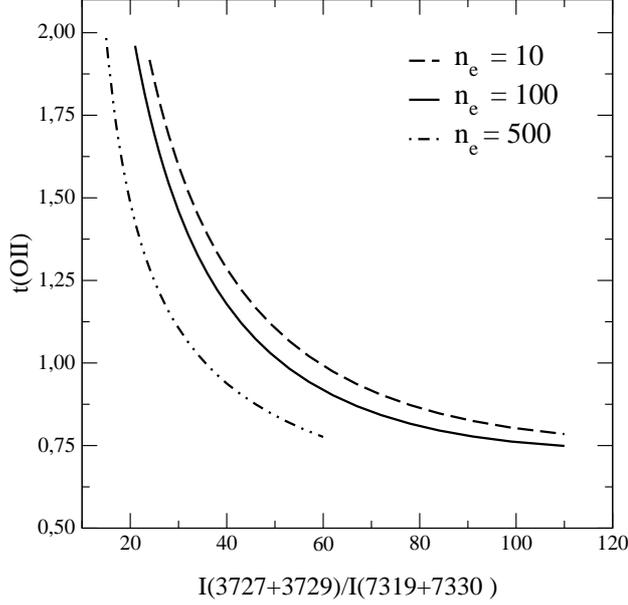
$$a_0(n) = 0.23 - 0.0005 \cdot n - \frac{0.17}{n}$$

$$a_1(n) = 0.0017 + 0.000009 \cdot n + \frac{0.0064}{n}$$

$$a_2(n) = 38.3 - 0.021 \cdot n - \frac{16.4}{n}$$

in this case, the atomic coefficients vary. PMD03 and PMD05 take the collisional strengths from McLaughlin & Bell (1998), as the default ones in CLOUDY 96. It is necessary to note, though, that Wang et al. (2004) underline that the coefficients fail to fit the observations (contrary to those of Pradhan, 1976). For the fitting presented here then, we have used Pradhan's.

Occasionally the auroral  $[\text{O II}]$  lines are not observed with good S/N or they are outside our observed spectral range. In that case it is practice to use some relation based on

FIGURE 3. Relation between  $R_{O2}$  and  $t[O\text{ III}]$ 

photoionization models in order to infer  $t[O\text{ II}]$  from  $t[O\text{ III}]$ . For instance, the relation

$$t([OII]) = \frac{2}{t([OIII])^{-1} + 0.8}$$

based on Stasińska (1990) models is frequently accepted. However, such expression neglects the dependence of  $t[O\text{ II}]$  on the density, consistent with the observed dispersion for the objects for which both temperatures have been derived from observations (see Figure 4). Also in Figure 4 we show a different set of models that do take this dependence into account (presented in PMD03). The fitting obtained in this case is:

$$t([OII]) = \frac{1.2 + 0.002 \cdot n + \frac{4.2}{n}}{t([OIII])^{-1} + 0.08 + 0.003 \cdot n + \frac{2.5}{n}}$$

## 2.2. Sulfur

The  $[SIII]$  lines ratio is commonly used to determine electron density for the low excitation zone. It is generally assumed that the nebula has constant density, although there is growing evidence for the existence of a density profile instead. Fortunately, the high excitation species have diagnostic ratios that are not density sensitive.

$$R_{S2} = \frac{I(6717\text{\AA})}{I(6731\text{\AA})}$$

Density is then calculated as:

$$n([SII]) = 10^3 \cdot \frac{R_{S2} \cdot a_0(t) + a_1(t)}{R_{S2} \cdot b_0(t) + b_1(t)}$$

where

$$\begin{aligned} a_0(t) &= 2.21 - 1.3/t - 1.25t + 0.23t^2 \\ a_1(t) &= -3.35 + 1.94/t + 1.93t - 0.36t^2 \end{aligned}$$

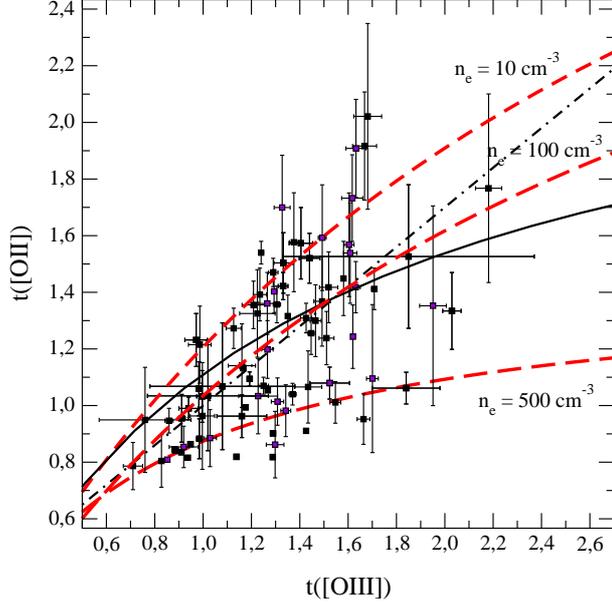


FIGURE 4. Relation between  $t[\text{O III}]$  and  $t[\text{O II}]$  for model sequences and for direct observations of H II galaxies

$$b_0(t) = -4.33 + 2.33/t + 2.72t - 0.57t^2$$

$$b_1(t) = 1.84 - 1/t - 1.14t + 0.24t^2$$

here  $t$  is generally  $t[\text{O III}]$ , although an iterative process could be used to calculate it with  $t[\text{S III}]$  given that this temperature, like  $t[\text{O II}]$ , a type  $np^3$  ion, is density dependent. The ratio used in this case is:

$$R'_{S2} = \frac{I(6717\text{\AA}) + I(6731\text{\AA})}{I(4068\text{\AA}) + I(4076\text{\AA})}$$

For the [S III] auroral lines it is enough to measure one of them, as they are related by a fixed theoretical ratio,  $I(4068\text{\AA}) \approx 3 \cdot I(4076\text{\AA})$ . We can calculate in this way the [S III] temperature.

$$t([\text{S II}]) = a_0(n) + a_1(n) \cdot R'_{S2} + \frac{a_2(n)}{R'_{S2}} + \frac{a_3(n)}{R'^2_{S2}}$$

where

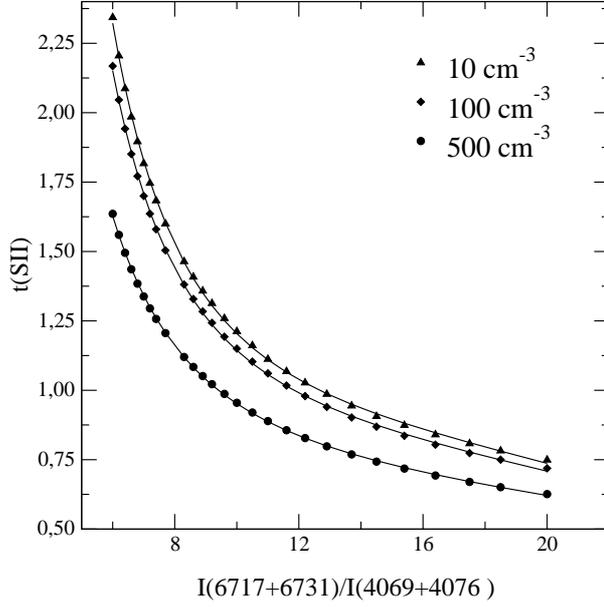
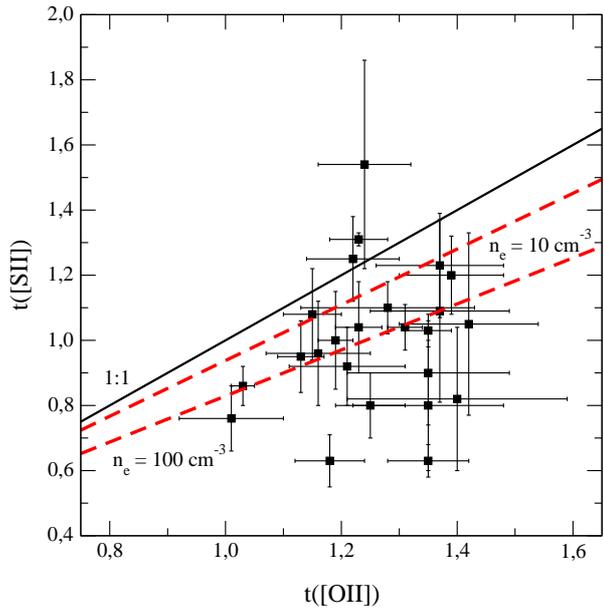
$$a_0(n) = 1.92 - 0.0017 \cdot n + \frac{0.848}{n}$$

$$a_1(n) = -0.0375 + 4.038 \cdot 10^{-5} \cdot n - \frac{0.0185}{n}$$

$$a_2(n) = -14.15 + 0.019 \cdot n - \frac{10.4}{n}$$

$$a_3(n) = 105.64 - 0.109 \cdot n + \frac{58.52}{n}$$

When the [S III] auroral lines are not available, it is usually assumed that  $t[\text{S III}] \approx t[\text{O II}]$ . There is evidence however, suggesting a somewhat lower value. From the models used

FIGURE 5. Relation between  $R_{S2}$  and  $t[\text{SII}]$ FIGURE 6. Relation between  $t[\text{O II}]$  and  $t[\text{SIII}]$  for model sequences and for a sample of HII galaxies.

for my PhD thesis, I obtained a lineal fitting:

$$t[\text{Siii}] = 0.71 \cdot t[\text{Oii}] + 0.12$$

for a  $100 \text{ cm}^{-3}$  number density. For lower densities, the first expression seems to be valid. In any case, for the few objects for which we have a simultaneous measurement of both temperatures, the dispersion is quite large as can be seen in Figure 6 for HII galaxies.

Direct measurements of the [S III] temperature have become possible with the availability of the collisional lines in the near IR.

$$R_{S3} = \frac{I(9069\text{\AA}) + I(9532\text{\AA})}{I(6312\text{\AA})}$$

expression that can be simplified in case of lacking one of the near IR lines, knowing that  $I(9532\text{\AA}) \approx 2.44 \cdot I(9069\text{\AA})$ . In that case, the [S III] temperature can be calculated by the fit:

$$t([SIII]) = \frac{R_{S3} + 36.4}{1.8 \cdot R_{S3} - 3.01}$$

There is also work relating the [O III] and the [S III] temperatures. As discussed in Garnett (1992),  $t[S III]$  is in between the temperatures of [O III] and of [O II] and allows to calculate the  $S^{2+}$  abundance from just the 6312 Å line. The importance of this relation has reversed recently, given that in very high metallicity objects one can more easily detect the [S III] auroral line than the [O III] one (e.g. S5 en M101; Kinkel & Rosa, 1994 or CDT1 in NGC1232; Castellanos et al., 2002). The relation given by Garnett is:

$$t([SIII]) = 0.83 \cdot t([OIII]) + 0.17$$

although this relation underestimates  $t[S III]$  given new atomic coefficients for sulphur (Tayal & Gupta, 1999). From my models, instead, I can derive:

$$t([SIII]) = 1.05 \cdot t([OIII]) - 0.08$$

plotted as a thin solid line in Figure 7. Since the relation between electron temperatures does not allow to obtain a realistic estimation of the associated uncertainties to the models and since there is growing number of observations in the spectral range where it is possible to obtain this temperatures, it is possible now to give an empirical fit:

$$t([SIII]) = (1.19 \pm 0.08) \cdot t([OIII]) - (0.32 \pm 0.10)$$

(Hägele et al. en prep.) that is plotted as thick solid line in Figure 7.

### 2.3. Nitrogen

The temperature of [NII] can be calculated from the quotient:

$$R_{N2} = \frac{I(6548\text{\AA}) + I(6584\text{\AA})}{I(5755\text{\AA})}$$

The nebular lines of [NII] are very close to  $H\alpha$  so they appear sometimes blended to this line and therefore it is not possible sometimes to measure them or, at least, to measure the weakest of them. In this case it is often used the theoretical relation between them, in such a way that  $I(6584) \approx 2.9 \cdot I(6548)$ . Besides, the auroral line of [NII] is affected by recombination emission, that can be corrected using the next expression proposed by Liu et al. (2000):

$$\frac{I_R(5755)}{I(H\beta)} = 3.19 \cdot t^{0.30} \cdot \frac{N^{2+}}{H^+}$$

in the range between 5000 and 20000 K.

In these conditions it is possible to calculate the temperature of [NII] by means of the expression:

$$t([NII]) = 0.537 + 0.000253R_{N2} + \frac{42.126}{R_{N2}}$$

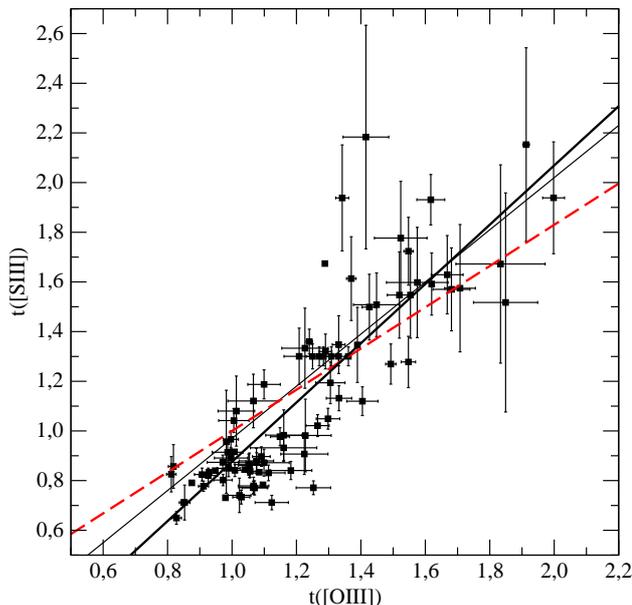


FIGURE 7. Relation between  $t[\text{O III}]$  and  $t[\text{S III}]$  for model sequences and for a sample of different objects.

that does not depend on density due to the  $np^2$  nature of  $\text{N}^+$ .

Unfortunately, the auroral line has very low signal-to-noise ratio, so it is usually considered the approximation  $t([\text{N II}]) \approx t([\text{O II}])$  as valid. This relation is confirmed by photoionization models but is quite sensitive to density and the inner ionization structure of the nebula, and it is possible to reach values closer to  $T([\text{S III}])$  in some cases, so in all case it can't be taken without any sort of uncertainty.

Other possibility is to calculate  $t([\text{N II}])$  directly from  $t([\text{O III}])$  using the expression derived using photoionization models by Pérez-Montero & Contini (2009):

$$t([\text{N II}]) = \frac{1.85}{t([\text{O III}])^{-1} + 0.72}$$

#### 2.4. Balmer temperature

The Balmer temperature depends on the value of the Balmer jump (BJ) in emission. In order to measure this value, it is necessary to fit the continuum in both sides of the discontinuity ( $\lambda_B = 3646 \text{ \AA}$ ). The contribution of the underlying population affects, between other factors, to the emission of hydrogen lines near from the BJ. The increment in the number of lines at shorter wavelengths produces blends between them that trend to reduce the level of the continuum at right of the discontinuity so it is necessary to take into the account all this in the final uncertainty. Once measured the BJ, the Balmer temperature ( $T(\text{Bac})$ ) is measured from the quotient of the flux of the jump and the emission of the line H11 by means of the expression:

$$T(\text{Bac}) = 368 \times (1 + 0.259y^+ + 3.409y^{++}) \left( \frac{\text{BJ}}{\text{H11}} \right)^{-3/2} K$$

where  $y^+$  and  $y^{++}$  are the ionic abundances of helium once and twice ionized, respectively, and BJ is in  $\text{ergs cm}^{-2} \text{ s}^{-1} \text{ \AA}^{-1}$ .

### 2.5. Other temperatures and densities

There are other line quotients for other ions, but they are not necessarily easily available. This is the case for [ArIV] for instance, for which we can deduce the density in the inner nebula, or the temperature of the rest of ions for which the corresponding auroral line is not available in the observed spectral range or there is not enough signal-to-noise ratio to detect the recombination lines, but whose nebular emission lines are detectable in the spectrum, allowing the determination of their ionic abundances. In this case, they are usually taken some assumptions about the gas structure and the position of each species is associated to another whose temperature is known. This is the case of the next ions:

$$t([NeIII]) \approx t([FeIII]) \approx t(HeII) \approx t([ArIV]) \approx t([OIII])$$

$$t([ArIII]) \approx t([SIII])$$

## 3. Ionic abundances and ICFs

In what follows I will describe the fittings derived for each ion once we know the available line fluxes and the appropriate electron densities and temperatures. The relations are listed in Table 1. Note however that the fittings have been obtained with the atomic coefficients as in *STSDAS* except those for  $O^+$ , and changes in them will produce changes in the equations. Most of the fittings follow equations as proposed by Pagel et al. (1992).

### 3.1. Helium

Helium lines, as equal as hydrogen ones in the visible spectrum have a recombination nature and they are generally strong and numerous, although many of them are usually blended with other lines. Besides they are affected by absorption of underlying stellar population, by fluorescence or by collisional contribution. They are generally used the lines of HeI  $\lambda\lambda$  4471, 5876, 6678 and 7065Å, and HeII  $\lambda$  4686 Å to estimate the helium abundances once and twice ionized respectively. In order to determinate the contributions to these lines, it exists a very deep study in Olive & Skillman (2001) who deduce the equations from the theoretical emissivities done by Smits (1996) In these conditions it is considered as valid to approximate that:

$$\frac{He}{H} = \frac{He^+ + He^{2+}}{H^+}$$

### 3.2. Oxygen

The chemical abundance of  $O^+$  can be determined from the intensity of the 3727 Å line, although recently Kniazev et al.(2003) suggest the use of the 7319,7330 Å doublet in the low redshift SLOAN spectra without 3727. For the  $O^{2+}$  abundance, both strong lines at 4959 and 5007 Å are used.

The total oxygen abundance can be approximated by

$$\frac{O}{H} = \frac{O^+ + O^{2+}}{H^+}$$

given that due to the charge exchange reaction the relative fractions of neutral oxygen and hydrogen are similar:

$$\frac{O^0}{O} = \frac{H^0}{H}$$

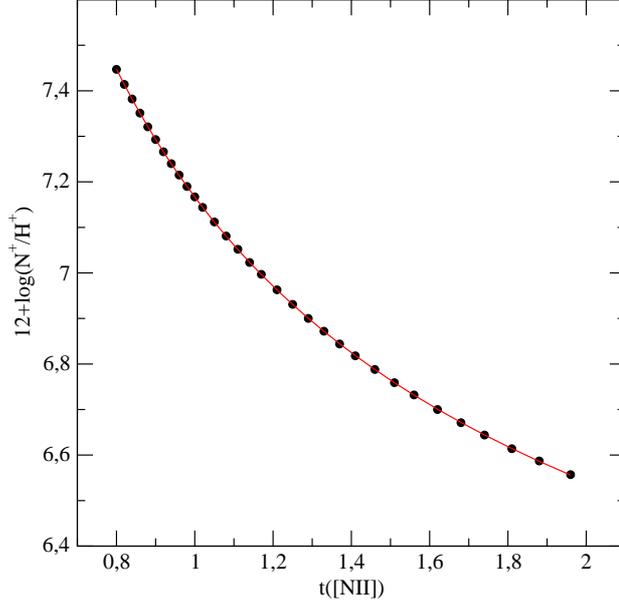
FIGURE 8.  $N^+$  abundance as a function of electron temperature, for a fixed line value.

TABLE 1. Equations for abundance calculations.

Ion†	Equation
$O^+$	$\log\left(\frac{I(3727)}{I(H\beta)}\right) + 5.992 + \frac{1.583}{t} - 0.681 \cdot \log t + \log(1 + 0.00023 \cdot n)$
	$\log\left(\frac{I(7325)}{I(H\beta)}\right) + 6.895 + \frac{2.44}{t} - 0.58 \cdot \log t - \log(1 + 0.0047 \cdot n)$
$O^{2+}$	$\log\left(\frac{I(4959)+I(5007)}{I(H\beta)}\right) + 6.144 + \frac{1.251}{t} - 0.55 \cdot \log t$
$S^+$	$\log\left(\frac{I(6717)+I(6731)}{I(H\beta)}\right) + 5.423 + \frac{0.929}{t} - 0.28 \cdot \log t + 0.0001 \cdot n$
$S^{2+}$	$\log\left(\frac{I(9069)+I(9532)}{I(H\beta)}\right) + 5.8 + \frac{0.771}{t} - 0.22 \cdot \log t$
	$\log\left(\frac{I(6312)}{I(H\beta)}\right) + 6.74 + \frac{1.672}{t} - 0.595 \cdot \log t$
$N^+$	$\log\left(\frac{I(6548)+I(6584)}{I(H\beta)}\right) + 6.273 + \frac{0.894}{t} - 0.592 \cdot \log t$
$Ne^{2+}$	$\log\left(\frac{I(3868)}{I(H\beta)}\right) + 6.486 + \frac{1.558}{t} - 0.504 \cdot \log t$
$Ar^{2+}$	$\log\left(\frac{I(7137)}{I(H\beta)}\right) + 6.157 + \frac{0.808}{t} - 0.508 \cdot \log t$
$Ar^{3+}$	$\log\left(\frac{I(4740)}{I(H\beta)}\right) + 4.705 + \frac{1.246}{t} - 0.156 \cdot \log t$

† As  $12+\log(X/H^+)$ 

### 3.3. Sulphur

The abundances are obtained from the 6717, 6731 Å lines for  $S^+$  and by the 9069, 9532 Å lines for  $S^{2+}$ , though for the latter, also the 6312 Å line can be used.

The ICF for sulphur takes into account the  $S^{3+}$  abundance which cannot be determined

in the optical range. A good approximation is given by Barker (1980) as:

$$ICF(S^+ + S^{2+}) = \left[ 1 - \left( \frac{O^{2+}}{O^+ + O^{2+}} \right)^\alpha \right]^{-1/\alpha}$$

Although it is customary to write Barker's expression as a function of the  $O^+/(O^+ + O^{2+})$  ionic fraction, we have reformulated it in terms of  $O^{2+}/(O^+ + O^{2+})$  since the errors associated to  $O^{2+}$  are considerably smaller than for  $O^+$ . and for a sample of objects with observed [SIV] line at  $10.5 \mu\text{m}$ , it was derived  $\alpha \approx 2.5$  (Pérez-Montero, Díaz, Vílchez & Kehrig, 2006).

### 3.4. Nitrogen

We can calculate  $N^+$  abundance from the 6548 and 6584 Å lines. Given their proximity to  $H\alpha$ , if one of them cannot be measured it doesn't matter as there is a theoretical relation between them as  $I(6584) \approx 2.9 \cdot I(6548)$ .

Starting from:

$$\frac{N^+}{N} = \frac{O^+}{O}$$

one can calculate quite precisely the total abundance of Nitrogen.

### 3.5. Neon and Argon

The [Ne III] line at 3868 Å is used for neon,

The ionisation correction factor for neon can be calculated according to the expression given by Pérez-Montero et al. (2007):

$$ICF(Ne^{2+}) = 0.142x + 0.753 + \frac{0.171}{x}$$

where  $x = O^{2+}/(O^+ + O^{2+})$ .

This expression deviates from the classical approximation  $Ne/O \approx Ne^{2+}/O^{2+}$  used to derive total neon abundances.

For argon, we use the [ArIII] 7137 Å line. It is possible to measure as well the lines of [ArIV] at 4713 and 4740 Å. Nevertheless, the first of them usually appears blenden with another line of HeII at 4711Å that is difficult to correct, so it is better to use the second and brighter to calculate the ionic abundance of  $Ar^{3+}$ .

As in the case of neon, the total abundance of argon has been calculated using the ionization correction factors ( $ICF(Ar^{2+})$  and the  $ICF(Ar^{2+} + Ar^{3+})$ ) given by Pérez-Montero et al. (2007). We use the first one only when we cannot derive a value for  $Ar^{3+}$ . The expressions for these ICFs are:

$$ICF(Ar^{2+}) = 0.507(1 - x) + 0.749 + \frac{0.064}{(1 - x)}$$

$$ICF(Ar^{2+} + Ar^{3+}) = 0.364(1 - x) + 0.928 + \frac{0.006}{(1 - x)}$$

where  $x = O^{2+}/(O^+ + O^{2+})$ .

## REFERENCES

- Barker, T. 1980, ApJ, 240, 99.  
 Castellanos, M., Díaz, A.I. & Terlevich, E. 2002, MNRAS, 329, 315.  
 De Robertis, M.M., Dufour, R.J. & Hunt, R.W. 1987, JRASC, 81, 195.

- Garnett, D.R. 1992, *AJ*, 103, 1330.
- Kingsburgh, R.L. & Barlow, M.J. 1994, *MNRAS*, 271, 257.
- Kinkel, U. & Rosa, M.R. 1994, *A&A*, 282, 37.
- Kniazev, A.Y., Grebel, E.K., Hao, L., Strauss, M.A., Brinkmann & Fukugita, M. 2003, *ApJ*, 593L, 73.
- Liu, X.-W., Storey, P.J., Barlow, M.J., Danzinger, L.J., Cohen, M. & Bryce, M. 2000, *MNRAS*, 312, 585.
- McLaughlin, B.M. & Bell, K.I. 1998, *J.Phys.B*, 31, 4317.
- Pagel, B.E.J., Simonson, E.A., Terlevich, R.J. & Edmunds, M.G. 1992, *MNRAS*, 255, 325.
- Pérez-Montero, E., Contini, T., Hägele, G.F., & Díaz, A.I., 2007, *MNRAS*, 381, 125
- Pérez-Montero, E. & Contini, T., 2009, *MNRAS*, 398, 949
- Pérez-Montero, E. PhD thesis. 2003. Universidad Autónoma de Madrid.
- Pérez-Montero, E. & Díaz, A.I. 2003, *MNRAS*, 346, 105. (PMD03)
- Pérez-Montero, E. & Díaz, A.I. 2005, *MNRAS*, 361, 1063 (PMD05)
- Pérez-Montero, E. Díaz, A.I., Vílchez, J.M. & Kehrig, C. 2006, *A&A*, 449, 163. bibitem[] Pradham, A.K. 1976, *MNRAS*, 177, 31.
- Shaw, R.A. & Dufour, R.J. 1995, *PASP*, 107, 896.
- Tayal, S.S. & Gupta, G.P. 1999, *ApJ*, 526, 544.
- Wang, W., Liu, X.-W., Zhang, Y. & Barlow, J. 2004, *A&A*, 427, 873.