Determination of the Primordial Helium Abundance Based on NGC 346, an H II Region of the Small Magellanic Cloud

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Abstract

To place meaningful constraints on Big Bang Nucleosynthesis models, the primordial helium abundance determination is crucial. Low-metallicity H II regions have been used to estimate it because their statistical uncertainties are relatively small. We present a new determination of the primordial helium abundance, based on long-slit spectra of the H II region NGC 346 in the small Magellanic cloud. We obtained spectra using three 4096 \times 512 slits divided into 97 subsets. They cover the range \(\lambda\lambda 3600–7400\) of the electromagnetic spectrum. We used PyNeb and standard reduction procedures to determine the physical conditions and chemical composition. We found that for NGC 346 \(X = 0.7465\), \(Y = 0.2505\), and \(Z = 0.0030\). By assuming \(\Delta Y/\Delta O = 3.3 \pm 0.7\) we found that the primordial helium abundance is \(Y_P = 0.2451 \pm 0.0026\) (1\(\sigma\)). Our \(Y_P\) value is in agreement with the value of neutrino families, \(N_{\nu}\), and with the neutron half-life time, \(\tau_n\), obtained in the laboratory.

Key words: galaxies: ISM – H II regions – ISM: abundances – Magellanic Clouds – primordial nucleosynthesis

1. Introduction

A highly accurate (\(\leq 1\%\)) primordial helium abundance (\(Y_P\)) determination plays an important role in understanding the Universe. In particular, it is extremely important to constrain Big Bang Nucleosynthesis models, elementary particle physics, and the study of galactic chemical evolution.

To obtain a value for \(Y_P\) using H II regions, it is necessary to use the relation between the helium mass fraction \(Y\) and the heavy element mass fraction \(Z\). The relation is then extrapolated to zero metallicity to estimate the primordial mass fraction of helium. The first determination of \(Y_P\) based on observations of H II regions was carried out by Peimbert & Torres-Peimbert (1974). Since then, many estimations have been done using this technique (e.g., Izotov et al. 2014; Aver et al. 2015). This method has been modified to use the O abundance instead of \(Z\), because measuring other chemical elements becomes impractical, although some groups seek alternatives to O, like using N or S (e.g., Pagel et al. 1992; Fernández et al. 2018).

In recent years, careful studies with high-quality determinations of \(Y_P\) have been done by Izotov et al. (2014), Aver et al. (2015), Peimbert et al. (2016), and Fernández et al. (2018). To diminish the uncertainties most determinations use many objects, but large samples only diminish the contribution of statistical errors, while systematic errors are not diminished; in fact, if one is not careful, systematic errors can be more significant for large samples, because low-quality objects/observations are often included (e.g., Peimbert et al. 2000, 2007; Ferland et al. 2010).

Low-metallicity H II regions have been permitted to estimate the primordial helium abundance, whose statistical uncertainties are very small (\(\leq 1\%\)) (e.g., Izotov et al. 2014); nevertheless, due to the numerous systematic uncertainties, obtaining better than 1\(\%\) precision for individual objects remains a challenge.

The Small Magellanic Cloud (SMC) hosts the region NGC 346, which is the best H II region to determine \(Y_P\). This is the brightest and largest H II region of the SMC. Its H\(\alpha\) intensity places it on the boundary between normal and giant extragalactic H II regions. In addition, its proximity, at a distance of 61 \(\pm 1\) Kpc (Hilditch et al. 2005), makes it possible to resolve individual stars and to avoid their contributions to the observed nebular spectra. We also note that NGC 346 hosts the largest sample of O-type stars throughout the SMC (Massey et al. 1989).

NGC 346 possesses some definite advantages over other H II regions. As it is a nearby object, the underlying absorption correction (due to the stellar absorption of the H and He lines) can be reduced significantly, at least by an order of magnitude; another advantage of its proximity is that it can be observed in different lines of sight. This feature has been used to rule out the presence of a significant helium ionization correction factor (ICF; e.g., Vilchez 1989; Mathis & Rosa 1991; Peimbert et al. 1992; Viegas et al. 2000; Zhang & Liu 2003). Another important characteristic is that the electronic temperature is smaller than that in the less metallic H II regions, which reduces the effect of collisional excitation of the ground level of H\(\alpha\); its metallicity (\(Z \sim 0.003\); \(\sim 20\%\) of solar) requires a small extrapolation to derive \(Y_P\). However, it has the disadvantage that the correction for its chemical evolution is larger than that in many other objects used for the determination of \(Y_P\) (Peimbert et al. 2000). Overall, there is no H II region that is simultaneously at least as close, as large, and with a metallicity as low as NGC 346.

In this paper, we present a very precise primordial helium abundance determination. For this we use a single H II region, NGC 346, which is probably the best object for this determination, for the reasons mentioned before (see the comment by Steigman in Ferland et al. 2010). Additionally, we include a detailed study of the sources of error involved in this determination, along with a justification for using a single H II region to determine \(Y_P\).

The paper is organized as follows. Section 2 presents the characteristics of the observations and the data reduction. The line intensity corrections due to extinction and the underlying absorption are presented in Section 3. Section 4 shows the
determination of physical conditions, using both recombination lines (RLs) and collisional excitation lines (CELs). Sections 5 and 6 show the determination of chemical abundances, the determination of ionic abundances, and of total abundances, respectively. Finally, the determinations of the primordial helium abundance and the conclusions are shown in Sections 7 and 8.

2. Spectroscopic Observations

We obtained long-slit spectra with the Focal Reducer Low Dispersion Spectrograph FORS1, the night of 2002 September 10 at the Very Large Telescope Facility, located at Melipal, Chile for the H II region NGC 346 ($\alpha = 00^h59^m05^s.0, \delta = -72^\circ10'38''$). We used three grism configurations: GRIS-600B +12, GRIS-600R +14 with the filter GG435, and GRIS-300V with the filter GG375.

Table 1 shows the resolution and wavelength coverage for the emission lines observed with each grism. We chose this resolution as a compromise: seeking high resolution without breaking the spectra into too many segments.

The slit width and length were 0.51 and 410, respectively. We have three different positions of the slit in the nebula; see Figure 1. From the first position A (slit position $\alpha = 00^h59^m06^s.0, \delta = -72^\circ10'29''/3$), we defined 32 extraction windows; from the second position B (slit position $\alpha = 00^h59^m06^s.0, \delta = -72^\circ10'37''/3$), we defined 31 extraction windows; and from the last position C (slit position $\alpha = 00^h59^m19^s.0, \delta = -72^\circ10'0''$), we defined 34 extraction windows. The limits for each window were selected to avoid regions with qualitatively different spectra (e.g., stars or clouds) or to break long regions into more homogeneous bits. To keep the same observed region within the slit, we used the LACD (linear atmospheric dispersion corrector) regardless of the air mass value.

We used IRAF to reduce the spectra, following the standard procedure: bias, dark, and flat correction, wavelength calibration, flux calibration, then spectra extraction. We use the low-resolution data to tie the calibration of the blue and red spectra by calibrating the brightest (blue and red) lines to the low-resolution spectra. For the flux calibration we used the standard stars LTT 2415, LTT 7389, LTT 7987, and EG 21 by Hamuy et al. (1992, 1994). We did an analysis of the full sample, for each spectrum, in the three slit positions A, B, and C.

We did a rough analysis of each of our 97 windows; in it we measured H$\beta$ (as a measure of the overall intensity), [O II] $\lambda 3727$ (as a measure of O$^+$), [O III] $\lambda 5007$ (as a measure of O$^{\text{III}}$), [O III] $\lambda 4363$ (as a measure of the quality of the temperature), and [S II] $\lambda 6717+31$ (as a measure of the quality of the density); see Figure 2. We also present the equivalent width (EW) of H$\beta$ in emission (to estimate the influence of the underlying absorption); see Figure 3. In a future paper, we will present the line intensities for each of the 97 windows.

![Figure 1. VLT images of NGC 346 in the SMC. This object is located at $\alpha = 00^h59^m05^s.0, \delta = -72^\circ10'38''$.](image1.png)

![Figure 2. Intensity of the five emission lines ($\lambda 3727, \lambda 4363, \lambda 4861, \lambda 5007,$ and $\lambda 6717+31$) in the 97 windows for slits A, B and C (up, middle, and down, respectively). On the x axis, the bar corresponds to the number of pixels that each window contains. The slits are centered on pixel zero.](image2.png)

We want to select the windows that cover the brightest parts while avoiding regions where there is too much stellar continuum (i.e., we remove the brightest stars to avoid the effects of underlying stellar absorption; see Figure 1).

To minimize the errors of the emission lines we define a new region, Region I, summing the windows with the best data. We
3. Line Intensities and Reddening Correction

Emission line flux measurements were made with the SPLOT routine of the IRAF package. These were measured by integrating all the flux in the line between two given limits and by subtracting the local continuum estimated by eye. Partially blended lines were deblended with two or three Gaussian profiles to measure the individual line fluxes and we forced the lines to have the same widths.

We derived the reddening by fitting the observed flux ratios of the brightest Balmer lines to the theoretical values normalized to the Hβ flux. We estimated the theoretical Balmer decrement using the data by Storey & Hummer (1995) assuming $T_e = 12500$ K, and $n_e = 100$ cm$^{-3}$. We used the extinction law of Seaton (1979). We assumed that the underlying absorption had the form $EW_{abs}(H\beta) = EW_{abs}(H\beta) \times g(H\beta)$, where $g(H\beta)$ was obtained from Table 2 of Peña-Guerrero et al. (2012). We produced a stellar template normalized to $EW_{abs}(H\beta)$ based in the low-metallicity instantaneous burst models from González Delgado et al. (1999), to correct for the underlying absorption EWs for the Balmer lines. The percent error for each line includes the uncertainties in the reddening correction and the uncertainties in underlying absorption.

Finally, we are left with regions that have $EQW > 150$, and that do not show Wolf–Rayet contamination in H II. Region I was composed of 20 windows (A10, A12, A13, A14, A21, A23, A24, B10, B11, B14, B22, B23, C3, C5, C6, C7, C9, C11, C14, and C15). In Figures 4, and 5, we show the blue and red spectra of Region I.

The underlying absorption, reddening correction, and a renormalization of Hβ were done simultaneously with the equation

$$I(\beta) \overline{I}(H\beta) = F(\beta) \overline{F}(H\beta) \times 10^{f(\beta)C(\beta)} \left(1 - {EW_{abs}(\beta) \over EW(\beta)}\right) \left(1 - {EW_{abs}(H\beta) \over EW(H\beta)}\right)^{-1} {100 \over \eta} ,$$

where $C(\beta)$ is the reddening coefficient, $F(\beta)$ is the absolute flux for each line, $EW(\beta)$ is the EW observed in $\lambda$, and $EW_{abs}(\beta) = EW_{abs}(H\beta) \times g(\beta)$ is the theoretical EW (Peña-Guerrero et al. 2012). Only the H and He lines were corrected for underlying absorption. This underlying absorption is due to the contribution of dust-scattered light; this contribution is expected to include in absorption the H and He lines present in the brightest stars.

Finally, we used a $\chi^2$ minimization routine to obtain the parameters $C(H\beta)$, $EW_{abs}(H\beta)$, and $\eta$; the fitted values for Region I were 0.159, 1.647, and 99.53, respectively. The values for our best spectrum are in good agreement with the ones derived from Peimbert et al. (2012). Note that Hβ is not normalized to 100 because with this value, the other lines of H, would be overestimated. In the same way that we want to average the helium lines, we want to do it with the hydrogen lines. We used the following Balmer lines Hα, Hγ, Hδ, Hε, H11, H12, and H13 to calculate the observations. This calculation implies that Hβ = 99.53. We decided to normalize the reddening-corrected fluxes using the seven brightest Balmer lines. The best fit to these seven lines yielded an Hβ different than 100. We took the unconventional decision of choosing this

**Figure 3.** EW(Hβ) for each window in three slit positions: sections A, B, and C (up, middle, and down, respectively). On the x-axis, the bar corresponds to the number of pixels that each window contains. The slits are centered on pixel zero.

**Figure 4.** Plot of a fraction of the observed wavelength blue range for the high-resolution spectrum of the chosen extractions.

**Figure 5.** Plot of a fraction of the observed wavelength red range for the high-resolution spectrum of the chosen extractions.
4. Physical Conditions

4.1. Temperatures and Densities from CELs

We calculated physical conditions through CELs, fitting the line ratios using PyNeb (Luridiana et al. 2015). Electron temperatures were determined using the [O III] λ4363/(λ4959 + λ5007), [O II] (λ7320 + 7330)/(λ3726 + λ3729), and [N II] λ5755/λ6584 auroral to nebular intensity ratios. Electronic densities were determined from the [O II] λ3726/λ3729, and [S II] λ6731/λ6716 ratios, which are strongly density-dependent.

In addition, we obtained the [Fe III] density from the computations by Keenan et al. (2001), using the f(λ4986)/I(λ4658) ratio. This ratio is strongly dependent on density, going from I(4986)/I(4658) ≈ 1.0 at n_e = 100 cm−3 to I(4986)/I(4658) ≈ 0.05 at n_e = 3000 cm−3. This ratio is a good density indicator, particularly for low densities, since it is more density-dependent than the [Cl III] and [Ar IV] determinations; moreover, the Fe++ density is much more representative of the whole object than the O++ and S+ densities, since Fe++/Fe ≈ 60% while O++/O ≈ 20% and S+2/S+ ≈ 15%. The temperatures and densities are presented in Table 3.

4.2. Temperatures and Densities from RLs

We used the program Helio14, which is an extension of the maximum likelihood method presented by Peimbert et al. (2000) to obtain values for the temperature and density from He I lines. Helio14 code determines the He+/He++ abundance, n_e(He I), T_e(He I), the optical depth of the line He I λ3889, and f(He I) through the information on the line intensities of He I, and the information of the [O II] (λ7320 + 7330)/(λ3726 + λ3729) and the [O III] λ4363/(λ4959 + λ5007) line ratios. For a full description of the method, see Peimbert et al. (2007).

Using 10 He I lines: λ3819, λ3889, λ4026, λ4388, λ4471, λ4922, λ5016, λ5876, λ6678, and λ7065, we obtained a temperature T_e(He I) = 11400 ± 550 K, and a density n_e(He I) < 10 cm−3.

4.3. Temperature Inhomogeneities

To derive the ionic abundance ratios, the average temperature, and the temperature standard deviation, we use the
Table 2
(Continued)

| \(\lambda\) | ID | \(f(\lambda)\) | \(F(\lambda)\) | \(R(\lambda)\) | \(\%\) Error |
|------------|----|----------------|----------------|----------------|----------------|
| 5755       | [N II] | −0.192 | 0.08 | 0.07 | 20 |
| 5876       | He I | −0.216 | 11.69 | 10.85 | 1.5 |
| 6312       | [S III] | −0.286 | 1.78 | 1.61 | 4 |
| 6548       | [N II] | −0.320 | 2.18 | 1.95 | 4 |
| 6563       | H I | −0.322 | 314.61 | 282.24 | 0.8 |
| 6583       | [N II] | −0.324 | 4.28 | 3.83 | 3 |
| 6678       | He I | −0.337 | 3.39 | 3.05 | 3 |
| 6716       | [S II] | −0.342 | 8.32 | 7.40 | 1.5 |
| 6731       | [S II] | −0.343 | 5.98 | 5.31 | 2 |
| 7065       | He I | −0.383 | 2.40 | 2.13 | 3 |
| 7136       | [Ar III] | −0.391 | 8.39 | 7.33 | 2 |
| 7281       | He I | −0.406 | 2.27 | 1.99 | 3 |
| 7320       | [O II] | −0.410 | 1.30 | 1.13 | 4 |
| 7330       | [O II] | −0.411 | 1.07 | 0.93 | 6 |

Notes.

* \(F(\lambda)\) in units of \(F(\text{H}\beta) = 100.0\); \(F(\text{H}\beta) = 1.57 \times 10^{-12}\) erg cm\(^{-2}\) s\(^{-1}\).

* \(R(\lambda)\) in units of \(R(\text{H}\beta) = 99.53; R(\text{H}\beta) = 2.26 \times 10^{-12}\) erg cm\(^{-2}\) s\(^{-1}\).

* Note that \(R(\text{H}\beta) = 99.53;\) see the text.

Table 3
Temperature and Density from CELs

| Temperature (K) | [O III] | [O II] | [N II] |
|-----------------|---------|--------|--------|
| NGC 346         | 12871 ± 98 | 12445 ± 464 | 10882 ± 767 |
| Density (cm\(^{-3}\)) | [O II] | [S II] | [Fe III] |
| NGC 346         | 23.7 ± 8.2 | 32 ± 21 | 101 ± 17 |

The formalism of \(\dot{r}^2\) (Peimbert 1967). This formalism takes into account the temperature structure inside the H II region. Each ionic species, \(X^{+i}\), has a unique value of average temperature and of thermal inhomogeneity given by

\[
T_0(X^{+i}) = \frac{\int T_e n_e n(X^{+i}) dV}{\int n_e n(X^{+i}) dV} \tag{2}
\]

and

\[
\dot{r}^2(X^{+i}) = \frac{\int [T_e - T_0(X^{+i})]^2 n_e n(X^{+i}) dV}{T_0(X^{+i})^2 \int n_e n(X^{+i}) dV}, \tag{3}
\]

where \(n_e\) and \(n(X^{+i})\) are the electron and ion densities, respectively, and \(V\) is the observed volume.

In general, RLs, emissivities are stronger when \(T_e\) is lower, while CELs, emissivities are stronger when \(T_e\) is higher; and also in general RL temperatures are lower than \(T_0\) while CEL temperatures are higher than \(T_0\). Overall, each temperature has a different \(\dot{r}^2\) dependence; for instance the [O III] \(4363/5007\) temperature, in the presence of thermal inhomogeneities, can be expressed as a function of \(T_0\) and \(\dot{r}^2\) as

\[
T_{4363/5007} = T_0 \left[ 1 + \frac{\dot{r}^2}{2} \left( \frac{91300}{T_0} - 3 \right) \right], \tag{4}
\]

equivalent equations can be derived for the [N II], [O II], and [S II] temperatures, as well as for He I and Balmer continuum temperatures.

Because temperatures can be presented as a function of \(\dot{r}^2\) and \(T_0\), when one measures two different temperatures in principle one can obtain \(T_0\) and \(\dot{r}^2\). In practice one prefers one temperature that weighs preferably the hot regions (e.g., \(T [O III]\)), and one that weighs preferentially the cold regions (e.g., \(T(\text{He} I)\) or \(T(\text{Bac})\)).

Combining results from CELs and values from He I and using the helium4 code we found that the maximum likelihood values are \(\dot{r}^2(\text{He}^+) = 0.033 ± 0.017\) and \(T_0 = 11900 ± 450\) K. These values can be used to determine abundances in the high-temperature regions, due to the similar ionization potentials of \(\text{He}^+\) and \(\text{O}^{++}\). Determinations for \(\dot{r}^2\) values from observations of H II regions range between 0.020 and 0.120 (Peimbert et al. 2012). The \(\dot{r}^2\) value depends on the specific characteristics of the thermal structure of each nebula (or even of each fraction of nebula).

5. Ionic Chemical Abundances

5.1. Heavy Element Ionic Abundances

We have followed the \(\dot{r}^2\) formalism for chemical abundances as presented by Peimbert & Costero (1969). The first step to do so is to derive the abundances using the so-called direct method and then to correct for the presence of thermal inhomogeneities.

We have used a two-zone direct method characterized by the zones where low- and high-ionization ions are present: \(T(\text{low})\) and \(T(\text{high})\). \(T(\text{low})\) is the mean of \(T_0[\text{N II}], T_0[\text{O II}]\) and \(T_0[\text{S II}],\) while \(T(\text{high})\) is \(T_0[\text{O III}];\) we used \(T(\text{low})\) for singly ionized heavy elements, and \(T(\text{high})\) for multiple ionized elements.

We used PyNeb to make the computations (Luridiana et al. 2015). In Table 4 we present two values for each ionic abundance. The first value is obtained directly from PyNeb (\(\dot{r}^2 = 0.00\)), and the second value is obtained using the formalism of \(\dot{r}^2\).

To obtain the ionic abundance using \(\dot{r}^2 = 0.00\), we used the equations by Peimbert et al. (2004):

\[
\left[ \frac{n_{\text{cel}}(X^{+i})}{n(\text{H}^+)} \right]_{\dot{r}^2=0.00} = T(\text{H}\beta)^{\alpha} T(\lambda_{\text{mm}})^{0.5} \quad \frac{n(\text{H}^+)}{T_{4363/5007}} \times \exp \left[ \frac{-\Delta E_n}{k T_{4363/5007}} + \frac{\Delta E_n}{k T(\lambda_{\text{mm}})} \right] \times \left[ \frac{n_{\text{cel}}(X^{+i})}{n(\text{H}^+)} \right]_{\dot{r}^2=0.00}, \tag{5}
\]

where \(\alpha = -0.89\) is the temperature dependence of H\(\beta\). \(\Delta E_n\) is the difference of energy between the ground and excited states.
levels of the CEL, and \( T(\lambda_{nm}) \) and \( T(H\beta) \) are line temperatures, as described by Peimbert (1967). The average emissivity of each line for the temperature distribution on the observed object can be written as

\[
T(\lambda_{nm}) = T_0 \left[ 1 + \frac{r^2}{2} \left( \frac{\Delta E_n/kT_0}{\Delta E_n/kT_0 - 1/2} \right)^2 - 3\Delta E_n/kT_0 + 3/4 \right],
\]

and

\[
T(H\beta) = T_0 \left[ 1 - \frac{r^2}{2} (1 - \alpha) \right].
\]

### 5.2. Helium Abundance

To determine the He abundance we also used the 10 lines mentioned in Section 4.2 as input for the Helio14 code. This code uses the effective recombination coefficients given by Storey & Hummer (1995) for \( H^+ \), and for the case of \( He^+ \) those given by Benjamin et al. (2002) and Porter et al. (2013). The collisional contribution was estimated from Sawey & Berrington (1993) and Kingdon & Ferland (1995), and we used the calculations made by Benjamin et al. (2002) for the optical depth effects in the triplets; see Table 5.

The correction of the underlying stellar absorption is important for \( H_1 \) lines, therefore we used the values from González Delgado et al. (1999) to correct lines with \( \lambda < 5000 \) Å, and for lines redder than 5876 Å, we used the values of Peimbert et al. (2005). The ionic abundance of \( He^+ \) was obtained from the \( \lambda 4686 \) line presented in Table 2, and the recombination coefficients given by Storey & Hummer (1995).

| Ion     | \( r^2 = 0.000 \) | \( r^2 = 0.033 \pm 0.017 \) |
|---------|------------------|-----------------------------|
| \( He^+ \) | 10.917 ± 0.004  | 10.915 ± 0.004             |
| \( He^{++} \) | 8.30 ± 0.04    | 8.30 ± 0.04                |

**Note.** In units of 12 + log \( n(X^+)/n(H) \).

For objects of low and medium degrees of ionization, the presence of neutral helium within the H II region is important and the ICF(\( He^{0} \)) > 1; for H II regions with a high degree of ionization, like NGC 346, we can consider the presence of neutral helium within the nebula to be negligible. However, one must be careful with objects with very high degrees of ionization, e.g., ICF(\( He^{0} \)) < 1.00 (even if the correction for helium is always ICF(\( He^{0} \)) \( \geq 0.99 \), when we are looking for a precise determination of \( Y_p \), one should try to avoid such objects).

To determine an accurate He/H ratio a precise value of the ICF(\( He^{0} \)) is needed. Previous computations by Peimbert et al. (2002, 2007), Relaño et al. (2002), and Luridiana et al. (2003) showed that the ICF(\( He^{0} \)) = 1.00 for NGC 346, because there is no transition zone from ionized to neutral He and H regions. We note that their results were calculated using tailor-made photoionization models with the code CLOUDY (Ferland et al. 2013).

For oxygen, to obtain the total abundance, we considered that

\[
\frac{N(O^{+3})}{N(O)} = \frac{N(He^{++})}{N(He^+)},
\]

and

\[
ICF(O^+ + O^{++}) = \frac{N(O^+) + N(O^{++}) + N(O^{+3})}{N(O^+) + N(O^{++})};
\]

therefore

\[
\frac{N(O)}{N(H)} = ICF(O^+ + O^{++}) \left[ \frac{N(O^+) + N(O^{++})}{N(H^+)} \right],
\]

where

\[
ICF(O^+ + O^{++}) = \left[ 1 - \frac{N(He^{++})}{N(He^+)} \right]^{-1}
\]

For nitrogen and neon we obtained the total abundances using the ICFs by Peimbert & Costero (1969):

\[
\frac{N(N)}{N(H)} = ICF(N^{++}) \left[ \frac{N(N^{++})}{N(H^+)} \right],
\]

where

\[
ICF(N^{++}) = \frac{N(O^+) + N(O^{++}) + N(O^{+3})}{N(O^+)},
\]

and

\[
\frac{N(Ne)}{N(H)} = ICF(Ne^{++}) \left[ \frac{N(Ne^{++})}{N(H^+)} \right],
\]

where

\[
ICF(Ne^{++}) = \frac{N(O^+) + N(O^{++}) + N(O^{+3})}{N(O^{++})}
\]

We observed the auroral lines of [S II] and [S III]. To obtain the total sulfur abundance we used the ICF proposed by Staśnińska (1978):

\[
\frac{N(S)}{N(H)} = ICF(S^+ + S^{++}) \left[ \frac{N(S^+) + N(S^{++})}{N(H^+)} \right],
\]

where

\[
ICF(S^+ + S^{++}) = \left[ 1 - \left( \frac{N(O^{++})}{N(O^+)} \right)^{1/3} \right]^{1/3}
\]

For the case of argon, we have measurements for lines of [Ar III] and [Ar IV]; we used the calculations by Pérez-Montero...
et al. (2007):

$$\frac{N(Ar)}{N(H)} = \text{ICF}(Ar^{++} + Ar^{+3}) \left[ \frac{N(Ar^{++}) + N(Ar^{+3})}{N(H^{+})} \right]. \quad (19)$$

where

$$\text{ICF}(Ar^{++} + Ar^{+3}) = 0.928 + 0.364(1 - x) + \frac{0.006}{1 - x}. \quad (20)$$

For chlorine, we have measurements for lines of $\text{Cl}III$, and to obtain the total abundance we used a reimplementation of the ICF proposed by Delgado-Inglada et al. (2014):

$$\frac{N(Cl)}{N(H)} = \text{ICF}(Cl^{++}) \left[ \frac{N(O^{++}) + N(O^{+})}{N(H^{+})} \right]. \quad (21)$$

where

$$\text{ICF}(Cl^{++}) = 2.914 \left[ 1 - \left( \frac{N(O^{++})}{N(O^{+}) + N(O^{++})} \right)^{0.21} \right]^{0.75}. \quad (22)$$

In Table 6 we present the results for $r^2 = 0.00$ as well as for $r^2 = 0.033 \pm 0.017$.

6.1. X, Y, and Z Values for NGC 346

To calculate the fraction of He by mass, we need to know the fraction of the heavy elements by mass. For low-metallicity objects, like NGC 346, the $O/Z$ ratio is expected to be $O/Z \approx 0.55$ (Peimbert et al. 2007); this ratio is slightly larger than that for high-metallicity objects where $O/Z \sim 0.40$, mainly due to the increase of the $C/O$ ratio (Carigi & Peimbert 2011). Using $O/Z = 0.55$, we can thus estimate the full value of $Z = 0.0030$ and from that $Y = 0.2490$.

One of the most subtle effects that we need to include in the determination of primordial helium or helium in very hot objects is the correction due to the collisional excitation of the Balmer lines. Because models in Peimbert et al. (2002) were later used in Peimbert et al. (2007, 2016), we know that for NGC 346 including this effect produces an increase of $0.0015 \pm 0.0005$, consequently $X = 0.7465$, $Y = 0.2505$, and $Z = 0.0030$. In this work, the contribution by collisions to the intensity of He and H lines is considerably smaller than that from objects with higher electron temperature like SBS 0335–052 or I Zw 18. To minimize this particular error it is important to study $O/H$ poor objects but not extremely poor ones. Therefore, if we wish to further minimize this error, we must observe colder objects, with heavy element abundances similar to those of NGC 346.

7. Primordial Helium Abundance

To obtain the $Y_P$ value, we need to compute the fraction of helium present in the interstellar medium produced by galactic chemical evolution. For this purpose it was assumed that

$$Y_P = Y - O \frac{\Delta Y}{\Delta O}, \quad (23)$$

where $Y$ and $O$ are the helium and oxygen abundances by mass. We adopted the determination $\Delta Y/\Delta O = 3.3 \pm 0.7$ by Peimbert et al. (2016). To obtain this value, they used the observations of brighter objects from Peimbert et al. (2007) and chemical evolution models for galaxies of low mass and metallicity by Carigi & Peimbert (2008) and Peimbert et al. (2010).

Most determinations are dominated by systematic errors rather than statistical errors. We can decrease the statistical errors by increasing the number of objects used to determine $Y_P$. Nevertheless, the systematic ones will not decrease.

In Table 7, we present the error budget of our determination. The error is derived from the same 13 sources considered in Peimbert et al. (2007). The most important source of error is due to the difficulty of correcting for the collisional excitation of the Balmer lines. In very hot objects, $T_P \geq 15,000$, this excitation produces a non-negligible increase in the Balmer line intensities relative to the case B recombination, and if ignored introduces a bias in the reddening correction deduced from the Balmer decrement, both effects affecting the calibration of all the lines in the spectra by up to a few percent. However, because this effect has not been studied extensively, the corrections that can be applied are uncertain at best, and thus such objects should be avoided. Another important source of error comes from the temperature structure, which arises because most $Y$ determinations are based on a single homogeneous temperature, $T(4363/5007)$. However, the $T(HeI)$ is systematically lower, showing the presence of temperature variations that should be included in the $Y_P$ determination. The error of $O(\Delta Y/\Delta O)$ correction implies the extrapolation to zero heavy element content. Based on chemical evolution models of galaxies of different types, it is found that $Y/O$ is practically constant for objects with $O < 4 \times 10^{-3}$. To estimate the error in the reddening correction, we used the statistical errors of the determinations in Carigi & Peimbert (2011) and the systematic errors of the determinations in Peimbert et al. (2007).
correction, we made comparisons among four classic extinction laws and a recent one. These laws are labeled S79 (Seaton 1979), W58 (Whitford 1958), CCM89 (Cardelli et al. 1989), and B07 (Blagrave et al. 2007). The systematic effect of ICF(He) has to be tested with tailor-made photoionization models for low-metallicity H II regions. A discussion of the remaining (smaller) sources errors is presented in Peimbert et al. (2007).

When compared with errors of the sample of 2007 we can see that some systematic errors go down because we are using the best-known object to determine $Y_P$. While a few of the statistical errors go up because we are using a single object with a notable increase in \( \Delta Y / \Delta O \). One of the advantages of NCG 346 is that the correction due to the collisional excitation of the Balmer lines is smaller 0.0015 \( \pm 0.0005 \) in comparison with 0.0144 \( \pm 0.0038 \) for SBS 0335–052, and the 0.0056 \( \pm 0.0015 \) for the sample by Peimbert et al. (2007, 2016). Overall, the total error goes down, and the $Y_P$ value amounts to 0.2451 \( \pm 0.0026 \).

We can also compare with previous determinations for $Y_P$, determined using only NCG 346: $Y_P = 0.2345 \pm 0.0026$ by Peimbert et al. (2000), $Y_P = 0.2384 \pm 0.0025$ by Peimbert et al. (2002), $Y_P = 0.2453 \pm 0.0033$ by Peimbert et al. (2007), and $Y_P = 0.2433 \pm 0.0034$ by Peimbert et al. (2016). The difference when comparing the values of 2000 and 2002 with those from 2007 and onward comes from a better model of NCG 346, newer atomic data, and a better understanding of the quantity and magnitude of the sources of error (particularly, the effects of the collisional excitation of the Balmer lines were ignored in the first two determinations).

To summarize, there are several advantages to observing a nearby H II region. The higher spatial resolution allows us to isolate the windows that are contaminated by the light of bright stars. This allows us to reduce the starlight emission and therefore obtain a greater signal-to-noise ratio in the emission lines. In this determination, the systematic errors due to collisional excitation of the H are smaller and some statistical errors (like the ICF(He\(^5\)) and the density structure) also become smaller.

It is useful to discuss the determination of $Y_P$ as a function time. Skillman et al. (2012) presents the plot with results from 1974 to 2012. Accordingly, in Figure 6 we present the $Y_P$ results from 2007 to the present. It is clear that the $Y_P$ values derived by different groups provide important constraints on the Big Bang theory. The main source of error is due to systematic uncertainties. In addition, the result by the Planck Collaboration (2016) corresponds to $Y_P = 0.24467 \pm 0.0002$.

8. Discussion and Conclusions

Low-metallicity H II regions are optimal objects to determine the primordial helium abundance. In this context, we studied the H II region NGC 346. By measuring its He I line intensities we determined a primordial helium abundance $Y_P = 0.2451 \pm 0.0026$.

Table 8 shows $Y_P$ measurements reported in the literature and in this work. Our $Y_P$ value is consistent with most of the previous estimates. We note that the Izotov et al. (2014) determination differs significantly from the other values. We consider that they have some systematic errors that they are ignoring. For example, part of this difference is due to the adopted temperature structure. Izotov et al. (2014) used the temperature from CLOUDY photoionization models (Ferland et al. 2013), which predict 0.000 < $T$ < 0.015, with a typical value of about 0.004 (Peimbert et al. 2017). Alternatively, the observed $T^2$ values for 28 H II regions are in the 0.19 to 0.120 range with an average value of 0.044 (Peimbert et al. 2012). A discussion of the possible reasons for the high $T^2$ values observed in H II regions is given by Peimbert et al. (2016).

Since the regions observed in all cases are not the same, we do not expect the values to be the same, e.g., there may be regions where the $T^2$ value may be higher due to stellar winds or other sources of energy input (e.g., Peimbert et al. 2017).

According to our error budget presented in Table 7, it is comforting that three of the other four independent $Y_P$ results are very similar to ours. The main points of our results are as follows. (i) The lower temperature of NGC 346 reduces considerably the collisional excitation of the hydrogen lines, as well as the uncertainty of its determination when compared to the other samples, all of which include a few very hot objects. (ii) The small error introduced by correcting for the underlying absorption of the He I lines in NGC 346 occurs because we were able to eliminate most of the contribution of the dust-scattered stellar light in our observations.

Peimbert et al. (2016) presents a $Y_P$ value determined only using NGC 346. That determination has its error divided into statistical and systematical components that amount to 0.0028, and 0.0019, respectively, while in this paper the statistical error has been diminished to 0.0018 and the systematical one remains at 0.0019. The systematical error was not expected to change, because we are using the same atomic data, the same reddening law, etc. On the other hand, the statistical error was indeed expected to diminish due to the greater quality of the data (the bigger telescope and the higher number of photons collected).

While this is a significant improvement over previous determinations, we would still like better determinations. To continue the quest to obtain a $Y_P$ value of higher precision, we need (a) helium atomic data of higher quality; (b) better
Table 9

| Yp source                     | Yp       | τn      |
|-------------------------------|----------|---------|
| Izotov et al. (2014)           | 0.2551 ± 0.0022 | 921 ± 11 |
| Aver et al. (2015)             | 0.2449 ± 0.0040 | 872 ± 19  |
| Peimbert et al. (2016)         | 0.2446 ± 0.0029 | 870 ± 14  |
| Fernández et al. (2018)        | 0.245 ± 0.007 | 872 ± 33  |
| This work                     | 0.2451 ± 0.0026 | 873 ± 13  |

Table 9 Values and the Neutron Mean Life, τn

Note. Assuming N_e = 3.046 (Mangano et al. 2005).

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