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Nuclear astrophysics: nucleosynthesis in the Universe

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Abstract: Nuclear astrophysics is a relatively young science; it is about half a century old. It is a multidisciplinary subject, since it combines nuclear physics with astrophysics and observations in astronomy. It also addresses fundamental issues in astrobiology through the formation of elements, in particular those required for a carbon-based life. In this paper, a rapid overview of nucleosynthesis is given, mainly from the point of view of nuclear physics. A short historical introduction is followed by the definition of the relevant nuclear parameters, such as nuclear reaction cross sections, astrophysical S-factors, the energy range defined by the Gamow peak and reaction rates. The different astrophysical scenarios that are the sites of nucleosynthesis, and different processes, cycles and chains that are responsible for the building of complex nuclei from the elementary hydrogen nuclei are then briefly described.

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Introduction

The role of nuclear reactions in our Universe is 2-fold: the production of energy and the formation of elements, a process that is called nucleosynthesis. In the following, we will give a short historical introduction of the main steps of this very fascinating journey. The idea of energy production in stars occurring through the nuclear fusion of H into ^{4}He was first raised by A.S. Eddington in 1920. Details of this fusion process were given by George Gamow in 1928, using the quantum mechanical concept of tunnelling. He proposed that due to the Coulomb repulsion between positively charged nuclei and due to the thermal energy that is much lower than the height of the Coulomb barrier, the fusion could only occur by tunnelling through the barrier.

Georges Lemaître, a Belgian priest and astrophysicist, proposed in 1931 the idea of a Big Bang (not the name, which was suggested later by Fred Hoyle), based on the evident expansion of the Universe: this expansion suggested, if projected backward, that everything began from a very small region in the past. After the Big Bang, the first generation of stars was made of hydrogen and helium only. Heavier elements, necessary for a carbon-based life, were produced by the nucleosynthesis in stars. Then the elements essential for life were made in supernova explosions of massive stars. This evolution proceeds on long timescales: it takes a few billion years for the interstellar medium in a galaxy to be seeded with astrochemical elements in the abundances needed for carbon-based life.

In 1939, Hans Bethe (Bethe 1939) established which could be the nuclear reactions responsible for the production of ^{4}He from H in the stars. He introduced the mechanism of the proton–proton (pp) chain and of the Carbon-Nitrogen-Oxygen (CNO) cycle, still valid nowadays. The next important contribution came from Fred Hoyle in 1946 (Hoyle 1946), who was the first to address the creation of heavier elements. He proposed that very hot nuclei could fuse into iron. A major contribution of Hoyle to astrophysics was his prediction, from astrobiological arguments, of a resonance in ^{12}C. This nuclear effect, which strongly enhances the ^{12}C production, is necessary to explain the observed ^{12}C abundance in the Universe. This resonance was then found experimentally, and is known as the 'Hoyle state'. It provides an excellent example of the interplay between astrophysics and nuclear physics. The scientific life of Fred Hoyle and his revolutionary theories are described in detail in Mitton (2011).

In 1948, Alpher, Bethe and Gamow (Alpher *et al.* 1948) proposed that all elements could be produced during the Big Bang, through successive neutron captures and photon emissions. The authors also predicted a relic background radiation, coming from the Big Bang to our days, with a temperature of a few Kelvin.

A very important paper published in 1957, by Burbidge, Burbidge, Fowler and Hoyle (also known as the B2FH paper) (Burbidge *et al.* 1957) has established the foundation of the theory of stellar nucleosynthesis. The authors proposed that all nuclei were produced in stars, they proposed the cycles and processes, and accounted for the relative abundances observed at that time. In the same year, Cameron (1957) also published a paper giving independently a description of the stellar nucleosynthesis. He introduced computers into time-dependent calculations of evolution of nuclear systems.

The observation in 1964 by Wilson & Penzias (1964) of the relic cosmic microwave background (CMB) radiation, corresponding to a blackbody spectrum with a temperature of 2.73 K, was of enormous importance, as it gave a strong support to the Big Bang theory for the model of our Universe. They received the Nobel Prize in 1978 for this discovery.

Recently, the Wilkinson Microwave Anisotropy Probe (WMAP) satellite of the NASA (Bennett *et al.* 2003) realized a precise mapping of tiny fluctuations of temperature density (anisotropy) on the microwave sky (Hinshaw *et al.* 2009; Komatsu *et al.* 2009). The observed anisotropies are closely related to specific cosmological parameters.

In this paper, a rapid overview of the origin of elements (nucleosynthesis) will be given, mainly from the point of view of nuclear physics. In particular, we discuss low-energy cross sections and the associated reaction rates. Hydrogen and helium burnings, which represent the most important part in the evolution of a star, are briefly described. We refer to recent review papers (José & Iliadis 2011; Coc *et al.* 2012) for more details.

Relevant nuclear information and parameters

Nuclear-binding energies

Let us consider a nucleus made of Z protons and N neutrons ($A=Z+N$), and with mass $M(Z,N)$. The binding energy $B(Z,N)$ of this nucleus is defined as the energy required to break it up into the A individual nucleons. Using Einstein relation of mass-energy equivalence, it is defined by

$$M(Z,N)c^2 = Nm_nc^2 + Zm_pc^2 - B(Z,N), \tag{1}$$

where m_n and m_p are the neutron and proton mass, respectively ($m_nc^2 = 939.57$ MeV and $m_pc^2 = 938.27$ MeV). The nuclear force makes the mass of the nucleus smaller than the sum of the individual nucleon masses. The difference is the binding energy that is positive for bound nuclei.

The binding energy per nucleon $B(Z,N)/A$ is displayed in Fig. 1. This graph illustrates many important properties of nuclei. Iron is the most tightly bound nucleus, as its binding energy is $B(Z,N)/A = 8.8$ MeV for the ^{56}Fe isotope. In the low-mass region, ^{4}He is strongly bound, and almost behaves as an elementary particle. It is often referred to as the α particle. This high binding energy also explains why the $\alpha+p$ and $\alpha+\alpha$ systems (^{5}Li and ^{8}Be) are unstable: they immediately breakup.

The behaviour of the nuclear binding energy with A in Fig. 1 shows that, for $A<56$, energy is released by increasing the mass or, in other words, by capturing a nucleon or an α particle. This is the origin of fusion reactions occurring in stars and in fusion reactors. In contrast, for masses $A>56$, nuclei increase their binding energy (or, equivalently, reduce their mass) by emitting particles. In this mass region, many nuclei are unstable by α emission. Spontaneous fission occurs in the uranium region ($A \approx 200$ and above).

In nature, there are 330 stable isotopes of 82 elements, from hydrogen (H, $Z=1$) to lead (Pb, $Z=82$), with the exceptions of Technetium (Tc, $Z=43$, $T_{1/2} \sim 10^6$ years for ^{98}Tc) and Promethium (Pm, $Z=61$, $T_{1/2} \sim 17$ years for ^{145}Pm) and 9 unstable isotopes with $Z>82$. There are no stable elements with $A=5$ and 8, due to the strong binding energy of the α particle ($A=4$).

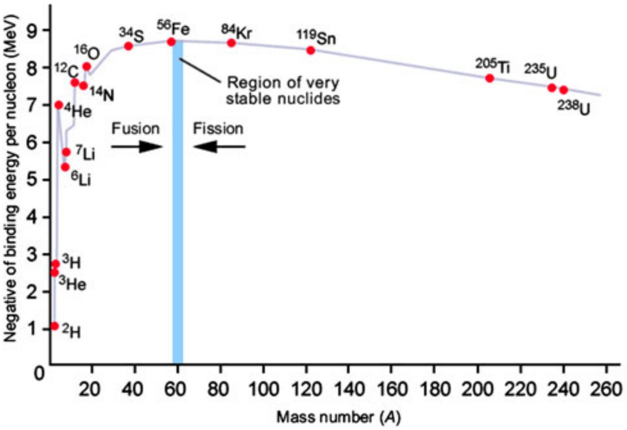

Fig. 1. Binding energy per nucleon $B(Z,N)/A$ as a function of the mass number (http://www.dlt.ncssm.edu/tiger).

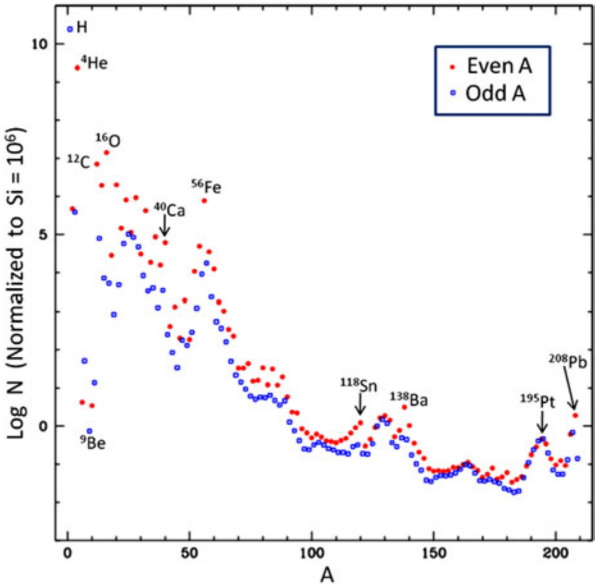

Fig. 2. Isotopic abundances of the solar system, as a function of the nuclear mass number A (reprinted from José & Iliadis 2011 with permission from IOP).

Most visible matter in Universe is made of atomic nuclei comprising protons and neutrons. Bound nuclei are those with mass smaller than the sum of masses of their free constituents. There are about 3000 known bound nuclei, which decay by weak interaction emitting electrons (e^-) or positrons (e^+) and neutrinos. This process is known as the β radioactivity. Heavy bound nuclei can decay by α emission and there are some examples of exotic proton decays.

In Fig. 2, we show the distribution of isotopic abundances in the solar system. The isotopic abundances vary on a scale of 10^{12}, H and He being the most abundant elements with 75% for H and 25% for ^{4}He. The most stable element, iron, corresponds to a peak in the abundance distribution. Other signatures of

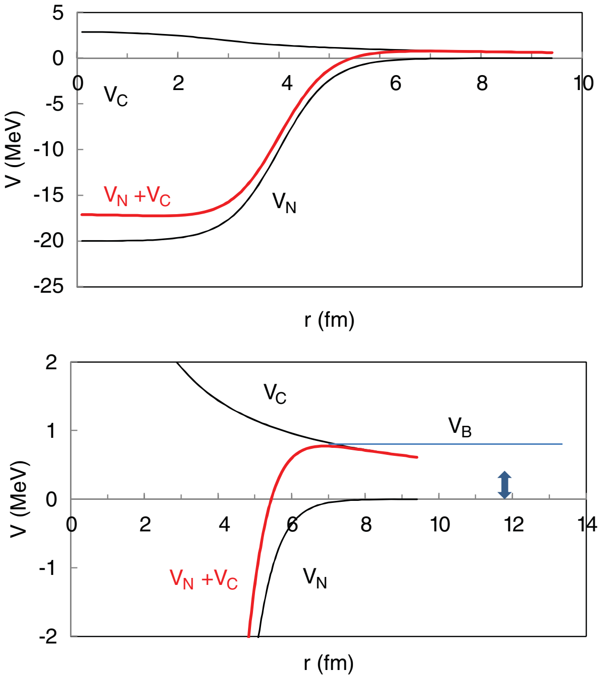

Fig. 3. Typical nuclear (V_N) and Coulomb (V_C) potentials. In the lower panel, V_B is the Coulomb barrier and the arrow shows the energy range of astrophysical interest.

nuclear physics properties are present in this diagram. Peaks above $A = 100$ are associated with magic numbers, corresponding to nuclei with high excitation and breakup energies. These nuclei are therefore more abundant in stars. The same comment can be given when comparing even and odd nuclei. It is well known in nuclear physics that odd nuclei are less bound than even nuclei, and are therefore more fragile.

The experimental data on chemical composition are collected not only from the composition of the Earth and of the Moon but also from meteorites and from absorption lines of the photosphere of the Sun. It is verified that the bulk isotopic composition of the solar system is homogeneous. The chemical composition of distant stars is obtained from the electromagnetic spectra emitted by their atoms and molecules.

Interaction potential between nuclei

The potential $V(r)$ between two nuclei depends on their relative coordinate r, and involves nuclear $V_N(r)$ and Coulomb $V_C(r)$ contributions. Both are presented schematically in Fig. 3, where we use standard units of nuclear physics (lengths are expressed in fm $= 10^{-15}$ m, and energies in MeV $= 10^6$ eV). The nuclear potential is not known precisely, but is attractive at short distances, and tends rapidly to zero when r increases. In contrast $V_C(r)$ is repulsive and extends to large distances.

A typical energy in nucleus–nucleus collisions is the Coulomb barrier. Its height V_B represents the maximum of the potential. At relative energies E lower than V_B, any reaction would be classically forbidden. However, quantum mechanics allows tunnelling effects, and the fusion of two nuclei is possible even for $E < V_B$. This probability, also called penetration factor $P(E)$, presents a fast energy dependence as

$$P(E) \sim \exp(-2\pi\eta), \tag{2}$$

where η is the Sommerfeld parameter, which provides a ‘measurement’ of Coulomb effects. It is defined by

$$\eta = Z_1 Z_2 e^2/\hbar v, \tag{3}$$

v being the relative velocity and Z_1, Z_2 the charges of the colliding nuclei. At low energies, reaction cross sections between charged particles follow this energy dependence, and therefore drop to very low values when $E \ll V_B$. The situation is different for neutron induced reactions, where $\eta = 0$. However, the short lifetime of the neutron (~ 15 min) limits their role in astrophysics to specific processes.

The nucleus–nucleus potential $V(r)$ also contains a centrifugal term $\hbar^2 L(L+1)/2\mu r^2$, depending on the angular momentum L. This additional repulsive contribution increases the height of the Coulomb barrier, and therefore still reduces the penetration factor. At low energies ($E \ll V_B$), $L = 0$ is dominant.

Astrophysical S-factor

The main characteristic of a reaction is the cross section $\sigma(E)$, which has the dimension of a surface (1 barn $= 10^{-28}$ m^2), and depends on energy. As mentioned before, the cross section at stellar energies is governed by Coulomb effects. The astrophysical S-factor is defined as

$$S(E) = \sigma(E) E \exp(2\pi\eta). \tag{4}$$

It removes the fast Coulomb dependence and is mainly sensitive to the nuclear effects.

An example is shown in Fig. 4 with the ^{3}He(α,γ)^{7}Be reaction. The cross section has a strong variation with energy (it follows the $\exp(-2\pi\eta)$ dependence) and it drops quickly to very low values ($\sim 10^{-10}$ barns), making impossible the extrapolation to lower energies. In contrast, the S-factor puts the nuclear effects in evidence and presents a weak energy dependence, which turns simple the extrapolation to low energies. The relevant astrophysical energy is around 23 keV for this reaction.

Reaction rates and stellar energies

The important quantity in stellar models is the reaction rate $\langle\sigma v\rangle$, which represents the averaged value of the cross section times the relative velocity. The study of stellar evolution requires a large number of reaction rates for different reactions (Clayton 1983; Rolfs & Rodney 1988; Iliadis 2007). The star is considered as a gas in equilibrium, where the energy distribution is given by the Maxwell–Boltzmann function $N(E, T)$. The reaction rate therefore depends on the temperature T of the star.

According to the energy dependence (2) of the non-resonant cross section, the reaction rate is obtained by integrating over energy

$$\sigma(E) v N(E) \sim \exp(-2\pi\eta) \times \exp(-E/k_B T), \tag{5}$$

where k_B is the Boltzmann constant. The former term stems from the penetration factor, and the latter from the

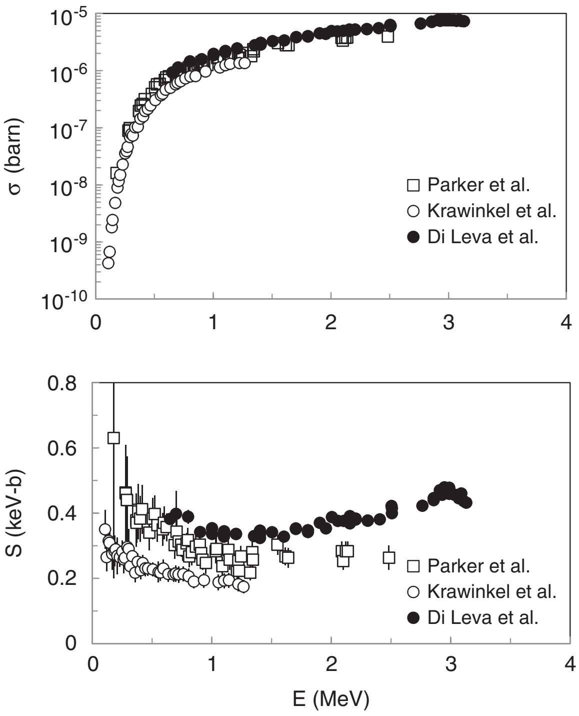

Fig. 4. $^{3}He(\alpha,\gamma)^{7}Be$ cross section (lower panel) and S-factor (upper panel), as a function of the energy.

Maxwell–Boltzmann distribution. Both terms present very different energy dependences and this product can be approximated by

$$\exp(-2\pi\eta)\times\exp(-E/k_{B}T)\sim\exp(-((E-E_{G})/2\Delta)^{2}), \qquad (6)$$

where E_G and Δ define the Gamow peak. These quantities can be easily determined from the properties of the system (masses and charges) and from the temperature (see Clayton 1983 for details). The Gamow peak defines the energy range where the cross section must be known to derive the reaction rate. In practice E_G is much lower than the Coulomb barrier, and typical stellar energies are very low at the nuclear scale. For example the Gamow energy of the important reaction $^{12}C(\alpha,\gamma)^{16}O$ is $E_G = 300$ keV (at the typical Helium burning temperature $T = 2\times10^{8}$ K), whereas the Coulomb barrier is approximately $V_B \sim 3$ MeV.

A general problem in nuclear astrophysics is that the cross sections at the Gamow energy are too small to be measured in the laboratory. Recently underground laboratories have been developed (as for example LUNA in Italy; Broggini *et al.* 2010) to reduce background effects coming from cosmic rays. This technique allows very low counting rates, but of course raises technological and practical problems. Until now the $^{3}He + {}^{3}He \rightarrow \alpha + 2p$ (important in the Sun) is the only reaction where the cross section has been measured in the Gamow peak. For many reactions, theoretical models are necessary to extrapolate the available data or to predict the cross sections.

Let us also mention that many reactions present resonances, in particular for heavy elements. In that case the reaction rate contains a resonant contribution, in addition to the non-resonant term discussed above. The resonant term depends on the properties of the resonances, such as their spin, energy and width. We refer the reader to Clayton (1983) for more details.

Types of nuclear reactions relevant in the nucleosynthesis

Essentially two types of nuclear reactions are important from the astrophysical point of view: radiative capture and transfer reactions. The former are determined by the electromagnetic interaction, whereas the latter arise from the nuclear force.

Radiative capture

In the radiative-capture process, two nuclei, A and B, fuse to the excited final nucleus C, with emission of γ-rays. This reaction is denoted as $A + B \rightarrow \gamma + C$, or $A(B,\gamma)C$. The (p,γ) reactions of the capture of protons, or the alpha capture (α,γ) are the most important, and are among the most common reactions occurring in stellar environments. The neutron capture (n,γ) is important in the formation of heavy nuclei (s- and r-processes).

The capture cross sections are usually small since they occur through the electromagnetic interaction, which is weaker than the nuclear force. Due to the small cross section they can put limit to the nuclear processing. The final nucleus C can be excited in several final states, necessary to be detected in order to obtain the total reaction cross section.

At energies where the cross section is not too small, capture cross sections can be studied in laboratory through a variety of methods: the detection of the recoiling C nucleus, the direct detection of the γ-rays and delayed decay measurements. Typical examples are $^{2}H(p,\gamma)^{3}He$ which is the first capture reaction in hydrogen burning, $^{7}Be(p,\gamma)^{8}B$ which determines the solar-neutrino spectrum at high energies, or $^{12}C(\alpha,\gamma)^{16}O$ which influences the $^{12}C/^{16}O$ ratio in stars. Many capture reactions, essentially involving protons or α particles, play a key role in stellar evolution.

Transfer reactions

In transfer reactions, nucleons are exchanged between the target and the projectile. They are denoted as $A + B \rightarrow C + D$ or $A(B,D)C$, and also play an important role in nuclear astrophysics. Typical examples are $^{3}H(d,n)^{4}He$, $^{6}Li(p,\alpha)^{3}He$ or $^{13}C(\alpha,n)^{16}O$. In this notation, A is the target, B the projectile and D the detected particle.

The process occurs through the nuclear interaction and the cross sections σ_t are much larger than those of radiative-capture reactions σ_C for the same entrance channel. This can be seen in Fig. 5, where the S-factors of the $^{6}Li(p,\alpha)^{3}He$ and $^{6}Li(p,\gamma)^{7}Be$ reactions are compared as a function of energy. They have the same entrance channel $^{6}Li + p$, and therefore present similar energy dependences. However, there is a factor of 10^{4} between the absolute values, in favour of the transfer process. This factor is typical of the ratio between the strengths of the nuclear (strong) and electromagnetic interactions. As a general statement, if the transfer channel is open, the radiative capture can be neglected since $\sigma_C \ll \sigma_t$. Another consequence is

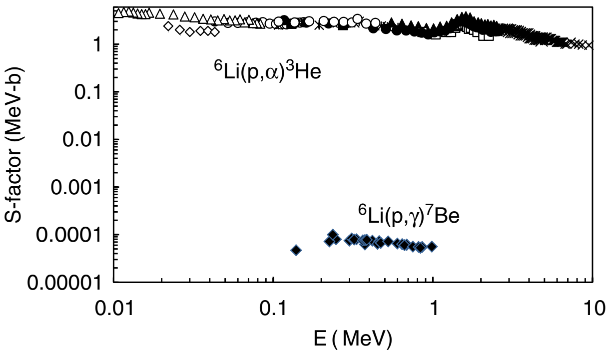

Fig. 5. Comparison of the S-factors for the ^{6}Li(p, α)^{3}He and ^{6}Li(p, γ)^{7}Be reactions.

that transfer cross sections can be measured at lower energies than capture cross sections.

Weak capture reactions

These reactions occur through the weak interaction, and the corresponding cross sections are far below the experimental possibilities. The p(p, e$^+$ν)d is, however, very important, since it is the first reaction of the hydrogen burning pp chain. It presents a tiny cross section and there are no experimental measurements. Stellar codes make use of theoretical calculations, which are expected to be very precise since the pp interaction at low energies is well understood. Owing to the very small cross section of this reaction, hydrogen burning is a slow process, our Sun has a long life time, and life could develop on Earth.

The weak-capture reaction ^{3}He(p, e$^+$ν)^{4}He plays a role in the neutrino spectrum, since it produces high-energy neutrinos that can be detected in terrestrial experiments. However, the cross section is made even smaller due to nuclear physics properties of this reaction. Only theoretical estimates are available.

Primordial nucleosynthesis

Timescale of events

The Big Bang occurred 13.7 billion years ago (13.7 Gyr). It was an extremely dense and hot initial state, followed by rapid expansion and cooling. In the first stage relativistic particle–antiparticle pairs were created and destroyed in collisions continuously, constituting what we call today quark–gluon plasma.

At $t = 10^{-6}$ s, quarks and gluons combine to form protons and neutrons. At the existing very high temperature particles (electrons, neutrinos, protons and neutrons) and photons were in equilibrium. At $t = 10$ s and $T \sim 3$ GK the decay of free neutrons into protons with half-life $T_{1/2} = 614$ s became the dominant weak interaction (José & Iliadis 2011), decreasing the neutron-to-proton number ratio n_n/n_p. At this temperature both the processes of creation ($p + n \rightarrow d + \gamma$) and destruction ($d + \gamma \rightarrow p + n$) of the deuteron were possible.

However, due to cooling, at $t = 250$ s and $T = 0.9$ GK, even the high energy tail of the Planck distribution has a photon energy that dropped below 2.2 MeV, the destruction of deuteron by photons could not occur anymore and the primordial nucleosynthesis, through successive nuclear reactions, started. At that time, the decay of free neutrons gave rise to a neutron-to-proton number ratio $n_n/n_p \sim 1/7$.

Due to cooling, at $t = 20$ min the Coulomb barrier between the existing nuclei suppressed all nuclear reactions and the Big Bang nucleosynthesis stopped. Afterwards, until the formation of the first stars, no new nuclides were created or destroyed (except ^{3}H and ^{7}Be that are unstable and decay by β emission).

At $t \sim 380\,000$ years and $T - 3000$ K, nuclei recombined with electrons to form neutral atoms. Consequently, the mean-free path of photons increased, and they could travel without interacting with matter. At that time, photons were decoupled from baryons and the Universe became transparent to electromagnetic radiation.

Today, after $t - 13.7$ Gyr, due to cooling and expansion, photons have continuously lost energy; photons have a black body spectrum of 2.725 K, isotropic with fluctuations of 10^{-5}. The WMAP has measured the temperature distribution of this microwave spectrum in all directions of the sky. The WMAP observation and the reliability of cosmological parameters deduced from these measurements is a key piece of evidence for the Big Bang cosmological model.

^{4}He production

After $t = 250$ s, the subsequent nuclear reactions were relatively fast and nearly all neutrons were incorporated into the tightly bound element ^{4}He, with a small fraction of other nuclides. A simple counting argument allows estimating the primordial ^{4}He abundance: for a neutron-to-proton ratio $n_n/n_p \sim 1/7$, in each eight nucleons, one proton and one neutron will end up bound in ^{4}He. Consequently the primordial ^{4}He mass fraction becomes $X_\alpha^{\text{pred}} \sim 2/8 = 0.25$.

Observations of ^{4}He (in clouds of ionized HII in dwarf galaxies), extrapolated to zero-metallicity, give the value $X_\alpha^{\text{obs}} = 0.248 \pm 0.003$ (Peimbert *et al.* 2007). Very recently, Izotov & Thuan (2010) presented a new determination of the primordial mass fraction of helium, based on 93 spectra of 86 low-metallicity extragalactic HII regions. The new value is $0.2565 \pm 0.0010(\text{stat}) \pm 0.0050(\text{syst})$. This value is higher, at the 2σ level, than calculations using the Standard Big Bang Nucleosynthesis (SBBN). The authors claim that this disagreement demands for modifications in the SBBN model.

From this reasoning, it becomes clear that the ^{4}He abundance depends on the free neutron half-life, on the weak interaction cross sections, on the neutron–proton mass difference and on the expansion rate, but it is insensitive to the baryon density and to the nuclear reaction cross sections. The destruction of deuterium occurred by the d(d, n)^{3}He, d(d, p)t and by the t(d, n)^{4}He reactions. Tritium was mainly produced by the d(d, p)t and ^{3}He(n, p)t reactions, and destroyed by t(d, n)^{4}He. The ^{4}He nuclide was mainly produced by the t(d, n)^{4}He and destroyed by the ^{4}He(t, γ)^{7}Li reactions. The ^{7}Be nuclide was mainly produced by ^{3}He(α, γ)^{7}Be, which decays to ^{7}Li, with a half-life of $T_{1/2} = 53.3$ days. The production of ^{7}Li was mainly via ^{7}Be, but also through ^{7}Be(n,p)^{7}Li

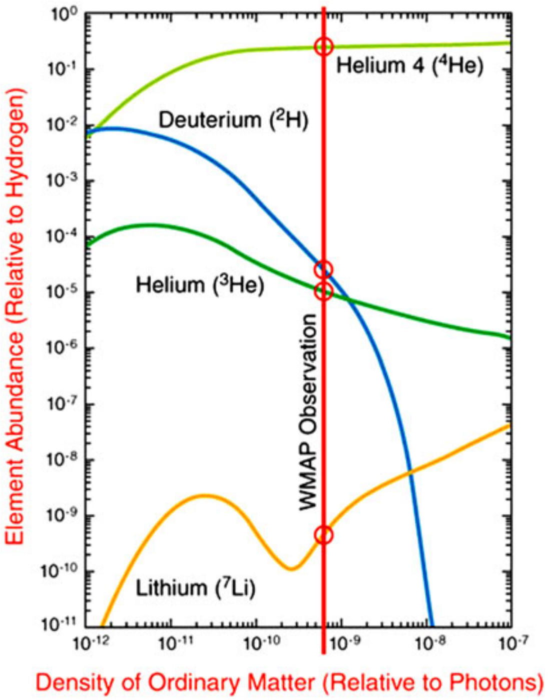

Fig. 6. Primordial abundances of light elements as a function of the baryon density η. Ref.: NASA/WMAP Science Team.

and the destruction through the ^{7}Li(p, α)^{4}He reaction. Direct cross section measurements exist for most of the above reactions at the relevant energies (Descouvemont *et al.* 2004).

In Fig. 6, the primordial abundances of elements produced shortly after the Big Bang (250 s to 20 min) are shown, as a function of the baryon density η, also called the baryon-to-photon ratio. The above-mentioned insensitivity of the ^{4}He abundance can be seen in the figure, as well as the fairly strong dependence of the other nuclides, as Deuterium, ^{3}He and ^{7}Li on η. The vertical line corresponds to the WMAP baryon density.

Primordial abundances are calculated using reaction network codes (Wagoner *et al.* 1967) and using all existing information on production and destruction reactions. The only free parameter for the primordial nucleosynthesis in the standard cosmological model is the baryon density η or baryon-to-photon ratio. The rates of nuclear reactions depend on this parameter and the final abundances also. If the WMAP value of $\eta = (6.2 \pm 0.2) \times 10^{-10}$ is adopted, the SBBN becomes a parameter-free model.

The primordial abundance of ^{4}He using the WMAP baryon density gives $X_{\alpha}^{\text{pred}} = 0.2486 \pm 0.0002$ (Cyburt *et al.* 2008), in very good agreement with older observational values given above, but at a 2σ level disagreement with more recent measurements.

Deuterium and lithium production

There is no alternative to the Big Bang for synthesizing deuterium: processes in stars destroy it rather than produce it. The predicted number abundance ratio of deuterium to hydrogen, using the WMAP η value is $(D/H)^{\text{pred}} = (2.5 \pm 0.2) \times 10^{-5}$ (Cyburt *et al.* 2008). Observation of absorption lines (Lyman-α) in low-metallicity gas clouds yields $(D/H)^{\text{obs}} = (2.8 \pm 0.2) \times 10^{-5}$ (Pettini *et al.* 2008). As the deuterium abundance depends strongly on η, this agreement is a key piece of evidence in favour of SBBN.

There is a large disagreement between the predicted and observed values and the primordial Lithium abundance is the central unresolved problem of the SBBN model. The predicted lithium to hydrogen abundance ratio is $(^7\text{Li/H})^{\text{pred}} = (5.2 \pm 0.7) \times 10^{-10}$ (Cyburt *et al.* 2008). The measurement of the ^{7}Li primordial abundance is a challenge, and carries a large systematic error; the most recent value is $(^7\text{Li/H})^{\text{obs}} = (1.1 - 2.3) \times 10^{-10}$ (Ryan *et al.* 2000; Asplund *et al.* 2006; Bonifacio *et al.* 2007).

Beryllium and boron production

The lithium, beryllium and boron elements have very low abundances as we can see in Fig. 2. Due to the non-existence of stable $A = 5$ and 8 isotopes, the Big Bang nucleosynthesis almost stopped at $A = 7$. In the nuclear burning in stars, they are by-passed by the triple α capture that goes directly from ^{4}He to ^{12}C. These fragile and rare nuclides are mostly destroyed in stellar interiors. Their production in an alternative model, called inhomogeneous BBN (IBBN) was proposed (Malaney & Fowler 1988; Kajino *et al.* 1990; Lara *et al.* 2006).

Stellar nucleosynthesis

Introduction

In contrast with the primordial nucleosynthesis, which occurred during a short period after the Big Bang, stellar nucleosynthesis determines the long-term, slow evolution of stars. After the Big Bang, the Universe is essentially made of protons (∼75%) and of α particles (∼25%). The birth of a star starts by the gravitational contraction of an hydrogen gas, and the first reaction in hydrogen burning is p(p, e$^+$ν)d. As mentioned before, this reaction occurs through the weak interaction and measurements are impossible owing to the smallness of the cross section. The p + p interaction is fairly well known, and accurate calculations of the cross section are available in the literature.

Hydrogen burning

Hydrogen is the most abundant element in the Universe. In the Hydrogen burning, four protons fuse through different reactions, with a ^{4}He nucleus as final product. The total energy release is 26.731 MeV. This process can occur in two different ways: the pp chains and the CNO cycle. Both are briefly discussed below.

The pp chains correspond to different paths to synthesize helium from hydrogen. These paths are given in Table 1. The first two reactions (p(p, e$^+$ν)d and d(p, γ)^{3}He) are common to the three chains. The pp1 chain is the most important since it involves the ^{3}He(^{3}He, 2p)α transfer reaction which has a high cross section. This cross section has been measured down to

Table 1. *Reactions (or α and β decays) involved in the pp chains*

pp1 chain	pp2 chain	pp3 chain
$p(p,e^+\nu)d$	$p(p,e^+\nu)d$	$p(p,e^+\nu)d$
$d(p,\gamma)^3He$	$d(p,\gamma)^3He$	$d(p,\gamma)^3He$
$^3He(^3He,2p)\alpha$	$^3He(\alpha,\gamma)^7Be$	$^3He(\alpha,\gamma)^7Be$
	$^7Be(e^-,\nu)^7Li$	$^7Be(p,\gamma)^8B$
	$^7Li(p,\alpha)\alpha$	$^8B(\beta^+\nu)^8Be$
		$^8Be(\alpha)\alpha$

Table 2. *Reactions (or β decays) involved in the CNO cycle*

CNO cycle
$^{12}C(p,\gamma)^{13}N$
$^{13}N(\beta^+\nu)^{13}C$
$^{13}C(p,\gamma)^{14}N$
$^{14}N(p,\gamma)^{15}O$
$^{15}O(\beta^+\nu)^{15}N$
$^{15}N(p,\alpha)^{12}C$

15 keV, which correspond to the central temperature of the Sun (15×10^6 K).

The pp2 and pp3 chains involve capture and β decays. The pp3 chain plays a minor role in the nucleosynthesis, but is important for the solar neutrino problem. The $^7Be(p,\gamma)^8B$ reaction produces high-energy neutrinos in the subsequent β decay of ^8B. This reaction has been recently studied by different groups.

If the star contains a small fraction of ^{12}C, this nucleus can act as a catalyst. This process is known as the CNO cycle, and also converts four protons into an α particle. The main CNO cycle is given in Table 2. All reactions involved in the CNO cycle are fairly well known. The last reaction $^{15}N(p,\alpha)^{12}C$ could in principle compete with the $^{15}N(p,\gamma)^{16}O$ reaction. However, the (p,α) cross section is much larger than the (p,γ) cross section (see Fig. 5), and the initial ^{12}C nucleus is reformed at the end of the cycle. As for the pp chains, the end result of each process is the transformation of four protons in an α particle, i.e. $4p\rightarrow{}^4He+2e^++2\nu$.

Other variants of the CNO cycle exist, but play a role in explosive burning only. For example, if the temperature is high enough, the $^{13}N(p,\gamma)^{14}O$ reaction is faster than the ^{13}N β decay. This leads to the 'hot' CNO cycle, and other reactions must be considered. This variant of the CNO cycle is initiated by the $^{13}N(p,\gamma)^{14}O$ reaction, which involves the radioactive nucleus ^{13}N ($T_{1/2}\approx10$ min). In those conditions, traditional experiments, using a proton beam, cannot be used, since the lifetime of ^{13}N is too short to be considered as a target. Significant progresses have been carried out with the availability of radioactive beams (see 'Primordial nucleosynthesis' section).

Helium burning

Hydrogen burning is the longest phase in the evolution of a star. When hydrogen is consumed in the core, the star contracts and its central temperature increases. In those conditions, helium burning can start. However, the $\alpha+p$ and $\alpha+\alpha$ reactions are impossible since ^5Li and ^8Be are unbound, and immediately break up. Helium burning therefore proceeds through the triple α process, where ^8Be can exist for a short time, in equilibrium with $\alpha+\alpha$, and captures a third α particle by the $^8Be(\alpha,\gamma)^{12}C$ reaction.

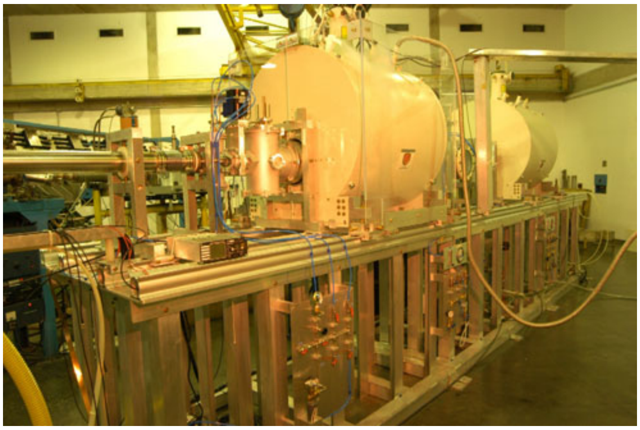

Fig. 7. The RIBRAS system installed in the experimental area of the Nuclear Physics Laboratory of the University of São Paulo (photo from O. Camargo).

The second step of the 3α process, i.e. the $^8Be(\alpha,\gamma)^{12}C$ reaction, provides a striking example of the interplay between astrophysics and nuclear physics. The observed abundance of ^{12}C in the Universe can only be explained if ^{12}C presents a 0^+ resonance in the Gamow energy range. This resonance was predicted by Hoyle in 1954 from anthropic arguments (a carbon-based life), and discovered experimentally a few years later. This 0^+ state of ^{12}C is known as the 'Hoyle state', and its properties are now firmly established.

The $^{12}C(\alpha,\gamma)^{16}O$ follows the 3α process. This reaction determines the $^{12}C/^{16}O$ ratio after helium burning, and is crucial in many stellar models. As for $^8Be(\alpha,\gamma)^{12}C$ the cross section is enhanced by specific nuclear properties of ^{16}O. Owing to the high Coulomb barrier, the cross section cannot be directly measured at stellar energies. A theoretical support is therefore necessary to extrapolate the available data to the Gamow energy range.

In principle, Helium burning could continue with the $^{16}O(\alpha,\gamma)^{20}Ne$ reaction. However, ^{20}Ne does not present any resonance at stellar energies, and the cross section is therefore quite small in standard conditions (temperature and density). In general, helium burning ends at ^{16}O. The formation of heavier elements involves various processes. Beyond iron, elements are essentially produced by neutron-capture reactions since the Coulomb barrier increases with the charge of the elements, and proton captures are impossible.

Influence of radioactive nuclides in the nucleosynthesis

We have seen that nuclear reactions are responsible for the formation of elements in the Universe. Most reactions have been studied experimentally, sometimes in a limited energy range. There are compilations of thermonuclear reaction rates initiated by the Caltech group (Caughlan & Fowler 1988) and more recent ones (Angulo *et al.* 1999; Descouvement *et al.* 2004; Longland *et al.* 2010).

However, many of these reactions involve radioactive nuclei, which can be formed during the Big Bang or in nuclear reactions in stars, such as ^{12,13}N, ^{8,9}Li, ^{10,11,12}Be, etc. The study of radioactive nuclei is one of the very important and active areas of nuclear physics nowadays. They can be produced in accelerators and usually they have half-lives long enough ($T_{1/2} \sim 1$ μs to s) to allow the production of secondary beams with them.

The use of radioactive beams to study nuclear reactions of astrophysical interest is a very challenging issue. As for other reactions, the cross sections in the relevant energy range are very small, but the beam intensity of the available radioactive beams is much lower than of stable beams. Their production and use in reactions of astrophysical interest is the subject of some review papers (Kubono 2001; Smith & Rehm 2001).

We have installed next to our 8MV Pelletron Tandem, in our Nuclear Physics Laboratory at the University of São Paulo, a facility to produce secondary radioactive ion beams, the Radioactive Ion Beams in Brazil (RIBRAS) system (see Fig. 7). It consists of a production target, where the radioactive nuclides are produced by means of transfer reactions, followed by a double superconducting solenoid system that selects and focuses the radioactive beam of interest.

We are producing beams of ^{6}He, ^{7}Be, ^{8}Li, ^{8}B, ^{10}Be and ^{12}B, and studying the reactions induced by beams of radioactive nuclides. The RIBRAS facility is the first and for the moment the only radioactive beam facility of the Southern hemisphere.

Conclusions

Astrophysics is a broad science that requires inputs from various fields: nuclear, atomic and particle physics, cosmology, thermodynamics, optics, etc. We have focused here on nuclear astrophysics and, more specifically, on nuclear reactions involved in stellar models. Nuclear physics aspects are not limited to reactions: nuclear masses and β-decay rates, for example, are also important nuclear ingredients to the star evolution.

Many nuclear reactions, essentially involving protons and α particles, play an important role in astrophysics. They are responsible for the energy released by the star and for its evolution. The interplay between astrophysics and nuclear physics is illustrated by many observations; the abundances of the elements, and therefore the conditions for a carbon-based life, are clearly associated with nuclear-physics properties.

The main problem in nuclear astrophysics is that, at the nuclear scale, stellar energies are low, and the corresponding cross sections between charged particles are often too small to be directly measured in the laboratory. Collaborations between experimentalists and theoreticians are therefore necessary to provide reliable reaction rates. Another issue is the need of reaction rates involving unstable nuclei. Significant progress has been made recently thanks to technological advances in the development of radioactive beams.

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